Very High Energy astronomy from H.E.S.S. to CTA. Opening of a new astronomical window on the non-thermal Universe
Mathieu de Naurois

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MÉMOIRE
d’HABILITATION À DIRIGER DES RECHERCHES de
l’UNIVERSITÉ PARIS VI

Spécialité:
Particules, Noyaux et Cosmos
présenté par
Mathieu de Naurois
Laboratoire Leprince-Ringuet
Ecole polytechnique, 91128 PALAISEAU Cedex

L’astronomie $\gamma$ de très haute énergie.
Ouverture d’une nouvelle fenêtre astronomique sur l’Univers non thermique.

Soutenu le 13 Mars 2012 devant le jury composé de:

MM. Pascal Vincent
Christian Stegman
Sylvie Rosier Lees
Manel Martinez Rapporteur
Stephane LeBohec Rapporteur
Fabian Zomer Rapporteur
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Part I

Toward precision $\gamma$-ray astronomy
Introduction

The energy spectrum of high energy cosmic rays$^{[1]}$ Fig. 1] is most probably one of the most famous plots of modern science. It summarizes, over more than 12 orders of magnitudes, the work of several thousands of physicists across the world, grouped in several dozens of experiments, and using a huge variety of techniques ranging from satellite observations, balloon-born experiments, ground-based particle detectors, huge arrays of water tanks or scintillators as well as ultra-fast telescopes for the detection of giant showers. It certainly represents one of the most collaborative efforts of the 20$^{th}$ century, deserving, according to some authors [e.g. 1], the title of “one of the seven wonders of the world of physics.”

The spectrum of cosmic rays exhibits a quite impressive regularity over 32 orders in magnitude, corresponding to about 12 orders in individual particle energy, and indicating that non-thermal, wide-band phenomena must be at its origin. The source of these energetic particles travelling across the cosmos remains largely unknown, and nearly one century of research has not yet allowed to pin-point the mechanism of acceleration nor the astrophysical objects where they are produced.

To achieve a full understanding of this wide-band phenomenon, very different detection techniques are in particular required in different energy bands. High fluxes at low energies allow direct detection and detailed, precise composition measurements, either from space or using high altitude balloon experiments. Around the knee, a break in the energy spectrum located at an energy of $\sim 5 \times 10^{15}$ eV, the already low fluxes implies to develop experiments with effective areas in the order of 1 km$^2$, that are rendered possible by the existence of extensive air showers in the atmosphere. At the very highest energies $\sim 10^{20}$ eV, very low fluxes of the order of 1 particle per square kilometer and per century impose the use of giant arrays covering surfaces of the order of $\geq 1000$ km$^2$.

In this wide field of research, very high energy (VHE) $\gamma$ ray astronomy plays a particular role: In contrast to charged particles, that are deflected by Galactic and extra-galactic magnetic fields and, therefore, do not allow us to identify their origin$^[2]$ $\gamma$ rays travel rectilinearly in the Universe, thus allowing the source of their emission to be pin-pointed.

Decades of technical and theoretical developments have been necessary to bring the field of VHE astronomy to the state of a mature, reliable, technique for the observation of the High Energy Universe. This thesis is the view of one insider on the evolution of this field. The first chapter aims to provide an introduction on the physics of cosmic rays, and on their production and radiation mechanisms. The second focuses on the basics of VHE astronomy with ground-based experiments. Starting from the first chapter, the H.E.S.S. experiment is explained in detail. Some results of particular interest are briefly explained in the second part.

$^1$Cosmic rays high energy particle continuously bombarding the earth, and produced in the distant Universe in violent, non-thermal phenomena.

$^2$except, at the very highest energies, for nearby sources
Figure 1: Energy Spectrum of cosmic rays. From [2].
Chapter 1

The Very High Energy Universe

1.1 A brief history of cosmic rays

History remembers the year of 1912 as the date of the discovery of cosmic rays by Victor Hess using balloon experiments. Measurements done at altitudes up to 5300 m showed a regular increase of ionizing radiations. Hess took advantage of the solar eclipse of 1912 to demonstrate, from the absence of shadow, that cosmic rays do not originate from the Sun and he was awarded the Nobel Prize in 1936 for this discovery. The photograph of Hess in his balloon (Fig. [11] is one of the most famous photographs in science.

The pioneering work of several physicists is somewhat forgotten in this statement, although of prime importance. A better time to start the history of cosmic ray physics is probably the year 1900 with the first discovery of a ionizing component in the atmosphere. Important milestones include in particular:

- 1900 : Charles Thomson Rees Wilson, when studying the formation of clouds in a closed chamber, and searching for the condensation nuclei, found that the air is weakly ionized. He experimented with the creation of cloud trails in his chamber caused by ions and radiation. Wilson noticed the reappearance of drops of condensation in expanded dust free gas. This led to the discovery of the cloud chamber, which he first presented in 1912, and for which he was awarded the Nobel Prize in 1927.

- 1910 : First measurements from Theodor Wulf at the top of Eiffel Tower using an electroscope. He found a higher radiation level at the top of the tower than at its base, but his paper, published in Physikalische Zeitschrift, was not widely accepted.

- 1912 : Historical balloon measurements by Victor Hess up to 5300 m, showing that the ionization level increases with altitude, and is therefore caused by extraterrestrial particles. These measurements were based on the spontaneous discharge of an electroscope (in the absence of ionizing radiation, the electroscope remains charged). The balloon carried three high accuracy electrometers. Hess wrote in his notebook “The results of the present observations are most easily explained by the assumption that radiation with very high penetrating power enters the atmosphere from above; even in its lower layers, this radiation produces part of the ionization observed in closed vessels... Since there was neither a decrease at night or during solar eclipse, the Sun can hardly be considered as the source...”
1.1. A BRIEF HISTORY OF COSMIC RAYS

Figure 1.1: Victor Hess in the balloon he used for his famous discovery of cosmic rays

- 1913-1914: Werner Kolhörster repeated the observations of Victor Hess up to 9 km and came to similar conclusions
- 1925: Robert Millikan introduced the term of cosmic rays and assumed they were neutral particle reaching the ground
- 1928-1929: Using Geiger counters instead of oscilloscopes, it was shown that Cosmic Rays are most probably charged particles (Science, 1930)
- 1928: Discovery of latitude effect by Jacob Clay: cosmic rays are bent by the Earth’s magnetic field, resulting into lower intensity of radiation at the equator than at the pole. This was a second hint in favour of a charged nature of cosmic rays.
- 1929: a Russian scientist, Dmitri Skobelzyn, discovered ghostly tracks made by cosmic rays in a cloud chamber. In the same year, Walther Bothe and Werner Kolhörster verified that the cloud chamber tracks were curved. Thus the cosmic radiation was composed of charged particles.
- 1932: Hot debate between Millikan and Compton. Millikan argued that primary cosmic rays are mainly gamma rays, resulting from the continuous creation ("birth
cries") of atoms by God to counteract entropy increase and prevent the heat death of the Universe. According to Millikan, the observed charged particles resulted from secondary electrons produced by Compton scattering of gamma rays. Sir Arthur Compton, in contrast, argued that cosmic rays are genuine charged particles. This debate made the front page of the New York Times.

- 1932: Discovery of the positron by Carl David Anderson in the cosmic rays. He obtained the Nobel Prize in 1936 for this discovery that started a new era of tight connection with particle physics. Many particles were discovered in the cosmic rays: \( \mu^\pm (1936) \), \( \pi^\pm (1947) \), strange particles (1947), ...

- 1934: Simultaneous discovery of the east-west effect by Bruno Rossi (Eritrea) and Thomas H. Johnson (Mexico): charged particles are bent by the Earth’s magnetic field depending on their signs. As a consequence, the Earth shadows particles originating from one direction. The ionization rate increased from east to west viewing angle indicating they were positively charged particles (protons). When making tests for his experiments about the east-west effect, Rossi also reported an observation of near-simultaneous discharges of two Geiger counters widely separated in a horizontal plane. In his report on the experiment, Rossi wrote “...it seems that once in a while the recording equipment is struck by very extensive showers of particles, which causes coincidences between the counters, even placed at large distances from one another.”. Unfortunately, he did not have the time to study this phenomenon more closely; although this was the first hint for extended air showers.

- 1934: Observation by Walter Baade & Fritz Zwicky of very intense stellar explosion (~galactic luminosity), which they called Super-nova. They immediately proposed supernova remnants as sources of cosmic rays and supposed to be the result of a star collapse into a neutron star (confirmed in 1937)

- 1934: Hans Bethe and Walter Heitler developed the electromagnetic cascade theory thus showing that ionizing particles reaching the ground are secondary particles produced in the atmospheric showers

- 1936: Discovery of the muon in cosmic rays by Carl David Anderson and Seth Neddermeyer

- 1938: Proposition from Fritz Zwicky to use supernovae as standard candles to estimate distances in deep space. This idea marks the birth of modern cosmology.

- 1937: Unaware of Rossi’s earlier report, Pierre Auger discovered giant showers by studying coincidences between detectors located several meters apart. He concluded that extensive particle showers are generated by high-energy primary cosmic-ray particles that interact with air nuclei high in the atmosphere, initiating a cascade of secondary interactions that ultimately yield a shower of electrons, photons, and muons that reach ground level, and estimated that the primary energy must exceed \( 10^{15} \text{ eV} \).

- 1949: Proposition of a cosmic ray acceleration mechanism by Enrico Fermi through a second-order scattering process
1.2 Acceleration of particles

Soon after their discovery in the early 1930’s, supernovae were identified to be the prime candidates for cosmic ray acceleration [3]. The first theory of particle acceleration in astrophysical shocks was published by Enrico Fermi in 1949 [4] and was extensively developed in the following years. For a review, see for instance [5,6]. This section summarizes the main results of diffusive shock-acceleration theory.

1.2.1 Collisionless shock

Due to the very low density of the medium, astrophysical shocks are almost always collisionless plasma shocks, meaning that two-body collisions can be neglected with respect to interactions between the charged particles constituting the plasma and collective modes of the plasma [e.g. 6]. This also means that the medium in which the shock is propagating is ionized and may contain a magnetic field.
1.2.2 Second order Fermi particle acceleration

Let’s consider particles of mass $m$, being randomly deflected by magnetic perturbations (Fig. 1.2). After an interaction at incidence angle $\theta$, the energy of the particle can simply be obtained by a Lorentz transformation into the perturbation frame, followed by an elastic reflection and a Lorentz transformation back into the laboratory frame:

$$E' = \gamma V E \left[ 1 + \frac{2vV \cos \theta}{c^2} + \left( \frac{V}{c} \right)^2 \right]$$

(1.1)

Figure 1.2: Second order Fermi acceleration

where $\gamma_V$ is the Lorentz factor associated with the speed of the magnetic perturbation $V$. For non relativistic perturbations, one has $\gamma_V \approx 1 + \left( \frac{V}{c} \right)^2$. After averaging over all incident angles, and accounting for the fact that head-on collisions are more frequent (by a factor $1 + \frac{V}{c} \cos \theta$) than back collisions, the net average energy gain is:

$$\left\langle \frac{\Delta E}{E} \right\rangle = \frac{\int_{-1}^{1} \Delta E \left( 1 + \frac{V}{c} \cos \theta \right) d \cos \theta}{\int_{-1}^{1} \left( 1 + \frac{V}{c} \cos \theta \right) d \cos \theta} = \frac{8}{2} \left( \frac{V}{c} \right)^2$$

(1.2)

This energy gain is of the second order with respect to the perturbation speed. Magnetic perturbations are usually Alfvén waves, whose speed, in case of low density and low magnetic field, can be approximated by:

$$v_A = \frac{B}{\sqrt{\mu_0 n_i m_i}} \approx (2.18 \times 10^9 \text{ m/s}) \left( \frac{m_i}{m_p} \right)^{-1/2} \left( \frac{n_i}{1 \text{ cm}^{-3}} \right)^{-1/2} \left( \frac{B}{1 \text{ G}} \right)$$

(1.3)

where $n_i$ is the ion number density and $m_i$ is the ion mass. For typical interstellar magnetic field of $B \sim 5 \mu \text{G}$, the order of magnitude of Alfvénic waves is $v_A \approx 10^4 \text{ m/s}$, leading to very small energy gains of the order of $10^{-8}$. Second order Fermi acceleration is therefore not efficient, except in the case of relativistic perturbations [e.g. 7].
1.2.3 First-order Fermi particle acceleration

One of the main limitation of the second order Fermi acceleration mechanism comes from its low efficiency. It was soon realized that strong shocks can provide a much more efficient acceleration process.

![Diagram of first order Fermi acceleration](image)

Figure 1.3: First order Fermi acceleration by reference change.

Rankine-Hugoniot conditions

The Rankine-Hugoniot conditions relate the thermodynamic quantities in the upstream and downstream media. They are obtained from several conservation laws (for non-relativistic shocks):

- Conservation of matter (continuity equation): \( \rho_1 v_1 = \rho_2 v_2 \)
- Conservation of momentum: \( p_1 + \rho_1 v_1^2 = p_2 + \rho_2 v_2^2 \)
- Conservation of enthalpy: \( \frac{\gamma}{\gamma - 1} \frac{p_1}{\rho_1} + \frac{v_1^2}{2} = \frac{\gamma}{\gamma - 1} \frac{p_2}{\rho_2} + \frac{v_2^2}{2} \)

The ratio of density (and pressure) between the two media can then easily be obtained:

\[
\frac{\rho_1}{\rho_2} = \frac{(\gamma + 1)p_1 + (\gamma - 1)p_2}{(\gamma - 1)p_1 + (\gamma + 1)p_2}, \quad \frac{p_1}{p_2} = \frac{(\gamma + 1)p_1 - (\gamma - 1)p_2}{(\gamma + 1)p_2 - (\gamma - 1)p_1}
\]

The Mach Number \( M_1 \) relates the shock speed with the sound speed \( c_1 \) in the unshocked (upstream) medium: \( M_1 = v_1/c_1 \), where \( c_1 = \sqrt{\gamma p_1/\rho_1} \). It can be used to rewrite the ratio of density (and pressure) between the two media:

\[
\frac{p_1}{p_2} = \frac{\gamma + 1}{2\gamma M_1^2 - (\gamma - 1)}, \quad \frac{\rho_1}{\rho_2} = \frac{v_2}{v_1} = \frac{2 + M_1^2(\gamma - 1)}{M_1^2(\gamma + 1)}
\]

Writing the compression ratio as \( r = v_1/v_2 = p_2/p_1 \), in the limit of a strong shock (\( M_1 \gg 1 \)), one obtains:

\[
r = \frac{\gamma + 1}{\gamma - 1}, \quad \frac{p_1}{p_2} \approx \frac{\gamma + 1}{2\gamma M_1^2} \to 0, \quad \frac{\rho_1}{\rho_2} \approx \frac{\gamma - 1}{\gamma + 1}
\]

For a mono-atomic gas, \( \gamma = 5/3 \) and the strong shock corresponds to \( r = 4, p_2 = 4p_1 \) and \( v_1 = 4v_2 \).
Energy gain

In first-order Fermi acceleration, we consider particles crossing the shock back and forth. At each shock crossing, due to the change of the reference frame, the particle moves from a fast medium into a slow one, the speed difference being \( v_1 - v_2 \). Each process of the cycle can be regarded as having no associated acceleration in some frame. The scattering in the upstream (resp. downstream) medium is at constant energy in that frame; the shock crossings are at constant energy in the shock frame. The combination of these reference frame changes leads to the fact that, after having completed a cycle of shock crossings and scatterings, a particle will have increased its energy if the initial and final energies are measured in the same local frame.

When taking into account the incident angle of the particle on the shock, the energy gain reads after each crossing (in the case of a non relativistic shock):

\[
\Delta E = E \frac{v_1 - v_2}{v} \cos \theta
\]

(1.7)

After averaging over the incident angle, and taking into account the fact that the crossing probability is proportional to \( \cos \theta \), the average energy gain is easily obtained :

\[
\left\langle \frac{\Delta E}{E} \right\rangle = \int_{0}^{1} \frac{v_1 - v_2}{v} \cos^2 \theta \cos \theta d \cos \theta
\]

\[
= \frac{2 v_1 - v_2}{3} \frac{2 v_1 - v_2}{c} \approx \frac{2 v_1 - v_2}{3 c}
\]

(1.8)

For a full cycle (crossing in both directions) the total energy gain is then:

\[
\left\langle \frac{\Delta E}{E} \right\rangle = \frac{4 v_1 - v_2}{3 c}
\]

(1.9)

Escape probability

Particles can escape the acceleration region due to advection into the downstream medium, which tends to tear particles away from the shock at the speed \( v_2 = v_1/r \). The advection rate away from the acceleration region is therefore \( nv_1/r \). The flux of particles crossing the shock is, for isotropic distribution and non-relativistic shock:

\[
\phi = \frac{nc}{\int_{0}^{1} \cos \theta d \cos \theta} = nc/4
\]

(1.10)

The escape probability over a full cycle is then \( P_{\text{esc}} = 4v_1/rc \), while \( P = 1 - 4v_1/rc \) is the probability for a particle to remain in the acceleration region.

Energy spectrum of accelerated particles

Let’s write \( \beta = \left\langle \frac{\Delta E}{E} \right\rangle \) the relative energy gain for each cycle. We have, in the case of first-order Fermi acceleration:

\[
\beta = \frac{4 v_1 - v_2}{3 c} = \frac{4 v_1 r - 1}{3 c r}
\]

(1.11)
1.2. ACCELERATION OF PARTICLES

In the limit of strong shocks, \( r = 4 \) and \( \beta = v_1/c \). After \( k \) cycles, \( N_k = N_0 P_k \) particles reached the energy \( E_k = E_0 (1 + \beta)^k \) (and can eventually be further accelerated). Solving for \( k \), one obtains the integral energy spectrum:

\[
\frac{N(E \geq E_k)}{N_0} = \left( \frac{E_k}{E_0} \right)^{\frac{\ln P}{\ln(1+\beta)}} = \left( \frac{E_k}{E_0} \right)^{\frac{\ln(1-P_{\beta,acc})}{\ln(1+\beta)}} \approx \left( \frac{E_k}{E_0} \right)^{-\beta/(r-1)} \tag{1.12}
\]

First order Fermi acceleration naturally produces a power-law spectrum, with integral index \( \alpha = 3/(r-1) \) (= 1 for a strong shock) and differential particle index \( \Gamma = (r + 2)/(r - 1) \).

**Acceleration time**

The acceleration time can be calculated from the diffusion time in the downstream and upstream media (with relative diffusion coefficients \( D_1 \) and \( D_2 \)). In the regime of Bohm diffusion, which corresponds to the slowest possible diffusion, the mean free path is equal to the Larmor radius of the particle \( R_L = c p/eB \approx E/eB \). In general, the diffusion coefficient can be expressed as \( D = rv/3 = \eta R_L c / 3 \), with \( \eta \geq 1 \) (the value \( \eta = 1 \) corresponding to the Bohm diffusion regime). During a time \( t \), a particle diffuses in the medium a typical distance \( L_d = \sqrt{D/t} \), while the shock moves during the same time by a distance \( L_S = v \times t \). Equating the two lengths gives the residence time scale: \( t = D/v^2 \) and the diffusion distance \( L_d = D/v \) that is travelled by the particle during a time \( \tau = L_d/c \). Using the upstream and the downstream media, one can show that the duration of a cycle is:

\[
\tau_{cycle} = \frac{4}{c} \left( \frac{D_1}{v_1} + \frac{D_2}{v_2} \right) \tag{1.13}
\]

and for a cycle,

\[
\beta = \left\langle \frac{\Delta E}{E} \right\rangle \approx \frac{\langle dp \rangle}{p} = \frac{4 v_1 - v_2}{3} \frac{1}{c} = \frac{4}{3} \frac{v_1}{c} \frac{r - 1}{r} \tag{1.14}
\]

The acceleration time is then:

\[
\tau_{acc} = \int_{p_i}^{p_f} \left( \frac{dt}{dp'} \right) dp' = \int_{p_i}^{p_f} \left( \frac{\tau_{cycle}}{\beta p'} \right) dp' \tag{1.15}
\]

For identical magnetic fields in both regions, \( D_1 = D_2 = \eta E_{c} / 3eB \). One then obtains [5]:

\[
\tau_{acc} = \frac{r(r + 1)}{v_1^2 (v_1^2 - 1)} \frac{E}{eB} \approx \frac{20 \eta}{3} \frac{E}{v_1^2 eB} = 2113 \text{ yr} \times \eta \times \left( \frac{E}{1 \text{ TeV}} \right) \left( \frac{1 \mu \text{ G}}{B} \right) \left( \frac{1000 \text{ km s}^{-1}}{v_1} \right)^2 \tag{1.16}
\]

For a typical young supernova remnant, \( v_1 = 10^4 \text{ km s}^{-1} \) and \( B = 2 \text{ nT} \). To reach 100 TeV, only 100 yrs are necessary, making shock acceleration a very efficient mechanism.

The maximum energy depends on the physical conditions of the system and is limited by several factors:

- Energy losses (Coulomb, radiative, inelastic collisions)
- Age of the system
- Particle escape, linked with the geometry of the system.
1.2.4 Maximum Energy – Hillas Criteria

A simple limitation on the energy an accelerator can provide comes from its size. In order to avoid the particles escaping the acceleration region, its size must be smaller than the Larmor radius of the particles. This criteria, known as Hillas criteria [9], simply reads:

\[
E_{\text{max}} = R ce Z B \approx 10^{12} Z \frac{B}{1 \text{ Gauss}} \frac{R}{1 \text{ pc}} \text{ GeV}
\]  

(1.17)

where \( R \) is the size of the acceleration region, \( Z \) the atomic number of the particles, and \( B \) the magnetic field. In the case of an acceleration region moving at relativistic speed, an additional Doppler factor has to be added: \( E_{\text{max}} = \Gamma R ce Z B \).

Potential astrophysical sources of high energy (and ultra high energy) cosmic rays are shown in Fig. 1.4 as function of the average size and magnetic fields. Diagonal lines correspond to the maximum energy than can be reached by a population of sources.

---

**Figure 1.4:** Hillas diagram showing possible sources of particle acceleration as function of their size and typical magnetic fields.
1.2.5 Maximum Energy: other criteria

In general, other criteria have to be taken into account to estimate the maximum energy a specific accelerator can reach. For that, the acceleration time scale has to be compared with:

- the age of the system
- the radiative losses time scale, mainly by synchrotron, inverse-Compton or Bremsstrahlung radiation
- the escape time scale, which decreases as the magnetic field in the system decreases (so that the diffusion coefficient increases)

1.2.6 Limit of test-particle approximation

These conditions are only valid in the test-particle approximation, when the cosmic rays have no retro-action on the shock structure.

![Figure 1.5: Density and temperature vs. radius for different values of the injection efficiency. As injection of particles in the main shock ($\eta_{\text{inj},p}$) becomes larger, retro-action of the particles on the shock becomes more and more important, leading to a shock precursor of increasing density on the right side (FS = Forward shock), and to a continuous decrease of the temperature at the main shock (Radius/contact discontinuity = 1). From [10].](image)
It was quickly realized [e.g. 11, 12, 10] that, since the efficient acceleration of particles leads to a rough energy equipartition between the relativistic particles and the thermal gas, the test particle assumption is in general not valid.

Charged particles (cosmic rays) scattering on the turbulent magnetic field back and forth over the shock front, can diffuse ahead of the shock into the upstream region and produce an adverse pressure gradient, called the shock precursor. This is illustrated in Fig. 1.3. This also leads to the fact that the ambient medium gets pre-heated.

Since the cosmic rays have, in general, a softer equation of state than the background plasma (for a fully relativistic gas $\gamma = 4/3$ whereas the background typically has an adiabatic exponent of $\gamma = 5/3$), this precursor decelerates and compresses the incoming plasma, leading to larger compression ratios [8, 10]. In particular, from Eq. 1.6, the compression ratio can reach 7 for cosmic-ray dominated shocks compared to 4 for thermal gas dominated shocks. The velocity of the plasma drops in the precursor and the temperature, and hence the sound speed, rises due to the compression, leading to a decrease of the local Mach number. The temperature behind the main shock becomes lower than one would expect in the test particle assumption.

One of the most interesting effects of cosmic ray retro-action on the shock structure is that it can lead, because of larger compression ratios, to more efficient acceleration of charged particles. Full relativistic treatment is required to fully understand the acceleration of particles in the case of significant retro-action.

### 1.3 Radiation processes

In this section we recall the main radiation processes, that convert energy in cosmic rays into $\gamma$ rays than can be detected from satellite or ground based instruments.

#### 1.3.1 Synchrotron Emission

Energy losses by synchrotron radiation are given by:

$$-\frac{dE}{dt} = 2\sigma_T c U_B \beta^2 \gamma^2 \sin^2 \theta$$

with $U_B = B^2/2\mu_0$ being the magnetic energy density and $\sigma_T$ the Thomson cross-section. For an isotropic electron population, averaging over angles yields:

$$-\frac{dE}{dt} = \frac{4}{3} \sigma_T c U_B \beta^2 \gamma^2$$

Important characteristics of Synchrotron emission are:

- The emission is proportional to $\gamma^2$
- The emission is proportional to $B^2$.
- The emission is bipolar in the particle frame. Due to relativistic motion of the particle, the emission in the laboratory frame is beamed in a cone of opening angle $1/\gamma$ in the direction of motion.
1.3. RADIATION PROCESSES

Peak of emission

The peak of emission occurs for the frequency:

\[ \nu_{\text{max}} = 0.29 \nu_c = 1.16 \left( \frac{E}{5 \text{ GeV}} \right)^2 \left( \frac{B}{1 \text{ nT}} \right) \text{ GHz} = 1.9 \left( \frac{E}{100 \text{ TeV}} \right)^2 \left( \frac{B}{1 \text{ nT}} \right) \text{ keV} \]

(1.20)

Emission from a particle population

When calculating the radiation from an electron power-law energy distribution, \( N(E) \text{d}E = KE^{-p} \text{d}E \), one can assume a delta-function approximation (where the radiation at frequency \( \nu \) is assumed to occur exclusively from electrons at energy \( E \) such as \( \nu_{\text{max}} = \nu \) from Eq. 1.20). Then, the spectral energy distribution of radiated emission reads:

\[ \nu \propto E^2 \times B \iff E \propto v^{1/2} B^{-1/2} \implies \text{d}E \propto v^{-1/2} B^{-1/2} \text{d}\nu \]

(1.21)

The energy flux per unit frequency reads:

\[ J(\nu) = \int -\frac{\text{d}E}{\text{dt}} \times N(E) \text{d}E \propto \sigma_T U_B \int -E^2 KE^{-p} v^{-1/2} B^{-1/2} \text{d}\nu \propto B^{(p+1)/2} \nu^{(1-p)/2} \]

(1.22)

The differential energy flux index is therefore \( \alpha = (p - 1)/2 \). The differential photon flux index is \( \Gamma = \alpha + 1 = (p + 1)/2 \). This is shown in Fig. 1.6 for different ranges of electron energies.

![Synchrotron Spectrum](image)

Figure 1.6: Energy spectrum of synchrotron radiation emitted by a power law population of electrons

The electron cooling time is:
\[ \tau = \left( \frac{1}{E} \times \frac{dE}{dt} \right)^{-1} = \frac{1}{bU_B E} \quad \text{with} \quad b = \frac{4}{3} \frac{\sigma_T c}{m_e^2 c^4} \]  

(1.23)

In weak magnetic fields, the cooling time is usually much longer than the supernova age, and can be safely ignored. Synchrotron emission at keV energies is usually produced by highly relativistic electrons having a much shorter cooling time-scale. Low energy synchrotron emission is the superposition of young and older electrons and is therefore suspected to last longer than high energy synchrotron emission.

### 1.3.2 Electron-photon interaction

Several electron-photon interactions involve two incoming particles and two outgoing particles, two of which are photons and the other two are electrons (or positrons). They all have the same Feynman diagram and therefore all have basically the same cross sections:

- Compton scattering: the incoming particles (and the outgoing particles) are an electron and a photon. The photon is scattered by the electron and eventually transfers a fraction of its energy. If the photon energy (in the electron rest frame) is small compared to the electron rest mass energy \( m_e c^2 \), the photon undergoes an elastic scattering (its energy is almost unchanged). The cross section is then the Thomson cross section. At higher energy, a significant fraction of the photon energy is transferred to the electron. The Klein-Nishina limit applies and the scattering cross section drops quickly with energy.

- Pair production: the incoming particles are two photons, and the outgoing particles are an electron and a positron. This process has an energy threshold which corresponds to the rest mass energy of the produced particles. This process is important in astrophysics: due to pair-creation on low energy photons (CMB and infra-red) the size of the observable Universe (horizon) shrinks as the energy increases.

- Likewise, an electron and a positron can undergo a pair annihilation, for which no threshold must be satisfied. The produced photons will have an energy of the order of \( m_e c^2 \) at least, so the pair annihilation will always occur in the Klein-Nishina regime (see below)

- Bremsstrahlung is a special case of Compton scattering where the incoming photon is replaced by a virtual photon representing the Coulomb field of a nucleus, leading to a reduction of the cross section by the fine structure constant \( \alpha \sim 1/137 \). In the same way, pair creation of a high energy photon on the Coulomb field of a nucleus will only be possible above the threshold of \( 2m_e c^2 \). These two processes are responsible for the development of atmospheric showers that shield the earth from high-energy cosmic rays.

### 1.3.3 Compton Scattering

The unpolarized Klein-Nishina cross section can be obtained from quantum electrodynamics:

\[ \frac{d\sigma}{d\cos \theta} = \frac{3}{4} \frac{\sigma_T}{\epsilon_i^2} \left( \frac{\epsilon_i}{\epsilon_f} + \frac{\epsilon_f}{\epsilon_i} - \sin^2 \theta \right) \]  

(1.24)
If the photon energy in the electron rest frame is comparable to, or larger than the electron rest mass energy, the photon energy is not conserved (due to electron recoil energy) and reads:

$$
\epsilon_f = \frac{\epsilon_i}{1 + \frac{\epsilon_i}{m_e c^2}(1 - \cos \theta)}
$$  \hspace{1cm} (1.25)

In the low energy region approximation, the photon energy is unchanged ($\epsilon_i = \epsilon_f$), but the photon is scattered at a different angle. The differential and integral cross section read:

$$
\frac{d\sigma}{d\cos \theta} = \frac{3}{4} \sigma_T(1 + \cos^2 \theta), \quad \sigma = \sigma_T
$$  \hspace{1cm} (1.26)

At high energy, relativistic effects make the cross section drop:

$$
\sigma_{KN} = \frac{3}{8} \sigma_T \frac{1}{x} \left( \ln 2x + \frac{1}{2} \right)
$$  \hspace{1cm} (1.27)

Figure 1.7: Compton scattering cross-section as function of photon energy scaled to the electron mass energy $x = \epsilon_i / m_e c^2$.

This is shown in Fig. 1.7 as function of photon energy scaled to the electron mass energy $x = \epsilon_i / m_e c^2$. The drop of the cross section at high energies has two main implications:

1. Inverse-Compton radiation will be less efficient at very high energies, resulting in a steepening of the spectral energy distribution

2. On the other hand, radiative cooling of very high electrons will be less efficient, thus allowing electrons to possibly reach higher energies.

**Inverse-Compton Scattering**

Inverse-Compton scattering is the opposite process, where the electron moves at relativistic speed and scatters a low energy photon up to high energies. We denote the Lorentz
factor of the electron in the laboratory frame as $\gamma$ (when the photon distribution is usually isotropic).

- In the electron rest frame, the incoming photon energy is of the order of $\gamma \epsilon$.
- In the same frame, when the incoming photon energy is small compared to the electron rest mass energy ($\gamma \epsilon < m_e c^2$), scattering occurs in the Thomson regime and the scattered photon energy remains close to $\gamma \epsilon$. In the high energy (Klein-Nishina) regime, the scattered photon energy is on the order of $m_e c^2$.
- After Lorentz transformation back into the laboratory frame, the scattered photon energy is on the order of $\gamma^2 \epsilon$ (in Thomson regime) and on the order of $\gamma m_e c^2$ in Klein-Nishina regime.

The maximum energy of the scattered photon occurs for head-on collision and is $4 \gamma^2 \epsilon$. The following orders of magnitude are obtained:

- On the CMB ($\epsilon = 0.61 \times 10^{-3}$ eV) a 1 TeV gamma ray is produced by an electron of energy $E_\gamma = \gamma m_e c^2 = 20$ TeV.
- On the infra-red background ($\epsilon = 0.01$ eV), a 100 keV X-ray photon is produced by an electron of energy 1.5 GeV.

**Inverse-Compton spectrum for a particle population**

Computation of the energy distribution of the scattered particles involves the following steps:

- Transformation of the (isotropic) photon distribution into the electron rest frame. In this frame the photon field is no longer isotropic.
- Integration over the photon energy distribution for every scattering angle.
- Integration over scattering angles and Lorentz transformation into the laboratory frame.

For a single accelerated electron, the energy loss by inverse Compton Scattering is found to be:

$$- \frac{dE}{dt} = \frac{4}{3} \sigma_T c U_R \beta^2 \gamma^2$$

(1.28)

where $U_R$ is the target photon energy density. For a power-law electron distribution $n(E) = \frac{dn}{dE} = KE^{-p}$, the radiated energy per unit frequency (or energy) i:

$$J(E_\nu) = \frac{2}{3} \sigma_T c U_R KE_\nu^{1-p} (m_e c^2)^{-\frac{1-p}{2}}$$

(1.29)

The differential index of specific energy is then $\alpha = (p - 1)/2$. The differential index of the photon distribution is $\Gamma = (p + 1)/2$: the inverse Compton spectrum has the same shape as the synchrotron spectrum. In the Thomson limit:

$$\frac{dN}{dt dV \epsilon_f} \propto \sigma_T c n_{\text{ph}} \epsilon_f^{-(p+1)/2}$$

(1.30)
Thus the radiated powers in inverse Compton and in synchrotron emission scale as the ratio of radiation field over magnetic field energy density:

\[
\frac{P_{IC}}{P_{\text{sync}}} = \frac{U_r}{U_B}
\]

(1.31)

**Electron Cooling**

In a similar way to synchrotron losses, the cooling time scale reads, in the Thomson approximation:

\[
\tau = \left( \frac{1}{E} \times \frac{dE}{dt} \right)^{-1} = \frac{1}{bU_B E} \quad \text{with} \quad b = \frac{4}{3} \frac{\sigma_T c}{m_e^2 c^3}
\]

(1.32)

In the Klein-Nishina regime, the electrons are less efficiently cooled and the time scale becomes almost independent of the energy:

\[
\tau_{KN} = \left( \frac{1}{E} \times \frac{dE}{dt} \right)^{-1} = \frac{1}{bU_B} \ln \left( \frac{2E}{m_e c^2} + \frac{1}{2} \right) \quad \text{with} \quad b = \frac{1}{2} \frac{\sigma_T c}{m_e c^2}
\]

(1.33)

### 1.3.4 Synchrotron – Self Compton processes

In a Synchrotron – Self Compton (SSC) process, the inverse-Compton scattering occurs on the photon field that is produced by the same electrons through synchrotron radiation. In this case, one has:

\[
- \frac{dE}{dt}_{\text{sync}} = \frac{4}{3} \sigma_T c U_B \beta^2 \gamma^2, \quad - \frac{dE}{dt}_{\text{IC}} = \frac{4}{3} \sigma_T c U_B \beta^2 \gamma^2
\]

(1.34)

The ratio of IC over synchrotron energy losses is therefore simply:

\[
- \frac{dE}{dt}_{\text{IC}} = - \left( \frac{dE}{dt}_{\text{sync}} \times \frac{U_R}{U_B} \right) \times \left( \frac{dE}{dt}_{\text{sync}} \right)^2 \times \frac{1}{U_B}
\]

(1.35)

As a consequence, the IC and synchrotron luminosity are expected to be related by a square relation. In the external Compton process (inverse Compton emission on an external radiation field), the relationship is expected to be almost linear.

An example of a SSC spectral energy distribution is show in Fig.1.8 using the full treatment (solid lines) and using the Thomson regime approximation (dashed lines). Inverse-Compton scattering on an external field is also shown (red lines). In the Thomson regime, the energies of the synchrotron and inverse Compton peaks differ by a factor of \( \gamma_{\text{max}}^2 \) (Fig.1.9) whereas in the Klein-Nishina regime, they differ by a factor of \( m_e c^2 \gamma_{\text{max}} \). In both cases, the distance between the two peaks provides a measurement of the maximum energy of the underlying electrons.

### 1.3.5 Bremsstrahlung

Bremsstrahlung emission in matter is due to the interaction between the electric field of an electron and the Coulomb field of a nuclei. The electron energy losses are almost proportional to the electron incident energy (Bethe-Heitler calculation)
Figure 1.8: Spectral Energy distribution for a synchrotron-self-Compton model (blue lines) and for external Compton (red lines).

Figure 1.9: Evolution of the position of the synchrotron and inverse Compton peak positions as function of the maximum energy of the underlying electrons, in the Thomson regime (dashed lines) and using the full Klein-Nishina treatment (solid lines).

\[ -\frac{dE}{dt} = \frac{Z(Z+1,3)e^6n}{16\pi^3\epsilon_0m_e^2c^4}\left[\ln\left(\frac{183}{X^{1/3}}\right) + \frac{1}{8}\right] \times E \]  

(1.36)

The energy density is almost constant up to a critical frequency \( \nu_c = (\gamma - 1)m_ec^2/h \). In terms of differential photon spectrum:
\[ I(\omega)d\omega = P(\omega)\hbar \omega N_{\gamma}d\omega \Rightarrow P(\omega) \propto 1/\omega \] \hspace{1cm} (1.37)

Spectrum from a population of particles

In a similar way, when assuming that the photon takes in average half of the kinetic energy of the parent particle, the gamma-ray spectrum shape reflects the parent population:

\[ \frac{dE}{dt} \propto E \Rightarrow I_{br}(E_{\gamma}) \propto E^{-s} \] \hspace{1cm} (1.38)

1.3.6 Hadronic Interaction

One of the primary reactions of high energy hadrons with interstellar medium results in the production of pions:

\[
\begin{align*}
p + p & \rightarrow p + p + \pi^0 \\
p + p & \rightarrow p + n + \pi^+ \\
p + p & \rightarrow p + p + \pi^+ + \pi^-
\end{align*}
\]

\(\pi^0, \pi^+ \text{ and } \pi^-\) are produced roughly in the same quantities. Neutral pions predominantly decay into a pair of \(\gamma\) rays, while charged pions decay into charged muons and neutrinos:

\[
\begin{align*}
\pi^0 & \rightarrow \gamma + \gamma \\
\pi^+ & \rightarrow \mu^+ + \nu_\mu \rightarrow e^+ + \bar{\nu}_\mu + \nu_e + \nu_\mu \\
\pi^- & \rightarrow \mu^- + \bar{\nu}_\mu \rightarrow e^- + \nu_\mu + \bar{\nu}_e + \bar{\nu}_\mu
\end{align*}
\]

Neutrinos therefore provide a smoking gun for identification of hadronic interactions. Two muonic neutrinos are produced on average per electronic neutrino, but due to neutrino oscillation the fluxes of neutrinos at the Earth are identical for the three flavours.

We consider the reaction \(p + p \rightarrow p + p + \pi^0\), where one proton is at rest while the other has a kinetic energy \(T\). The reaction is only possible above the threshold

\[ T_{th} = \frac{(2m_p c^2 + m_\pi c^2)^2 - 2m_p^2 c^4}{2m_p c^2} = 2m_\pi c^2 + \frac{m_\pi^2 c^2}{2m_p c^2} = 280 \text{ MeV} \text{ (with } m_\pi = 135 \text{ MeV)} \] \hspace{1cm} (1.39)

\(E_{\gamma} = \frac{1}{2} m_\pi c^2 = 67.5 \text{ MeV}\) \hspace{1cm} (1.40)
After Lorentz transformation and integration over angles, the produced $\gamma$ ray spectrum has the following characteristics:

- At high energy, the shape of the $\gamma$ ray spectrum is almost identical to the shape of the parent proton spectrum.
- The $\gamma$ ray spectrum exhibits a bump around 67.5 MeV
- The mean energy of the photons is of the order of $E_p/10$

1.3.7 Summary

The following tables summarize the spectral shape of the emitted radiation for several processes, assuming a parent particle power-law spectrum $N(E_p) \propto E_p^{-s}$ up to a cutoff energy $E_C$ (or Lorentz factor $\gamma_C$):

<table>
<thead>
<tr>
<th>Process</th>
<th>Energy Losses</th>
<th>Particle Energy</th>
<th>Radiated Spectrum</th>
</tr>
</thead>
<tbody>
<tr>
<td>Synchrotron</td>
<td>$dE/dt \propto E^2$</td>
<td>$E_\gamma \approx \frac{3eB}{4\pi m_e^2 c^4} E_b^2$</td>
<td>$I(E_\gamma) \propto E_\gamma^{(1-s)/2}$ up to $\frac{3eB}{4\pi m_e^2 c^4} E_C^2$</td>
</tr>
<tr>
<td>Inverse Compton (Thomson)</td>
<td>$dE/dt \propto E^2$</td>
<td>$E_\gamma \approx \frac{E_e^2 E_0}{m_e^2 c^3}$</td>
<td>$I(E_\gamma) \propto E_\gamma^{(1-s)/2}$ up to $\gamma_C^2 E_0$</td>
</tr>
<tr>
<td>Inverse Compton (Klein-Nishina)</td>
<td>$dE/dt \propto E^2$</td>
<td>$E_\gamma \approx \frac{E_e^2 E_0}{m_e^2 c^3}$</td>
<td>$I(E_\gamma) \propto E_\gamma^{(1-s)/2}$ up to $\gamma_C^2 m_e c^2$</td>
</tr>
<tr>
<td>Bremsstrahlung</td>
<td>$dE/dt \propto E$</td>
<td>$E_\gamma \approx 0.5 E_e$</td>
<td>$I(E_\gamma) \propto E_\gamma^{-s}$</td>
</tr>
<tr>
<td>Hadronic Interaction</td>
<td>$dE/dt \propto E$</td>
<td>$E_\gamma \approx 0.1 E_p$</td>
<td>$I(E_\gamma) \propto E_\gamma^{-s}$ above 67.5 MeV</td>
</tr>
</tbody>
</table>

1.4 A quick visit to the zoo

Supernova remnants

Supernova, resulting either from the thermonuclear explosion of an accreting white dwarf (type Ia supernova) or from the collapse of a massive star having exhausted its nuclear fuel, result into the propagation of an almost spherical blast wave into the interstellar medium. Particles (electrons and ions) can be accelerated in this blast wave up to $10^{14}$ eV (at least) through the Fermi mechanism. Supernova remnants usually exhibit an intense radio signal attributed to synchrotron emission of accelerated electrons. X-ray emission originates from thermal emission of hot ejecta superimposed with non-thermal emission from the same accelerated particles.
1.4. A QUICK VISIT TO THE ZOO

Although supernova remnants are most probably effective accelerators only during a relatively short time (a few thousand years), they might release a large amount of accelerated particles that could subsequently be detectable through their interaction with clouds of interstellar matter. This could, in particular, lead to diffuse emission in the neighbourhood of accelerators.

Pulsar Wind Nebula

Pulsar resulting from the collapse of massive stars are characterized by very intense magnetic fields (up to $10^8$ Y) and fast rotation with periods ranging from a few seconds to milliseconds. They are known to act as efficient cosmic accelerators although several models (polar cap, outer gap and slot gap) are still in competition to explain the details of their emission. Pairs of particles and antiparticles produced in the intense electromagnetic field spiral along the magnetic field lines and escape the system along the open lines reaching the light cylinder (where co-rotating lines would rotate around the pulsar at the speed of light).

The radiated power can be easily estimated from the neutron star period $P$ and its time derivative $\dot{P}$:

\[
W = 4\pi^2I \frac{\dot{P}}{P^3}
\]

where $I$ is the moment of inertia of the neutron star. The interaction of the intense wind from the pulsar with the ejecta of the supernova explosion can lead to the formation of a synchrotron nebula, or plerion, in the form of a stationary shock wave, where particles might be accelerated up to VHE energies. At the shock, the ram pressure of the wind is balanced by the pressure of the surrounding nebula and the kinetic energy of the wind is transformed into random motion.

Since the discovery of the Crab nebula [13], such plerions or pulsar wind nebula (PWN) have been firmly established as a major population of VHE sources [e.g. 14]. Radiative cooling of electrons away from the pulsar has been observed in several cases [e.g. 15].

Binary systems

Binary systems are very common throughout the Galaxy, as about 70% of stars live in binary or more complex systems. If one of the two members of the system is a compact object, acceleration of particles can take place up to very high energies.

The physical environment inside a close binary system is characterized by very high radiation densities (up to 1 erg cm$^{-3}$ of rather high frequency photons) and high magnetic fields, and is therefore radically different from the interstellar medium.

In the case when the compact object is a neutron star, interaction of the pulsar and the massive star winds can lead to the formation of a bow shock suitable for particle acceleration. For a black hole compact object, accretion can provide the needed energy for VHE particle acceleration in a scaled-down version of an AGN.

Stellar cluster, stellar winds

Stellar clusters can harbour a large number of massive stars in a relatively small volume, as well as pulsar wind nebula and supernova remnants. Interaction of the dense winds
from the massive stars could result into the formation of shocks where particle acceleration could take place [e.g. 16].

Collective effects from individual stellar winds can also lead to the formation of super-bubbles where acceleration can also take place [e.g. 17].

Active galactic nuclei

Active Galactic Nuclei are galaxies comprised of a super-massive black hole with a mass from millions to billions of solar masses surrounded by a large accretion disk. In about 10% of the cases they also exhibit Mpc-scale jets of plasma in the form of collimated highly relativistic outflows. Since the discovery of the galaxy Mrk 421 at VHE energies [18] AGNs now form a major class of VHE sources, with nearly 20 objects known, with red-shifts up to 0.536. Emission of AGNs is thought to arise inside the jets from blob of relativistic particles moving inside the jet, and is characterized by a double-humped spectral shape, with a low-frequency peak in the optical to X-ray regime, attributed to synchrotron emission of accelerated particles and a high-frequency peak caused by inverse Compton scattering on either the synchrotron photons or external photons from the accretion disk.

γ ray bursts

γ ray bursts (GRB) are events associated with very intense and brief emission of low energy γ ray photons, lasting from a few milliseconds to a few hundred of seconds, detected up to cosmological distances and usually followed by an afterglow that can last up to several months. Short GRBs, lasting less than a few seconds, could be associated with merger events of compact objects (spiralling binaries), while long GRBs may be produced by the gravitational collapse of a supermassive star ($\gtrsim 30M_\odot$). In both cases, the formation of a fireball would lead to the formation relativistic shocks from which VHE emission could arise. Despite several models, no GRB has been detected at VHE energies to date.

Star-burst galaxies

Starburst galaxies are essentially ordinary galaxies in which a episodes of massive star formation occur, usually in a very localized region (referred as start-burst region) of a few thousand light years in diameter. This star-burst activity is believed to be either triggered by external events, such as a galaxy merger, or induced by instabilities in the bar of the galaxy.

Due to the large concentration of gas and dust and to an large population of young and very massive stars, star-burst galaxies could naturally be efficient accelerators of VHE particles through the collision of winds from massive stars.

Clusters of galaxies

Clusters of galaxies are the largest gravitationally bound systems in the Universe, and therefore act as cosmological cosmic-ray storehouses. High energy particles produced by supernova activity in cluster galaxies and by particle acceleration in AGNs have been accumulating inside the cluster since the beginning of the Universe. Very high energy emission would unavoidably arise from the interaction of high energy cosmic rays with the inter-cluster medium.
Non thermal emission from high-energy particles in clusters has been detected mainly in radio, but no VHE emission has been detected so far. Emission from galaxy clusters is predicted at levels just below the sensitivity of current instruments.

Since the escape time of high-energy particles in clusters of galaxies vastly exceed the age of the Universe, probing the density of cosmic rays in clusters via their gamma-ray emission would provides a calorimetric measure of the total integrated non-thermal energy output of galaxies.
Chapter 2

Very High Energy $\gamma$-ray astronomy

Introduction

The last decade saw the emergence of Very High Energy (VHE) gamma-ray astronomy as a robust observational discipline [e.g. 19], mainly opened up through the work of the H.E.S.S., MAGIC and VERITAS experiments. Very high energy photons are produced mostly by energy transfer from accelerated electrons through Compton scattering off ambient light, or by interaction of protons on the interstellar medium though neutral-pion production.

More exotic processes such as annihilation of still undiscovered Dark Matter particles, such as neutralinos, are also predicted in some super-symmetric theories beyond the Standard Model. In contrast to charged particles deflected by magnetic fields, high energy photons travel in straight lines in the Universe, and can emerge even from very dense environments such as the Galactic Centre, thus allowing the very central engine to be identified.

At 1 TeV, photon fluxes are low and their measurement requires large collection areas, which are incompatible with space-borne experiments. Ground-based experiments rely on the detection of the extended shower of secondary particles initiated by the gamma-ray impinging in the upper atmosphere. In the Atmospheric Cherenkov Telescope technique, these secondary particles are detected through the very short duration (a few nanoseconds) light flash they emit as they travel faster than the speed of light in the atmosphere.

Fast cameras at the focal plane of each telescope record the shower images. The light distribution in the camera is used to reconstruct the parameters of the impinging gamma-ray (energy, direction, impact,...) and also to discriminate the showers initiated by gamma rays from those initiated by the much more numerous hadrons (mainly protons, alpha particles and nuclei).

Early reconstruction techniques were based on the so-called Hillas Parametrization [20], in which shower images in the camera are parametrized by an elliptical shape and reduced to a few numbers. Discrimination is then based on the comparison of image width with expectations from simulations.

Since HEGRA, stereoscopy [21] has been used to provide a geometrical reconstruction of shower direction and impact and an improved background rejection. More recently, a new reconstruction technique [22] based on the log-likelihood comparison of the actual, raw image with a pre-calculated model, resulted in an improvement of the sensitivity by a factor of $\sim 2$. This method is described in detail in chapter [6]. Effective areas of the order of $10^5$ m$^2$, which corresponds to the size of the light pool on the ground, can be
easily achieved, even with modest-sized telescopes.

The H.E.S.S. experiment is an array of four Atmospheric Cherenkov Telescopes (ACT) of 107 m² each, located in the Khomas Highlands, Namibia, and equipped with fast cameras comprised of 960 individual photo-multiplier pixels. H.E.S.S. achievements include the first VHE scan of the inner parts of the Galaxy [23, 24], which revealed a rich harvest of new sources belonging to the categories of shell-type supernovae remnants (SNR), pulsar wind nebulae (PWN), star forming regions and binary systems. Some of these sources, called "dark accelerators", have no known counterparts in other energy domains, thus strengthening the discovery capabilities of VHE astronomy. Key results also include the first VHE image of an SNR ever produced [25], which showed strong correlations between X-ray and VHE features, the detection of diffuse gamma-ray emission around the Galactic Centre [26], the detection of VHE orbital modulation in a binary system [27] and tight constraints on the diffuse intergalactic infra-red background obtained from the observation of distant active galactic nuclei [28], as well as the first observation of a star-burst galaxy [29].

2.1 Atmospheric Showers

2.1.1 Ingredients

When a high-energy gamma ray enters the atmosphere, it can interact with the Coulomb field of the atmospheric nuclei and produce an electron-positron pair. These undergo Bremsstrahlung emission in the same Coulomb field, thus resulting in new gamma-rays of smaller energy, which will again produce pairs of electron-positrons and so on. A shower of particles, consisting of sometimes billions of particles, develops in the atmosphere, until the particles drop into the energy regime where they are quickly stopped by ionization losses. For showers induced by hadronic particles, such as protons, additional reactions (spallation, ...) result into the production of nuclear fragments, π and K mesons, which decay into gamma rays (for π⁰), long-lived muons which reach the ground (for π⁺ and π⁻) and neutrinos.

The lateral spread of the shower is governed by multiple scattering and, for hadronic interactions, by large transverse momentum transfer. Showers induced by gamma rays have a typical width of a few meters. High-energy particles in the shower produce Cherenkov emission, which, for a typical Cherenkov angle of ∼ 1°, illuminates an area of > 10⁵ m², thus providing a very large effective area in which to detect the primary gamma ray.

The atmosphere behaves as an inhomogeneous calorimeter, where the primary energy is deposited. Only incomplete and degraded information is recorded at the bottom of the shower. Observables can be Cherenkov emission, the charged secondary particles that reach the ground (for sufficient high energy showers) and, for very high energy showers, nitrogen fluorescence light in the atmosphere and radio emission from the decelerated electrons. The goal of the atmospheric Cherenkov technique is to identify the incident particle (γ ray or hadron), and measure its direction and energy.

2.1.2 Processes

The elementary processes playing a role in the shower development are:

- Bremsstrahlung of e^± in the Coulomb field of nuclei,
Figure 2.1: Illustration of the development of atmospheric showers. **Left**: shower initiated by a γ ray of 300 GeV. **Right**: Shower initiated by a proton of the same energy. From [30].

- Conversion of high-energy γ rays into pairs of \( e^+ e^- \) in the Coulomb field of nuclei,
- Coulomb multiple scattering of \( e^\pm \)
- Energy losses of \( e^\pm \) by ionization and atomic excitation, leading to rapid extinction
of the shower at low energy. Energy losses become dominant below the critical energy, defined as the energy where Bremsstrahlung and ionization losses are equal. The critical energy is 83 MeV in air.

In addition to these processes, several processes are of less importance:

- Compton scattering and positron annihilation leads to a negative charge excess of about 10% (Askaryan effect).

- At high energy, photo-production or electro-production of hadrons (pions) can induce a hadronic component in electromagnetic showers. However, the photo-production cross section is typically a factor $10^{-3}$ less than the pair creation cross section.
At ultra-high energy, coherent interaction of a single $e^\pm$ with several atoms (due to contraction of length) leads to the Landau-Pomerantchuk-Migdal (LPM) effect which can drastically change the Bremsstrahlung and multiple scattering due to interference.

Hadronic showers comprise several components:

- Hadronic component: nuclear fragments, nucleons, $\pi$ & K mesons, etc.
- Electromagnetic component: from $\pi_0 \rightarrow \gamma\gamma$ and other radiative decays
- Muonic component: from decay of charged mesons ($\pi^\pm$ & $K^\pm$)
- Atmospheric neutrinos from decay of $\pi^\pm$, $K^\pm$ & $\mu^\pm$

### 2.1.3 Heitler model

Above the critical energy, two processes dominate in the longitudinal shower development: Bremsstrahlung and pair creation. Both depend on the same characteristic length: the electromagnetic radiation length, defined as:

$$\frac{1}{X_0} = 4\alpha r_e^2 \frac{N_A}{A} Z^2 \ln \left(183 Z^{-1/3}\right) \left[ \text{g cm}^{-2} \right]$$

(2.1)

In the atmosphere (dry air), the radiation length corresponds to 36.7 g cm$^{-2}$. The atmosphere is therefore a thick calorimeter of $\sim 27$ radiation lengths. The Bremsstrahlung emission leads to energy losses as a function of traverse depth $X = \int_{z}^{\infty} \rho(z)dz$:

$$E(X) = E_0 \exp \left(\frac{-X}{X_0} (1 + b)\right)$$

(2.2)

where $b = \frac{1}{18 \ln(183/Z^{1/3})} = 0.0122$ (in air).

Each electron loses half of its energy after a depth $R = X_0 \ln 2$. In the Heitler model, we assume that this energy is transferred to a single gamma ray. The integrated pair creation probability is given by:

$$\mu(X) = 1 - e^{\frac{-X}{X_0} (\frac{7}{9} + \frac{1}{2})}$$

(2.3)

In the limit $b \ll 1$ and neglecting the factor $7/9$, a gamma-ray photon undergoes a pair creation after the same distance of $R = X_0 \ln 2$. This leads to the simple model illustrated in Fig. [2.3].

- Each electron loses half of its energy to one gamma-ray after a depth $R = X_0 \ln 2$
- Each gamma-ray photon undergoes a pair creation after the same distance $R = X_0 \ln 2$. The electron energy is assumed to be distributed equally between the electron and positron.

This simple model leads to the following conclusions:

- The number of particles is $2^t$ after a normalized depth $t = X/R$
2.1. ATOMIC SHOWERS

![Diagram of shower development](image)

- The depth of maximum of shower development is $X_{\text{max}} = X_0 \ln \frac{E_0}{E_c}$, which corresponds to $t_{\text{max}} = \ln \frac{E_0}{E_c} / \ln 2$, where $E_c$ is the critical energy at which ionization losses become important.

For an hydrostatic atmosphere, the pressure and density as a function of altitude are $P = P_0 \exp(-z/z_0)$ and $\rho = \rho_0 \exp(-z/z_0)$ with $z_0 = RT/gM = 8.4$ km. The depth as a function of altitude is therefore:

$$X(z) = \int_{z_0}^{\infty} \rho(z) dz = \rho_0 z_0 \exp(-z/z_0), \quad \frac{dX}{dz} = -\rho(z) = \rho_0 \exp(-z/z_0)$$

(2.4)

with $\rho_0 = 1.2 \text{ kg/m}^3$. The altitude of maximum of shower development is therefore:

$$z_{\text{max}} = z_0 \ln \left( \frac{\rho_0 z_0}{X_0 \ln E_0 / E_c} \right)$$

(2.5)

leading to a slow evolution with energy. At 1 TeV, the depth of maximum is $t_{\text{max}} = 13.5$ leading to

$$z = -z_0 \ln \left( \frac{t_{\text{max}} X_0 \ln 2}{\rho_0 z_0} \right) \approx 9 \text{ km}$$

(2.6)

2.1.4 Cherenkov Emission

The wavelength distribution of the Cherenkov radiation is given by:

$$\frac{d^2 N}{dxd\lambda} = 2\pi \alpha \frac{\sin^2 \theta}{\lambda^2} \quad \text{with} \quad \cos \theta = \frac{c}{nv} \approx \frac{1}{n}$$

(2.7)
and the index of refraction of air is primarily a function of pressure:

\[(n - 1) = 2.92 \times 10^{-4} \times \frac{P}{P_0} \times \frac{288.15 \text{ K}}{T}\]  \hspace{1cm} (2.8)

From that the energy spectrum of produced photons follows:

\[\frac{d^2N}{dxd\lambda} \approx A \frac{\exp(-z/z_0)}{\lambda^2}, \quad \text{with} \quad A = 2.68 \times 10^{-5} \text{ photons}\]  \hspace{1cm} (2.9)

After integration over the altitude of the emission, the number of Cherenkov photons above the wavelength \(\lambda_0\) is found to be directly proportional to the primary energy.

\[N(\lambda > \lambda_0) = \frac{2X_0 A \ln 2}{\rho_0 \lambda_0} \int N(t)dt = \frac{2X_0 A}{\rho_0 \lambda_0} \exp(t_{\text{max}} \ln 2) = \frac{2X_0 A}{\rho_0 \lambda_0} \times \frac{E_0}{E_c}\]  \hspace{1cm} (2.10)

For \(\lambda_0 = 300\text{ nm}\), one finds \(N(\lambda > 300\text{ nm}) \approx 5450 \left(\frac{E_0}{E_c}\right)\).

The Heitler model is of course oversimplified, but reproduces the basic features of shower development. It’s interesting to note that the increase of Cherenkov emission at low altitudes is exactly compensated by the larger density, so that the Cherenkov yield is constant per unit depth.

Cherenkov radiation suffers significant absorption in the atmosphere. Ultraviolet emission below 300 nm is almost completely absorbed. The low altitude layers of the atmosphere contribute the most to the absorption. Using a simple atmospheric model based on Rayleigh diffusion theory, one can estimate the transmission of the atmosphere as a function of wavelength. Above 300 nm, the average transmission is of the order of 75%, but the ultraviolet emission is fully absorbed.

**Radius of emission**

The radius of Cherenkov emission (intersection of the Cherenkov light pool produced by a particle at height \(z\) with the ground) is simply:

\[r \approx z\theta = z\sqrt{5.84 \times 10^{-4} \exp(-z/z_0)} = 2.4 \times 10^{-2} \times z \exp(-z/2z_0)\]  \hspace{1cm} (2.11)

This radius is shown in Fig. 2.4 as a function of altitude of emission. In a radius range of \(\sim 130 - 170\text{ m}\), the relation between radius and altitude of emission is degenerated: each radial distance corresponds to the superposition of an emission at high altitude with an emission at low altitude. Close to the shower axis, only the low altitude emission contributes to the light pool. The maximum extension of the light-pool is obtained for \(z = 2z_0\), corresponding to a radius:

\[r_{\text{max}} = 2.4 \times 10^{-2} \times 2z_0/e \approx 150\text{ m}\]  \hspace{1cm} (2.12)

**Lateral Profile**

Radii around 150 m correspond to the emission in a large range of altitudes, close to the shower maximum (from roughly 10 to 22 km) thus leading to an intense emission. Lateral intensity profiles are shown at the sea level for various primary energies in Fig. 2.5.
Figure 2.4: Radius of Cherenkov light pool on the ground as a function of altitude of emission.

Figure 2.5: Lateral profile of showers of different energies at sea level (left) and at an altitude of 1600 m (right).

Lateral profiles are characterized by an intense plateau emission up to distances of the order of 150 m, followed by a slowly decreasing tail that is caused by multiple scattering of electrons in the shower development (thus leading to a angular widening of the shower). At higher energies, or for instruments at higher altitudes (Fig. 2.5 right), charged particles reaching the ground produce a very peaked emission close to the shower axis.

Multiple Coulomb scattering, responsible for the widening of the shower, is modelled in Molière theory by a average angle:

$$\theta_0 = \frac{13.6 \text{ MeV}}{\beta c p} z \sqrt{x/X_0} \left[1 + 0.038 \ln(x/X_0)\right] \quad (2.13)$$

where $z$ is the particle charge. For GeV energies, the mean scattering angle is $1.7^\circ$ after 5 radiation lengths, compared to a Cherenkov angle $\theta_c \approx 1.3^\circ \exp(-z/2z_0)$. Multiple Coulomb scattering significantly spreads the Cherenkov emission, but do not destroy the spatial structures.
Temporal aspects

The shower development is illustrated in Fig. 2.7. The Cherenkov angle reads \( \theta_c \approx 1.3^\circ \exp(-z/2z_0) \). At an impact distance \( r \), two different effects compete:

- A delay related to the refraction index of the air, which makes the light propagate slower than relativistic charged particles (reason for Cherenkov emission)

- The geometric delay due to the inclination of the track.

The time difference between the projected arrival time of the primary on the ground and the photons can be estimated simply as a function of altitude of emission \( z_e \) and angle \( \alpha = x/z_e \):

\[
\Delta t \approx \frac{x^2}{2cz_e} + 2.92 \times 10^{-4} \frac{z_0}{c} (1 - \exp(-z_e/z_0))
\]

(2.14)

![Figure 2.6: Time delay as a function of lateral distance for various altitudes of emission.](image)

The time delay as a function of impact distance, for different altitudes of emission, is shown in Fig. 2.6. Close to the shower axis, photons emitted at low altitude reach the ground first, whereas at distances larger than 150 m, geometric effects dominate and photons emitted at high altitude arrive first. The time spread of the signal is minimal at a distance of the order of 150 m where the total signal duration is less than 2 ns.

2.1.5 Approximation A

The Approximation A is a further, and more realistic development of the initial Heitler model. The processes considered are Bremsstrahlung and pair creation, as in Heitler model, but using coupled integral-differential equations (Bethe-Heitler Formulas) instead of a simple assumption. As in the Heitler model, the lack of any energy scale in microscopic processes leads to simple steady state solutions with separated variables which can be used as a base of solutions on which any particular initial condition can be developed.
Figure 2.7: Generation of Cherenkov Image.
2.1.6 Semi-empirical model (Greisen)

Following the development of the Approximation A, Greisen proposed a semi-empirical model that, in particular, takes into account ionization losses which where neglected in the previous models.

One defines a *shower age* parameter, independent of the particle energy, which depends on the critical energy $E_c$:

$$ s = \frac{3t}{t + 2y} \quad \text{with} \quad y = \ln \left( \frac{E_0}{E_c} \right) $$

(2.15)

The age is $s = 1$ at the depth of shower maximum $t_{max} = y = \ln(E_0/E_c)$ and $s > 1$ in the following extinction phase. The semi-empirical Greisen formula gives average number of electrons at depth $t$:

$$ \overline{N_e}(t) = \frac{0.31}{\sqrt{y}} \exp \left[ t \left( 1 - \frac{3}{2} \ln s \right) \right] $$

(2.16)

At the shower maximum, the number of particles is almost proportional to energy:

$$ \overline{N_e}(t_{max}) = \frac{0.31}{\sqrt{y}} \left( \frac{E_0}{E_c} \right) $$

(2.17)

![Graph](image)

Figure 2.8: Shower development versus primary energy. Maximum of development corresponds to $s = 1$.

Orders of magnitude for the number of particles and altitude of maximum developments are given in table 2.1 and shown in Fig. 2.8.
2.2 Detection techniques

The Cherenkov emission for electromagnetic showers illuminates an area of the order of $10^5 \text{ m}^2$, thus allowing huge effective areas even with modest size detectors. The main experimental challenges are, on the one hand, the very low intensity and short duration of the signal, leading to the requirement of very fast acquisition systems (with typical integration times of $\sim 1 \text{ ns}$ and, on the other hand, the huge background, both from night sky background light and from air showers produced by charged cosmic rays. In addition, both the spatial and angular distributions of Cherenkov photons reveal some information about the shower development, but it is impossible to measure both at the same time with a single detector. Rejections factors of the order of $10^2 - 10^4$ are required to suppress the hadronic background. Fine pixelization and stereoscopy proved to be the the most efficient way to reduce the cosmic-ray background, but this has only been possible by using small field-of-view instruments.

Several approaches have been followed to investigate very high energy $\gamma$ ray emission from astrophysical sources.

2.2.1 Sampling Experiments

Sampling experiments aim to make a dense spatial and temporal sampling of the shower front. The spatial distribution of Cherenkov photons as a function of time is measured, and is used to discriminate between $\gamma$ rays and hadronic cosmic rays. A simple trigonometric approach is in general used to reconstruct the shower direction. Different kind of sampling experiment have been tested:

Solar farms

Following first sampling experiments such as ASGAT [31] and THEMISTOCLE [32], several projects based on the re-use of former solar farms have been proposed [33, 34, 35, 36].

Using existing infrastructure, experiments such as CELESTE (Fig. 2.1) managed to develop experiments with very large collection areas in the range $1000 - 4000 \text{ m}^2$. Using a distributed trigger over several heliostats¹ a trigger threshold as low as $\sim 20 \text{ GeV}$ was achieved. Secondary optics were used to focus the image of each heliostats onto a single photo-multiplier, equipped with a glass Winston Cone to limit the field of view.

These resulted into a few detections in the domain between $\sim 10 \text{ GeV}$ and $100 \text{ GeV}$ [e.g. 37, 38, 39, 40, 41] that was almost unexplored at that time. However, the constrain of secondary optics located at a distance of $\sim 100 \text{ m}$ implied a very small field of view, of

<table>
<thead>
<tr>
<th>$E_0$</th>
<th>$T_{max}(\text{gU/cm}^2)$</th>
<th>Altitude (m)</th>
<th>$N_e(t_{max})$</th>
</tr>
</thead>
<tbody>
<tr>
<td>30 GeV</td>
<td>216</td>
<td>12000</td>
<td>50</td>
</tr>
<tr>
<td>1 TeV</td>
<td>345</td>
<td>8000</td>
<td>1200</td>
</tr>
<tr>
<td>1000 TeV</td>
<td>600</td>
<td>4400</td>
<td>0.9 $\times 10^6$</td>
</tr>
<tr>
<td>$10^{19}$ eV</td>
<td>936</td>
<td>1200</td>
<td>7.4 $\times 10^6$</td>
</tr>
</tbody>
</table>

Table 2.1: Orders of magnitude of shower development for different primary energies

¹Heliostats are large mirrors of $\sim 50 \text{ m}^2$ which were originally used to focus the light from the sun on the top of the tower, where a boiler was installed
the order of $\sim 0.5^\circ$, on each heliostat. Although this was a major help in reducing the threshold, it also made it almost impossible to discriminate between $\gamma$ ray showers and hadronic induced ones, the small field of view making the electromagnetic sub-showers in hadronic shower resemble genuine $\gamma$ ray-induced showers.

**Wide-field sampling experiments**

Several approaches have been used to develop wide-field sampling experiments, based on the detection of the charged particles in the shower tail instead of the emitted Cherenkov light. The advantages are two-fold:

1. a 24/7 duty cycle, compared to about 10% of the time for Cherenkov experiments, due to the fact that charged particles can be detected on any weather and luminosity conditions.

2. a $\sim 2\pi$ sr field of view, compared to $\sim 0.05$ sr for solar farms, making these technology very suitable for survey instruments.

Two approaches have been proved to be successful so far:
2.2. DETECTION TECHNIQUES

Water Cherenkov The Milagro [12] experiment was a water Cherenkov detector (Fig. 2.10) at an altitude of 2630 m, and composed of a central 60 m × 80 m pond with a sparse 200 m × 200 m array of 175 “outrigger” water tanks surrounding it. The pond is instrumented with two layers of photo-multiplier tubes, the top one being dedicated to the measurement of the electromagnetic component in showers, and the bottom, located 6 m below the surface, dedicated to the identification of hadronic showers through their muonic content.

Figure 2.10: Sketch of the central Milagro water pond detector.

The top layer allows the accurate measurement of shower particle arrival times which are then used for direction reconstruction and triggering. An outrigger array, added in 2003, improved the angular resolution of the detector by providing a longer lever arm with which to reconstruct events. Milagro’s large field of view (2 sr) and high duty cycle (> 90%) allow it to monitor the entire overhead sky continuously, making it well suited to measuring diffuse emission.

This is illustrated by the view of the Galactic Plane by Milagro (Fig. 2.11) above ∼ 20 TeV, representing 2358 days of data collected by Milagro between July 2000 and January 2007.

Four sources (including the Crab) are detected with post-trial significances greater than 4 σ, and four additional lower significance candidates are identified. Six of these eight TeV excesses are coincident with the locations of EGRET sources.

The next generation water pond experiment, HAWC [44, 45], under construction at Sierra Negra, Puebla, Mexico, is expected to gain more than one order of magnitude in sensitivity compared to Milagro.

Other particle detectors Another possible approach is to cover a large area with cheap and easy to handle particle detectors, such as resistive plate chambers. This approach has been followed by the ARGO-YBJ experiment [46], an extensive air shower detector which combines a high altitude location (Tibet) and a fully instrumented coverage using resistive plate counters.

This results in an energy threshold of a few hundred GeV, for a large field of view (∼ 2 sr) and high duty cycle, thus allowing continuous monitoring of the Northern sky.

Here again, precise timing is used to reconstruct the shower direction. The distribution of charged particles on the ground provides hadronic discrimination, however, at a level that still severely limits the sensitivity of the instrument. Results of ARGO-YBJ include in particular the measurement of the moon shadow in cosmic rays [47] (allowing a validation of the reconstruction algorithms) and the monitoring of strong VHE sources such as the blazer Markarian 421 [48].
Figure 2.11: Significance map of the galactic emission as seen by the Milagro Water Cherenkov experiment. The color code shows the pretrial significance, smoothed according to the instrument PSF. From [13].

2.2.2 Atmospheric Imaging Cherenkov Telescopes

In the Imaging Atmospheric Cherenkov Telescope technique (Fig. 2.12), initiated by the Whipple group [13], the short duration Cherenkov flash is collected by large optical telescopes (of the order of 100 m²) and detected by fast cameras at the focal plane of each. This approach provides effective area of the order of 10⁵ m² even with modest-sized telescopes. The energy threshold of the instrument depends on the collection area of the telescope \( A \), the solid angle of the photo-detectors \( \Omega \), the integration time \( \Delta t \), the average luminosity of the night \( \phi_{NSB} \) (expressed in terms of photo-electron rate per unit area and solid angle) and the global efficiency of the system \( \epsilon \), and can be expressed according to the scaling law:

\[
E_{th} \propto \sqrt{\frac{\Omega \Delta t \phi_{NSB}}{A \epsilon}}
\]  

(2.18)

The light distribution ("image") in the camera is used to reconstruct the parameters of the impinging gamma ray (energy, direction, impact,...) and also to discriminate the showers initiated by gamma rays from those initiated by the much more numerous hadrons (mainly protons, nuclei and alpha particles). In the reconstruction and analysis
Figure 2.12: Imaging Atmospheric Cherenkov technique. **Left:** the Cherenkov light emitted by the charged particles in the shower is collected by several dishes. **Right:** The shower angular image is projected into the camera focal plane.

techniques based on the so-called “Hillas parametrization”, shower images in the camera are parametrized by an elliptical shape and reduced to a few numbers. Discrimination is then based on the comparison of image width with expectations from simulations.

The CAT experiment [49], in the French Pyrenees, introduced in 1998 the concept of fine camera pixilation assorted with a new, more powerful, analysis technique [50], based on the adjustment of a semi-analytic model on the recorded images.

At roughly the same time the HEGRA collaboration [21] used stereoscopic observation to provide a geometrical reconstruction of shower direction and impact together with improved background rejection.

These developments, done in parallel, have then been merged in third generation instruments such as H.E.S.S., MAGIC and VERITAS.

Tables 2.2 compares the parameters of the early, second generation instruments (Whipple, HEGRA and CAT) with the third generation of instruments currently in operation.

**Conclusion**

Very-high-energy γ-ray astronomy is a very recent, and very active field of astronomy, that has led to the opening of a new window on the non-thermal Universe. Recent improvements are largely illustrated by the rate of new detections, as illustrated in Fig. 2.13, which indicates the accumulated number of detected sources as a function of time.

The current VHE sky is shown in Fig. 2.14 and now comprises more than 100 sources, belonging to the classes of pulsar wind nebulae, active galactic nuclei, supernova remnant, binary systems, starburst galaxies and unidentified sources.

H.E.S.S. was awarded the prestigious Descartes Prize for 2006, in recognition of the
Table 2.2: Comparison of existing Atmospheric Cherenkov Telescopes. From left to right, columns of the table respectively give the instrument name, its location, its altitude, the number of telescopes, the surface area of the telescope reflector, the total reflecting surface area, the number of pixels per telescope, the field of view and the threshold energy.

<table>
<thead>
<tr>
<th>Instrument</th>
<th>Site</th>
<th>Alt. (m)</th>
<th>$N_{Tel}$</th>
<th>$S_{Tel}$ (m$^2$)</th>
<th>$S_{Tot}$ (m$^2$)</th>
<th>$N_{pix}$</th>
<th>FOV (°)</th>
<th>$E_{th}$ (GeV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>H.E.S.S.</td>
<td>Namibia (S)</td>
<td>1800</td>
<td>4</td>
<td>107</td>
<td>428</td>
<td>960</td>
<td>5</td>
<td>0.1</td>
</tr>
<tr>
<td>VERITAS</td>
<td>Mont Hopkins (N)</td>
<td>1275</td>
<td>4</td>
<td>106</td>
<td>424</td>
<td>299</td>
<td>3.5</td>
<td>0.1</td>
</tr>
<tr>
<td>MAGIC</td>
<td>La Palma (N)</td>
<td>2225</td>
<td>2</td>
<td>234</td>
<td>468</td>
<td>574</td>
<td>3.5</td>
<td>0.06</td>
</tr>
<tr>
<td>Whipple</td>
<td>Mont Hopkins (N)</td>
<td>2300</td>
<td>1</td>
<td>75</td>
<td>75</td>
<td>379</td>
<td>2.3</td>
<td>0.3</td>
</tr>
<tr>
<td>HEGRA</td>
<td>La Palma (N)</td>
<td>2200</td>
<td>5</td>
<td>9</td>
<td>43</td>
<td>271</td>
<td>4.3</td>
<td>0.5</td>
</tr>
<tr>
<td>CAT</td>
<td>Targassone (N)</td>
<td>1650</td>
<td>1</td>
<td>18</td>
<td>18</td>
<td>600</td>
<td>4.8</td>
<td>0.25</td>
</tr>
</tbody>
</table>

Figure 2.13: Number of detected VHE Galactic sources as a function of publication time. The number of Galactic sources has increased tremendously by steps over the last years, particularly through the Galactic Plane Surveys (GPS, EGPS standing for the extension of the GPS) conducted with H.E.S.S.. From [31].

quality of results concerning the “non-thermal universe” that established a new field of astronomy and was awarded the Rossi Prize in 2010.
Figure 2.14: Known VHE sources as of July 2011 from the TeVCat catalog, plotted in Galactic coordinates. The colored regions show the accessible sky from the H.E.S.S. (red) and VERITAS/MAGIC (blue) sites. Point color identifies sources type: pulsar wind nebulae (magenta), AGN (red), SNR (green), binaries (yellow), star-burst (orange), other identified (blue), unidentified (grey). The proposed CTA southern sites (Argentina, Namibia) cover almost the same region of the sky as H.E.S.S.. The blind spots correspond to zenith angles > 50°. From [52].
Chapter 3

The H.E.S.S. experiment

The High Energy Stereoscopic System (H.E.S.S.) is an array of four imaging atmospheric Cherenkov telescopes located in the Khomas Highland of Namibia, 1800 m above sea level, and dedicated to very high energy $\gamma$-ray astronomy between 100 GeV and 100 TeV. The telescopes are identical in construction and arranged in a square of 120 m length. Each telescope is comprised of a 107 m$^2$ optical reflector composed of segmented spherical mirrors. These focus the incident light into a camera built of 960 photo-multiplier tubes (PMTs). H.E.S.S. was built by an international collaboration of 35 different institutions from 12 countries. Observations were started in 2001 with the completion of the first telescope. The full array was completed in December 2003 and is taking data on a regular basis since then.

The site location, in the Khomas Highland of Namibia, was chosen because of the quality of the sky (low back-scattered background luminosity thanks to the absence of any nearby city, dry atmosphere and low cloud coverage), but also because of its position close to the tropic of Capricorn allowing optimal observation of the inner regions of the Galactic Plane.

![Photograph of the array of H.E.S.S. telescopes.](image)
3.1 Telescopes

The structure of the telescopes (Fig. 3.2) is made of steel with alt-azimuthal mount allowing observation in any position of the sky. Pointing of the telescopes in any direction of the sky can be achieved in less than 2 minutes in order to allow the observation of transient phenomena.

![Figure 3.2: Left: Structure of the H.E.S.S. Telescopes. Right: Principle of a Davis-Cotton mount (see text).](image)

The dish, of 13 m diameter and 107 m² area, is composed of 380 panels of 60 cm diameter each, arranged on a Davis-Cotton mount: panels of focal length 15 m are arranged on a spherical surface of focal length of 7.5 m, half of the focal length of the facets (Fig. 3.2 right). This mount reduces the coma aberration of the telescope at the expense of an anisochronism that varies quadratically with the distance to the dish center, and reaches 5 ns on the edge of the dish. The time difference between rays falling at the center of the dish and at a distance \( r \) is simply given by

\[
c\Delta t = f(1 - \cos \alpha) \quad \text{with} \quad \sin \alpha = r/f
\]

This can be simplified into

\[
\Delta t = \frac{r^2}{2cf} - \frac{1}{8 cf^3} + O\left(\frac{r^6}{cf^5}\right)
\]

In the case of H.E.S.S., \( r_{\text{max}} = 6.5 \text{ m} \), \( f = 15 \text{ m} \) and thus \( \Delta t_{\text{max}} = 4.5 \text{ ns} \).

Focusing for on-axis rays (blue lines) are unchanged with respect to pure parabolic optics. The focal point for inclined rays (red lines) is given by the focal length of the facets,
not the focal point of the structure. There are always facets that are on-axis with respect to the inclined rays, which reduces the amount of comatic aberrations. Davis Cotton mounts are not possible with larger telescopes as the induced anisochronism would widen significantly the signal from the showers and thus reduce the signal over background ratio.

The mirrors facets are made of glass aluminized in front surface, and covered by a protective quartz layer. There reflectivity has been measured to be around 70% in the wavelength domain of the Cherenkov emission, with an intrinsic point spread function of 0.45 mrad. Facets are bound to the telescope dish with a mechanical structure comprising three actuators, which allow a precise alignment of each of the facets. The alignment has been performed once after the installation of the telescopes, and was redone after mirror refurbishment. The procedure involves the tracking of a bright star and measurement of the image of the star on the focal plane with a CCD placed at the center of the dish.

3.2 Camera

3.2.1 General description

The cameras of the H.E.S.S. telescopes [53] (Fig. 3.3) combine a fine pixelization (for optimal imaging of the shower image) with fast electronics (for optimal background suppression). They comprise 960 photo-multipliers arranged in drawers of 16 PMT each, power supplies and the full trigger, readout and slow control electronics. The cameras are connected to the ground only with power cables and optical fibers, weigh 900 kg and have a width of 1.6 m.

Figure 3.3: Photograph of one of the H.E.S.S. cameras taken during the installation of the Winston cones.

Each PMT has a field of view of 0.16° on the sky. The H.E.S.S. cameras are characterized by a wide field of view of 5° diameter compared to other Atmospheric Cherenkov Telescopes. Winston cones, placed in front of each PMT, limit their field of view to the actual size of the mirror and therefore reduce the albedo noise from the ground. They also reduce the effect of the dead space between the PMTs by guiding the light onto the photo-cathode.
3.2.2 Front End Electronics

The camera comprises 60 structures called Drawers, which are the base units of the front-end electronics. A Drawer (Fig. 3.4) contains the acquisition and control electronics for a cluster of 16 PMTs, as well as a fraction of the trigger and high voltage generation electronics. Drawers are connected to the main acquisition and control crates (located in the back-plane of the camera) through a custom designed data bus (Box-bus). Each drawer is comprised of one slow control board and two analogue boards, dedicated to readout, and connected to eight PMTs each, as depicted on Fig. 3.5.

![Figure 3.4: Photograph of one of the H.E.S.S. drawers](image)

![Figure 3.5: Sketch of the H.E.S.S. camera electronics. Adapted from [54]](image)

The signal from each PMT is split into one trigger channel and two sampling channels. The two sampling channels have different amplification factors in order to increase the dynamical range of the instrument. The so called high gain channel has a dynamical range of $[1 - 150]$ photo-electrons with single photo-electron resolution. The other channel (low gain) has a gain 13.5 times lower than the high gain and covers the range between 20 and $\sim 2000$ photo-electrons. After amplification, the signal is sent to an analogue memory,
which acts as a circular buffer able to continuously sample the signal at the frequency of 1 GHz and store the last 128 samples of signal (Fig. 3.6). Each analogue memory comprises 5 independent channels, out of which only 4 are used.

![Photograph of the ARS chip](image)

**Figure 3.6: left:** Photograph of the ARS chip. **right:** Principle of the ARS Readout

The signal is also sent, through the third channel, to the trigger system, which is designed to identify spatial structures in the camera that could correspond to image of showers, and to reduce the rate of accidental events induced by the night sky background. When a trigger decision occurs, the analogue memories are stopped, and the signal from read in the analogue memories and converted to ADC counts. The position of the signal in the analogue memories (with respect to the stopping position) corresponds to the time needed by the trigger system to make its decision, plus the transit time of the signal. This time, called \( Nd \), is calibrated precisely using pulsed signals.

![Data flow in the camera of the H.E.S.S. experiment](image)

**Figure 3.7: Data flow in the camera of the H.E.S.S. experiment.**

As depicted in Fig. 3.7, each drawer contains two analogue cards, each of them handling the signal from a sub-group of eight PMTs. Four analogue memories are installed on each
analogue card, each of them handling either the high or the low gains for a line of four pixels. High and low gains are handled with different analogue memories in order to reduce the risk of cross talk.

Once a camera trigger decision is made, the signal in the analogue memories is read out and transferred into a memory buffer inside a FPGA located on the board. The conversion in the ARS takes 275 μs. Data are then transferred to the camera CPU using the box bus which works using a token mechanism, each drawer sending their data one after the other. Among other tasks, the slow control handles the communication with the box bus, as well as configuration and monitoring of the drawer. Temperatures are measured on a regular basis at three different positions in the drawer, high voltage and anode current are regularly monitored as well. The slow control also controls the programming of pixel high voltage and trigger threshold.

3.2.3 Trigger

The trigger system of the H.E.S.S. array of telescopes is a dual level system. The first level, or local level, is internal to each camera, and exploits the correlation between neighbouring pixels to reduce the influence of the night sky background. The cameras are divided into 38 overlapping sectors of 64 pixels each. The trigger decision is made when the signal in a number \( S_2 \) of pixels inside the same sector exceeds a threshold \( S_1 \), within an gate of \( \sim 1.3 \) ns. In normal operations, the thresholds are respectively \( S_1 = 4 \) p.e. and \( S_2 = 3 \) pixels, leading to a single camera trigger rate of \( \sim 1.4 \) kHz.

In a second stage, coincidences between telescopes are taken into account to reduce the triggering rate cause by isolated muons or localized showers, but also to ensure stereoscopic view of the showers that greatly improves the reconstruction precision and the rejection capabilities. When a camera triggers, it sends a signal to a central facility (central trigger) which delays the signal from the different telescopes according to the pointing conditions and makes a trigger decision based on the number of telescopes having triggered on the same shower. The details of the central trigger operations are described in section 4.2.

3.3 Environment

In addition to the major components of the telescope array, several smaller instruments are installed in the H.E.S.S. array to allow for proper monitoring of the atmosphere and of the instrument behaviour.

3.3.1 Flatfielding LED

A flat-fielding system composed of a flashing LED and an isotropic diffuser, is installed at the center of the dish. It is used to cross-calibrate the response (quantum efficiency multiplied by collection efficiency) of the PMTs.

3.3.2 Atmospheric monitoring

A proper monitoring is required for array operation as well as for reduction of systematic effects at the analysis level. The role of the atmospheric monitoring are:
• Provide a decision help for the shift crew regarding the quality of the sky and the safety of observation. Monitoring of cloud coverage and wind speed allows observer to decide to stop observations and put the camera in a safe state before rain occurs. In order to be efficient, this requires a full sky monitoring.

• Give quantitative criteria to estimate the quality of the sky and to decide which data are hampered by bad quality weather and which one can be safely used. This includes the continuous monitoring of the transparency of the atmosphere (using a ceilometer) and the monitoring of the cloud coverage in the field of view of the telescopes.

• Provide quantitative measurement of the atmospheric properties that could be used offline, possibly with simulations, to reduce the reconstruction systematics and improve the precision of the measurement. In this regard, a precise measurement of the atmosphere composition as a function of altitude (with a Raman Lidar) could be very useful, although measurements have not been used in the data analysis to date.

The atmosphere monitoring equipment includes:

• One radiometer per telescope, installed on the telescope mount, and pointing at the same field of view as the telescope. These radiometers measure the temperature of the sky, which depends on the humidity level and provide monitoring of clouds crossing the telescope field of view, however they provide no information on the altitude of these clouds. Cloud coverage in the field of view is also very well correlated with the trigger rate.

• A scanning radiometer, making full sky images of the cloud coverage

• A weather station, making regular measurements of temperature, pressure, humidity, wind speed and direction.

• A Lidar, making measurements of the aerosol composition of the atmosphere as a function of altitude.

• A setup consisting of LEDs installed on the Gamsberg mountain, 30 km away from the site and 500 m higher in altitude and a detector installed on the site is used to monitor the transparency of the ~ 500 m lower part of the atmosphere.
Chapter 4

Data Acquisition Software

H.E.S.S. is an heterogeneous system, consisting in particular of four telescopes with several subsystems that must be controlled and read out.

The telescope subsystems comprise a camera with 960 individual photo-multiplier tubes, light pulser systems for calibration purposes, a source tracking system, an infra-red radiometer for atmosphere monitoring, and a CCD system for measuring pointing offsets. Common to the whole array is a set of devices for the monitoring of atmospheric conditions, including a weather station, a ceilometer and an all-sky radiometer.

In addition to its primary task of controlling and coordinating these subsystem to read out event data, the data acquisition system must also provide high level capabilities such as online data calibration and analysis, exhaustive monitoring of the instrument behaviour and health as well as react to worldwide scientific alerts such as those sent for instance for Gamma-Ray bursts.

The data acquisition must therefore be based on a flexible framework that allows communication between very different systems, running on different hardware architectures and running different operating system. The system must also allow easy inclusion and removal of new hardware without major reconfiguration.

4.1 General Software Architecture

The data acquisition software is based on a multi-processor, multi-process, multi-threaded architecture. More than 100 different processes spread over more than 20 machines running different operating systems and written with different languages cooperate to achieve the data acquisition strategy. All processes are based on the same state-machine structure, and are configured using a central database.

This ensures the flexibility of the system, as new processes corresponding to new hardware or new software tasks can be added to the system without modification of the existing structure. The relations between the processes, and in particular the dependencies (i.e. which processes are required, for instance, to perform a given task, and in which order they need to be started and configured) are dynamically configured using databases. The central database also specifies which components are required for each run type (for instance, light pulsers are required for calibration purpose but not for standard acquisition). The runs that can be performed are selected depending on the available resources.

\[1\text{Currently, the systems comprises 5 servers, 10 nodes in the DAQ cluster, 5 PCs in the control room plus the PCs in the electronic huts and cameras}\]
4.1.1 Inter-process communication

The inter-process communication (IPC) is based on the CORBA (Common Object Request Broker Architecture) protocol which was created in 1992 by the Object Management Group [23], a group initiated by several computer manufacturers and software developers including Sun, Oracle, and IBM.

CORBA is based on a fully object oriented client-server architecture which allows to call methods of objects in remote processes and to pass data objects.

The operations are schematically depicted in Fig. 4.1. On the client side, a calling object, written in a programming language \( A \), obtains a pointer to a stub object (also sometimes called proxy) and calls the corresponding function \( foo \) using its natural language. The function call and eventually the arguments are converted by the CORBA layer into a CORBA message that is propagated to the server side. On the server side, the CORBA layer translates the message back to a function call of the skeleton (using language \( B \)), which in turn actually calls the function of the instance object\(^2\). The result of the function call (with possible read-write parameters) are then propagated back to the stub, and thus to the calling object.

![Figure 4.1: CORBA Inter-Process communication layer principles.](image)

Each object can expose some functions using a dedicated programming language. A first compilation pass generates the stub (shadow class, located on the client side), and the skeleton from which the actual object instance \( \text{interface} \) must derive. A call to a function \( foo \) on the client side is propagated by the CORBA layer to the server side where the actual function is executed. The answer of the call is propagated back (as well as, for instance, exception calls) to the client side.

A central directory \( \text{name service} \) permits object registration (using a hierarchical naming tree) and allows any object to easily obtain pointer to stub objects.

4.1.2 State machine

In order to standardize the operations on the controller, a state machine structure has been set up: each process, whether it is a hardware controller, a display manager, an element of the calibration software or any other element, is build upon a machine state. The states and corresponding transitions between them are shown in Fig. 4.2.

\(^2\)The foo function of the skeleton is made purely virtual
Figure 4.2: Machine states and transitions for the H.E.S.S. acquisition building blocks.

- **Safe**: the hardware is not configured and is in a secured state (for instance, high voltage turned off on the camera, telescope parked, ...)
- **Ready**: the hardware is online, and a minimal configuration has been performed. For instance, for the camera, this state corresponds to power switched on, cooling in operation and slow control running.
- **Configured**: the hardware has been configured appropriately for the next data taking task. For instance, for the camera, the high voltage is set on the photo-multipliers, the trigger is configured, and all electronic components are ready
- **Running**: the hardware is taking data

The corresponding transitions are:

- **GetReady**: Initialize the hardware
- **Configure**: Configure the hardware according to the parameters of the next data taking
- **Start**: Run data taking and processing
- **Stop**: Stop data taking and processing
- **UnConfigure**: Terminate the data taking and go to a state where a new acquisition (using eventually different parameters) can be configured.
- **GotoSafe**: Send the hardware to a safe state and disconnect

The **GetReady** transition is typically executed only once per night. Normal run operation usually implies the following transitions: **Configure, Start**, wait 28 mn, then **Stop, UnConfigure**. The next run would then follow the same sequence: **Configure, Start, ...**

### 4.1.3 Messaging system

A centralized, standardized messaging system has been set up to allow any process to communicate with any other processes, based on their name in the central naming service. A registration mechanism allows any process to request the message from any other processes. The messages are then send using a push paradigm. Several message types exist:
<table>
<thead>
<tr>
<th>Process</th>
<th>Safe</th>
<th>Ready</th>
<th>Configured</th>
<th>Running</th>
</tr>
</thead>
<tbody>
<tr>
<td>Tracking</td>
<td>Telescope Parked</td>
<td>Telescope Parked Out</td>
<td>On Target</td>
<td>On Target</td>
</tr>
<tr>
<td>Camera</td>
<td>Power OFF</td>
<td>Power ON, cooling ON</td>
<td>Drawer Configured</td>
<td>Reading Data</td>
</tr>
<tr>
<td>Camera HV</td>
<td>HV OFF</td>
<td>HV ON, 800 V</td>
<td>HV ON, Nominal Values</td>
<td>HV ON, Nominal Values</td>
</tr>
<tr>
<td>Camera LID</td>
<td>Lid Closed</td>
<td>Lid Opened</td>
<td>Lid Opened</td>
<td>Lid Opened</td>
</tr>
<tr>
<td>Camera Trigger</td>
<td>Trigger OFF</td>
<td>Trigger OFF</td>
<td>Trigger input configured, not running</td>
<td>Trigger Running</td>
</tr>
<tr>
<td>Node Receiver</td>
<td>Input buffer</td>
<td>Camera configuration saved in memory</td>
<td>Receiving data, saving to file</td>
<td></td>
</tr>
</tbody>
</table>

Table 4.1: Matrix of machine states for some important processes in the data acquisition

![Diagram](image)

Figure 4.3: Messaging system in H.E.S.S.

- **Print**: Simple message, printed in the log files of the process. Not propagated to any other process.
- **Information**: Displayed on the central display, forwarded to other processes.
- **Caution**: Forwarded to other processes. Generates a short click noise audible to the observers.
- **Warning**: Forwarded to other processes and displayed on the central display, requesting an action from the observers.
- **Error**: Forwarded to other processes and displayed on the central display, requesting an action from the observers. The acquisition is immediately stopped (if one was running) or forbidden (if not running) until the problem is solved, and the DAQ initiates a transition towards the Ready state.
- **Fatal Error**: Emergency procedure. Forwarded to other processes and displayed on the central display. The acquisition is immediately stopped, and the DAQ initiates a transition towards the Safe state. Telescopes are blocked for security reasons,
high voltages are switched off immediately, lid closed, and so on, to ensure maximal
security.

4.1.4 Data storage and access

The transport and storage of objects needs a serialization mechanism which is provided by
the ROOT Data Analysis Framework [36]. Both the ROOT-based H.E.S.S. data format
and the ROOT graphics and histogram classes are used online and offline allowing a
seamless integration of data analysis code and hence fast feedback.

Data is stored in a collection of ROOT files, using data classes that are specific to the
experiment. Data based on the ROOT format can also be transported from one process
to another using the CORBA IPC. The same software architecture is used for on-line
and off-line software. As a consequence, any algorithm (reconstruction, calibration ...) developed for off-line analysis can easily be incorporated into the online data acquisition
software, thus allowing on-the-fly data calibration or analysis. Moreover, the introspection
capabilities of ROOT allows to access any member of any object without any modification
to the code, since any member function of any object can be called with the help of the
C++ interpreter. The online display system of H.E.S.S. is therefore able to display any
combination of any data available in the data stream, using several display types such
as histograms, bar charts, pie charts, sky view, ... The list of displayed objects and
parameters is fully customizable purely from database, without the need for any change
to the code.

4.1.5 Controller Structure

The standard software architecture of a typical hardware controller is depicted in Fig. 4.3.
This architecture is substantially standardized across all hardware elements in the H.E.S.S.
experiment.

- A dedicated controller communicates with the hardware potentially using proprietary
protocols defined by the hardware manufacturer. This controller is based on
the state machine structure, to ensure standard communication with the central
DAQ. The role of the controller is to configure and run the hardware, and also to
receive the data from the hardware and convert it into the H.E.S.S. data format.

- A data receiver binds to the controller via a subscription mechanism, and receives
the data during data taking. In some cases, several data receivers can run in parallel.
This is the case for instance for the weather station, for which a central receiver runs
all the time and stores the data into a day-wise file. Another receiver related to the
data taking stores a second copy of the the data in run-wise files that are more easily
analyzed. In order to ensure that no data is lost, the data receivers work in a push
mode with significant buffers on both sides of the communication.

- Optional readers bind to the receiver and receive either a fraction of the data (in a
sampling mode) for display, or to the full data for any analysis. These readers are
optional for data taking, any problems they could experiment would not interrupt
data taking. Readers that only need a subset of the data (such as event displayer)

---

3meaning that the data are sent by the controller whatever the status of the receiver is
can request data using a pull mode, requesting data only at a certain state or when they are available for further treatment.

- A manager acts as an orchestra chief, establishing the communication between the different CORBA controllers, and coordinating transitions of the state machines.

![Software architecture diagram](image)

**Figure 4.4**: Software architecture of a typical hardware controller. Data are sent using a push mode from the hardware to the receiver, which store it on disk. Optional readers, such as event display, can access a subset of the data using a pull mode.

### 4.2 Camera trigger & Central trigger operation

The main data flow of the H.E.S.S. experiment is produced by the cameras which generate events with a size of 1.5 kB. At the design trigger rate of 1 kHz, this yields a maximum data rate of 6 MB/s for four telescopes, resulting in roughly 100 GB of data per observation night. The data rates from the other subsystems are significantly smaller.

The general data flow of the H.E.S.S. camera and central trigger operation are depicted on Fig. 4.5. The trigger of H.E.S.S. is a two-level system:

- At the first level, telescopes trigger independently from local information. The trigger of the H.E.S.S. cameras is derived from a multiplicity trigger within overlapping sectors, each containing 64 pixels. A camera trigger occurs if the signals in M pixels within a sector (sector threshold) exceed a threshold of N photo-electrons (pixel threshold). The time-window for the multiplicity trigger is dictated by the minimum integrated charge over a programmable threshold.

  For a typical PMT pulse shape the effective trigger window is 1.3 ns (with a jitter of 0.14 ns). This rather narrow gate is possible due to the sorting of PMTs by required high voltage within the camera. This sorting minimizes the time dispersion introduced by different PMT transit times. The narrow gate guarantees maximum NSB suppression.

- At the second level a multi-telescope coincidence decision is made by the central trigger system in order to make optimum use of the stereoscopic approach.
Simultaneous observation of an air shower in multiple telescopes is required at the hardware level in order to reduce the rate of background events (for example single muons) significantly, enabling a reduction of the energy threshold of the system by a factor \(\sim 2\) compared to single telescope operation.

![Diagram of Central Trigger Operation](image)

**Figure 4.5: Central Trigger Operation.**

The central trigger operates in the following way:

- The triggered camera sends, through an optical fiber, a TTL trigger signal to a CPU located on the central farm together with a TTL level indicating the current readout status. At the same time, in the case of an active trigger, the analogue memories are stopped and the camera starts to convert the photo-multiplier data and transfers it on a specifically designed data bus, called the box bus.

- In the central hardware, the trigger signal from the different cameras are corrected for a variable delay that depends on the actual pointing of the telescopes (up to 1 ms by steps of 1 \(\mu s\)), and enter a coincidence module. The trigger condition is checked using a programmable look-up table in which all allowed coincidence patterns are stored. In normal operations, the coincidence window is set to 80 ns.

- When a trigger decision is made (when the number of triggered telescope exceeds the usual threshold of 2), a confirm signal is sent to all triggered cameras, together with a unique event number and a node ID (sent through a serial line), which identifies the machine on the farm where the data should be sent to.

- If no confirm signal is sent by the central trigger to the camera within a gate of \(\sim 5 \, \mu s\), a fast clear signal is generated within the local trigger module of the camera, and the data transfer is immediately interrupted. The analogue memories are restarted.

To enable the measurement of the system dead-time, cameras provide their current readout status together with the trigger signal (i.e. whether the camera is busy acquiring a
previous event or active and able to read out this event). The camera trigger and readout status information is stored for each telescope on an event by event basis in a FIFO. From these data the dead-time of the system can be derived. Fig. 4.6 shows the distribution of time differences between consecutive events for a single telescope. This distribution for all triggers is well described by an exponential function as expected for uncorrelated events. The distribution of events where the telescope was read out (light grey) is also exponential but with a sharp cut-off at the camera readout dead-time of 446 µs.

![Distribution of time differences between consecutive events for a single telescope.](image)

Figure 4.6: Distribution of time differences between consecutive events for a single telescope. The dark grey histogram shows all events triggering the system, the light grey histogram shows only events for which the telescope was read out. The inset shows the same distributions with a logarithmic scale to highlight the exponential behaviour. From [ref].

## 4.3 Camera Control

The cameras are one of the most complex components of the H.E.S.S. experiment, because of the complexity of the involved components on one hand, and, on the other, of the time constraints implied by shower duration and real time acquisition capabilities, which require to minimize the amount of communication between the camera and the central DAQ during data taking.

The camera DAQ runs on Motorola embedded systems inside the camera, controlled by a H.E.S.S. controller in the farm (Fig. 4.7 and 4.8). A specific configuration syntax is used to communicate between the central DAQ and the camera, and in particular to extract the full camera configuration from the database and send it to the hardware.

The camera control interface (GUI) allows full access to the manual configuration of each element inside the camera. For instance, the pixel configuration page (Fig. 4.8)
allows access to the high voltage status and value as well as trigger participation of each pixel inside the camera. This can be used in some cases to switch off a mis-behaving pixels that would generate additional noise in the trigger. However it is usually operated in a database-driven mode where the configuration parameters are extracted from the central database according to the run type.

### 4.4 Camera Readout

The camera readout scheme is depicted in Fig. 4.9.

When a system trigger occurs, the central trigger system provides a unique, system-wide event identification, composed of a bunch number and an event number in the bunch, which are distributed via the trigger hardware to all telescopes. These two numbers are read by the camera in a *local module* together with the address of the farm node they should sent the data to. For a given event, the data from all read-out camera are sent simultaneously to the same data receiver to allow for event building.

In order to avoid generating any dead time during the acquisition, the farm nodes only buffer the data they receive, using one memory buffer per telescope.

A dedicated controller, called the *connector*, periodically selects a new idle farm node where the data should be sent to. The usual pace is 4 s. When a new node is selected, the central trigger software sends the trigger data (central trigger mask, event number together with an unique bunch identifier and GPS time stamp for each event in the bunch) to the previous active node, which enters the *Running* state and starts to build the events it has previously buffered together with the central trigger data. Each event is identified with a bunch number and an event number: the bunch number is increased at each node
switch and the event number is reset to zero.

Figure 4.8: Camera Control Interface - Pixel control page.

Figure 4.9: Camera Readout scheme

4.5 Run Control and Naming Scheme

Setting up the required processes for a specific run type is simplified by combining related processes in groups that are called contexts. An example for a context is CT1 which comprises all Controllers accessing the hardware of telescope number 1. Every context
contains one Manager that controls the other processes in the context. At start up, the Manager reads the processes in its context from database tables and launches them. The Manager serves as an intervention point to all processes in the context, passes on state transitions, and takes over error handling in predefined ways.

![Diagram of Run Control in H.E.S.S.]

Figure 4.10: Run Control in H.E.S.S. For the sake for clarity, not all processes are shown.

The overall coordination of data-taking is handled by a dedicated run manager (Fig. 4.10), that can handle several data taking tasks in parallel (using different resources). Controllers corresponding to different hardware elements or software tasks are organized in a hierarchical structure, as described in Table 4.2. Each context can host several controllers, and is managed by a specific manager. Managers take care of the state transition of each of the controllers belonging to the context.

<table>
<thead>
<tr>
<th>Context</th>
<th>Signification</th>
</tr>
</thead>
<tbody>
<tr>
<td>Array</td>
<td>Central array control</td>
</tr>
<tr>
<td>Atmosphere</td>
<td>Weather monitoring</td>
</tr>
<tr>
<td>Atmosphere/Meteo</td>
<td>Meteo station controller</td>
</tr>
<tr>
<td>Atmosphere/Ceilometer</td>
<td>Lidar</td>
</tr>
<tr>
<td>Atmosphere/ScanningRad</td>
<td>Infrared radiometer for detection of clouds</td>
</tr>
</tbody>
</table>

Table 4.2 – continued on next page
<table>
<thead>
<tr>
<th>Context</th>
<th>Signification</th>
</tr>
</thead>
<tbody>
<tr>
<td>CentralTrigger/Controller</td>
<td>Central Trigger controller</td>
</tr>
<tr>
<td>SubArray0x/</td>
<td>Run control using a fraction of all resources</td>
</tr>
<tr>
<td>SubArray0x/Receiver/xxxxx</td>
<td>Data receivers associated to run</td>
</tr>
<tr>
<td>SubArray0x/Display/xxxxx</td>
<td>Event and monitoring display associated to run</td>
</tr>
<tr>
<td>SubArray0x/Services/</td>
<td>Utilities associated to run, such as database configuration reader or online analysis</td>
</tr>
<tr>
<td>SubArray0x/Trigger/Connector</td>
<td>Central trigger control</td>
</tr>
<tr>
<td>CTx/</td>
<td>Telescope control (camera, tracking, ...)</td>
</tr>
<tr>
<td>CTx/Camera</td>
<td>Camera control (power, slow control)</td>
</tr>
<tr>
<td>CTx/CameraHV</td>
<td>Camera High Voltage Control</td>
</tr>
<tr>
<td>CTx/CameraTrigger</td>
<td>Camera Trigger Control</td>
</tr>
<tr>
<td>CTx/CameraLid</td>
<td>Camera LID control</td>
</tr>
<tr>
<td>CTx/CameraCalibPMT</td>
<td>Control of external PMT used for calibration purposes</td>
</tr>
<tr>
<td>CTx/Tracking</td>
<td>Telescope tracking control</td>
</tr>
<tr>
<td>CTx/Radiometer</td>
<td>Par-axial Radiometer</td>
</tr>
<tr>
<td>CTx/StarController</td>
<td>Security process, checking trajectory of stars in the camera, sun rise and moon position</td>
</tr>
<tr>
<td>CTx/SkyCCD</td>
<td>CCD looking at the sky for tracking deviations</td>
</tr>
<tr>
<td>CTx/FlatField</td>
<td>Flatfielding calibration unit</td>
</tr>
<tr>
<td>CTx/LidCCD</td>
<td>CCD used for measurement of telescope deformations</td>
</tr>
<tr>
<td>Node0x/</td>
<td>Control of a data receiver node</td>
</tr>
<tr>
<td>Node0x/Receiver</td>
<td>Data receiver node</td>
</tr>
<tr>
<td>Services/</td>
<td>Central services, running all the time (data merger, DST production, sound control, ...)</td>
</tr>
<tr>
<td>Services/DataMerger</td>
<td>off-line data merger</td>
</tr>
<tr>
<td>Services/Summary</td>
<td>generates summary of runs.</td>
</tr>
<tr>
<td>Services/GCNAlert</td>
<td>Receiver for gamma-ray burst alerts</td>
</tr>
<tr>
<td>Services/Sound</td>
<td>Central service for sound alerts</td>
</tr>
<tr>
<td>SlowControl/</td>
<td>Slow control of the entire array, running all the time</td>
</tr>
<tr>
<td>SlowControl/Receiver/xxxxx</td>
<td>Data receivers receiving a copy of the slow control data and storing it in day-file files</td>
</tr>
<tr>
<td>SlowControl/Display/xxxxx</td>
<td>Global monitoring displays</td>
</tr>
</tbody>
</table>

Table 4.2: List of main contexts used in the H.E.S.S. data acquisition.

In addition, a specific database table specifies dependencies between controllers, thus imposing a specific configuration order: the transition of any controller is blocked until the controllers it depends on reached the correct state. The dependencies are handled in opposite direction during configuration and start up phases on one hand, and stopping and un-configuring phases on the other hand. For instance, the camera trigger depends on the camera high voltage: the trigger is only switched on once the high voltages reach their nominal values. In contrast, the trigger is stopped before the high voltages are lowered.
4.6  EVENT BUILDING

For data-taking a central DAQ Manager reads the observation schedule from the
data base and determines the actual sequence of runs according to the availability of
contexts. Runs that require different contexts can be processed in parallel. The central
DAQ Manager launches the required context Managers and initiates the run preparation.
After the preparation of a run the starting and stopping of the data-taking is taken over
by a dedicated Manager which controls the participating contexts.

![Central control GUI and monitoring displays](image)

Figure 4.11: Central control GUI (left) and examples of monitoring displays (right).
From [58].

The shift crew interacts with the DAQ system via a central control GUI providing an
access point to the system and direct monitoring of the states of the different processes (cf.
Fig. 4.11 (left)). Data monitoring information is shown by a variety of different displays.

The displays are generated by instances of a generic Reader which are configurable
via database tables allowing a simple and flexible setup of the quantities to be displayed.
The introspection mechanism of ROOT is used to access the relevant members of data
classes dynamically. Fig. 4.11 (right) shows some examples of monitoring displays.

4.6  Event Building

At the end of a run, the data consist typically of 5 different data files\(^4\) (one for each
farm node), each of them containing a certain number of data bunches. A dedicated
DataMerger program merges the data offline, sorting the events based on the bunch
and event numbers. This is depicted in Fig. 4.12.

\(^4\)in the early days of H.E.S.S., due to small computing power per node, the data were split in 20
different files
4.7 Redundancy and reliability

Many hardware elements in the camera sometimes suffer from software or hardware problems. Among the problems, one can cite:

- Incorrect event or bunch numbers read from the local modules, usually caused by some stuck bits in the registers,
- Free-running camera GPS clocks (meaning that the clock is not locked on the GPS signal),
- Duplicated event header without actual data, generated by self-triggering of the the camera induced by electromagnetic noise occurring during data transfer,
- ...
For instance Fig. [4.13] shows the distribution of time difference between consecutive events in a normal observation runs, for all events (red) and read-out events (blue). In addition to the (expected) exponential distribution a prominent peak occurs at a time difference of 350 μs. This peaks is caused by the start of the data transfer on the camera bus, which in turns generates an additional electromagnetic noise inside the camera and provokes a self-triggering of the camera. Theses events thus generate busy-trigger events which lead to a wrong estimation of the system dead-time (since they do not originate from a real shower), and are called ghost triggers.

The effect of these problems is usually a non chronological data file, with events mixed up or improperly merged events. A dedicated program, called **StickyBitFixer**, tries to correct for most of the problems based on a probabilistic approach, and using the redundancy available in the system. It is based on the following information:

- GPS clocks from each of the 4 camera, plus the GPS clock of the central trigger.
- CPU clocks from the camera CPU, which is used to check whether the GPS clock is flickering
- Hardware camera trigger counter, which is monotonically increasing, but with gaps due to single telescope trigger
- Event and Bunch number sent by the central trigger (i.e. software trigger counters)

It has to be noted that none of the aforementioned information is fully reliable and only a statistical approach can successfully recover the data.

### 4.8 Conclusion

We have designed a flexible, evolutive and easily scalable DAQ system for the H.E.S.S. experiment. This system, based on industrial technologies such as the CORBA inter-process communication protocol, and implemented with free software, has worked smoothly since the beginning of the experiment in 2001. Incorporation of new hardware has been effected on several occasions and has been done in a smooth and effective manner. The DAQ system is ready to accept the coming H.E.S.S.-II telescope, and this incorporation is expected to be done in a seamless way.

Such flexible and evolutive systems are clearly required for Atmospheric Cherenkov arrays of telescopes which incorporate many independent devices that, however, need to cooperate closely during data acquisition. Similar approaches are foreseen for next-generation instruments such as CTA.
Chapter 5

Calibration

5.1 General principles

The signal from the Cherenkov light produced by the showers is recorded by the cameras in analogue memories and then converted to ADC counts which do not give a direct measure of the light actually emitted. The goal of the calibration is to convert these electronic signals into a physical unit, the density of Cherenkov photons. The calibration involves several steps:

- Determination of the electronic baseline (Pedestal determination)
- Conversion of the electrical signal into a number of detected photo-electrons. This step involves the gain calibration of the photo-multipliers using a flashing LED in front of the camera
- Relative calibration of the high and low gain electronic chains, using the overlapping band in amplitude on real events
- Conversion of detected photo-electrons into a number of Cherenkov photons using the optical transmission efficiency of the system (mirror, Winston Cones, but also atmosphere). In general the reflection coefficient of each element cannot be measured individually, but a global transmission coefficient can be obtained relatively easily from the measurement of the signal generated by muons crossing the telescope. Indeed, the signal generated by those muons involves only Cherenkov emission by ultra-relativistic particles in the lower atmosphere (which can be analytically calculated) and transmission by the detection system. Since the muon measurement only provides an average efficiency measurement across the camera, a flat-fielding procedure is needed to inter-calibrate the quantum and collection efficiencies of the different PMTs.

The calibration coefficients required in the full system calibration are described in Eq. (5.1)

\[
S_{HG} = \frac{ADC_{HG} - P_{HG}}{\gamma_{HG}} \times FF
\]

\[
S_{LG} = \frac{ADC_{LG} - P_{LG}}{\gamma_{HG}} \times (HG/LG) \times FF
\]

where:

\[HG\]

\[LG\]
5.2. GAIN CALIBRATION

- $\text{ADC}_{HG}$ and $\text{ADC}_{LG}$ are the measured charges, in ADC counts, respectively in high and low gains
- $P_{HG}$ and $P_{LG}$ are the *pedestals*, or baseline of the electronics, respectively in high and low gains
- $\gamma_{HG}$ is the conversion factor between ADC counts and photo-electrons, also named *gain*
- $FF$, or *flat-fielding coefficient*, characterizes the relative collection efficiency (and quantum efficiency) of one specific PMT with respect to the camera average
- $HG/LG$ is the ratio of high to low gain electronic amplification factors

### 5.2 Gain calibration

The gains of the PMTs are calibrated using a flashing LED pulsed at 70 Hz, placed in the camera shelter (to avoid contamination from the night sky background), $\sim 2$ m in front of the camera. The intensity of the LED is adjusted to enable the measurement of the single photo-electron peak on the PMTs. Dedicated runs, using a specific trigger synchronized on the LED, are done every two nights. Only the high gain has enough resolution to allow for gain calibration.

The charge distribution is adjusted using the expression \[5.2\] where the first term represents the pedestal distribution and the second one a Poisson distribution for the number of photo-electrons convoluted with Gaussian of increasing width, representing the resolution of the PMT for a signal of $n$ photo-electrons:

\[
G(x) = \frac{e^{-\mu}}{\sqrt{2\pi}\sigma_p} \sum_{n=1}^{m} \frac{e^{-\mu} \mu^n}{\sqrt{2\pi n} \sigma_{\gamma e} n!} \left[ \frac{(x - P_{HG})^2}{2\sigma_p^2} \right] + \frac{e^{-\mu} \mu^n}{\sqrt{2\pi n} \sigma_{\gamma e} n!} \left[ \frac{(x - (P_{HG} + n\gamma_{HG}))^2}{2n \sigma_{\gamma e}^2} \right]
\]

where

- $N$ is the total number of events
- $\mu$ is the average luminosity, also called the *single photo-electron fraction*. This corresponds to the average number of photo-electrons per event
- $P_{HG}$ is the pedestal position in the high-gain channel (expressed in ADC counts)
- $\sigma_p$ is the pedestal width
- $\sigma_{\gamma e}$ is the single photo-electron peak width
- $\gamma_{HG}$ is the gain of the PMT
- $N_S$ is a normalization constant that characterize the compliance of the full distribution with a Poisson distribution. In the case of a true Poisson distribution, it should read $N_S = 1$. 

\[5.2\]
Figure 5.1: Single photo-electron spectrum. The high voltage is determined to get the peak of the single photo-electron at 80 ADC counts above the pedestal.

An example of a single photo-electron distribution with such an adjustment is shown in Fig. 5.1. In this example, the gain is measured to be 90.25 ADC counts per photo-electron (DiffGamPed) and the normalization constant reads $N_S = 1.17$ for a single photo-electron fraction of 75%.

The resolution of the PMT can be expressed as the single photo-electron width normalized to the gain: $Res = \sigma_{\gamma_e}/\gamma_{HG} \sim 45\%$.

The LED that is used for the single photo-electron calibration is placed in the camera shelter, two meters in front of the camera. The distribution of the single-photo-electron fraction across the camera is shown in Fig. 5.2 for one single photo-electron run. The average illumination is 49% with a fluctuation of ±14% from the center of the camera (where the illumination is maximum) to the edges. Although the illumination is not really homogeneous, the gains are still determined with a good accuracy on all pixels.

Figure 5.2: Homogeneity of the LED luminosity across the camera (left) and corresponding distribution (right).

All single photo-electron runs of one observation period are merged together to derive more precise, average calibration coefficients. The average gains are around 80 ADC counts
per photo-electron for an uncertainty of usually less than 1 ADC count and are homogeneous up to a few ADC counts across the camera field (Fig. 5.3).

![Average Gain](image1.png)  ![Average Gain Uncertainty](image2.png)

Figure 5.3: Distribution of pixel average gains for the telescope CT1 and the period of September 2011 (left) and corresponding uncertainties (right).

### 5.3 Pedestal determination

Pedestals are the baseline of the electronics. They need to be calibrated with precision as the Cherenkov signals are calculated with respect to this baseline, so the resolution on the actual signals directly depends on the precision with which the pedestal is obtained.

Dedicated runs with the camera lid closed are used to determine the electronic pedestal, in the absence of night-sky background noise. The width of these electronic pedestals (typically 15 ADC counts in high gain) constitute a measurement of the electronic noise in the camera, which, for the nominal gain, is well below the signal from a single photo-electron ($16/80 \approx 0.3\gamma_c$). Examples of such measurements are shown in Fig. 5.4. These measurements also allow various instrumental effects to be investigated and calibrated, such as the dependence of the pedestal position on the temperature (the pedestal typically moves by 10 ADC count per degree Celsius).

Because the pedestal position on each pixel depends on various instrumental effects, including the atmospheric temperature and the level of night sky background, calibration of the pedestal must be done for each observation run.

### 5.3.1 Pedestal shape in observation runs

In the presence of night-sky background, the shape of the pedestal distribution is significantly altered: for a small rate of photo-electrons from the sky, it is similar to the shape of the photo-electron spectrum, with the difference that the time of the photo-electron pulse is not correlated with the trigger so that a random fraction of the charge is recorded. This leads to a degradation of the peak-to-valley ratio in the charge spectrum due to the accumulation of fractional photo-electrons (Fig. 5.5 upper.). When the acquisition
Figure 5.4: Shape of electronic pedestal in high and low gains. Adapted from [34].

window is large compared with the duration of the single photo-electron pulse, the single photo-electron spectrum is reproduced almost unaltered.

For increasing level of NSB, the pedestal becomes more and more symmetrical as the average number of photo-electrons per acquisition window approaches one (Fig. 5.5 lower.).

The relation between the level of NSB and the pedestal shape has been extensively tested using a dedicated test bench [39].

5.3.2 Cherenkov light filtering

The determination of pedestal from the data is relatively tricky as the trigger is specifically designed to trigger on showers and not on pure noise. Two strategies can be adopted to circumvent this problem:

- Interleave actual triggers on shower with random triggers. This ensures the acquisition of pure noise events, but increases the amount of data and thus the average dead time of the acquisition

- Use the data recorded on showers. This solution requires the part of the camera that has been illuminated by the shower to be identified and excluded on a event by event basis, so that only the data from the remaining pixels are accumulated. This is the solution actually used by H.E.S.S.. The algorithm is the following:

  - The data are accumulated without any filtering using a sliding window of $\sim 10$ s to provide a first estimate of the pedestal position on each pixel
  - The charge on each pixel is then estimated using this preliminary pedestal, and the intensities on the neighbouring and next-neighbouring pixels are averaged, taking into account the uncertainty on each charge and excluding the
Figure 5.5: Determination of pedestal (charge distribution in absence of signal) in high gain for some pixels with various NSB levels: \textbf{up} for low NSB levels of \( \sim 30 \) MHz, \textbf{down} for high NSB levels of \( \sim 100 \) MHz.

pixel being worked on. In parallel, the number of neighbouring pixels with an intensity above a certain threshold is calculated. A pixel is considered to be contaminated by Cherenkov signal if the average intensity of the neighbours exceed a certain threshold (usually \( \sim 0.4 \, \gamma_e \)) or if more than 6 neighbouring or next-neighbouring pixels see a significant contribution to the signal caused by Cherenkov light, typically above 4 standard deviations from the noise.

– Charge histograms are filled only for those pixels that are not identified using the previous algorithm as having a significant contribution from Cherenkov signal. To avoid any bias on the pedestal determination, the intensity seen by a particular pixel is never used to determine if it has seen some Cherenkov light. Only the neighbouring and next-neighbouring pixels are used.

The effect of this selection based on the intensity seen by neighbouring and next-neighbouring pixels is shown in Fig. 5.6, where the white histogram corresponds to the charge distribution seen by a particular pixel without any filtering and the blue filled one with the Cherenkov filtering procedure applied. About 20\% of events are filtered out, and as expected, the high intensities are properly removed without alteration of the low charge part.
5.3.3 Pedestal evolution and stability

The shape of the pedestal distribution depends, among others, on the temperature of the ADC converter chips which can vary significantly during the night. Due to the capacitive coupling behind the PMTs, its average level also depends on the actual level of night sky background. As the camera is rotating onto the sky during observation, the later is subject to significant variations during an observation, as faint stars might enter or leave a pixel. For these reasons, the pedestal needs to be recalculated on a regular basis.

In the calibration procedure, the pedestal over the full camera is recomputed as soon as the average position of the pedestal of at least four pixels in the camera has moved by more than 10 ADC counts compared to the previous calculation. In addition, the time gradient of the pedestal is calculated each time to provide a linear interpolation of the pedestal level with time.

Different patterns of pedestal evolution with time are shown in Fig. 5.7. Similar evolution in high and low gains as in the upper panel are characteristic of a pedestal shift caused by a temperature variation. Evolution with similar pattern although with very different amplitude such as in the middle panel are usually produced by a change of the night sky luminosity in a specific pixel, in particular due to the passage of a star.

Despite this careful treatment of pedestal evolution, detailed studies of the pedestal position as a function of time resulted in the identification of several instabilities. Fig. 5.8 shows for instance the evolution of the pedestal level for 16 PMTs belonging to the same analogue card. The completely parallel evolution in all PMTs (except one disconnected due to a HV problem) indicates that the shift cannot be attributed to a problem in the analogue memories since the 16 PMTs are grouped by four on different chips. The origin of such brutal jumps is not yet understood and must be somehow related to one of the two analogue memory cards (grouping several analogue memories) in the drawer or the slow control board. Detection for such events is implemented in the calibration chain and the pedestal calibration is re-initialized after such after an important jump.
Figure 5.7: Evolution of the pedestal position as a function of time for a specific pixel in low gain (left) and high gain (right). **Up:** As the amplitude of variation is similar in both gains, the reason of this shift is most probably an instrumental effect such as variation of the temperature of the chip. **Middle:** As the amplitude of variation is a factor of $\sim 6$ higher in high gain than in low gain, however with a similar evolution pattern, the reason of this shift is most probably a variation of the luminosity of the sky in this specific pixel. **Bottom:** Pixel showing no significant variation of pedestal position.
Figure 5.8: Unstable pedestals seen in 16 PMTs belonging to the same drawer, in high gain. Low gain data show an identical pattern.

5.3.4 Common modes

Detailed investigation of the property of the pedestal on an event-by-event basis led to the observation of coherent shift of the pedestal on the upper or the lower part of the camera, as shown for example in Fig. 5.9. In the lower part of the camera, the pedestal is shifted up by $\sim 2 \gamma_e$, thus leading to a complete overestimation of the image amplitude. The red part in the center of the camera corresponds to the image of the shower. Since the separation between the upper and the lower part of the camera actually corresponds to two different power supplies, an instrumental problem was immediately identified as the cause.

Statistical analysis of this effect have been performed. We define the common mode as the coherent pedestal shift on one of the two half of the camera, expressed in photo-electrons. It is determined using a similar procedure to the one used for pedestal determination: the pixels affected by Cherenkov light are excluded from the determination and the average value of the charge over all remaining pixels is determined and converted in photo-electrons. The distribution of the common mode for one observation run and for one camera half is shown in Fig. 5.10. About 5% of the events are affected by a significant shift of the pedestal, between 0.5 and 2 photo-electrons. For a typical shower illuminating about 20 pixels in the camera, the total effect can be of the order of $40 \gamma_e$, thus leading
Figure 5.9: Display of the intensity over one camera, for which the common mode oscillation is observed. From [60].

to large systematic effect in the energy determination of the affected events.

Figure 5.10: Distribution of Common Mode for the lower part of the camera.

Fig. 5.11 shows the distribution of the common mode as a function of the time difference between the affected event and the previous event. The observed damped oscillation is a clear indication of the presence of a capacitive coupling transfer chain: During the acquisition and the transfer of one event on the data bus, the power supplies are significantly drained and need some time to return to equilibrium. The recovery time scale has been studied in details in [60] and ranges between 300 and 400 μs. Due to its much smaller dead-time, this effect might be more pronounced in the case of the soon-to-be installed H.E.S.S.-II telescope.
Figure 5.11: Distribution of Common Mode for the upper part of the camera as a function of the time difference to the previous event [00]. The lack of events below 500 µs is due to the acquisition dead-time.

5.3.5 NSB Measurement with PMTs

Pedestal width constitutes the most reliable observable of NSB level, as it does not depend on additional calibration constant such as the temperature. Other observables, like the high voltage current or the anode current of in the PMTs need a much more precise calibration [20, 01].

The total width of the pedestal results from three different components:

- the width of the electronic pedestal \( \sigma_{p0} \)
- the resolution of the PMT (width of the single photo-electron peak) \( \sigma_{\gamma} \)
- the statistical fluctuations in the NSB

For an acquisition window of \( \Delta t = 16 \) ns, and a gain \( \gamma_{HG} = 80 \) ADC \( \gamma^{-1} \), the NSB rate (in units photo-electrons per second) can be expressed as [e.g. 54]:

\[
\frac{r}{\gamma_{HG} \times \Delta t} = \frac{\sigma_{p}^2 - \sigma_{p0}^2 - \sigma_{\gamma}^2}{
\gamma_{HG} \times \Delta t}
\] (5.3)

Examples of the variation of the pedestal width caused by transit of stars are shown in Fig. 5.12. The transit of two stars on to different pixels is clearly seen, both in the high and low gains.

This NSB estimator is used in the H.E.S.S. analysis to investigate potential systematic effects related to the NSB variation across the sky. Fig. 5.13 shows the comparison of the NSB determined using this method with measurements from optical telescopes. The very obvious correlation stresses the capabilities of the method. Typical NSB rate in H.E.S.S. observations ranges from \( \sim 40 \) MHz outside the Galactic Plane up to \( \sim 300 \) MHz in the brightest regions of the plane, such as one degree below, in latitude, the Galactic Centre.
5.4 Cross-calibration of high and low gains

The single photo-electron peak can only be resolved in the high-gain channel. Proper calibration of the low gain channel, required to extend the dynamical range of the instrument above 200 p.e., is done relative to the high-gain channel. Fig. 5.14 shows the charge...
collected in the high-gain channel, in ADC counts, as a function of the charge measured in the low gain channel. Saturation occurs in high gain for signal of ~ 200 p.e. and above. At low charges, below ~ 40 p.e., the resolution of the low-gain channel deteriorates significantly. The overlapping dynamical ranges of the two channels, between about 40 and 200 p.e., is used to inter-calibrate the high and low gain channels, on a run-by-run basis.

Figure 5.14: High gain versus low gain charge from pedestal.

Inter-calibration of the high and low gain channels is done individually for each pixel, on a run-by-run basis. Calibration coefficients are then merged by observation periods to reduce the statistical fluctuations on the individual pixel values per period below 1%.

The obtained inter-calibration coefficient is of the order of 13.5, with a dispersion of about 10% from one pixel to the other, clearly dominated by the intrinsic properties of the front-end amplifiers and not by statistical fluctuations.

Long-term evolution of the average calibration coefficient (averaged over all pixels of one camera) is shown in Fig. 5.15 for data taken between 2006 and end of 2011. The observed seasonal variation of 0.1 in amplitude has not been fully understood yet, but could be attributed to the effect of seasonal temperature variation on the amplifiers. However, this effect would then also be noticeable at the time scale of one night, which is not the case yet. Effects of other parameters, such as average humidity, are under investigation.

5.5 Flat-fielding

Precise measurement of the collection and quantum efficiency of the individual PMTs is not an easy task, although this information is required to convert photo-electrons into incident photons. Quantum efficiency calibration requires in particular a stable and controlled light source. Relative calibration of pixel response is much easier to achieve, using dedicated observation runs with a flashing LED installed at the center of the dish, in front
5.6 Absolute calibration using atmospheric muons

The different calibration coefficients investigated so far, and described in Eq. [5.1] only allow the electrical signals recorded in the camera to be converted into a number of photo-electrons, averaged over the spatial efficiency of the camera. The final step towards an absolute calibration of the camera requires the conversion between photo-electrons and actual Cherenkov photons. This conversion coefficient depends, theoretically, on:

- the reflectivity or transmittivity, as a function of wavelength, of the various optical elements of the system: mirrors and Winston cones mainly, but also entrance window of the PMTs
- the collection efficiency, which depends on the actual placement and orientation of the various elements
- the collection and quantum efficiency, as a function of wavelength, of the PMTs

Individual calibration of the various elements is almost impossible since, due to the continuous nature of the Cherenkov spectrum, it would require a precise calibration at all wavelengths.

Thus the approach chosen by operators of all the major atmospheric Cherenkov telescopes is to rely on the Cherenkov signal emitted by atmospheric muons to calibrate the instrument. This *natural* calibration beam indeed offers some advantages:

- the radiation spectrum is very similar to the Cherenkov emission of electrons in gamma-ray induced showers. Actually the main difference comes from the fact that muons are more penetrating, leading to an emission occurring in average at lower altitudes, and thus suffering less from ultra-violet absorption.
• muons reaching the ground are usually ultra-relativistic and saturate the Cherenkov angle \((\cos \theta \approx 1/n)\). As a consequence, the light yield only depends on the track length and not anymore on the muon energy.

• the muon signal is very short in time, and very similar to the signal induced by showers.

The main difficulty in this absolute calibration with muons comes from the design of the trigger strategy: muons usually trigger only one telescope; usage of stereoscopy reduces the number of observable muons to an handful per telescope and per hour. Moreover, muons going through a mirror produced a complete ring, easily identifiable (Fig. 5.16 left), although pretty rare, whereas muon falling at larger distances produce a small ring segment that resemble a \(\gamma\) ray image (Fig. 5.16 right).

![Figure 5.16: Examples of muon images recorded by the H.E.S.S. camera.](image)

Detailed modeling of the Cherenkov emission by muons has been done extensively [e.g. 61] and is used on a regular basis to follow the overall optical efficiency of the system. Fig. 5.17 shows the evolution of the overall efficiency of the H.E.S.S. telescopes from 2003 to 2011. On average less than 10\% of Cherenkov photons reaching the telescopes are detected. The continuous drop has been attributed mainly to degradation of the reflectivity of the mirrors. Re-aluminization of the mirrors of CT3 in 2011 led to a sudden increase of optical efficiency.

### 5.7 Non-operating pixels

Before any proper analysis of the data, non-operational or mis-operating pixels must be identified and removed from the charge analysis. Problems affecting pixels can originate from different elements:

• High voltage problems, such as unstable or switched off HV supply. These problems can be easily identified using the available monitoring data.
• Missing or abnormal data, usually due to corruption of data sent on the data bus. Missing data are automatically identified and flagged by the acquisition software. Corrupted data can be identified because these events show a very different charge distribution.

• De-synchronization of analogue memories, which causes four pixels to have random charge readout in high or low gains, thus having effectively only pedestal data. These pixels can be identified by a detailed comparison of the signals measured in high and low gains.

5.7.1 Unlocked ARS

Sampling of the signal by the ARS analogue memories, at a frequency of 1 GHz, is performed using a chain of 128 elements with a transition time of 1 ns between each element. Transition from one element to the next one is triggered by the propagation of an electronic edge signal. An intermittent problem exists in which two or more fronts propagate simultaneously, resulting into several distinct locations in analogues memories being written simultaneously. This mis-behaviour is triggered by an incorrect initialization of the chip when the power is switched on, and therefore remains present until the power is switch off again. When camera power supplies are switched on, a random fraction of the order of 5 – 10% of the analogue memories are improperly initialized. The memories affected therefore change randomly from camera reconfiguration to camera reconfiguration.

Each analogue card, connected to eight pixel, comprises four analogue memories. Pixels are grouped by four in each analogue memory, and the high and low gains are handled by different chips to reduced the risk of cross-talk. The result of this electronic configuration is that the so called unlocked ARS behaviour always affects a group four pixels aligned vertically in the camera.

Identification of affected pixels is done through a detailed comparison of the signal recorded in high and low gains, for events with significant signal in either of the two
channels. Only events with charge in the overlapping dynamical range ([40 – 200] p.e.) are used in the comparison. We use the *ARS Lock Ratio* defined from the signals $S_{HG}$ and $S_{LG}$ recorded in high and low gains respectively, and expressed in photo-electrons after application of the full calibration:

$$ARS_L = \log_{10} \left( \frac{S_{HG}}{S_{LG}} \right)$$  \hspace{1cm} (5.4)

In the case of normal electronic behaviour, the distribution of the *Ars Lock Ratio* should be centered on zero with a small width dominated by the charge resolution of the low gain channel. Such an example is shown in Fig. 5.18 top left.

![Figure 5.18](image)

**Figure 5.18:** Determination of unlocked ARS using pedestal subtracted charge. **Top left:** distribution for a correct pixel. **Top right:** distribution for a pixel with bad charge readout. **Bottom left:** distribution for a pixel for which the low gain ARS is unlocked. **Bottom right:** distribution for a pixel for which the high gain ARS is unlocked.

When one of the two channels is connected to an unlocked ARS, the corresponding signal is read out at an incorrect location in the analogue memory, resulting into a smaller charge, essentially compatible with pedestal, whereas the other channel keeps a high charge. The *Ars Lock Ratio* distribution then shows an asymmetrical distribution, with a tail extending on the positive (resp. negative) values in the case of low (resp. high) gain unlocked ARS. Examples of this behaviour is shown in Fig. 5.18 bottom. The central peak on these distributions corresponds to the cases when the readout window in the affected channels corresponds, by chance, to the actual location where the signal was written. Some pixels, such as that shown in Fig. 5.18 top right, show an incorrect behaviour that does not correspond to unlocked ARS.
Figure 5.19: Identification of mis-operating PMTs using the comparison of high-gain and low-gain signals.

5.7.2 PMTs with incorrect charge behaviour

Some pixels show strange behaviour, that can also be identified through the ARS Lock Ratio variable. Examples are given in Fig. 5.19. Although the detailed reasons of these pathological distributions is not fully understood yet, the ARS Lock Ratio provides an efficient way to identify and reject them. The origin of these problems are most probably due to an incorrect behaviour of the ARS analogue memories. In the case of unlocked ARS, only the affected channel (high or low gain) is excluded in the analysis. In the case of pathological distributions such as those shown here, both channels are rejected.

5.7.3 Automatic calibration check

In order to check the quality of the calibration procedure, we developed and implemented a precise statistical test, based on a detailed log-likelihood comparison of the charge distribution in pedestal events with expectations from pure noise.

The probability to observe a signal \(x\) when an intensity \(\mu\) is expected is given by:

\[
P(x|\mu) = \sum_n \frac{\mu^n e^{-\mu}}{n! \sqrt{2\pi(\sigma_p^2 + n\sigma_s^2)}} \exp\left(-\frac{(x - n)^2}{2(\sigma_p^2 + n\sigma_s^2)}\right)
\]

where \(\sigma_p\) is the width of the pedestal (NSB + electronic) and \(\sigma_s\) the resolution of the PMT. The log-likelihood and the goodness of fit read:

\[
\ln L = -2 \times \ln(P(x, \mu))
\]

\[
G = \langle \ln L \rangle - \langle \ln L \rangle
\]

where \(\langle \ln L \rangle\) is the expectation value of the log-likelihood. In the case of pure noise \((\mu = 0)\) this simplifies to:

\[
P(x|\mu = 0) = \exp\left(-\frac{x^2}{2\sigma_p^2}\right)
\]

and

\[
G = \frac{x^2}{\sigma_p^2} - 1
\]
Figure 5.20: Test of calibration using goodness estimator for properly calibrated data. **Left:** Evolution of Goodness $G$ with intensity. The filled area depicts the intensity distribution. **Right:** expected goodness distribution for pure pedestal events (convolution of left distribution with pedestal intensity distribution).

If the intensity distribution is consistent with pure noise, the distribution of $G$ should have a mean value of 0 and a RMS of $\sqrt{2}$. This is confirmed by simulations as shown in Fig. 5.20. Further simulations of incorrect calibration, such those show in Fig. 5.22 confirm that this statistical test is sensitive to an incorrect determination of the pedestal and/or the gain.

Figure 5.21: Sensitivity of the likelihood test to wrong estimation of the pedestal position (X axis) or width (Y-axis). **Left:** Evolution of the average goodness as a function of pedestal position and width; **Right:** Evolution of the goodness RMS. The black lines corresponds to a mean goodness within $[-0.1,0.1]$ and a RMS within $\pm 10\%$ from the average value of $\sqrt{2}$.

The sensitivity of the goodness test is shown in Fig. 5.21 as a function of estimated pedestal position with respect to true one (X axis), and as a function of estimated pedestal width for a true width of 1.6 p.e. (Y axis). The black lines correspond to a mean goodness within $[-0.1,0.1]$ and a RMS within $\pm 10\%$ from the average value of $\sqrt{2}$, values that are
Figure 5.22: Test of calibration using goodness estimator obtained from simulations. **Top:** the intensity is shifted up due to an incorrect determination of the pedestal. **Middle:** the intensity distribution is too narrow because of a wrong gain estimation. **Bottom:** the intensity distribution is too wide because of a wrong gain estimation. In each case, the mean value or RMS significantly deviates from the expected values.
used to estimate that a pixel is mis-operating. These plots show that the test is very sensitive to a mis-estimation of the pedestal width, but can accept a significant pedestal shift if it is accompanied with a mis-estimation of the width in a precise way. This combination is however very unlikely.

This test is implemented in the calibration procedure and helps identifying mis-operating pixels. An example of this is shown in Fig. 5.23 for one observation run. The left panel shows the distribution over the camera of the average value of the goodness for each pixel. The distribution of the pixel average goodness is centered on 0.06 with an RMS of 0.06. In this example, two pixels show a significant deviation of their average goodness value, around 0.64 and 1.2. This is confirmed by an analysis of the actual goodness distribution, shown on the right panel. Further investigation on this particular pixel showed that it was affected by a sudden rise of the night sky background luminosity, a sudden rise that was too rapid to be properly taken into account in the pedestal evolution.

![Figure 5.23: Calibration test using the goodness estimator.](image)

**Left:** The distribution of the pixel average goodness is centered on 0.06 with an RMS of 0.06. Two pixels show a significant deviation. **Middle:** Likelihood distribution for a correct pixel. **Right:** Likelihood distribution for one of the two pixels that show a significant deviation.

Further tests done on data from the blazar PKS 2155-304 are shown in Fig. 5.24. Pixels with incorrect average value of $G$ are displayed in red in the left panel, in high (resp. low) gain for the upper (resp. lower) panel. Two arcs are clearly visible, corresponding to a bright star rotating into the camera. The goodness procedure described here allows the time periods when the pedestal calculation is not sufficiently accurate to be easily identified and rejected from the data.

### 5.7.4 Summary on non-operating pixels

All algorithms described here operate on sliding windows of $\sim 1$ mn duration. When a particular problem is detected, the pixel is rejected from the analysis (only in the affected gain) for the duration of this sliding window extended with a security margin of 30 s before and after the event. Calibration files collecting the information about mis-operating pixels as a function of time are written for each observation run and are used in the reconstruction and analysis procedure.
5.8 Extensive tests of PMT Aging

The observation in the early years of H.E.S.S. operation of a continuous degradation of the overall efficiency (as shown in Fig. 5.17) led to extensive investigations of the possible causes of this loss. The fraction of photons detected depends on a number of factors including the collection coefficient of the Winston cone attached to the front of each photo-multiplier tube, the reflectivity of the mirrors covering the area of each dish and the quantum efficiency of the PMTs themselves. An attempt to clean the Winston cones showed a slight increase in the efficiency but could not account for the observed loss.

For that reason a project was set up, in the form of an Erasmus internship, to determine whether or not the quantum efficiency of H.E.S.S. phase 1 PMTs does drop due to long term exposure to bright light and if this drop can account for the observed decrease in efficiency.

The tests were conducted in a dark room with controlled light sources which allows to expose the PMTs twenty-four hours a day. Five test PMTs were exposed for the equivalent of a few observing years in Namibia and four control PMTs were used as a reference. These PMTs were not exposed to the high luminosity light source but had their gains and measured luminosity monitored carefully in the same way as the aging PMTs, in order to provide a reference measurement. The aging process consisted of exposing the test PMTs to a bright light of continuously monitored intensity for a given period of time and then measuring the evolution of the gain and the luminosity seen by both the test and the control PMTs.

The anode current was used as a measure of the efficiency evolution; For a given PMT it is calculated as:

\[ I = \frac{dN}{dt} \times \epsilon \times \gamma_{HG} \]  

(5.11)
where:

- $\frac{dN_\gamma}{dt}$ is the rate of incident photons
- $\epsilon$ is the overall PMT efficiency (product of quantum efficiency with collection efficiency)
- $\gamma_{HG}$ is the PMT gain

Thus, a photo-electron rate of 100 MHz with the nominal gain of $2 \times 10^5$, corresponds to an intensity of 3.2 $\mu$A. For an average annual observation time of 1200 h, the charge delivered by each PMT during one year of operation in Namibia is

$$L_{\text{year}} = 1200 \times 3600 \times 3.2 \times 10^{-6} = 13.8 \text{ C year}^{-1}$$

The comparison of the anode currents between a test PMT and a reference PMT provides a measurement of the evolution of their relative efficiencies, assuming that the ratio of luminosities is stable with time.

### 5.8.1 Experimental setup

A automated test-bench for PMT characterization were set up in LPNHE, based on the H.E.S.S. electronics, to allow for:

- measurement of gains and characterization of the relation between gain and high voltage
- pedestal characterization, including evolution with temperature and luminosity
- bi-dimensional characterization of photo-cathode response
- measurement of after-pulsing rate

![Figure 5.25: A schematic of the test-bench.](image-url)
5.8. **EXTENSIVE TESTS OF PMT AGING**

- ... 

A sketch of the test-bench is shown on Fig. 5.25 with a functional block diagram shown in Fig. 5.26. It consists of 4 H.E.S.S. drawers, each comprising 16 PMTs and associated electronics, and a full acquisition system based on the H.E.S.S. DAQ system.

A pulsed LED light source and a white light, associated with computer-controlled filter wheels with filters of varying densities and diffusers, illuminate the 64 PMTs approximately uniformly (within $\sim 10\%$). Computer controlled X-Y arms with a pulsed LED light source allow a precise mapping of the response of the photo-cathode with sub-millimeter resolution. The maximum intensity of the white light source corresponds to a photon rate of $\sim 1.5 \text{ GHz}$, corresponding to roughly ten to fifteen times the night sky brightness in Namibia.

![Block diagram for the test-bench.](image)

The test-bench is connected via a TCP-IP socket connection to the control room and all functions carried out by the test-bench are controlled by a network of Linux-PC's communicating with one another via the CORBA protocol. A central graphical user interface (GUI), developed in Python with the Gtk toolkit, controls all involved processes. The tests have been fully automatized, allowing a round-the-clock sequence of PMT illumination and performance measurements.

### 5.8.2 Experimental protocol

The following tests are done successively:

- LED stability tests, consisting of a pedestal run followed by one single photo-electron run (to measure the PMT gain) and a LED run with high intensity. This sequence of
runs allows a precise measurement of the LED luminosity multiplied by the efficiency (collection and quantum efficiency) on each PMT, using the ratio of the single photoelectron and pedestal peaks. Control PMTs are then used to derive a relative value of the LED luminosity, assuming that the efficiency of those PMTs is stable with time. These runs also give a continuous monitoring of the gains of the PMTs.

- Aging runs, consisting of a pedestal run followed by a long acquisition run with the white light switched on at an average photon rate of \( \sim 1.5 \) GHz. Because the white light takes about \( \sim 2 \) min to stabilize in intensity, the light is switched on before the run is actually started, with a black filter in front of it. The filter, mounted on a wheel, is then removed at run start. The aging runs typically last \( \sim 2 \) h, corresponding to \( \sim 30 \) h of exposure in Namibia.

After each aging run, two independent stability sequences are done at \( \sim 10 \) min interval, in order to measure the recovery capabilities of the PMT.

In addition to this sequence of runs, half of the PMTs used for aging tests had their gains readjusted to \( 2 \times 10^5 \) as soon as they showed a drop of about 20%.

### 5.8.3 Evolution of gain

Fig. 5.27 shows the evolution of the gain for two PMTs exposed to aging. The first one had its high voltage kept constant during the course of the tests while the second one had its gain periodically readjusted to the nominal value of \( 2 \times 10^5 \). A total integrated

![Figure 5.27: Gain against the integrated luminosity plots for drawer 304. Pixel 5 (down) had its gain readjusted during the experiment. The red dashed lines indicate the passing of an equivalent year. For the PMTs with an initial ADC separation of 200 ADC counts (pixel 4) we have applied the fits from equation 5.13 (blue line).](image)
luminosity of more than 100 C has been delivered of those PMTs, corresponding to several years of exposure in Namibia.

The evolution of gain with integrated luminosity \( L_I \) can be adjusted with an expression of the form:

\[
g = g_0 \exp \left( - \left( \frac{L_I}{L_0^p} \right)^p \right)
\]  

with typical characteristic luminosities of \( L_0^p \) ranging from 100 to 400 C depending on PMT and index \( p \) of the order of 0.2 to 0.5. The corresponding fit is shown as a blue line on Fig. 5.27 top.

It has to be noted as well that after significant illumination, PMTs recover a fraction (5 to 10%) of the lost gain with typical recovery time scales of \( \sim 100 \) h, but this recovery is only transient: this fraction is immediately lost again after new exposure to intense light.

### 5.8.4 Evolution of PMT efficiency

The evolution of PMT collection and quantum efficiency can be monitored using the ratio of the integrals under the single photo-electron peak and in the pedestal peak. This value, called *single pe fraction* is the average value of photo-electrons detected by event, and corresponds to the value of \( \mu \) in Eq. 5.5. It depends, among others, on the intensity of the flashing LED. The later was monitored using a control PMTs and did not show any significant variation (within 5%) during the course of the experiment.

![Figure 5.28](image)

Figure 5.28: Single photo-electron fraction against the integrated luminosity plots for drawer 304. The stability of the single photo-electron fraction indicates that the aging of the PMTs is not accompanied with a degradation of the quantum efficiency nor the collection efficiency.
The *single pe fraction* is shown in Fig. 5.28 as a function of the integrated luminosity for the same PMTs exposed to aging process. It does not show any significant variation during the course of the experiment, thus indicating that the loss of gain is not accompanied with a change of the quantum efficiency of the PMTs. The resolution on the *single pe fraction* is of the order of or better than 4% for all PMTs used in the test.

The loss of gain observed in the aging process is therefore most probably not due to degradation of the photo-cathode itself, but rather to dynode aging where the multiplication factor induces significant current.

### 5.8.5 Evolution of PMT resolution

The evolution of the PMT resolution is measured through the comparison of the single photo-electron peak width and the gain. This value is shown in Fig. 5.29 and shows no degradation with time; a slight improvement of a few % is even noticeable for most PMTs.

![Pixel Resolution vs Integrated Luminosity](image)

Figure 5.29: PMT resolution as a function of integrated luminosity for drawer 304.

### 5.8.6 Conclusion

Extensive aging test on H.E.S.S. PMTs showed that aging due to large integrated luminosities was not responsible for the degradation of the optical performances of the instruments. Much more likely are the degradation of the reflectivity of the mirrors and the Winston cones as well as degradation of the transparency of the PMT front window due to dust and sand abasing.

Cleaning test attempts on the Winston cones did not result into any significant improvement whereas replacement of mirrors during the year 2011 led to the recovery of about half of the lost performances, as shown in Fig. 5.17, the second half remaining still unexplained.
Conclusion on calibration

Atmospheric Cherenkov Telescopes are complicated devices, using cutting-edge technologies such as very fast digitization (at GHz frequencies), in a dusty and not very well controlled environment. Relying on a calorimetric measurement of the impinging $\gamma$ ray, the properties of the atmosphere (which is part of the detector) have to be monitored with the best possible precision. The calibration of atmospheric Cherenkov telescopes is a fundamental task that must be done with very great care, as the final resolution of the instrument depends strongly on the quality of the calibration.

Systematic effects due to variation of atmospheric conditions are estimated to be in the $10 - 20\%$ range for current-generation instruments. Following for instance the work done by the Auger collaboration to characterize the composition and layout of the atmosphere, next generation of instruments (such as CTA) aim at reducing significantly the systematic uncertainties and will therefore have to make a significant improvement in the quality of atmospheric monitoring and interpretation.
Chapter 6
Reconstruction

From the beginning of ground-based gamma-ray astronomy, data analysis techniques have been mostly based on the “Hillas parametrization” \[20\] (section 6.1) of the shower images, relying on the fact that the gamma-ray images in the camera focal plane are, to a good approximation, elliptical in shape. This is shown in Fig. 6.1 where the shower image on the left can probably be associated to a $\gamma$-ray induced shower (due to its regular shape), whereas the heterogeneity of the image on the right, with several clusters, points toward a shower induced by a hadronic cosmic ray.

![Image](image.png)

Figure 6.1: Examples of images from atmospheric showers as recorded by one of the H.E.S.S. camera. **Left:** the regular shape of the image indicates that it has most probably been produced by a $\gamma$-ray. **Right:** the irregular shape with several isolated cluster points toward an hadronic shower

The goals of the reconstruction are two fold:

1. reconstruct the parameters of the primary particle: direction, impact location and energy

2. provide discriminating variables to reduce the enormous amount of background events induced by isotropic, charged cosmic rays. Recent reconstruction techniques reach rejection factors of the order of $10^4$, where a factor of $\sim 10^2$ is provided by the discriminating parameters constructed from the shower image shape, and an additional factor of $\sim 10^2$ is provided by the precise direction reconstruction from
which allows the number of events from the direction of the observed source can be compared with events from the neighbouring, empty fields.

A major step, stereoscopy, was introduced in 1997 by the HEGRA collaboration [21]. The same event is recorded by several telescopes simultaneously, thus allowing a simple, geometric reconstruction of the shower direction and impact parameter (Section 6.1.2). More elaborate analysis techniques were pioneered by the work of the CAT collaboration [50] on a model analysis technique, where the shower images are compared to a realistic pre-calculated model. This has been further developed with a more precise model of showers, which depends on the altitude of the first interaction, and the introduction of a complete modelling of the background using a log-likelihood approach [22] (Section 6.3). Other analysis techniques, such as the 3D Model analysis [64, 65] (Section 6.2) were developed recently with the start of the third-generation telescopes. The 3D Model analysis is, for instance, based on the assumption of a 3 dimensional elliptical shape of the photo-sphere, that is adjusted on the observed images simultaneously on all telescopes.

These analysis techniques are complementary in many senses. They are sensitive to different properties of the shower, and can therefore be used to cross-check the analysis results or be combined together to improve the sensitivity.

More recently, multi-variate techniques have been introduced to improve the rejection capabilities of the various methods; they also led to significant improvement in sensitivity [66, 67, 68, 69].

6.1 Hillas Parameters

In a famous paper of 1985 [20], M. Hillas proposed to reduce the recorded images to a few parameters, reflecting the modelling of the image by a two-dimensional ellipse. These parameters, shown on figure 6.2 are usually:

- length $L$ and width $w$ of the ellipse
- size (total image amplitude)
- nominal distance $d$ (angular distance between the centre of the camera and the image centre of gravity)
- azimuthal angle of the image main axis $\phi$
- orientation angle $\alpha$

![Figure 6.2: Geometrical definition of the Hillas Parameters.]

This parametrization can be expressed analytically from the measured intensity $q_i$ and the position $(x_i, y_i)$ of each pixel, which makes it fast and easy to implement. Starting from the first and second moments:

$$\langle x \rangle = \frac{\sum_i x_i q_i}{\sum_i q_i}, \quad \langle y \rangle = \frac{\sum_i y_i q_i}{\sum_i q_i}$$

(6.1)
\[
\langle x^2 \rangle = \frac{\sum_i x_i^2 q_i}{\sum_i q_i}, \quad \langle y^2 \rangle = \frac{\sum_i y_i^2 q_i}{\sum_i q_i}, \quad \langle xy \rangle = \frac{\sum_i x_i y_i q_i}{\sum_i q_i}
\]  \hspace{1cm} (6.2)

and the variances and covariances:

\[
\sigma_{x^2} = \langle x^2 \rangle - \langle x \rangle^2, \quad \sigma_{y^2} = \langle y^2 \rangle - \langle y \rangle^2, \quad \sigma_{xy} = \langle xy \rangle - \langle x \rangle \langle y \rangle
\]  \hspace{1cm} (6.3)

we define the intermediate variables:

\[
\chi = \sigma_{x^2} - \sigma_{y^2}, \quad z = \sqrt{\chi^2 + 4\sigma_{xy}} \hspace{1cm} (6.4)
\]

\[
b = \sqrt{(1 + \chi/z) \langle x \rangle^2 + (1 - \chi/z) \langle y \rangle^2 - 2\sigma_{xy} \langle x \rangle \langle y \rangle}/2
\]

and the Hillas parameters simply read:

\[
\begin{align*}
  d &= \sqrt{\langle x \rangle^2 + \langle y \rangle^2} \\
  L &= \sigma_{x^2} + \sigma_{y^2} + z \\
  w &= \sigma_{x^2} + \sigma_{y^2} - z \\
  \alpha &= \arcsin \left( \frac{b}{d} \right) \hspace{1cm} (6.5)
\end{align*}
\]

In the determination of the Hillas parameters, only the pixels participating in the shower image need to be taken into account. This reconstruction method therefore relies on an image cleaning procedure, which is usually based on a dual-threshold algorithm: pixels with intensity above a threshold \(q_1\) are kept only if they have at least a neighbour with intensity above a threshold \(q_2\), and vice-versa. In H.E.S.S., the cleaning threshold are respectively set to \(q_1 = 7\ \gamma_e\) and \(q_2 = 10\ \gamma_e\). Moreover, this parametrization is subject to biases caused by missing (or mis-operating) pixels, which significantly degrade the estimation of the parameters. In particular images located close to the edge of the camera are often truncated and are then mis-reconstructed because of mis-identification of main axis (Fig. 6.3). Therefore, Hillas-parametrization-based reconstruction methods reject these images by applying a selection on the distance of the image barycentre to the centre of the camera.

### 6.1.1 Single telescope reconstruction

In single telescope observations, the shower direction is estimated from the Hillas parameters themselves (and in particular from the image length and size), either with look-up tables or with ad-hoc analytic functions. But the choice of a symmetrical parametrization of the shower leads to degenerate solutions, on each side of the image centre of gravity along the main axis.

In order to break this degeneracy, other parameters — based in particular on the third order moments (skewness and kurtosis) — were added later.

The shower energy is usually estimated with a similar technique, from the image size and nominal distance.
6.1. HILLAS PARAMETERS

6.1.2 Stereoscopic reconstruction

The stereoscopic imaging technique, pioneered by HEGRA [21], provides a simple geometric reconstruction of the shower: the source direction is given by the intersection of the major axes of the shower images in the camera (Fig. 6.4), and the shower impact point is obtained in a similar manner, using a geometrical intersection of the planes in the sky containing the telescopes and the shower track (Fig. 6.5). The energy is then estimated from a weighted average of each single telescope energy reconstruction.

Figure 6.4: Geometric reconstruction of shower direction in stereoscopic mode. In the camera frame, the main axis of the shower corresponds to a plane on the sky that contains the actual shower track. The primary particle direction corresponds to a point on this main axis. The intersection of the main axis of the images recorded by the different telescopes immediately provides the primary particle direction.
6.1.3 Gamma-hadron separation

The Hillas parameters not only allow the reconstruction of the shower parameters, but also can provide some discrimination between γ ray candidates and the much more numerous hadrons. Several techniques have been developed, exploiting to an increased extent the existing correlation between the different parameters (e.g. 

Supercuts \cite{70}, Scaled Cuts \cite{21} and Extended Supercuts \cite{71}). We will use here the Scaled Cuts technique, in which the actual image width \( w \) and length \( l \) are compared to the expectation value and variance obtained from simulation as a function of the image charge \( q \) and reconstructed impact distance \( \rho \), expressed by two normalized parameters Scaled Width(SW) and Scaled Length(SL):

\[
SW = \frac{w(q, \rho) - \langle w(q, \rho) \rangle}{\sigma_w(q, \rho)}, \quad SL = \frac{l(q, \rho) - \langle l(q, \rho) \rangle}{\sigma_l(q, \rho)}
\]  

These parameters have the noticeable advantage of being easily combined in stereoscopic observations in Mean Scaled Width and Mean Scaled Length:

\[
MSW = \frac{\sum_{\text{tels}} SW}{\sqrt{\text{ntels}}}, \quad MSL = \frac{\sum_{\text{tels}} SL}{\sqrt{\text{ntels}}}
\]  

From simulations, one can show that the Mean Scaled Width and Mean Scaled Length are almost uncorrelated for γ candidates \( (\rho = 0.15 \pm 0.01) \) and can therefore be combined in a single variable Mean Scaled Sum (MSS): \( MSS = (MSW + MSL)/\sqrt{2} \). Distributions
Figure 6.6: Distribution of the Mean Scaled Width and Mean Scaled Length parameters for simulated $\gamma$-ray spectra (red), real background events (blue) and real $\gamma$-rays originating from the Crab Nebula (green).

of the Mean Scaled Width and Mean Scaled Length parameters are shown in Fig. 6.6 for simulated $\gamma$-ray spectra in red, real background events in blue and real $\gamma$-rays originating from the Crab Nebula in green. This figure shows that the distributions are well understood.

The main interest of the Mean Scaled Parameters comes from the fact that they provide a uniform selection, because they are scaled by the variance of the parameters in the denominator. In particular, the selection efficiency is almost independent of energy, and will therefore not bias the energy spectrum determination.

### 6.1.4 Energy Reconstruction

In a similar manner, look-up tables are used to reconstruct the energy of the primary particle as a function of image amplitude and distance to the impact parameter. These tables allow the relations that exists between the true primary energy, the impact parameter and measured signal intensity to be inverted numerically.

These tables are constructed from simulations using the following procedure: for each simulated shower, the shower impact point is first reconstructed, as well as the image intensity in each telescope. Events are then binned by reconstructed distance to the telescope and image intensity, and for the corresponding data set, the average value and variance of the true energy is computed. This procedure is repeated for 20 zenith angle bands, 7 optical efficiency bands and 6 off-axis angle bands. An interpolation procedure is then used to estimate the energy of the primary particle independently for each telescope, but using the impact parameter reconstructed using the full array. An averaging over all telescopes is then used, weighted by the uncertainty on the individual energy measurement, to provide the final energy estimation. An example of such a look-up table is shown in Fig. 6.7.

This look-up approach permits to reach typical energy resolutions of the order of 20%, with however significant biases at low and high energies.
### 6.1.5 Possible improvements

Several authors discussed possible improvement of the Hillas stereoscopic reconstruction. Four algorithms were in particular proposed [72]:

(a) Intersecting pairs of image axes, followed by an averaging over intersection points.
(b) Intersecting image axes taking into account the errors on image location and image orientation, resulting in an error ellipse for the image of the source.

(c) Using in addition the width/length-ratio to constrain the source image to two regions on either side of an image.

(d) Optimizing the shower geometry such that the predicted image axes best match the observed images.

However, these algorithms result in only modest improvements of the sensitivity, of the order of 10 – 20%, compared to the simple approach.

6.2 3D Model analysis

6.2.1 Introduction

The second analysis presented here, the 3D Model Analysis [64, 65], is a kind of 3 dimensional generalization of the Hillas parameters: the shower is modelled as a Gaussian photo-sphere in the atmosphere (with anisotropic light angular distribution), which is then used to predict — with a line-of-sight path integral — the light collected in each pixel. A comparison of the actual image to the predicted one (with a log-likelihood function) allows eight shower parameters to be reconstructed: mean altitude, impact point and direction, 3D width and length and luminosity.

![Figure 6.9: Geometrical definition of the 3D Model Parameters.](image)

6.2.2 Energy reconstruction

One of the major improvements of the 3D Model reconstruction over the Hillas reconstruction is that the relation between the actual image amplitude in the telescopes and the total shower luminosity is encoded in an unequivocal way. The altitude of the photo-sphere
is another estimator of the energy of the shower, as high energy showers penetrate deeper into the atmosphere than low energy ones. The best estimator of the primary energy is thus the luminosity of the Gaussian photo-sphere. This is illustrated in Fig. 6.10 which shows, for different primary energies, the average photo-sphere depth as a function of its luminosity.

![Diagram](image)

Figure 6.10: Depth of development (Y) as a function of photo-sphere luminosity (X) for different primary particle energies: 200 GeV, 500 GeV, 1 TeV and 5 TeV at zenith.

6.2.3 Gamma-Hadron separation

The 3D-Model analysis is based on the strong assumption of a rotational symmetry, which is used to reject about 70% of the hadrons during the fit procedure. For the remaining events, the most powerful parameters for discriminating between the \( \gamma \) ray candidates and the hadrons is found to be the 3D shower width, as hadronic showers are typically much wider than the electromagnetic ones. The shower width, expressed in units of radiation length, is found from simulations to be proportional to the slant thickness. This is used to define a zenith-angle-independent Reduced-width parameter:

\[
\omega = \frac{w \times \rho(z_{\text{max}})}{D}, \quad (6.8)
\]

where the slant thickness \( D \) reads:

\[
D = \frac{1}{\cos \theta} \int_{z_{\text{max}}}^{\infty} \rho(z) dz \quad (6.9)
\]
6.3 Model Analysis

A new, sophisticated gamma-ray likelihood reconstruction technique has been developed [22] for Imaging Atmospheric Cherenkov Telescopes. The technique is based on the comparison of the raw images of showers with the predictions from a semi-analytic model. The approach was initiated by the CAT experiment in the 1990's [54], and has been developed further to include a new fit algorithm based on a log-likelihood minimization using all pixels in the camera, a precise treatment of night sky background noise, the use of stereoscopy and the introduction of first interaction depth as a parameter of the model.

The reconstruction technique provides a more precise direction and energy reconstruction of the photon induced shower compared to other techniques in use, together with a better $\gamma$ ray efficiency, especially at low energies, as well as an improved background rejection. For data taken with the H.E.S.S. experiment, the reconstruction technique yields a factor of $\sim$2 better sensitivity compared to the standard H.E.S.S. reconstruction techniques based on second moments of the camera images (Hillas Parameter technique).

6.3.1 Principles

The shower images generated from the semianalytical model are derived from the Cherenkov light distribution of charged particles in electromagnetic showers taking into account light collection efficiency, atmospheric absorption etc. The Cherenkov light distribution of a shower is determined by the longitudinal, lateral, and angular distributions of charged particles in the shower. These distributions are derived from Monte Carlo simulations and parametrized to yield an analytic description of the shower, i.e. the shower model, including the depth of the first interaction as a new parameter in the parametrization. Additionally, the contribution of the night sky background noise in the camera in every pixel is modelled on the basis of a detailed statistical analysis.

The light density due to a shower in the camera can be calculated from an eight-dimensional integral:

- integral over altitude $z$ or depth $t$ (longitudinal development of shower),
- integral over energy of the electron/positron in the shower,
- integral over electron/positron direction with respect to the telescope ($u$ and $\phi$).
- integral over electron/positron position with respect to its direction and the shower main axis ($X_r$ and $Y_r$).
- integral over Cherenkov photon wavelengths,
- integral over Cherenkov photon azimuthal angle around the electron (the angle between the electron and the Cherenkov photon being fixed for a given electron energy).

\[
I(x, y) = \int dz \int dE \frac{dN_e}{dE} (t, E) \frac{dt}{dz}(y) \\
\int du \times F_u(u(E, s)) \int \frac{d\phi}{2\pi}
\]

(6.10)
\[
\int dX_r \int dY_r F_{XY}(X_r, Y_r, E, s, u) \\
\int d\phi_{ph} \int \frac{d\lambda}{\lambda^2} \frac{d^2 n_{\gamma}}{d \lambda d \gamma} \times \exp(-\tau(z, \lambda)) \times Q_{\text{eff}}(\lambda) \\
\times \text{Col}(z, X_r, Y_r, u, \phi, \phi_{ph})
\]

Where:

- \(dN_c/dE(t, E)\) is the energy dependant longitudinal distribution of charged particles in the shower as a function of atmospheric depth from first interaction,
- \(s\) is the shower age, expressed from depth and energy,
- \(u\) is the angle of the particle with respect to shower main axis and normalized to the average angle for the corresponding age and energy,
- \(F_u(u(E, s))\) is the normalized angular distribution of particles,
- \(F_{XY}(E, s, u)\) is the normalized lateral distribution of particles,
- \(1/\lambda^2 \times d^2 n_{\gamma}/(\cos \theta d z d \lambda)\) is the Cherenkov photon production rate (per unit of vertical track length and emitted photon wavelength) for an electron angle \(\theta\) with respect to the shower axis,
- \(\exp(-\tau(z, \lambda))\) is the atmospheric absorption (\(\tau\) being the optical depth),
- \(Q_{\text{eff}}(\lambda)\) is the detector quantum efficiency (multiplied by mirror reflectivity and other wavelength-dependent transmission coefficients in the detector),
- \(\text{Col}(z, X_r, Y_r, u, \phi, \phi_{ph})\) is the average geometrical collection efficiency for photons emitted by a electron at position \((X_r, Y_r, z)\) with direction defined by \((u, \phi)\), and with an azimuthal photon angle \(\phi_{ph}\) around the electron.

The corresponding distributions have been calibrated from simulations, and are described in [22]. In addition to the aforementioned ingredients, instrumental effects such as the point spread function and the electronic response of the camera, including in particular trigger response and integration duration, have to be taken into account in the calculation.

These effects, as well as the geometric light collection efficiency are obtained from a detailed simulation of the telescope response and parametrized in look-up tables. Atmospheric absorption of light, wavelength-dependent quantum efficiency and reflectivity used in the model generation are also implemented as look-up tables.

### 6.3.2 Example

The output of the model generation procedure is a bank of two-dimensional shower images in the frame of a perfect camera, with very small \((0.01^\circ)\) pixel size. For a given set of primary particle parameters, a predicted image is computed using an interpolation procedure on the generated images in a 4 dimensional space (energy, impact distance, primary interaction depth and zenith angle). Shower direction and azimuthal angle are then taken into account as a rotation and a translation in the camera frame to compute the final predicted image. Examples of such two dimensional shower images are shown in Fig. 6.11.
Figure 6.11: Model of a 1 TeV shower started at one radiation length and falling 20 m (top-left), 100 m (top-right) and 250 m (bottom-left) away from the telescope. X and Y axes are in units of degrees in the camera frame. Bottom-right: same as top-right but with a first interaction point located deeper in the atmosphere (at 3 $X_0$). Note that the vertical scale (image amplitude) differs.
Figure 6.12: Comparison between 1 TeV simulated shower images (black) at zenith and model prediction (red). **Left:** image amplitude as a function of impact distance for primary interaction point of one radiation length. **Right:** image length (top) and width (bottom), in units of milliradians, estimated with the standard Hillas parametrization technique, and as a function of impact distance.

### 6.3.3 Comparison with simulation

A comparison between the image shapes from simulation and model prediction is shown in Fig. 6.12 for 1 TeV γ-ray showers. The images were calculated for a H.E.S.S. camera, with pixels of 0.16° diameter. The image length and width were estimated using the standard Hillas parametrization applied to the predicted images. In each plot, the average value of the simulation is drawn as a black histogram, with error bars indicating the shower-to-shower fluctuations (spread). The model prediction (image of average shower) is represented by a solid thick red line. The agreement between the model and the average values of the simulation is excellent up to core distance of about 300 m, where trigger effects explain the differences: at this distance, the total image amplitude does not exceed 100 photo-electrons, distributed over many pixels. Showers fluctuating up to higher intensities are more likely to trigger the telescope, thus resulting in a higher average image amplitude in the simulation compared to the model. The bars in the simulation histograms (in black) are due to shower-to-shower fluctuations, which are not taken into account in the model generation. At small core distances, and when taking into account the evolution with primary interaction depth, the shower intensity fluctuates by about 20% from shower-to-shower in addition to the fluctuation related to the depth of first interaction.

### 6.3.4 Fit procedure

Once a shower model has been obtained, actual images on the camera can be compared to those predicted by the shower image model for a given set of primary parameters. A minimization procedure is then used to obtain the most likely parameters of the incoming particle (energy, direction, impact, depth of the first interaction) under the hypothesis that the primary particle is a γ ray. The minimization procedure involves a precise comparison of the intensity in each pixel of the camera with the prediction from the model (interpolated between grid points to the actual set of parameters). In order to take into account the Poisson nature of the detected number of photons in each pixel, a log-
likelihood approach has been chosen. The fit procedure uses all pixels in the camera and does not require a dedicated image cleaning procedure to extract the pixels illuminated by the shower. The parameters of the calculated shower that best fit the measured shower image are determined in a minimization procedure which yields a selection criteria to discriminate $\gamma$ ray induced showers from the hadronic background.

The probability density (likelihood) to observe a signal of $s$ (in units of photo-electrons) in a pixel for an expectation value $\mu$ is given by the convolution of the Poisson distribution of the photo-electron number $n$ with the photo-multiplier resolution. The latter is well represented by a Gaussian of width $\sqrt{\sigma_p^2 + n\sigma_\gamma^2}$ where $\sigma_p$ is the width of the pedestal (charge distribution under pure noise, including night sky background) and $\sigma_\gamma$ the width of the single photo-electron peak (photo-multiplier resolution):

$$P(s|\mu, \sigma_p, \sigma_\gamma) = \sum_n \frac{\mu^n e^{-\mu}}{n!} \frac{1}{\sqrt{2\pi(\sigma_p^2 + n\sigma_\gamma^2)}} \exp \left( -\frac{(s - n)^2}{2(\sigma_p^2 + n\sigma_\gamma^2)} \right)$$  \hspace{1cm} (6.11)

If the observed signal is only due to noise (night sky background), the probability density functions simplifies to a Gaussian of width $\sigma_p$. At high $\mu$, the Poisson distribution can be replaced by a Gaussian of width $\sqrt{\mu}$, and the probability density function can be well approximated by the convolution of two Gaussians:

$$P(s|\mu \gg 0, \sigma_p, \sigma_\gamma) \approx \frac{1}{\sqrt{2\pi \sigma_p^2 + \mu(1 + \sigma_\gamma^2)}} \exp \left( -\frac{(s - \mu)^2}{2(\sigma_p^2 + \mu(1 + \sigma_\gamma^2))} \right)$$ \hspace{1cm} (6.12)

In both case, for null intensity and large $\mu$, the average value of the pixel log-likelihood, $\ln L = -2\ln P$, reads

$$\langle \ln L \rangle |_\mu = \int ds \ln L(s|\mu = 0, \sigma_p) \times P(s|\mu = 0, \sigma_p)$$
$$= \int ds \left( \ln(2\pi) + \ln \sigma_p^2 + \frac{s^2}{\sigma_p^2} \right) \frac{1}{\sqrt{2\pi\sigma_p^2}} \exp \left( -\frac{s^2}{2\sigma_p^2} \right)$$ \hspace{1cm} (6.13)

$$= 1 + \ln(2\pi) + \ln \sigma_p^2$$

Pixels in a camera are assumed to be independent. We define the telescope log-likelihood as the sum over all pixels of the pixel log-likelihood:

$$\ln L_{tel} = \sum_{\text{pixel } i} \ln L_i = \sum_{\text{pixel } i} -2 \times \ln P(s_i|\mu_i, \sigma_p, \sigma_\gamma).$$  \hspace{1cm} (6.14)

### 6.3.5 Fit procedure and Goodness-of-fit

The model photon reconstruction relies on the pixel-per-pixel comparison of the actual shower images with those that are predicted for a given set of parameters. A minimization procedure is used to find the best parameters (direction, impact, depth of the first interaction and energy). In contrast to Hillas-parameter based reconstruction techniques, the raw image amplitudes are directly used, without any image cleaning procedure. All pixels are used in the fit, not only those close to the actual image.
A Levenberg-Marquardt fit algorithm \([73,74]\) is used to minimize the telescope log-likelihood (eq.\([6.14]\)). This algorithm is very efficient in the case that the first and second derivative of the minimized function (log-likelihood or \(\chi^2\)) can be expressed analytically and depend mostly on the first derivative of the model (the second derivative being negligible). This is in general valid when the minimized function is a quadratic form (\(\chi^2\) or similar) and when the model function exhibits a smooth behaviour when varying the model parameters.

The output of the minimization procedure are:

- Best guess of the 6 shower parameters: direction (2 parameters), impact (2 parameters), depth of the first interaction and energy
- Correlation matrix and therefore uncertainty on these best fit parameters
- Final log-likelihood value

Since Atmospheric Čerenkov Telescopes are background dominated systems, the performance of any analysis is mainly driven by its ability to discriminate between gamma-ray induced showers and the much more numerous hadronic background. In the Model Analysis, a goodness-of-fit approach is chosen to compare the model prediction and the actual shower images, in order to check the compatibility of the recorded events with a pure \(\gamma\)-ray hypothesis. The goodness-of-fit is defined as a normalized sum over all pixels of the difference between the actual pixel log-likelihood and its expectation value, properly normalized:

\[
G = \sum_{\text{pixel } i} \left[ \ln L(s_i|\mu_i) - \langle \ln L \rangle |_{\mu_i} \right] \over \sqrt{2 \times \text{NDoF}}
\]  

(6.15)

where NDoF is the number of degrees of freedom (number of pixels - 6). The goodness of fit behaves asymptotically like a \(\chi^2\) and therefore provides a measure of the fit quality. This is used later in the hadronic discrimination. By construction, if the pixel likelihoods behave like independent random variables, \(G\) is expected to behave like a normal variable.

### 6.3.6 ShowerGoodness and BackgroundGoodness

The discrimination between \(\gamma\) ray-induced showers and hadron-induced showers can use several distinct features:

- Hadron-induced showers are more irregular, and contain several electromagnetic sub-showers initiated by disintegration of neutral pions. As a consequence, the image in a Čerenkov camera often exhibits several well separated clusters.

- In addition, the hadronic component of the shower itself emits a low intensity Čerenkov light spread over a large fraction of the camera. This emission, denoted as Hadronic min, is in general eliminated by cleaning procedures used in standard reconstruction techniques.

- Finally, the charged nucleus entering the atmosphere can emit Čerenkov light before interacting. This emission is produced very high in the atmosphere and therefore generates a faint spot in the camera, close to the shower direction.
6.3. MODEL ANALYSIS

In order to fully exploit the differences between γ ray and hadron-induced showers, individual pixels contributing to the goodness-of-fit are classified into two different groups at the end of the fit:

- Pixels belonging to the shower core, defined as pixels whose predicted amplitude is above 0.01 p.e., are grouped together with three rows of neighbours around them to construct a variable named ShowerGoodness (SG) in a similar way to eq. [6.15]. These pixels are selected at the end of the fit procedure to avoid changes of the number of degrees of freedom during the fit. Due to the large reduction of the number of degrees freedom, this variable is more sensitive than the Goodness to discrepancies between the model prediction and the actual shower images.

- The remaining pixels, denoted as background pixels, are grouped together to construct a variable named BackgroundGoodness (BG), which is sensitive to hadronic clusters outside the main image, hadronic rain and other irregularities.

6.3.7 Performance

Several configurations have been defined for the Model Analysis, and are summarized in Tab. 6.1.

<table>
<thead>
<tr>
<th>Name</th>
<th>Min. Charge (p.e.)</th>
<th>Max. Nom. Distance (deg.)</th>
<th>#Tels</th>
<th>(SG_{\text{max}})</th>
<th>(t_0 (X_0))</th>
<th>(\theta_{\text{max}}^2) (deg²)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Standard</td>
<td>60</td>
<td>2</td>
<td>2</td>
<td>0.6</td>
<td>[-1, 4]</td>
<td>0.01</td>
</tr>
<tr>
<td>Faint Source</td>
<td>120</td>
<td>2</td>
<td>2</td>
<td>0.4</td>
<td>[-1, 4]</td>
<td>0.005</td>
</tr>
<tr>
<td>Loose Cuts</td>
<td>40</td>
<td>2</td>
<td>2</td>
<td>0.9</td>
<td>N/A</td>
<td>0.0125</td>
</tr>
</tbody>
</table>

Table 6.1: Reconstruction configurations: Minimum charge, maximum nominal distance, maximum ShowerGoodness, primary interaction depth range and maximum squared angular distance from the reconstructed shower position to the source. A minimum of two telescopes passing the per-telescope cuts on image amplitude and distance from the centre of the field of view, are also required.

The effective area as a function of energy for the Model Analysis at zenith is shown in Fig. 6.13 compared to areas obtained with Hillas-parameter-based reconstructions. The reconstruction efficiency for the standard configuration is similar to the values obtained with Hillas reconstruction with a minimum image amplitude of 60 p.e.. As expected, the loose configuration has a larger effective area and a lower threshold. In most of the H.E.S.S. publications to date, a minimum image amplitude of 200 p.e. was used, yielding a much larger threshold of ∼ 300 GeV at zenith.

The energy resolution (after cuts) as a function of energy is shown in Fig. 6.14 for zenith, where the energy resolution is defined as the RMS of the \(\Delta E/E\) distribution. The energy resolution is better than 15% for the whole energy range (from 80 GeV up to more than 20 TeV), with biases not exceeding 5% in the central range. In the very central energy range (500 GeV to more than 10 TeV), the energy resolution is better than 10% and reaches values as low as 7 to 8%. Larger energy biases appear at very low energy (up to 20% at 80 GeV), due to trigger selection effects. These biases are however smaller than those obtained with an Hillas parameters based reconstruction. In a similar manner, negative energy biases appear at very high energies, due to very distant
Figure 6.13: Effective area as a function of energy, at zenith, compared with the values obtained for the standard and hard cuts Hillas parameters based analyses.

Figure 6.14: Energy resolution (main plot) and bias (inset) as a function of energy, at zenith, compared with the values obtained for the standard and hard cuts Hillas parameters based analyses (resp. blue circles and triangles).
high energy showers reconstructed closer to the telescope and with an underestimated energy. In the medium energy range (500 GeV to a few 10 TeV), the Model Analysis largely outperforms standard reconstruction techniques.

The Model Analysis provides an uncertainty estimation for each reconstructed parameter, as part of the correlation matrix. This can be used to select in a simple and natural way a sub-sample of the events with improved angular or energy resolution (for more precise morphological or spectral analysis).

The depth of the first interaction being a parameter of the model, is also a direct product of the reconstruction procedure (instead of being calculated, as in the Hillas parameters based reconstruction, from the shower maximum).

Fig. 6.15 shows the ability of the model analysis to reconstruct the depth of the first interaction with almost no bias and a resolution of 0.7X₀ (slightly worse at large zenith angles). This is better than Hillas parameters based analyses, which obtain resolution not better than 1 X₀ in the best cases.

![Reconstructed Primary Depth](image)

Figure 6.15: Distribution of reconstructed primary depth for ON events, OFF events and γ-rays from the Crab Nebula (green). The black line represents the convolution of a exponential distribution (for the true primary depth) with a Gaussian distribution of width 0.7X₀. The reconstructed distribution is compatible with a resolution of 0.7X₀ over the whole energy range.

The precise reconstruction of the primary depth not only improves the energy resolution, but also provides a new discriminating parameter that allows to separate the γ-ray induced showers from the much more numerous hadronic induced one; it could even, on a statistical basis, help to reject the electronic background at low energies, since showers induced by electrons start in average 0.5X₀ higher in the atmosphere than those induced by γ rays.

---

1Very distant shower produce almost parallel images in the camera, which introduces a degeneracy in the reconstruction.
6.4 Comparison of reconstruction techniques

6.4.1 Correlations between discriminating variables

The three reconstruction techniques presented here exploit different aspects of the showers. In particular, the Hillas reconstruction, through the use of extensive simulations that are used to construct the Scaled Parameters, take into account the shower-to-shower fluctuations, but does not properly take into account the correlations between the telescopes. In contrast, the 3D Model Analysis, assuming that different telescopes have just a different view on the same photo-sphere, naturally takes into account the correlations between the different telescopes, but does not treat properly the shower-to-shower fluctuations. The Model analysis, relying on a detailed model of shower development, takes into account a fraction of the shower-to-shower fluctuations (through the incorporation of the depth of first interaction) as well as correlations between telescopes.

Figure 6.16 shows the distribution of Model goodness-of-fit versus Hillas Mean Scaled Goodness and 3D Model rescaled width for simulated γ's (left) and real OFF data (right). Surprisingly, the discriminating parameters appear to be almost uncorrelated for γ-rays although some clear correlation appears for background events.

This clearly indicates that no analysis so far exploits the complete available information, and that further improvement might be possible. At the first level, combination of discriminating variables, using multi-variate techniques, provides a first approach to improve the sensitivity of the instruments, at the cost of having a larger sensitivity to systematic effects.

6.4.2 Sensitivity to varying level of NSB

The variation of the efficiency to γ-rays with respect to the Night Sky Background level (NSB) is shown for the three analyses considered in figure 6.17. The efficiency of the Hillas and the 3D Model analyses both start to drop quickly above 200 MHz, whereas the efficiency of the Model analysis is much flatter, due to its complete treatment of the NSB level in the goodness-of-fit parameter. The bulk of the H.E.S.S. observations corresponds to an average NBS level of 100 MHz, varying between 40 MHz outside the Galaxy to 300 MHz in the most luminous parts of the Galactic Plane.

6.4.3 Off-axis observations

Figure 6.18 shows the relative evolution of the efficiency for γ-rays as a function of the OFF-axis angle (distance of the shower axis to the centre of the camera) for the three analyses. As expected, the efficiency of the Hillas analysis starts to fall off before the others: the efficiency of the Mean Scaled Width/Length parameters degrades quickly due to truncated showers. Neither the Model nor the 3D Model analysis relies on the actual — truncated — images but rather extrapolate the available information and have therefore a flatter efficiency.

6.5 Conclusions

As was the case for the calibration, the quality of the reconstruction is essential to improve the sensitivity of Atmospheric Cherenkov Telescopes. The level of background discrimi-
Figure 6.16: Correlations between different discriminating variables.
Figure 6.17: Sensitivity of the different reconstruction techniques to varying NSB. From [73]

Figure 6.18: Efficiency as a function of OffAxisAngle. From [73]
nation, and thus the ability to extract the signal, depends directly on the precision of the reconstruction. Recent work, such as that which we did on the Model Analysis, resulted in an improvement by a factor of $\sim 2$ of the sensitivity of the instrument. To illustrate the importance of this improvement, we recall the fact that a gain of a factor of 2 in sensitivity is more or less equivalent to a multiplication by four of the number of telescopes. In addition, stability of the discriminating parameters against variation of external parameters, such as variation of the NSB level across the field of view, is absolutely essential to allow a proper background subtraction, as will be shown in the next chapter.

However, it appears that the story is yet not over. Detailed comparison of reconstruction techniques shows that the discriminating variables they provide are only poorly correlated. Recent work on combining these variables using numerical approaches (Neural networks, boosted decision trees, etc.) \cite{67,68,69} or analytic parametrization \cite{70} led to further steps in sensitivity improvement, however at the expense of a larger analysis complexity.

The definitive reconstruction method most probably remains to be invented, and significant improvements in the fields are to be expected in the next years. The use of timing, pioneered by the MAGIC and VERITAS although with little success, could be another handle, and will be discussed in chapter \ref{ch:timing}.
Chapter 7

Analysis

7.1 Introduction

Imaging Atmospheric Cherenkov Telescopes are highly background dominated instruments. Moreover, observations are split in runs of 28 mm each (in the case of H.E.S.S.), performed in different pointing or weather conditions. The instrument response varies notably with observing conditions, and in particular with zenith angles (as the depth of atmosphere increases quickly with zenith angle). Last but not least, a large number of runs usually need to be accumulated, spread over several months or even years, to be able to extract a significant signal.

The role of the analysis software is to combine observations taken under different conditions, accumulate them, and provide a precise background estimation. This requires in particular to be able to predict the expected level of background in any observation; in the absence of a calibration beam, this must be achieved from the data themselves.

Once the signal has been detected definitively, further stages in the analysis consist in:

- extraction of an energy spectrum (or energy distribution of the $\gamma$-rays), which often constitutes a valuable hint of the underlying acceleration mechanism,
- study of the evolution of the source luminosity at different time scales ($light curves$) and eventually search for repetitive processes,
- study of the spatial distribution of the events ($morphology$) which also often constrains the source emission mechanisms,
- and in some cases, combination of these possibilities in the form, for example, of an energy resolved morphology. This requires more statistics, but has proved to be particularly valuable.

Due to the large level of background, special care must be given on the precise statistical treatment of the collections of events all along the process.

A flexible analysis chain, named $ParisAnalysis$, has been designed to automatize the analysis steps in a simple and powerful way. It is able to work with any kind of reconstruction technique, cross-correlate them in an easy manner and provide tools for diagnostics and plotting. The structure is depicted in Fig. 7.1.

$ParisAnalysis$ is based on the concept of a $Data Bus$ owning basic parameters such as shower direction, impact and energy, but also an arbitrary number of $properties$ identified.
by names. There is one instance of the Data Bus for each reconstruction technique in use, and a specific part of the code is designed to fill that Data Bus. The part of the code that is specific to a reconstruction technique is implemented in a separate shared library. New reconstructions can be added this way to the analysis, as soon as they provide the needed interface, without any recompilation of the analysis software. In the same way, a registry is used internally to organize the list of algorithms to apply. Any algorithm provided by the system can be replaced in a easy an transparent manner by a user defined version.

All plotting, background subtraction and projection mechanisms use only the Data Bus and are therefore completely decoupled from the actual analysis implementation. They work equally well with the Hillas analysis as with the Model or any other analysis. This allows the user to cross-check the different analyses, by plotting for instance a variable of one analysis versus a variable of another, in the cuts of a third.

The same applies for the spectrum and morphology modules, described in sections 7.5 and 7.6 which come as separate plug-ins connected to ParisAnalysis.

7.2 Extraction of signal

7.2.1 Event classes

The reconstruction techniques usually provide discriminating parameters, that allow to separate, to some level, the γ-ray candidates from the background events. The distribution of a typical discriminating variable is shown in Fig. 7.2 Events are usually classified as:

- \(\gamma\)-candidate events (green band in Fig. 7.2) : these events have the value of the discriminating variable located in the signal peak. In real data however most such events are of hadronic origin via, for example, the decay of a \(\pi^0\). The response of the instrument on the tail of the hadronic distribution is particularly difficult to establish as they correspond to only a small \((\sim 10^{-2})\) fraction of the cosmic-ray induced showers.
Figure 7.2: Classification of events.

- background events (brown band in Fig. 7.2): these events are clearly identified as being cosmic ray events. Since their distribution is isotropic on the sky and independent of time, they can be used to monitor the response of the instrument.

- un-categorizable events lie in between. They are usually not used at all in the analysis, as the instrument response to them is particularly difficult to establish.

7.2.2 Statistical significance

The extraction of the signal, on a run-by-run basis, is based on the comparison of the number of events between the region of interest ($N_{on}$), and a control region which is presumed to contain only background ($N_{off}$) events (i.e. charged cosmic-ray events). In the case where the response of the instrument is not the same between the two regions, an additional factor $\alpha$ has to be taken into account, which gives the expected number of background events in the region of interest given the number $N_{off}$ of background events in the control region: $N_{on} = \alpha \times N_{off}$.

According to Li & Ma statistics [76] the significance of an excess $N_{on} - \alpha \times N_{off}$ is given by the formula

$$S = \sqrt{-2 \ln \lambda}$$

where $\lambda$ is the likelihood-ratio between the two hypotheses:

1. the observed number of events $N_{on}$ results from background events (expectation number $\alpha \times B$) plus signal ($S$), and the number of events in the control region only results from background ($B$). We then have the combined probability

$$P(N_{on}, N_{off}|S, B) = \frac{e^{-(S+\alpha B)}(S + \alpha B)^{N_{on}}}{N_{on}!} \times \frac{e^{-B}B^{N_{off}}}{N_{off}!}$$
2. the observed number of events $N_{on}$ results only from background events $\alpha B$. This is the so-called null hypothesis, for which the probability reads:

$$
P_0(N_{on}, N_{off}|B) = \frac{e^{-(\alpha B)(\alpha B)^{N_{on}}}}{N_{on}!} \times \frac{e^{-B}B^{N_{off}}}{N_{off}!}
$$

(7.3)

In the two cases, the optimal number of signal and background events $\bar{S}$ and $\bar{B}$ (or $\bar{B}_0$ in the null hypothesis) are found by maximizing the likelihood:

$$
\left\{ \begin{array}{l}
\frac{\partial \log P(N_{on}, N_{off}|S, B)}{\partial S} = 0 \\
\frac{\partial \log P(N_{on}, N_{off}|S, B)}{\partial B} = 0
\end{array} \right.
$$

which solves to give:

$$
\begin{align*}
\bar{S} &= N_{on} - \alpha N_{off} \\
\bar{B} &= N_{off} \\
\bar{B}_0 &= \frac{N_{on} + N_{off}}{\alpha + 1}
\end{align*}
$$

(7.5)

The final likelihood ratio reads:

$$
\lambda = \frac{P_0(N_{on}, N_{off}|\bar{B}_0)}{P(N_{on}, N_{off}|S, \bar{B})} = \left[ \frac{\alpha}{1 + \alpha \left( \frac{N_{on} + N_{off}}{N_{on}} \right)} \right]^{N_{on}} \times \left[ \frac{1}{1 + \alpha \left( \frac{N_{on} + N_{off}}{N_{off}} \right)} \right]^{N_{off}}
$$

(7.7)

In the case of pure Poisson statistical fluctuations on the number of events $N_{on}$ and $N_{off}$, the distribution of $S = \sqrt{-2\ln \lambda}$ follows a Gaussian distribution of width 1 and average 0.

In general, the complexity of the signal evaluation task is hidden in the factor $\alpha$: this number characterizes the ratio of the expected number of events in the two regions. Its evaluation requires a detailed and precise knowledge of the instrument response. There are however a few cases where this number can be estimated easily.

### 7.2.3 Accumulation of runs

Usually many runs need to be accumulated to be able to establish the emission of a source. The statistics need therefore to be properly handled.

Assuming we have $i$ runs, with respective number of events $N_{on}^{(i)}$ and $N_{off}^{(i)}$ and normalization factors $\alpha^{(i)}$, the number of excess events reads in each run is $N^{(i)}_\gamma = N^{(i)}_{on} - \alpha^{(i)} N^{(i)}_{off}$. Obviously, one wants to have:

$$
N_{on} = \sum_i N_{on}^{(i)}, \quad N_{off} = \sum_i N_{off}^{(i)}, \quad N_\gamma = \sum_i N^{(i)}_\gamma
$$

(7.8)

The unique solution to this system is to write

$$
\alpha = \frac{\sum_i \alpha^{(i)} N_{off}^{(i)}}{\sum_i N_{off}^{(i)}}
$$

(7.9)
7.3 Background subtraction

7.3.1 Excluded regions

In any background determination technique, regions of the sky where a significant γ ray emission has been detected must be excluded from the background determination. Such regions are called excluded regions and are either determined automatically with an iterative procedure, or registered in a database which needs periodic updating. A map of the currently defined excluded regions in the Galactic Plane is shown in Fig. 7.3. In some parts of the sky, the Galactic Plane is almost completely excluded, leading to potential difficulties to find suitable background regions.

Figure 7.3: Significance map across the Galactic Plane, with the excluded regions shown in yellow. From [60].

7.3.2 Reflected Background

The main inhomogeneity of the instrument response results from the dependency of the efficiency on the off-axis angle (angular distance to the center of the camera). This is
illustrated in Fig. 7.4 which shows the variation of the response with off-axis angle as a function of zenith angle left and energy right.

![Graph showing radial acceptance as a function of zenith angle and energy](image)

**Figure 7.4:** Radial acceptance as a function of zenith angle (left) and as a function of energy (right), for the Hillas reconstruction and for a specific set of cuts. From [74].

It is possible to calibrate this response using large samples of data, or in principle from simulations, but since this response must be calculated for $\gamma$-like cosmic-ray events, calibrating it from simulations is a complicated task. Moreover, if one wants to be able to derive the energy spectrum of the source, the calibration needs to be done for every energy and every zenith angle band. It is therefore easier to arrange the observations to allow the use of control regions at the same distance to the centre of the camera. There are two ways of doing this:

- **On-Off observations**: Two runs are taken consecutively, with the observation position shifted in right ascension so that the trajectory on the sky is exactly the same. The control region is taken at the same position in the camera as the region of interest, but in the OFF run. This ensures the same acceptance as a function of time but the observation time is doubled. Moreover, variations of the sky conditions between the on-source and control observations (such as inhomogeneities of the night sky background) are not taken into account.

- **Wobble observations**: The source position is deliberately shifted with respect to the centre of the camera, and the control region is taken symmetrically (Reflected background) with respect to the centre. In comparison to the On-Off background technique, the effective time on source is almost doubled and the sky conditions are usually much more similar (except for very strong gradients). But the control region has a slightly different position on the sky, and in particular a different zenith angle, thus resulting in a trigger rate that can differ by about 1%. Another advantage with respect to the On-Off background method is that several control regions can be taken all around the centre of the field of view. This is illustrated in Fig. 7.3 where the observation positions are shown in pink and the corresponding background regions in gray. Note that, however, regions intersecting already known sources must be rejected from the background estimation. The analysis is therefore
often iterative, newly discovered sources being constantly added to a database of regions that cannot be safely used as background control regions.

Figure 7.5: Reflected background subtraction technique, illustrated for two different runs. The green circle is the region of interest, corresponding to the potential source. The pink stars indicate the pointing positions of the centre of the camera in the two runs. For each of the two observations, many control regions can be derived (grey circles), at the same distance to the camera centre. However, some of these regions need to be excluded because of already known sources (blue circles) that would lead to an incorrect estimation of the background.

The main advantage of the reflected background method is that, to the f, the acceptances are similar in the test and control regions. This is also true in every energy band, thus allowing the extraction of the signal as a function of energy. Of course, converting that excess per energy into a spectrum requires a precise knowledge of the instrument response to γ rays, but at least, the energy-dependency of the response to background events does not need to be calibrated precisely.

The reflected background background technique requires however the source to be significantly offset from the camera centre, otherwise no background region can be found. Therefore, other background subtraction techniques have been developed to circumvent this problem.

7.3.3 Ring background

The idea of the Ring Background subtraction technique is to estimate the background in an annular region around the test position (Fig. 7.6). This is not done on a run-by-run basis, but when all events have been accumulated. Here again, regions known to have significant γ-ray emission must be excluded from the background determination (regions in red in Fig. 7.6).
Because the control region is not generally symmetrical with respect to the camera centre, a precise estimation of the number of expected background events in both regions (α factor) needs to be done. This determination is explained in section 7.4.

In principle, the α factor depends on the energy; its precise determination as a function of the energy should also allow, as in the reflected background technique, to derive the energy spectrum of the source. This has not, however, been done so far due to the to large complexity of the involved acceptance determination. Nevertheless, the Ring background subtraction technique is widely used to generate sky maps, such as those published by the H.E.S.S. collaboration in the Galactic Plane survey.

### 7.3.4 Template Background

The Template Background [78] uses a completely different approach to estimate the expected number of background events in the region of interest: instead of using γ ray-like events from different regions in the camera or on the sky, events from the same region are used, but belonging to the class of background events. This method is particularly adapted to the analysis of very extended sources, or very crowded fields of view where no suitable control region without significant γ ray emission can be found.

A more elaborated technique, called the Weighting Method [79], attributes a weight to each event, corresponding to its probability of being a γ ray event. The background subtraction, similar in spirit to the Template Background method, uses the distribution of the discriminating variable to exploit the full statistics available. Although this results in an optimal use of the data sample, this technique requires a precise calibration of the distribution of the discriminating variable and its evolution across the field of view.

Note that, similarly to the Ring Background method, it would in principle be possible to do the background modelling and subtraction in energy bands, and therefore to derive the energy spectrum of the source, although this has not yet been done.
7.4 Acceptance determination and predicted background

Several background determination techniques (Ring Background and Template Background methods) require a precise knowledge of the expected number of events in either two different regions of the sky, or for two different classes of events at the same position. This expected number of events is calculated from the acceptance of the detector, which characterizes the probability that an event of a given energy and type triggers the system, is reconstructed, and is selected according to the considered event class.

The acceptance depends on the event type (γ ray or background event), the reconstruction method and the associated selection cuts, but also on the observation conditions: zenith angle, off-axis angle, overall optical efficiency of the telescopes, ... Three different acceptances are used:

- Acceptance to γ rays: this is required to compute the energy spectrum of the source, and is obtained from simulation
- Acceptance to γ-like background events: this is required to estimate the contamination by cosmic rays in the region of interest, and to properly subtract the background in the ring background method. It is in general estimated directly from the data.
- Acceptance to hadron-like events: this is required, in the template model, to estimate the number of events belonging to the background event class. As in the previous case, it is in general estimated directly from the data.

7.4.1 Principles of determination

Calculating the acceptance from the data itself requires the events to be redistributed, assuming some underlying symmetry, and excluding the regions with significant γ ray emission. In particular, to avoid any bias and ensure statistical independence, one must take care not to use the data from one particular region to estimate the acceptance in that region. The data is then redistributed across the field of view using the identified symmetry.

7.4.2 Radial acceptance

The most frequently used assumption is that the response of the system is invariant against rotation of the camera around the observation position. This allows the derivation of the radial response of the instrument on a run-by-run basis. Excluded regions are taken into account either by allocating a weight to the events that properly corrects for the covered area as a function of angular distance to the centre of the field of view, or by using as a measure of covered area the fraction of background events in excluded regions as a function of the angular distance.

This technique allows the acceptance to be determined on a run-by-run basis, and therefore the complete dependency on, for instance, the zenith angle to be investigated. It’s only when an excluded region appears to be in the center of the field of view that this technique does not work any more. Averaging over runs therefore appears to be necessary to provide a complete coverage. Radial acceptances can also be provided as look-up tables, calculated from different observations, and are then just renormalized to the number of events in each run.
The simple assumption of rotational symmetry is however not completely sufficient. In particular, zenith angle variation across the field of view results in a variation of the trigger rate, and of the event rate in cuts, of the order of $2 - 3\%$. This gradient needs to be corrected on a run-by-run basis as well.

### 7.4.3 2D Acceptance model

The assumption of rotational symmetry in the field of view is only valid to the first order of approximation, and only if events close to the edges of the camera are rejected (since the shape of the camera is not circular). Moreover, inhomogeneities of the response across the camera are not taken into account. It is possible to use an alternate acceptance model, derived on the assumption that the acceptance in the camera field of view does not vary significantly in shape from one run to an other.

The first stage in the determination of a this 2D acceptance model is the calculation of an exposure map, which determines, for each position in the camera, the fraction of the time that this position was in an *excluded region*. This is depicted in Fig. [7.7] in each run, the trajectory of each excluded region in the camera has to be taken into account for proper determination of the exposure. This is usually done by using the fraction of *background* events in excluded regions as a function of position in the camera.

![Figure 7.7: Determination of exposure maps from accumulation of runs.](image)

In the second stage, the exposure maps of all runs are summed up, weighted by the number of events in each run. At the same time, the raw distributions of events in the camera field of view, outside any excluded regions, are summed up. The acceptance map (Fig. [7.8]) is then obtained by simply dividing the actual distribution of events by the exposure map.

![Figure 7.8: Determination of acceptance from event and exposure maps.](image)
In order to account for the evolution of the detector response with zenith angle, acceptance maps are determined in zenith angle bands (typically 20).

### 7.4.4 Expected background maps

The expected background maps (for \(\gamma\)-like or hadron like events) are obtained from the acceptance maps and the list of runs. For each run in the list, the actual observation conditions are computed, and the corresponding acceptance map (or radial acceptance) is re-projected on the sky, normalized by the number of events in the run.

### 7.4.5 Precision of acceptance calculation

Using very deep field observations, such as those on PKS 2155-304, it is possible to compare, in each sky bin, the actual number of observed events with the prediction from

![Graphs showing Rel. Event Rate](image)

**Figure 7.9:** Number of events, normalized to the calculated acceptance, for \(\sim 340\) h of data taken on the blazar PKS 2155-304, as a function of right ascension (left) and declination (right). Fluctuations are well within the 1% level and are, for the \(\gamma\) rays, still dominated by statistical fluctuations. No significant gradient is noticeable.
the acceptance determination. This is shown in Fig. 7.9 for γ-like events (upper) and background like events (lower), as a function of right ascension (left) and declination on the sky (right). The acceptance calculation is controlled at a level of better than ~ 1%, without any noticeable gradient.

Fig. 7.10 shows the same analysis, but in a two dimensional manner. No structure in the ratio between the observed and expected number of events is seen, neither for the γ-like events (left) nor for the background events (right). The red peak in the γ-like events map is due to the very intense emission (~ 60000 excess events) in the direction of PKS 2155-304.

Figure 7.10: Number of events, normalized to the calculated acceptance, for ~ 340 h of data taken on the blazar PKS 2155-304, as a function of position on the sky, respectively for γ ray candidates (left) and for background events (right). For each position, events within a circle of 0.25° radius have been accumulated to reduce fluctuations. The bottom panels show the distribution of the ratio of actual number of events over expected number across the sky. No significant gradient and no structure are noticeable.

Table 7.1 summarizes the ratio between the actual number of events integrated in a circle of 0.25° radius around each position and the expected number of events in the same area, for the deepest fields of observation with H.E.S.S.. The precision of the acceptance
calculation, defined as the RMS of this ratio across the field of view, is given in columns 3 (γ-like) and 4 (hadron-like). It ranges from 3% (resp. 1%) in the least exposed fields of view to ~1% (resp. 0.5%) in the deepest fields of observation, respectively for γ-like and background events.

Comparison between galactic and extra-galactic sources tends to indicate that the precision is slightly worse in galactic fields of view with large number of sources or significant variation of the night sky luminosity.

### 7.4.6 Sensitivity to NSB

As explained in section 5.3.5, the width of the pedestals in high gain provides a precise measure of the night sky background in each pixel. Proper reprojection on the sky (taking into account rotation of the field of view in the camera) and averaging over many runs allows maps of the night sky background across the sky to be constructed. This is shown for instance in Fig. [7.11] for the field around the Galactic center.

In this field of view, the noise level varies between less than 100 MHz in the darkest part (the Galactic Center itself is obscured by dust) to about 400 MHz in the brightest parts, south of the Galactic Center. Bright stars are also clearly identified.

Such fields can be used to further investigate the precision of the acceptance determination, and its sensitivity to NSB variations. Fig. [7.12] shows the evolution of the event rate (normalized to the prediction from the acceptance calculation) with the noise level. Variations of the order of ±3% are observed in the brightest regions for γ-like events, whereas the rate of background events remains stable. This evolution is probably caused by the fact that the Goodness, calculated in the Model Analysis, incorporates the pedestal variations.
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Figure 7.11: Distribution of average night sky background in the field of view of the Galactic Center, as estimated from the width of the pedestal distribution in high gain.

Figure 7.12: Evolution of the event rate in the field of view of the Galactic Center as a function of the average night sky background rate. **Left:** for $\gamma$-like events. **Right:** for hadron-like events.
width in its calculation and is therefore slightly more tolerant for showers observed from bright sky regions. Deep galactic fields might require a precise calibration of the event rate variation with NSB level.

### 7.4.7 Sensitivity to the magnetic field

The development of electromagnetic showers in the atmosphere is affected by the geomagnetic field, which tends to separate apart particles of opposite charge, particularly in the shower tails. This charge separation induces in particular a larger spread of the shower in the direction orthogonal to the magnetic field. Several authors studied the effect of the magnetic field on the image shape in Atmospheric Cherenkov Telescopes, both on the theoretical and experimental levels [e.g. 80][81]. The interaction of the field and the cascade electrons produces a broadening of the atmospheric Cherenkov light image resulting in a reduction of the density of light sampled by the telescope; so the energy threshold for the telescope increases.

The images are slightly distorted and rotated compared to showers developing in the absence of magnetic field; this effect induces a systematic bias in the direction estimation and results in inhomogeneities in the background level, as shown in Fig. 7.13.

\[
\begin{align*}
\text{Az} = 0^\circ & \quad \theta = 12^\circ \\
\text{Az} = 180^\circ & \quad \theta = 87^\circ \\
\end{align*}
\]

Figure 7.13: Expected distortion in the reconstruction of the shower direction due to the effect of the magnetic field. **Left:** when the magnetic field is almost parallel to the shower development, the effect is minimal. **Right:** when the magnetic field is inclined with respect to the shower direction, rotation of the images induce a systematic bias in the shower direction. From [80].

A study on this effect in the H.E.S.S. data has been conducted. The geometry for this analysis is given in Fig. 7.14

- B is the direction of the magnetic field in the camera field of view,
- S is the reconstructed shower direction,
- \( \phi \) is the azimuthal angle of shower direction with respect to the magnetic field,
7.4. ACCEPTANCE DETERMINATION AND PREDICTED BACKGROUND

- $D_S$ is the angular distance between the shower direction and the magnetic field.

![Diagram showing $D_S$ and $D_0$](image)

Figure 7.14: Geometry used for the analysis of the magnetic field effect

The distribution of azimuthal angle $\phi$ is given, in Fig. 7.13, for all reconstructed events (left) and for $\gamma$-like events (right) for the field of view of the start-burst galaxy NGC 253, one of the deepest fields of H.E.S.S.:

![Graphs showing event rate vs. azimuth](image)

Figure 7.15: Evolution of the event rate in the field of view of NGC 253 as a function of the azimuth with respect to the direction of the magnetic field. **Left:** for all reconstructed events. **Right:** for $\gamma$-like events.

An oscillation of the event rate of $\pm 2\%$ as a function of the azimuthal angle is observed, with a minimum around $90^\circ$. Fig. 7.16 shows the amplitude of the observed oscillation in 13 deep fields of view of H.E.S.S., as a function of the angular distance to the magnetic field.

It appears that the effect is maximal for average angles of $45^\circ$ and not $90^\circ$ as would naively be expected. More detailed studies based, in particular, on simulations are required to fully understand the relation between image rotation and azimuthal distribution of events. The use of stereoscopic observation most probably changes the picture notably compared to the study done by the MAGIC experiment. The fact that the amplitude of the oscillation is compatible with zero for small angular distances is expected, and confirm the analysis according to which this oscillation is indeed related to the magnetic field.

Although this effect was identified more than 10 years ago, and some corrections were proposed (based on the correction of the Hillas parameters) [82], these corrections only...
apply to analyses based on the Hillas parametrization. Recent studies applied to the H.E.S.S. case [53] estimate a variation of the order of 10\% of the event rate (at trigger level) as a function of the orientation of the telescopes with respect to the direction of the magnetic field, and a disorientation of shower images of the order of 2\degree, whereas attempts to propose a de-rotation in MAGIC turned out to be not very efficient [80].

Because the amplitude of this effect is comparable to the precision at which the acceptance is obtained, further improvement in the reconstruction and analysis would require the inclusion of the effect of the magnetic field directly in the reconstruction scheme. Concerning the Model Analysis, a direct comparison between the actual images and a rotated model, incorporating the effect of the magnetic field, should completely correct for this effect at the reconstruction level.

Conclusions

The precise determination of the expected response of the instrument (acceptance) is a crucial part of data analysis for Atmospheric Cherenkov Telescopes, especially when dealing with faint sources. Much progress has been made in the proper understanding of the background, which is now properly modelled at a precision level of \(\sim 1 - 2\%\). Reduction of systematics in the very deep fields of view (\(\sim 200\ h\)) will now require the incorporation of additional effects such as the image distortion caused by the geomagnetic fields. This issue might be problematic for the coming H.E.S.S.-II telescope as the lower threshold will certainly induce larger distortion by the magnetic field.
7.5 Spectra and Light Curves

The response of Imaging Atmospheric Cherenkov Telescopes depends on many factors, among which are in particular the zenith and the off-axis angles. The threshold and the effective area vary significantly from one observation to the other, thus requiring accurate statistical techniques to properly extract the intrinsic energy distribution of the observed source.

Spectral determination techniques, in a similar manner to signal extraction techniques, rely on the comparison of the number of events between a source (ON) region and one or several background (OFF) regions. In general, to avoid introducing systematic effects, the Reflected background method is used, ensuring a similar energy-dependant response in the ON and OFF regions. However the techniques described below could, in principle, be extended to other background subtraction techniques. This would require in particular an energy-dependant $\alpha$ normalization factor which is not easy to derive.

The method described here relies on a precise comparison, using a log-likelihood technique, between the expected and the observed number of events in four dimensional bins of energy, zenith angle, off-axis angles and telescope optical efficiency. The basics of the method were described in details in [34][35], but further enhancements have been done in order to improve its precision.

7.5.1 Forward folding technique

The forward folding technique relies on the assumption of a spectral shape $\Phi(E_{\text{true}}|\bar{\alpha})$, which can be a function of some parameters $\bar{\alpha}$. Quite a large variety of shapes are implemented:

- Pure power-law spectrum
  \[
  \frac{dN}{dE} = \Phi_0 \left( \frac{E}{E_0} \right)^{\alpha} \quad (7.10)
  \]

- Curved power-law spectrum
  \[
  \frac{dN}{dE} = \Phi_0 \left( \frac{E}{E_0} \right)^{\alpha - \beta \log(E/E_0)} \quad (7.11)
  \]

- Broken power-law. This is the combination of two different power-law spectra:
  \[
  \frac{dN}{dE} = \Phi_0 \times \left\{ \begin{array}{ll}
  \left( \frac{E}{E_0} \right)^{\alpha_0} & \text{for} \quad E \leq E_{\text{cut}} \\
  \left( \frac{E_{\text{cut}}}{E_0} \right)^{\alpha_0 - \alpha_1} \left( \frac{E}{E_0} \right)^{\alpha_1} & \text{for} \quad E \geq E_{\text{cut}}
  \end{array} \right.
  \quad (7.12)
  \]

- Power law with exponential cutoff
  \[
  \frac{dN}{dE} = \Phi_0 \left( \frac{E}{E_0} \right) \times \exp \left( -\frac{E}{E_{\text{cut}}} \right) \quad (7.13)
  \]

- Power law spectrum with super-exponential cutoff
  \[
  \frac{dN}{dE} = \Phi_0 \left( \frac{E}{E_0} \right) \times \exp \left( -\left( \frac{E}{E_{\text{cut}}} \right)^{\gamma} \right) \quad (7.14)
  \]
Figure 7.17: Example of energy resolution table, giving the probability density of measuring a reconstructed energy \( E_{\text{rec}} \) for a given true energy \( E_{\text{true}} \), for a given set of observation parameters. From [10].

- Dark matter Annihilation Lines

\[
\frac{dN}{dE} = \sum_i \Phi_i \times \delta(E - E_i) \quad (7.15)
\]

- Dark Matter particle annihilation spectrum, calculated from dark matter simulation codes

In addition, spectral shapes can easily be superimposed. It is, for example, possible to adjust the superposition of a power law with annihilation lines. Spectral shapes absorbed by pair creation on the extragalactic background light are also available.

### Acceptance and resolution tables

The acceptance \( A(E_{\text{true}} | \delta, \Psi, \epsilon) \) is calculated from simulation and is tabulated as a function of true energy \( E_{\text{true}} \), zenith angle \( \delta \), off-axis angle \( \Psi \) and optical efficiency \( \epsilon \). In a similar fashion, the resolution function \( R(E_{\text{rec}}, E_{\text{true}} | \delta, \Psi, \epsilon) \) is defined as the probability density function of measuring a reconstructed energy \( E_{\text{rec}} \) for a given true energy \( E_{\text{true}} \), and is obtained from simulations. An example of such an energy resolution table is given in Fig. [7.17]. Obviously, energy resolution depends on the selection parameters so that tables have to be generated for each reconstruction technique and each set of cuts.

The expected number of \( \gamma \)-ray events in a given reconstructed energy bin \([E_{\text{rec},1}, E_{\text{rec},2}]\) is obtained by an integration over true energies and reconstructed energies:

\[
n_\gamma = \int_{E_{\text{rec},1}}^{E_{\text{rec},2}} dE_{\text{rec}} \int_0^\infty dE_{\text{true}} R(E_{\text{rec}}, E_{\text{true}} | \delta, \Psi, \epsilon) \times A(E_{\text{true}} | \delta, \Psi, \epsilon) \times \Phi(E_{\text{true}} | \bar{\alpha}) \quad (7.16)
\]

In practice, data are divided in bands of zenith angle \( \delta \), off-axis angle \( \Psi \) and optical efficiency \( \epsilon \) and the double integral is performed in each band individually.
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Log-likelihood comparison

In a given bin of zenith angle $\delta$, off-axis angle $\Psi$ and optical efficiency $\epsilon$, we denote:

- $N_{ON}$ the measured number of events in the ON data set,
- $N_{OFF}$ the measured number of events in the OFF data set,
- $T_{ON}$ and $T_{OFF}$ the respective live times for the two data sets,
- $\beta = T_{ON}/T_{OFF}$ the live time normalization,
- $n_{\gamma}$ and $n_{h}$ the expected number of gamma and background events in the bin (from Eq. 7.16 for $n_{\gamma}$).

The probability of observing $N_{ON}$ and $N_{OFF}$ events when we expect $n_{\gamma}$ gamma events and $n_{h}$ hadrons is given, in Poisson statistics, by the formula

$$P(N_{ON}, N_{OFF}|n_{\gamma}, n_{h}) = \frac{n_{\gamma} + \beta n_{h}^{N_{ON}}}{N_{ON}!} e^{-(n_{\gamma} + \beta n_{h})} \times \frac{n_{h}^{N_{OFF}}}{N_{OFF}!} e^{-n_{h}}$$

(7.17)

The best background estimate is given by the value of $n_{h}$ which maximizes the log-likelihood $\log L = \log(P)$. It’s the solution of quadratic equation in $n_{h}$:

$$\frac{\beta N_{ON}}{n_{\gamma} + \beta n_{h}} + \frac{N_{OFF}}{n_{h}} - (\beta + 1) = 0$$

(7.18)

whose solution is given by:

$$C = \beta \times (N_{ON} + N_{OFF}) - (1 + \beta) \times n_{\gamma}$$

$$\Delta^2 = C^2 + 4\beta(\beta + 1) \times N_{OFF} \times n_{\gamma}$$

(7.19)

$$n_{h} = \frac{C + \Delta}{2\beta(\beta + 1)}$$

(7.20)

The uncertainties on $n_{h}$ can be calculated from the second derivative of the log-likelihood:

$$\Delta n_{h} = \sqrt{\left(\frac{\beta^2 N_{ON}}{(n_{\gamma} + \beta n_{h})^2} + \frac{N_{OFF}}{n_{h}^2}\right)^{-1}}$$

(7.21)

Log-likelihood maximization

The log-likelihood described above is then maximized against the parameters $\hat{\alpha}$ of the assumed spectral shape. A Levenberg-Marquardt [23, 24] maximization algorithm is used for maximal efficiency. The results of the maximization procedure are:

- the best fit parameters $\hat{\alpha}^0$ and their uncertainties,
- the covariance matrix between the parameters,
- the expected number of events in each bin,
• the final value of the log-likelihood, which can, under some assumption, be converted into a goodness-of-fit parameter.

The flux uncertainty at each energy is obtained from the covariance matrix and the spectral parameters. Writing $W_{ij}$ the error matrix (inverse of covariance matrix), and $\Phi(E)$ the differential flux at energy $E$, the flux uncertainty reads

$$\Delta \Phi(E) = \sqrt{\sum_{i,j} W_{ij} \frac{\partial \Phi}{\partial \alpha_i} \frac{\partial \Phi}{\partial \alpha_j}}$$  \hspace{1cm} (7.22)$$

In general, the uncertainty on the logarithm of the flux is computed instead of the uncertainty on the flux, because it’s more symmetrical. Residuals in each energy bin are then used to derive spectral points for the sake of clarity. It should be noted however that the spectrum points appearing in publication are only a sub-product of the spectrum determination technique.

### 7.5.2 Light curve models

Light curves are the primary ingredient for following the evolution of a source activity with time, in general assuming a fixed spectral shape. Keeping the same notation ($N_{ON}$, $N_{OFF}$, $N_{\gamma}$ and $\beta$) and $\Phi_0$ the corresponding flux (integral or differential), there are several different methods for finding the flux normalization or the integrated flux above some energy:

#### Log-likelihood determination

We can use a log-likelihood approach, similar to that used in the spectrum shape determination. In each bin the optimal value of the number of background events is determined, and the likelihood is then maximized against the number of signal events.

There is however a caveat in this method, in the case when the number of ON or OFF events is zero. Indeed, if the number of ON events is zero, the likelihood reads

$$logL = log(P) = N_{OFF} \times \log n_h - n_{\gamma} - (\beta + 1)n_h + C_{ste}$$  \hspace{1cm} (7.23)$$

The optimal number of background events is then found to be $n_h = N_{OFF}/(\beta + 1)$ and the likelihood is a linear function of $n_{\gamma}$, which is maximal for $n_{\gamma} = -\infty$. This problem is related to the Bayes theorem and the Bayesian approach: we actually want to maximize the probability $P(n_{\gamma}, n_h|N_{ON}, N_{OFF})$ that the signal and the background consist respectively of $n_{\gamma}$ and $n_h$ background events when the observed number of events are $N_{ON}$ and $N_{OFF}$. This probability can be expressed as a function of the inverse probability $P(N_{ON}, N_{OFF}|n_{\gamma}, n_h)$ of observing $N_{ON}$ and $N_{OFF}$ events when the signal and the background consist respectively of $n_{\gamma}$ and $n_h$ using the Bayes theorem:

$$P(n_{\gamma}, n_h|N_{ON}, N_{OFF}) = \frac{P(N_{ON}, N_{OFF}|n_{\gamma}, n_h) \times P(n_{\gamma}, n_h)}{P(N_{ON}, N_{OFF})}$$  \hspace{1cm} (7.24)$$

The prior $P(n_{\gamma}, n_h)$ is not known and is assumed, in the likelihood technique, to be uniform, although it is obviously not: in particular, for physical sources, it cannot be negative. Refinement of the technique might be needed to incorporate more realistic priors.
7.5. SPECTRA AND LIGHT CURVES

Simple statistics determination

To circumvent the problems of the "log-likelihood" method, a simpler method has been developed:

In this approach, the excess in each bin is simply \( S = nON - \beta \times nOFF \). The uncertainty on the excess is in general asymmetric and is determined using Li&Ma statistics, and the flux reads:

\[
\Phi = \Phi_0 \times \frac{N_{ON} - \beta \times N_{OFF}}{N_{\gamma}}
\]  
(7.25)

Event-wise integrated flux determination

This method doesn’t assume a spectral shape and directly determines the integrated flux above some threshold. But, since no spectral shape is assumed, it is impossible to compute the effect of finite energy resolution, and in addition only events above the energy threshold are taken into account.

Since the energy distribution of events can be expressed by:

\[
\frac{dN}{dE \, dt} = \Phi(E) \times A(E)
\]  
(7.26)

we have

\[
\Delta t \times \int_{E_1}^{E_2} \Phi(E) \, dE = \int_{E_1}^{E_2} \frac{dN}{A(E)} \approx \sum \frac{\alpha_i}{A(E_i)}
\]  
(7.27)

where \( \alpha_i = 1 \) for events in the ON region and \( \alpha_i = -\beta \) for events in the OFF regions. The integrated flux is simply estimated from an weighted sum of all events

\[
\Phi(E > E_{th}) = \sum \frac{\alpha_i}{A(E_i)}
\]  
(7.28)

The uncertainty is calculated in a similar way:

\[
\Delta^2 \Phi(E > E_{th}) = \sum \frac{\alpha_i^2}{A^2(E_i)}
\]  
(7.29)

Comparison

None of the above methods is absolutely perfect. The log-likelihood technique makes optimal use of the available statistics and properly deals with low number of events, using a full Poisson statistical treatment. It provides reliable error estimate (at least for non-zero number of events), but assumes on a spectral shape. The Event-wise integrated flux determination does not assume any spectral shape and is therefore able to provide a true integrated flux above some energy. But the finite resolution of the instrument is ignored and only events above the maximum threshold energy can be used. In the case of very extended data-sets with significant variation of the threshold energy from one run to the other, this can result into the drop of a large fraction of events. The simple statistics approach is robust but provides slightly incorrect estimation of uncertainties.
7.6 Morphological studies

7.6.1 Maximum likelihood approach

As in the spectral determination, the basic principles for a precise and quantitative morphological determination is to compare the observed event distribution (in some parameter space) with what is expected given a source morphology assumption, using Poisson statistics and a forward-folding technique: Given a source morphology shape and instrument angular response, one can compute the expected number of gamma-ray events in a given space domain. The difference between the expected and observed number of events can then be minimized against the morphological parameters. The method is described in details in [54].

Let’s denote:

- $\theta$ the angular distance between the reconstructed and the true event direction
- $PSF(\theta)$ the instrument point spread function (PSF) for the current set of cuts
- $\delta$ the zenith angle
- $\Psi$ the off-axis angle (distance of the event to the center of the camera)

For a given set of cuts, the system angular resolution is defined as the probability density function of reconstructing an event at a squared angular distance $\theta^2$ from its actual direction.

It is computed for a power-law energy spectrum, and depends also on the zenith angle and off-axis angle.

**Expected number of events in $\theta^2$ bins**

To compute the expected number of events at a square distance $\theta^2$ from the source origin, we need to integrate over the possible shower true direction (given a source morphology) and the reconstructed shower direction given the true direction.

Let’s note (Fig. 7.18):

- $P(x', y')$ the reconstructed shower direction at a distance $\theta$ from the source origin,

![Diagram](figure)

**Figure 7.18:** Geometry for the $\theta^2$ distribution. $O$ denotes the center of the source, $M$ the true shower direction and $P$ the reconstructed one.
7.6. MORPHOLOGICAL STUDIES

- $M(x,y)$ the true shower direction, at a distance $r'$ from P,
- $O$ the centre of the source,
- $\phi$ the angle between OP and PM.

Since the PSF is invariant under rotation, the probability of having a shower reconstructed at P when its true direction is M is simply $PSF(r')$. Using simple trigonometric relations, the distance between the source origin and the true shower direction is given by:

$$r^2 = \theta^2 + r'^2 + 2 \times \theta \times r' \times \cos \phi'$$

(7.30)

If we denote $\Phi(r)$ the radial source luminosity, the $\theta^2$ distribution is given by:

$$\frac{dN}{d\theta^2} = \int_0^{2\phi} d\phi \times \int_0^{\infty} dr'^2 PSF(r') \times \Phi(r', \phi)$$

(7.31)

The expected number $n_\gamma$ of gamma events in a reconstructed $\theta^2$ bin $[\theta_1^2, \theta_2^2]$ (and for a given zenith angle and off-axis angle) is then calculated by the formula:

$$n_\gamma = \int_{\theta_i}^{\theta_2} d\theta^2 \frac{dN}{d\theta^2}$$

(7.32)

**Expected number of events in a sky map bin**

The calculation is very similar. By integrating over the true shower direction, we have

$$\frac{d^2N}{dx'dy'} = \iint dx \, dy \, PSF(r') \times \Phi(x, y)$$

(7.33)

And then:

$$n_\gamma = \iint dx' \, dy' \, \frac{d^2N}{dx'dy'}$$

(7.34)

Using a 2D sky map involves one integral more than using the $\theta^2$ distribution, and is therefore more time consuming. It is therefore used when one wants to accurately determine the source position or the morphology of a non-radially-symmetric source.

**Shapes**

- Point like source
  $$\mathcal{L}(x, y) = \sum_i L_i \delta(x - x_i) \delta(y - y_i)$$
  (7.35)

- Symmetrical Gaussian source
  $$\mathcal{L}(x, y) = \frac{L_0}{2\pi\sigma^2} \exp \left[ -\frac{(x - x_0)^2 \cos^2 \beta + (y - y_0)^2}{2\sigma^2} \right]$$
  (7.36)
• Asymmetrical Gaussian source: This is an elliptical shape, with the major axis orientation given by the angle $\phi$:

$$
L(x, y) = \frac{L_0}{2\pi\sigma^2} \times \exp \left[ -\frac{((x - x_0)\cos \beta \cos \phi + (y - y_0)\sin \phi)^2}{2\sigma^2_a} \right] \\
\times \exp \left[ -\frac{(-(x - x_0)\cos \beta \sin \phi + (y - y_0)\cos \phi)^2}{2\sigma^2_b} \right]
$$

(7.37)

(7.38)

• Ring source: A ring shape, with Gaussian profile; this is not the projection of a 3D shell like a supernovae remnant shell, but already a quite good approximation

$$
L(x, y) = L_0 \times \exp \left[ \frac{(x - x_0)^2 + (y - y_0)^2}{2\sigma^2} \right]
$$

(7.39)

• 3D shell source: Shape obtained by a line-of-sight integral of a 3D emitting shell
• Navarro Frank Wright dark matter halo

Log-likelihood maximization

A log-likelihood maximization technique is then used, in a similar way to the energy spectrum determination, to derive the best-fit parameters. The main advantage of this forward folding technique, compared to a simple fit of the event distribution, is that it provides estimation of the true source shape, not the shape convoluted by the point spread function.

Using this toolbox, many analysis tasks can be performed in a reliable and efficient way:

• Fit of morphology shape, leading to confidence contours (or constraints) on source parameters such as extension and position
• Source subtraction with proper statistical treatment
• Smoothing with actual instrument Point Spread Function
• Superposition of instrument Point Spread Function corresponding to the actual data set
  • ...

An example of a source subtraction is shown in Fig. 7.19. The image on the left shows the distribution of events on the sky, smoothed by the instrument PSF. The center image is the result of a fit of two point like sources in the field of view. The right image is the result of the subtraction of the best fit from the actual event distribution, smoothed again by the instrument PSF (Note that the subtraction is done with the un-correlated distribution of events, before the smoothen).

A second example of morphology determination is shown in Fig. 7.20, where a shell-like shape is adjusted on the data, properly taking into account the projection effects. The resulting goodness of fit can be used to estimate the quality of the model, and precise information on the parameters can be derived.
7.7. CONCLUSION: TOWARD A SPECTRO-IMAGER

Figure 7.19: Example of subtraction of two sources. **Left:** excess map. **Middle:** fitted model. **Right:** Residuals after subtraction. A additional source near $l \sim 314.4^\circ$ becomes apparent after subtraction.

![Smoothed excess map](image1)

Figure 7.20: Example of source morphology fit on the supernova remnant RCW 86. **Left:** excess map. **Right:** fitted ring model.

![Smoothed excess map](image2)

7.7 Conclusion: Toward a spectro-imager

A complete analysis package has been developed, with special emphasis on the precision and reliability of the various implemented algorithms. Acceptance determination, required for proper background subtraction and signal estimation, reaches the precision of $1 - 2\%$ even for very deep fields. Remaining effects affecting the stability of the instrument response, such as the deterioration of images by the magnetic field, have been identified and will have to be handled in the next generation of instruments. Precise spectrum and morphology derivation methods have been developed, thus allowing a quantitative comparison between the observed data and a source model.

We believe that the next step in Cherenkov data analysis will be to able to provide spectro-imaging capabilities. Some preliminary attempts have already been successful, such as the generation of flux maps in energy bands. Proper statistical treatment of hadronic background should make the use of template models for spectrum determina-
tion possible, thus turning Cherenkov telescopes into genuine spectro-imagers, allowing spectral determination all across the field of view. Such tools will be very useful for next generation of instruments such as CTA.
Part II

A new view on the very high energy sky
Chapter 8

The Galactic zoo

The H.E.S.S. Galactic Plane survey (GPS) has been a core component of the observation program since 2004. The original GPS \cite{23,24}, consisting of $\sim 230$ h of observations after standard run-quality selection, covered the inner part of the Galaxy, from the Norma to the Scutum-Crux spiral-arm tangents ($l \pm 30^\circ$ in longitude and $b \pm 3^\circ$ in latitude). It resulted in the firm discovery of eight previously known sources of VHE $\gamma$ rays with a statistical significance above $6 \sigma$ (post-trials) and six likely sources above $4 \sigma$, all of them confirmed by subsequent deeper observations.

Since the GPS contains a large number of test positions, the significance has to be corrected for the "trial factor." This trial factor accounts for the increased probability of finding a fake signal with an increased number of test positions for which a significance is calculated. A rough estimate is given by the area of the GPS divided by the PSF size, and a more precise derivation of the number of trials can be obtained with a full Monte Carlo simulation.

Between 2005 and 2009, the GPS was extended significantly in longitude, from $l \sim -60^\circ$ to $l \sim 275^\circ$ \cite{30}. In addition, the overall exposure along the Galactic Plane was significantly increased with more than 1400 hours of accumulated data (representing roughly one third of the total H.E.S.S. data set). The H.E.S.S. exposure inside the Galactic Plane varies from a few hours on the less-observed area to more than 100 hours in the deep exposure regions centered around targets of particular interest, such as Sgr A*, RX J1713.7-3946, or LS 5039, leading to a sensitivity varying between less than 1% to about 10% of the Crab Nebula flux.

The pre-trials significance map of the Galactic Plane, reproduced from \cite{30}, and calculated using the ring-background subtraction technique \cite{87} and hard cuts, is shown in Fig. \ref{fig81}. A total of 56 Galactic sources are detected in the GPS. The major population consists of pulsar wind nebula (PWN) (29 identified sources) followed by supernova remnants (SNR) (9 associations) and binary systems (3 systems).

Most of the Galactic VHE sources are found to be significantly extended, with sizes greater than the $\sim 0.1^\circ$ H.E.S.S. PSF. The few sources in the Galactic Plane that appear point-like are associated with young pulsar wind nebulae, including the Crab Nebula \cite{87}, or with VHE $\gamma$ ray emitting high-mass X-ray binaries (HMXB) which include the very well established binaries PSR B1259-63 \cite{88} and LS 5039 \cite{27,89}. The point-like VHE source HESS J0632+057 \cite{90} is now a strong candidate for a HMXB system following a recent multi-wavelength campaign \cite{91}.

After excluding five sources well off the Galactic Plane, with $|b| > 2^\circ$ (HESS J1356-645, HESS J1442-624, HESS J1507-662, HESS J1514-591, and SN 1006), the latitude
distribution of the Galactic sources is very narrow ($\langle b \rangle = -0.26^\circ$ with an RMS of $0.40^\circ$). This scale is significantly smaller than the width of the region of significant H.E.S.S. exposure (of the order of $\sim 2^\circ$ in RMS), and similar to the scale of the molecular gas distribution. The latitude distribution (Fig. 8.2) is, at the first glance, compatible with what is presumably the parent populations of SNRs [92] and high spin-down luminosity pulsars ($\dot{E} > 10^{34}$ ergs s$^{-1}$) from [93].

Figure 8.2: Latitude distribution of sources in the GPS. Figure reproduced from [86].
Figure 8.3: Spin-down energy loss rate of radio pulsars listed in the ATNF catalog versus their characteristic age. The blue dots overlaid indicate the pulsars detected at high energies with the Fermi-LAT telescope whereas the red ones mark the pulsars associated to TeV PWNe. Figure reproduced from [94].

As shown in Fig. 8.3, most young pulsars with large spin-down power ($\dot{E} > 10^{33}$ erg s$^{-1}$) produce prominent PWNe detectable at VHE energies. Pulsars detected by the Fermi-LAT satellite in the energy range $100$ MeV $< E < 100$ GeV are shown in blue and comprise an additional population from re-accelerated, old pulsars (millisecond pulsars, on the right side of the plot).
Chapter 9

LS 5039: shedding light on binary systems

9.1 $\gamma$ ray binaries

X-ray binaries (XRBs) comprise a compact object, such as a neutron star or black hole, orbiting around a companion star. They are one of several types of astrophysical systems that provide an environment in which the acceleration of particles and subsequent production of radiation might be periodic. Modulation of this radiation, linked to the orbital motion of the binary system, provides key insights into the nature and location of the particle acceleration and emission processes.

The physical environment inside a close binary system (or an eccentric binary close to periastron) is characterized by very high radiation densities of the order of $1 \ erg \ cm^{-3}$ and high magnetic fields (from mG to G), and is therefore radically different from the environments in which supernova remnants and pulsar wind nebula exist, which are characterized by the much lower densities of the interstellar medium.

Due to this high radiation density, relativistic electrons suffer rapid radiative cooling (occurring mainly in the deep Klein-Nishina regime) and cannot escape the system. In addition, the relevant timescales in such systems scale as the inverse of the mass of the compact object and are therefore much shorter than those observed in other systems such as active galactic nuclei, thus allowing the mechanisms of accretion and ejection to be tested on human time scales.

The acceleration and cooling times for relativistic electrons, as well as the orbital period, are relatively short compared to observation timescales (of the order of years), and lead to the expectation of variable VHE emission (if it is of leptonic origin). Due to the lack of target matter, hadrons could, in some cases, escape the production region without significant energy losses. This assumption, which is usually valid for emission inside the binary system, would result into a steady and extended $\gamma$ ray emission as the escaping hadrons interact with the surrounding interstellar medium.

The population of $\gamma$-ray binaries now comprises five objects: PSR B1259-63/SS [88, 95, 96], LS 5039 [27, 89, 97], LS I +61° 303 [98, 99, 100, 101, 102, 103, 104, 105, 106, 107], HESS J0632+057 [90, 91, 104, 105, 106, 107] and recently 1FGL J1018.6-5856. [108].

The most detailed TeV measurements thus far have been made of LS 5039. These measurements are described in section 9.2. A summary of TeV binaries is given in Tab. 9.1.
<table>
<thead>
<tr>
<th>Name</th>
<th>D (kpc)</th>
<th>$L_r$</th>
<th>$L_{\nu}$</th>
<th>$L_{\nu \alpha}$</th>
<th>$L_{\nu \alpha}$</th>
<th>$L_{\nu \alpha}$</th>
<th>$L_{\nu \alpha}$</th>
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<td>LS 5039</td>
<td>2.5</td>
<td>1.3</td>
<td>5-50</td>
<td>70</td>
<td>4-11</td>
<td>0.46</td>
<td>0.45 - 0.6</td>
</tr>
<tr>
<td>LS I +61 303</td>
<td>2.0</td>
<td>1-17</td>
<td>3-9</td>
<td>60</td>
<td>8</td>
<td>-0.6 - 0.45</td>
<td>0.53</td>
</tr>
<tr>
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<td>0.02-0.3</td>
<td>0.3-6</td>
<td>...</td>
<td>2.3</td>
<td>-2.2 - 0.3</td>
<td>0.78</td>
</tr>
<tr>
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<td>0.3</td>
<td>$10^4$</td>
<td>...</td>
<td>12</td>
<td>0.1</td>
<td>0.8</td>
</tr>
<tr>
<td>HESS J0532+057</td>
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<td>0.13</td>
<td>&lt;9</td>
<td>0.9</td>
<td>0.6</td>
<td>0.26</td>
</tr>
</tbody>
</table>

Table 9.1: Properties of the TeV binaries, adapted from [102]. Note that the spectral indices are defined by $F_{\nu} \propto \nu^{-\alpha}$ or equivalently $dN/dE \propto E^{-(1+\alpha)}$. All luminosities are in units of $10^{33}$ erg s$^{-1}$ except the radio luminosities which are in units of $10^{31}$ erg s$^{-1}$. Luminosities are given for the following ranges: $L_r$: 0.1 – 100 GHz, $L_{\nu}$: 1 – 10 keV, $L_{\nu \alpha}$: 1 – 10 GeV, $L_{\nu \alpha}$: 0.2 – 10 TeV. Reproduced from [110].

### 9.1.1 Emission models

Two major models aim at explaining VHE emission from close binary systems. For a review, see [e.g. 111].

![Microquasar and Binary Pulsar](image)

Figure 9.1: Models for very high energy emission from micro-quasars. **Left**: Micro-quasars are powered by compact objects (neutron stars or stellar-mass black holes) via mass accretion from a companion star. This produces collimated jets that, if aligned with our line of sight, appear as micro-blazars. The jets boost the energy of stellar photons to the range of very energetic $\gamma$ rays. **Right**: Pulsar winds are powered by the rotation of neutron stars; the wind flows away to large distances in a comet-shaped tail. Interaction of this wind with the companion-star outflow may produce very energetic $\gamma$ rays. Credit: P. Huey/Science. From [111].

### 9.1.2 Micro-quasar scenario

The most obvious energy source inside a binary system is accretion onto a compact object. In a similar way to the giant active galactic nuclei (AGN), where this mechanism is well established, particle acceleration is expected to occur in the collimated jets emerging on both sides of the accretion disk. Galactic analogues to AGN were discovered in the 1990’s
with evidence of superluminal motion \cite{112}, and were therefore dubbed Micro-quasars. The internal and external shocks associated with such jets would then provide potential sites for particle acceleration up to TeV energies. In this view, emission could arise from the relativistic outflow of leptonic \cite{113,114,115} and/or hadronic plasma \cite{116} and some correlation, although with some delay, is expected between variation of the accretion rate and high energy emission \cite{117}.

### 9.1.3 Pulsar wind scenario

The principal alternative energy source in close binary system powered by a neutron star is the collision of stellar and neutron-star winds. This model \cite{113,114,115} requires neutron stars young enough to provide large spin down luminosities as the primary energy source. A \textit{standard} pulsar of period \(\sim 0.1\) s, with energy loss rate of \(\sim 10^{35}\) erg s\(^{-1}\), generates the pair plasma that inflates a pulsar wind nebula of sub-AU size. The PWN expands and produces shocks as it interacts with the outflows from the two-component wind from the massive star (consisting of a fast polar and a slow and denser equatorial component). This results in the formation of a bow shock in a scaled-down version of a pulsar wind nebula (Fig. 9.1 right).

![Image of LS I +61° 303 at 3.6 cm arranged by orbital phase \(\phi\). Contours begin at \(\pm 0.2\) mJy, and step by a factor \(\sqrt{2}\). The resolution is \(1.5 \times 1.1\) mas = \(3 \times 2.2\) AU. The image locations have been adjusted for non-overlap, and are hence approximate and illustrative only. Orbit size is greatly exaggerated. The time progression is counterclockwise. From \cite{121}.](image)

The acceleration of particles is then possible at the termination shock. A detailed review of these systems can be found in e.g. \cite{122}. Non-thermal emission from radio to keV energies arises from synchrotron emission of accelerated electrons and inverse-
Compton scattering on the stellar (UV) photons and provides the input for emission up to TeV energies.

In a variation of this model \[123\], the highest TeV energies arise from hadronic interactions at the shock.

At least one system of this kind is known, PSR B1259-63/SS 2883, where the pulsar is clearly identified through its radio pulses around apastron. PSR B1259-63 and its B2e companion SS 2883 form a highly eccentric binary with an orbital period of \( \sim 3.4 \) years \[124\]. From the relatively large spin-down luminosity \( \dot{E} \sim 10^{36} \) erg s\(^{-1}\) and its small distance \( d \sim 2.3 \) kpc, TeV emission from the system close to its periastron passage had been predicted \[125\] based on inverse-Compton scattering of the electrons and positrons in the shocked pulsar wind on the target photons from the Be star. A precise prediction of the variation of the emission with orbital phase was made in the paper, allowing the pulsar wind scenario to be disentangled from other models. TeV emission close to the periastron passage \[88, 95\] was indeed observed by the H.E.S.S. experiment with similar variation, thus strengthening the case for the pulsar-wind scenario.

LS I +61\(^\circ\) 303 is a Be-Pulsar binary system, comprising a radio pulsar orbiting at a period of 26.496 days around a Be star. The pulsation of the pulsar is suppressed at periastron by the dense wind. Detection of TeV emission from this system by the MAGIC collaboration \[98\] with orbital modulation of the VHE flux \[99\], shed some light on the pulsar-wind scenario. VLBI observations of LS I +61\(^\circ\) 303 \[121\] (Fig. 9.2) over a full orbit resulted into the detection of a clear cometary shaped emission, at scales of 1 mas (\( \sim 2 \) AU), that is always pointing away from the massive star. No relativistic motion nor large-scale structure, are detectable at any phase of the orbit, which clearly favours the pulsar wind nebula hypothesis over the jet model. Drastic change of the morphology around periastron is attributed to the large outflow velocity of \( V \sim 7500 \) km/s, which reduces gradually to \( V \sim 1000 \) km/s near apastron. Modulation of the radio flux (flares a few days after periastron) could be explained by the geometry of the system, the flux increasing during the near side of the orbit, when the tail is seen un-absorbed by the head.

### 9.2 Periodicity measurement in LS 5039

LS 5039 (distance \( d \sim 2.5 \) kpc) is a High Mass X-ray Binary comprising a compact object in a \( \sim 3.9 \) days eccentric (\( e \sim 0.35 \)) orbit around a massive O6.5V star \[126\]. Persistent radio outflows (observed with extension in the range 2 to \( \sim 1000 \) AU) have been attributed to a mildly relativistic (\( v \sim 0.2 \) c) jet \[127, 128\] which would have placed LS 5039 in the \underline{micro-quasar} class.

The detection of radio \[129, 130\] and X-ray \[117\] emission and their possible association with the MeV to GeV \( \gamma \)-ray sources GRO J1823—12 suggested the presence of multi-GeV particles.

LS 5039 was detected by H.E.S.S. in 2004 \[27\] in \( \sim 11 \) h hours of observation. The best fit position was within 30 arcsec of the radio position of LS 5039 (Fig. 9.3), revealing for the first time that LS 5039, and hence XRBs, are capable of multi-TeV particle acceleration. The limited statistics did not allow for detailed timing or variability analyses.

Deeper observation of \( \sim 70 \) h distributed over many orbital cycles between 2004 and 2006 yielded a modulation of the VHE \( \gamma \) ray flux (\( > 100 \) GeV) with a period of 3.9078 ± 0.0015 days \[89\], consistent with the orbital period reported by \[126\].

The corresponding H.E.S.S. Lomb-Scargle periodogram is shown in Fig. 9.4. The source HESS J1825—137, observed simultaneously in the same field of view \[13\], did not
Figure 9.3: Map of excess $\gamma$-ray emission in units of counts for the region around LS 5039, smoothed by the point spread function. The white ellipse shows the 3$\sigma$ confidence region for HESS J1826-148. The radio emission from the SNR G16.8-1.1 is represented by gray contours obtained from the Parkes-MIT-NRAO 6 cm radio survey map [131, 132]. The yellow contours show the 68%, 95% and 99% confidence level region of the EGRET source 3EG J1824-1514. The green star marks the position of the radio source associated with the micro-quasar LS 5039. HESS J1825-137 is discussed in [23].
exhibit similar variation and therefore served as a cross-check of the timing analysis.

In this section, we describe how this result, which is one of the best examples of precision measurement in VHE $\gamma$ ray astronomy so far, was obtained.

### 9.2.1 Lomb Scargle Periodogram

When data are unevenly spaced, as it is the case with H.E.S.S., a simple Fourier analysis is not the best tool to establish a periodic behaviour of a source. An alternate approach, based on the least-square fitting of a sinusoidal function, has been developed by Lomb and Scargle [133, 134]. This method, however, assumes similar uncertainties for each measurement and is not directly applicable to analysis of VHE light-curves, thus requiring further developments.

We assume a set of data \( \{t_j, X_j\} (j = 1 \ldots N) \) with null average. The original Lomb-Scargle periodogram consists of the least-square fitting of the sinusoidal function \( f(t_j) = A \cos \omega t_j + B \sin \omega t_j \) on the data with a \( \chi^2 \): \( \chi^2(A, B) = \sum (X_j - f(t_j))^2 \). The minimization problem can be solved analytically into:

\[
\begin{pmatrix}
\cos \omega t_j & \sum_j \cos \omega t_j \sin \omega t_j \\
\sum_j \cos \omega t_j \sin \omega t_j & \sin^2 \omega t_j
\end{pmatrix}
\begin{pmatrix}
A \\
B
\end{pmatrix} = 
\begin{pmatrix}
\sum_j X_j \cos \omega t_j \\
\sum_j X_j \sin \omega t_j
\end{pmatrix}
\]

(9.1)

A further simplification is obtained by translating the times with an offset \( \tau \) that is chosen in a way that the matrix is diagonal. The solution to the problem is then simply:

\[
\tan(2\omega \tau) = \frac{\sum_j \sin 2\omega t_j}{\sum_j \cos 2\omega t_j}
\]

\[
\begin{align*}
A &= \frac{\sum_j X_j \cos \omega(t_j - \tau)}{\sum_j \cos^2 \omega(t_j - \tau)} \\
B &= \frac{\sum_j X_j \sin \omega(t_j - \tau)}{\sum_j \sin^2 \omega(t_j - \tau)}
\end{align*}
\]

(9.2)

The un-normalized Lomb Scargle Periodogram then reads:

\[
\chi^2(\omega) = \sum_j \left( X_j^2 - A^2 \cos^2 \omega(t_j - \tau) - B^2 \sin^2 \omega(t_j - \tau) \right)
\]

\[
\hat{\chi}^2(\omega) = \sum_j X_j^2 - \chi^2(\omega)
\]

(9.3)

\[
= \left( \frac{\sum_j X_j \cos \omega(t_j - \tau)}{\sum_j \cos^2 \omega(t_j - \tau)} \right)^2 + \left( \frac{\sum_j X_j \sin \omega(t_j - \tau)}{\sum_j \sin^2 \omega(t_j - \tau)} \right)^2
\]

Proper normalization of the periodogram is essential for the calculation of the cumulative distribution function, and thus for the estimation of false-alarm, or confidence level, probability. Two normalizations are possible:
9.2. PERIODICITY MEASUREMENT IN LS 5039

Figure 9.4: Top: Lomb-Scargle (LS) periodogram of the VHE run-wise flux for LS 5039 (chance probability to obtain the LS power vs. frequency). Inset: zoom around the highest peak (pre-trial probability $\sim 10^{-20}$), which corresponds to a period of $3.9078 \pm 0.0015$ days, compatible with the ephemeris value of $3.90603 \pm 0.00017$ days (vertical red line at 0.2560 days$^{-1}$ on the inset) from [126]. Middle: LS periodogram of the same data after subtraction of a pure sinusoidal component at the orbital period of 3.90603 days. Bottom: LS periodogram of the H.E.S.S. source HESS J1825—137 observed simultaneously in the same field of view. From [89].
• traditional normalization, in which the periodogram is normalized to the empirical data variance:

$$z(\omega) = \frac{N}{2} \frac{\sum X_j^2 - \chi^2(\omega)}{\sum X_j^2}$$

(9.4)

• residual normalization, in which the periodogram is normalized by the actual $\chi^2$ at each frequency:

$$z(\omega) = \frac{N - 2}{2} \frac{\sum X_j^2 - \chi^2(\omega)}{\chi^2(\omega)}$$

(9.5)

![Figure 9.5: Comparison of traditional normalization (black) and residual normalization (red) on the VHE run-wise flux for LS 5039. inset: zoom around the orbital period of the system.](image)

It has been found that the residual normalization is more powerful for very unevenly spaced data, as it increases the peaks corresponding to the best adjustments due to lower $\chi^2$. This is illustrated in Fig. 9.5 for most of the periodogram, the two normalizations yield consistent results whereas around the significant peaks at the orbital period (inset), the relative normalizations differ by a factor of $\sim 2$. For a Gaussian noise, the cumulative distribution function of the normalized periodogram can be calculated analytically. In the case of residual normalization, it reads:

$$P(z > z_0) = \left(1 + \frac{2z_0}{N - 2}\right)^{-(N-2)/2}$$

(9.6)

### 9.2.2 Floating Lomb Scargle Periodogram

The Lomb-Scargle in its original form makes two simplifications that are not very well suitable to the H.E.S.S. experiment:
9.2. PERIODICITY MEASUREMENT IN LS 5039

- Uncertainties on individual measurements are not taken into account, thus giving equal weight to every observation. In real H.E.S.S. light-curves, uncertainties can vary by a factor of several units from one point to another one and need to be taken into account.

- The average value of the signal has been assumed to be zero. The experimental average value is subtracted from the data to make it zero average, but this is only an estimate of the average value. A cleaner approach consist in letting the average value free in the adjustment.

For these reasons, further developments were done, following the ideas of Cumming et al [135]. A statistical weight \( w_j = 1/\sigma_j^2 \) is applied to each data point and the average value is let free in the fit function \( f(t_j) = A \cos \omega t_j + B \sin \omega t_j + C \). The corresponding \( \chi^2 \) reads:

\[
\chi^2(\omega) = \sum_j w_j [(X_j - C)^2 - A^2 \cos^2 \omega(t_j - \tau) - B^2 \sin^2 \omega(t_j - \tau)]
\]

\[
= \chi^2_{N-1} - \sum_j w_j a^2 \cos^2 \omega(t_j - \tau) - \sum_j w_j B^2 \sin^2 \omega(t_j - \tau)
\]

(9.7)

where \( \chi^2_{N-1} = \sum_j w_j (X_j - C)^2 \) and \( \tau \) is defined as:

\[
tan(2\omega \tau) = \frac{\sum_j w_j \sin 2\omega t_j}{\sum_j w_j \cos 2\omega t_j}
\]

(9.8)

Analytic expressions of \( A \) and \( B \) are possible, although unnecessarily complicated. A Gauss-Jordan algorithm is used to solve the linear system. The un-normalized and normalized periodograms are respectively (using a residual normalization):

\[
\hat{\varepsilon}(\omega) = \chi^2_{N-1} - \chi^2(\omega), \quad z(\omega) = \frac{(N - 3)}{2} \frac{\chi^2_{N-1} - \chi^2(\omega)}{\chi^2(\omega)}
\]

(9.9)

9.2.3 False alarm probability

Let \( P(z > z_0) \) be the probability that a power \( z \), calculated from Gaussian noise, exceeds a reference value \( z_0 \) at a given frequency. For \( M \) independent frequencies, the false alarm probability, i.e. the probability to have one power exceeding \( z_0 \) by chance at any frequency, reads:

\[
F = 1 - [1 - P(z > z_0)]
\]

(9.10)

\( M \) is in principle of the same order of magnitude as the number of points in the data sample, but depends on the spacing of the data. It can be estimated more accurately using two different techniques:

- the bootstrap technique, where the reference time of the sample points are kept, but the fluxes are randomly exchanged between points,
• the simulation technique, where random light-curves are generated with the same reference time as the data, and random fluxes centered on the measured value with Gaussian fluctuation having the same RMS as the data.

In any case, a large sample of light-curves are generated, and the corresponding Floating Lomb Scargle periodogram is calculated for each of them. The characteristics (power and frequency) of the peak with highest power are then extracted and used to derive chance probability. This is illustrated in Fig. 9.6 and 9.7 where $10^6$ different light-curves have been generated from the LS 5039 observation using the bootstrap method. Fig. 9.6 shows the distribution of Floating Lomb Scargle periodogram power on the right, and the distribution of the power of the highest peak in each periodogram on the left. The power distribution is very well described by an exponential distribution of slope $-1$, thus confirming that the residual normalization is correct. The maximum power distribution on the left is described with a distribution from Eq. 9.10 yielding a measurement of the effective number of independent frequencies $M \sim 58$.

![Highest peak power](image1)

![Power Distribution](image2)

Figure 9.6: Power distribution using bootstrapping technique. **Left**: distribution of highest peak in periodogram. **Right**: distribution of power accumulated over all generated periodograms.

Fig. 9.7 shows the distribution of the frequency of the highest peak found in the periodogram of each generated light-curve (left) and the distribution of power versus frequency (right). The frequency distribution is almost flat, thus indicating that no characteristic frequency should emerge from the data by chance.

The two techniques are used to estimate the chance probability of the highest peak in the periodogram. $10^6$ periodograms have been generated using each of the two methods, yielding in both cases maximum power of the order of 20, well below the maximum power of $P \sim 60$ measured in the true LS 5039 periodogram. Extrapolation of the cumulative distribution functions allow the chance probability of having a peak at $P \sim 60$ to be estimated to be less than $10^{-20}$ (taking into account only statistical uncertainties), thus confirming the robustness of the result (Fig. 9.8).
Figure 9.7: Frequency distribution using bootstrapping technique. **Left:** distribution of frequency of the highest peak in generated periodogram. **Right:** distribution of power versus frequency for all entries in all periodograms.

Figure 9.8: Derivation of the normalization of the LS periodogram using the bootstrapping (blue) and simulation (red) techniques.
9.2.4 Precision on the measured period

The precision on the measured period arising from statistical uncertainties can be estimated by generating a large sample of simulated light-curves, starting from the measured LS 5039 light-curve and moving fluxes measured in each run randomly within the statistical uncertainties (using a Gaussian distribution).

![Frequency Distribution](image)

Figure 9.9: Determination of central period and uncertainty using simulated light-curves within statistical uncertainties of the individual points.

This is illustrated in Fig. 9.9 for 1000 generated light-curves within statistical uncertainties of the individual points. The distribution of maximum peak period yields a period of $3.9067 \pm 0.002$ days fully consistent with the orbital period determined by Casares et al. [126] ($3.90603 \pm 0.00017$ days) from radial velocity measurements of the stellar companion. This undoubtedly constitutes one of the first precision measurement from VHE astronomy.

9.2.5 Folded light-curve

The phasogram (light-curve of integral fluxes at energies $E > 1$ TeV, folded at the orbital period) of LS 5039 is shown in Fig. 9.10. It clearly indicates that the bulk of the VHE $\gamma$-ray emission is confined to roughly half of the orbital period, covering the phase interval $\phi \in [0.45, 0.9]$. The VHE flux maximum appears to lag somewhat behind the apanastron epoch, and aligns better with inferior conjunction ($\phi = 0.716$) of the compact object. Inferior conjunction of the compact object occurs when it is aligned along our line-of-sight in front of the stellar companion. The VHE flux minimum occurs at a phase $\phi \sim 0.2$, slightly further along the orbit than superior conjunction ($\phi = 0.058$), which is when the compact object is lined up behind the stellar companion. Note that the inclination upper limit $i < 64^\circ$ (obtained from the lack of X-ray eclipse) implies that direct views of both compact object and stellar companion are always available.
Figure 9.10: Integral $\gamma$-ray flux ($F > 1$ TeV) light-curve (phasogram) of LS 5039 from H.E.S.S. data (2004 to 2005) on a run-by-run basis folded with the orbital ephemeris of Casares et al. (126). Each run is $\sim$28 minutes. Two full phase ($\phi$) periods are shown for clarity. The blue solid arrows depict periastron and apastron. The thin red dashed lines represent the superior and inferior conjunctions of the compact object, and the thick red dashed line depicts the Lomb-Scargle Sine coefficients for the period giving the highest Lomb-Scargle power. From [89].

9.3 Orbital spectral modulation

9.3.1 Phase-resolved H.E.S.S. energy spectra

We define two broad phase intervals for further study: INFC ($0.45 < \phi \leq 0.9$) encompassing the inferior conjunction, and SUPC ($\phi \leq 0.45$ and $\phi > 0.9$) for the superior conjunction (Fig. 9.11).

The differential energy spectrum of the VHE $\gamma$-ray emission for the two phase intervals, between 0.2 to 10.0 TeV, are shown in Fig. 9.12 and constitute an important diagnostic-probe of the emission mechanism powering this object. For the INFC, the differential photon energy spectrum is consistent with a hard power law where $\Gamma = 1.85 \pm 0.06_{\text{stat}} \pm 0.1_{\text{syst}}$ with exponential cutoff at $E_\gamma = 8.7 \pm 2.0$ TeV (for fitted function $dN/dE \sim E^{-\Gamma} \exp(-E/E_\gamma)$). In contrast, the spectrum for SUPC is consistent with a relatively steep ($\Gamma = 2.53 \pm 0.07_{\text{stat}} \pm 0.1_{\text{syst}}$) pure power law (0.2 to 10 TeV).

The spectra from these phase intervals are mutually incompatible, with the probability that the same spectral shape would fit both simultaneously being $\sim 2 \times 10^{-6}$.

Fitting a pure power-law (which is statistically sufficient at present) to narrower phase intervals of width $\Delta\phi = 0.1$, and restricting the fit to energies $E \leq 5$ TeV to reduce the effect of any cutoff, also demonstrates that a harder spectrum occurs when the flux is higher (Fig. 9.13). Notably, the VHE flux at $E \sim 0.2$ TeV appears to be quite stable over phases and the strongest modulation occurs at a few TeV.

The results constitute the very first observation of spectral modulation at VHE energies and provide, as will be shown in section 9.4, strong constraints on the understanding of the underlying mechanisms.
Figure 9.11: Same as Fig.9.10 with indication of phase intervals.

Figure 9.12: H.E.S.S. spectra of LS 5039 for around inferior and superior conjunction. From [89].
9.3. ORBITAL SPECTRAL MODULATION

Figure 9.13: Phase resolved spectra of LS 5039 in phase intervals of width \( \Delta \phi = 0.1 \). **Top:** Power-law normalization (at 1 TeV) vs. phase interval. **Bottom:** Fitted pure power-law photon index (for energies 0.2 to 5 TeV) vs. phase interval. From \[59\].

### 9.3.2 Multi-wavelength picture

Significant variability of LS 5039 was detected by the Fermi-LAT instrument \[97\] at the period of \( 3.903 \pm 0.005 \) days, consistent with the orbital period. The light curve (Fig. 9.14) is characterized by a broad peak around superior conjunction, in phase opposition with the results from H.E.S.S. and, as will be shown later, in agreement with inverse Compton scattering models.

Using the same INFC and SUPC phase intervals, the Fermi spectrum (Fig. 9.15) is compatible with a pure power law of index \( \Gamma = 2.25 \pm 0.11 \) at inferior conjunction. At superior conjunction, a harder power law with an exponential cutoff is preferred with \( \Gamma = 1.91 \pm 0.16 \) and a cutoff energy of \( 1.9 \pm 0.5 \) GeV.

A detailed modelling is required to understand the multi-wavelength picture of the emission of LS 5039. Competition between phase-dependant inverse-Compton emission and pair-creation on the stellar photon field are the basic ingredients required in the modelling. They are described in detail in the following section.
Figure 9.14: Light curve of LS 5039 as seen by Fermi-LAT. **Top:** flux vs. orbital phase for $0.1 - 10$ GeV. **Bottom:** variations with orbital phase in the hardness ratio of $1 - 100$ MeV to $0.1 - 1$ GeV. From [97].

Figure 9.15: Phase resolved SED of LS 5039 with Fermi and H.E.S.S.. The black points (dotted line) represent the phase-averaged Fermi/LAT spectrum. The red data points (dotted line) represent the spectrum (overall fit) at inferior conjunction (Phase $\phi \in [0.45 - 0.9]$); blue data points (dotted line) represent the spectrum (overall fit) at superior conjunction (Phases $\phi < 0.45$ and $\phi > 0.9$). From [97].
9.4 Pair creation in the LS 5039 system

9.4.1 Geometry of LS 5039

The geometry of LS 5039 is described in Fig. 9.16. LS 5039 is a compact binary system in which the separation between the compact object and the companion star varies between 2 and 10 times the radius of the massive star \([126]\), leading to drastic changes of the photon radiation field of average energy \(kT \approx 3 \text{ eV}\).

Gamma rays emitted in the vicinity of the compact object with energies above the energy threshold \(E_{\text{th}} \approx (2m_e^2c^4)/(\hbar\nu) \approx 30 \text{ GeV}\) will inevitably suffer strong losses by pair creation on the stellar photons [e.g. \([136, 137, 120]\) and references therein], whereas photons emitted in the Fermi energy range would be largely unaffected by absorption.

The energy threshold for pair creation depends on the relative direction of the high \((E_{\text{th}})\) and low \((\hbar\nu)\) energy photons according to:

\[
E_{\text{th}} \times \hbar\nu \geq \frac{2m_e^2c^4}{1 - \cos \theta}
\]  

Around inferior conjunction, VHE photons emitted from behind the star see the stellar photons face on and suffer maximal losses by pair creation, whereas photons emitted around superior conjunction will remain mostly un-absorbed. In the following section, a quantitative estimate of absorption with phase (originally done in \([120]\)) is summarized.

9.4.2 Absorption vs phase

The geometry used in the calculation is described in Fig. 9.17.

The optical depth by pair creation can be written as a path-integral on the photon density:

\[
\tau_{\gamma\gamma} = \int d\epsilon \int dl \int d\Omega \times \frac{d\tau_{\gamma\gamma}}{d\Omega d\epsilon}(\epsilon, \Omega, E)
\]  

where

- \(\Omega\) describes the direction of stellar photons
- \(l\) is the gamma ray trajectory
- \(n_{\epsilon}(\epsilon, \Omega)\) is the stellar photon number distribution
- \(\frac{d\tau_{\gamma\gamma}}{d\Omega d\epsilon}(\epsilon, \Omega, E)\) is the elementary optical depth.

The calculation is usually easier to perform in the photon pair center-of-mass frame. The elementary optical depth indeed simply reads, as a function of pair creation cross-section \(\sigma_{\gamma\gamma}\):

\[
d\tau_{\gamma\gamma} = 2dr' \int d\epsilon' d\Omega' \sigma_{\gamma\gamma} n'_{\epsilon'}
\]  

where

- \(dr'\) is the elementary path in the photon pair center-of-mass frame
Figure 9.16: The orbital geometry (Casares et al. [126]) viewed from directly above LS 5039. Shown are: phases (φ) of minimum (periastron) and maximum (apastron) separation between the two components; epochs of superior and inferior conjunctions of the compact object representing phases of co-alignment along our line-of-sight of the compact object and stellar companion. The orbit is actually inclined at an angle in the range $13^\circ < i < 64^\circ$ with respect to the view above. VHE $\gamma$-rays (straight black lines with arrows) can be absorbed by optical photons of energy $h\nu_\gamma$, when their scattering angle $\theta$ exceeds zero.
Figure 9.17: Orientation of the various frames used in the calculation of the absorption by pair creation.

- $\epsilon'$ is the target photon center-of-mass energy
- $\sigma'_{\gamma\gamma}$ is the pair creation cross-section
- $n'_\epsilon$ is the photon density in the photon pair center-of-mass frame

The elementary path $dr'$ transforms under the Lorentz transformation as $dr' = dr \times \gamma_b(1 - \beta_b \cos \chi)$, which is the same as the energy transformation of the incident photon energy. Using also the fact that $\frac{1}{\epsilon} n_\gamma \, d\epsilon \, d\Omega$ is Lorentz invariant [e.g. [225]], the absorption depth can be written as:

$$d\tau = 2dr \times \frac{\epsilon'}{E} \int d\epsilon \, d\Omega \, \sigma_{\gamma\gamma} \frac{\epsilon'}{\epsilon} \, n'_\epsilon$$  \hspace{1cm} (9.14)

Using then the center-of-mass energy determination $4\epsilon'^2 = 2 \times (1 - \cos \theta) \epsilon E$, we obtain

$$d\tau = dr \times (1 - \cos \theta) \int d\epsilon \, d\Omega \, \sigma_{\gamma\gamma} \, n'_{\epsilon}$$  \hspace{1cm} (9.15)

For a star of finite size, the angular distribution of photons at the interaction point is uniform within a cone of opening angle given by the size of the star, and the distribution can be written as:

$$n_\epsilon(\epsilon, \Omega) = \frac{d^2N}{d\epsilon d\Omega} = \frac{1}{\epsilon} \frac{d^2U}{d\epsilon d\Omega} = \frac{B_\nu}{\epsilon \hbar c}$$  \hspace{1cm} (9.16)

For every position on the orbit, we work in the plane that contains the gamma-ray trajectory and the star. The integration along the photon path can easily be replaced by an angular integral:

$$dl = -\frac{d}{\sin^2 \alpha} \, d\alpha$$  \hspace{1cm} (9.17)

The integration over the stellar photon distribution can be expressed as a double integral.
Figure 9.18: Geometry of stellar photon distribution.

\[
\int d\Omega = \int_{0}^{\omega_{\text{max}}} \sin \omega \int_{0}^{2\pi} d\phi
\]

where the value of \( \omega_{\text{max}} \) is obtained with a tangency condition: \( \sin \omega_{\text{max}} = \frac{R_*}{r} = \frac{R_* \sin \alpha}{d} \).

The stellar photons are first expressed in a frame linked to the star (with \( y' \) being the normal to the star), before being rotated back around the \( z \) axis to the two-photon frame.

Finally, the optical depth can be written as a four-fold integral

\[
\tau_{\gamma\gamma} = \int_{\alpha_{\text{min}}}^{\pi} \frac{d}{\sin^2 \alpha} \int_{0}^{\omega_{\text{max}}} \sin \omega \, d\omega \int d\phi \int \frac{d\varepsilon}{\varepsilon \hbar c} \times (1 - \cos \Psi) \sigma_{\gamma\gamma}
\]

with \( \Psi \) being the angle between the two photons.

For 60 degrees inclination, the obtained results are shown in Fig. [9.19]. Since the system is nearly seen edge-on, the modulation is important. Complete extinction of the signal between \( \sim 20 \) GeV and a few TeV is expected, which is not strictly compatible with observations.

For 20 degrees inclination, the system is seen more from above and the modulation is less pronounced (Fig. [9.20]) but significant absorption of the VHE emission is expected from all phases, thus making low inclinations very unlikely. Comparison between VHE light-curves and prediction arising from pair creation clearly favour large inclinations.

### 9.4.3 Other Orbital-dependant processes

Emission in the Fermi range is largely unaffected by phase-dependant absorption and should allow easier identification of the intrinsic spectrum and variability of the source.
9.4. PAIR CREATION IN THE LS 5039 SYSTEM

Figure 9.19: Absorption probability at different phases for an inclination of 60 degrees. **Left:** Phase as a function of position on orbit. **Right:** Absorption as a function of phase. Around inferior conjunction ($\phi = 0.716$), absorption remains negligible and appears only above 1 TeV. Around superior conjunction ($\phi = 0.058$) the absorption is in contrast maximal with a threshold around 20 GeV.

Figure 9.20: Same as 9.19 but for a system inclination of 20°.
Modulation of the GeV light-curve indicates that other processes have a dependency on the phase.

Several phase-dependent processes have been identified:

- The efficiency of inverse-Compton scattering also depends on the relative angle between the high energy electron and the target stellar photon. The efficiency of the process and the energy of the scattered photon are expected to be maximal for head-on collisions, around superior conjunction, and minimal around inferior conjunction. This effect goes into the opposite direction from the phase-dependent absorption by pair-creation and could be responsible for the enhancement of the GeV flux, where absorption is inefficient, around superior conjunction. Detailed treatment of angle-dependant inverse-Compton scattering has been done by several authors [e.g.138, 139]. Fig. 9.21 shows an example of such a calculation for inverse-Compton scattering occurring in the Thomson or in the Klein-Nishina regimes.

- Cascading of higher-energy photons on the winds from the massive star could reprocess the absorbed VHE photons towards lower energies (until the pair creation threshold is reached), thus giving another reason for anti-correlation between H.E.S.S. and Fermi observations.

Consistent treatment of phase-dependant inverse Compton scattering, cascading and phase dependant absorption by pair creation manage to reproduce the H.E.S.S. light-curve in a very satisfactory manner (Fig. 9.22).

In this picture, anti-correlation between Fermi (with HE modulation peaking at $\phi \approx 0.0 - 0.1$, close to superior conjunction) and H.E.S.S. (with VHE modulation peaking close to inferior conjunction, that is $\phi = 0.72$) would arise from competition between IC scattering on high-energy electrons and VHE pair production.

An additional pulsed component from the pulsar magnetosphere itself cannot however be excluded in the GeV domain, as seen in the dozens of pulsars that have now been detected by Fermi [141] and that usually show hard spectra $\Gamma \sim 1.5$ with exponential cutoff at a few GeV. This component would then be modulated at the pulsar period. In the absence of prior knowledge of this period, identification of the magnetospheric component remains however very challenging.

### 9.5 Future Prospects

Although the understanding of $\gamma$ ray binary systems in general, and in LS 5039 in particular, have made some tremendous progresses in the last years, some points are still subject to detailed investigations:

- The long term stability of the emission has not been assessed. It could help to rule out definitively the hypothesis of accretion powered systems where eruptive events should be expected. Unpublished data from 2011 (Fig. 9.23) point toward long term stability of the VHE emission.

- In the case of LS 5039, significant VHE emission is observed close to orbital minimum, although predictions from pair creation indicate that the VHE flux should unavoidably be completely absorbed. Contamination from the extended, neighbouring pulsar wind nebula HESS J1825-137, needs to be better quantified. Fig. 9.24
Figure 9.21: Dependence of the viewing angle on the inverse Compton spectrum, for target photons originating from the star with $kT = 1$ eV and for interaction occurring in the Thomson regime (left) and in the Klein-Nishina regime (right). In each case, the spectrum is shown at viewing angles $\psi = 15^\circ$ (bottom), $30^\circ$, $60^\circ$, $90^\circ$, $120^\circ$ and $180^\circ$ (top). From [139].

Figure 9.22: Predicted VHE light-curve of LS 5039 using a consistent model taken into account phase-dependant inverse Compton scattering, cascading and phase dependant absorption by pair creation, and for two different inclinations. From [140].
Figure 9.23: H.E.S.S. integral light-curve of LS 5039 using data up to 2011. Recent data (in red) superimpose nicely with published data. Unpublished data.

shows the results of H.E.S.S. observation in the phase range $\phi \in [0.1, 0.2]$, where no emission is expected due to pair absorption. In this phase range, LS 5039 is detected at the $5.3\sigma$ significance level in 13 h of data, however with possible spill-over from HESS J1825-137. A new observation campaign in 2012 focusing on this phase range should help estimating the contamination fraction, thus bringing new observational constrains on the emission models.

Figure 9.24: H.E.S.S. sky map of LS 5039 for orbital phase $\phi \in [0.1, 0.2]$ where no significant emission is expected. Unpublished data.

- There are some indications of a peak of emission around phase $\phi \sim 0.85$. More sensitive instruments such as CTA are required to further investigate fine structures in the light-curve.
9.5. FUTURE PROSPECTS

Conclusions

LS 5039 is undoubtedly one of the best examples of precision measurements in modern γ ray astronomy. Precise measurements of orbital modulation with H.E.S.S. and additional multi-wavelength data from, in particular, the Fermi-LAT instrument, provide very valuable information allowing to precisely constrain the emission mechanisms and understand the underlying phenomena. The story of LS 5039 also illustrates the continuous exchange between observation and modelling, each taking benefits from the other. It illustrates the fact that only precise, multi-wavelength measurement really allow significant progresses in the understanding of the emission mechanisms to be made.
Chapter 10
Expanding Young Supernovae Remnants

Expanding shock waves in SNRs are believed to be able to accelerate cosmic rays up to multi-TeV energies through the mechanism of diffusive shock acceleration (DSA) [e.g. 6]. It was realized very early that, given the rate of ~ 3 explosive events per century, SNRs are capable of maintaining the galactic CR flux at the observed level [3] if a fraction of about 10% of their explosion energy is converted into cosmic rays. About 300 SNRs are currently known in the Milky Way, mostly from their non thermal radio emission [92, 142]. Most shell-type SNRs are non-thermal radio emitters, confirming that electrons are accelerated up to at least GeV energies. For a recent review of diffusive shock acceleration in the context of SNRs, see e.g. [143].

Several classes of SNRs are potential sites of VHE cosmic-ray acceleration. Composite SNR host an energetic pulsar at their center, whose non-thermal outflow leads to the formation of a pulsar wind nebula, capable of accelerating electrons up to VHE energies. Young shell-type SNRs are expected to release a significant fraction of their non thermal output at TeV energies, whereas older SNRs, through interactions with clouds of interstellar matter, are more likely to shine at GeV energies, thus allowing the different stages of SNRs evolution to be probed using multi-wavelength observation.

Four SNRs with clear shell-type morphology resolved in VHE γ rays have been detected by H.E.S.S., allowing direct investigation of the SNRs as sources of cosmic rays. They are all remnants of recent supernovae (less than a few kyr): RX J1713.7-3946 [25, 144], RX J0852.0+4622 - also known as Vela Junior [145], SN 1006 [146] and HESS J1731-347 [147]. A fifth case, RCW 86 [148], might be added to this list although the TeV shell morphology has not yet been clearly proven. All of them show a very clear correlation between non-thermal X-ray emission and VHE γ ray emission.

10.1 Evolution phases of SNRs

In this first section, we review the current knowledge of the evolution of supernova remnants.

A sketch of a supernova remnant is shown in Fig. [10.1]. During the explosion of a supernova, the major fraction (~ 99%) of the available energy is emitted in the form of prompt neutrinos. The remaining kinetic energy of the order of $E_k = 10^{51}$ erg = $10^{44}$ J is released into the interstellar medium in the form of a shell of ejected matter (with mass $M_{ej} \approx M_\odot$) expanding into a uniform low density medium.
The initial supersonic blast waves expands in the interstellar medium and gradually compresses, heats up and accelerates the up-stream medium. The swept-up medium leads to the formation of a reverse shock, propagating inwards in the direction of the center of the remnant, compressing further the ejecta. The surface separating the expanding shell of shocked ISM and the inward propagating shocked ejecta is called the surface discontinuity.

The description of the evolutionary phases of a supernova remnant has been investigated by many authors, using in particular self-similar solutions [e.g. [149] [150] [151] [152]]. Several phases are identified, as outlined below:

### 10.1.1 Free expansion

During the first phase, the amount of swept-up matter is negligible compared to the mass of ejecta. To first approximation, the speed of shock is constant. The kinetic energy also remains constant, and the equations of evolution are:

\[
V_S = C_{ste} \\
R_S = V_S \times t \\
E_k = \frac{1}{2} M_{ej} V_S^2
\]

For \( E_k = 10^{51} \) erg and \( M_{ej} \approx M_\odot \), one obtains a supersonic shock speed in the order of \( V_S \approx 10000 \) km s\(^{-1}\).
For a more realistic ejecta profile described by the expression $\rho_{\text{ejecta}} \propto r^{-n}$, a slightly different solution is obtained. For a type Ia supernovae, one has $n \approx 7$. For type II supernovae, $n$ is approximately 3 in the centre and between 5 and 14 at larger distances [153, 150]. Truelove & McKee [150] derived a more general solution, yielding in particular:

- For $n = 7$, $R \propto t^{4/7}$, $v \propto r^{-3/7}$
- For $n = 9$, $R \propto t^{2/3}$, $v \propto r^{-1/3}$
- For a profile with index $n > 5$, $R \propto t^{n-3/n}$, $v \propto r^{-3/n}$

The free expansion phase stops when the swept-up mass $\frac{4}{3} \pi \rho_0 R_S^3 = \frac{4}{3} \pi \rho_0 V_S^3 t^3$ equals the mass of the ejecta. This corresponds to an approximate time:

$$t_{ST} \approx 209 \left( \frac{E_{SNR}}{10^{51} \text{ erg}} \right)^{-1/2} \left( \frac{n_{ISM}}{1 \text{ cm}^{-3}} \right)^{-1/3} \left( \frac{M_{\text{ejecta}}}{M_{\text{sun}}} \right)^{5/6} \text{ yr} \quad (10.1)$$

The only nearby supernova that is in the free-expansion phase is SN 1987A.

### 10.1.2 Sedov-Taylor phase

In the Sedov-Taylor phase, the swept-up mass dominates over the ejected mass, and introduces a gradual slowing of the shock. In this phase, the total momentum is conserved. Auto-similar solutions exist, that describe the adiabatic evolution of the expanding shock wave.

Dimensional arguments can be used to derive the shape of the solution: the ejected mass does not play a role any more. The interstellar medium pressure is still too low compared to the shock speed to play any role either. Only the swept-up mass, the radius, and the time count. The only way to form an energy from these variables is through:

$$[E] = [M][L^2][T^{-2}] \quad \Rightarrow \quad E \propto M \times R_S^2 \times t^{-2} \quad (10.2)$$

The swept-up mass is $M = \frac{4}{3} \pi R^3 n_{ISM}$, thus $E \propto R^5 t^{-2} n_{ISM}$. In the first approximation, the fraction of the energy in the form of kinetic energy remains constant and the solution can be expressed as:

$$R_S \propto \left( \frac{E}{n_{ISM}} \right)^{1/5} (t - t_{ST})^{2/5}$$

$$V_S = \frac{dR_S}{dt} \propto (t - t_{ST})^{-3/5}$$

In this phase, the main shell of the SNR suffers Rayleigh-Taylor instabilities and the SNR ejecta are mixed up with the gas that was just shocked by the initial shock wave. This mixing also enhances the magnetic field inside the SNR shell. The Sedov-Taylor phase typically lasts 10,000 - 20,000 years.
10.1.3 Pressure-driven snow-plough phase

When the temperature of the shock wave decreases below $10^6$ K, radiative processes start to cool the medium efficiently. In this phase, the energy is not conserved any longer, and the expansion remains driven by the conservation of momentum.

This phase is actually subdivided in two sub-phases. When the medium inside the SNR is still hot, the thermal pressure inside the remnant governs the evolution and the energy losses are negligible (adiabatic pressure-driven snow-plough phase). When the medium cools down, the thermal pressure inside the remnant becomes negligible and the evolution is governed by momentum conservation. This phase starts when the time-scale of radiative losses becomes similar to the remnant age, at:

$$t_R \approx 19 \times 10^3 \left( \frac{E_{SNR}}{10^{51} \text{ erg}} \right)^{-4/17} \left( \frac{n_{ISM}}{1 \text{ cm}^{-3}} \right)^{-9/17} \text{ yr} \quad (10.3)$$

Using momentum conservation:

$$\frac{d}{dt} \left( \frac{4}{3} \pi R_S^3 n_{ISM} V_S \right) = 4\pi^2 R_S^2 P_{int}$$

(10.4)

When the evolution is pressure driven and the losses negligible (adiabatic evolution) one has $P_{int} V_S^2 = C_{ste}$, thus $P_{int} \propto R_S^{-35} \propto R_S^{-5}$. From Eq. (10.4) we obtain:

$$\frac{d}{dt} \left( R_S^3 \dot{R}_S \right) \propto R_S^2 R_S^{-5}$$

(10.5)

Searching for a solution in the form $(t - t_P)^{-n}$ gives $n = 2/7$. The evolution equation then reads

$$R_S \propto (t - t_R)^{2/7}$$

$$V_S \propto (t - t_R)^{-5/7}$$

Using Rankine-Hugoniot conditions, the downstream temperature decreases with time according to:

$$T_2 \propto \frac{P_2}{\rho^2} \propto (t - t_R)^{-6/5}$$

(10.6)

10.1.4 Momentum-driven snow-plough phase

When the temperature has dropped enough, the pressure becomes negligible and the shell moves with constant radial momentum. At this stage, electrons begin recombining with the heavier atoms (like oxygen) so the shell can more efficiently radiate energy. This, in turn, cools the shell faster, making it shrink and become more dense. The more the shell cools, the more atoms can recombine, creating a snowball effect. Because of this snowball effect, the SNR quickly develops a thin shell and radiates most of its energy away as optical light.

Momentum conservation then reads:

$$\frac{d}{dt} \left( \frac{4}{3} \pi R_S^3 n_{ISM} V_S \right) = 0 \Rightarrow \frac{d}{dt} \left( R_S^3 \dot{R}_S \right) = 0 \Rightarrow R_S^3 \dot{R}_S = R_c^3 \dot{R}_c$$

(10.7)

where $R_c$ is a characteristic radius. This can be solved into:
\[ R_S = R_c \left[ 1 + \frac{4\dot{R}_c}{R_c} (t - t_c) \right]^{1/4} \propto (t - t_c)^{1/4} \]
\[ V_S = \dot{R}_S = \dot{R}_c \left[ 1 + \frac{4\dot{R}_c}{R_c} (t - t_c) \right]^{-3/4} \propto (t - t_c)^{-3/4} \]

### 10.1.5 Dilution phase

After millions of years, the SNR will be absorbed into the interstellar medium due to Rayleigh-Taylor instabilities breaking material away from the SNR’s outer shell. The shell speed is in this phase smaller than the sound speed.

Fig. [10.2](#) summarizes the shell radius evolution as function of time for the various phases.

![Figure 10.2](image_url)

**Figure 10.2:** Summary of evolution phases of a supernova remnant.

### 10.1.6 Maximum energy of accelerated particles

The maximum particle energy is governed by three different constraints:

- the acceleration time to reach that energy must be smaller than the supernova age,
- the shell thickness must be larger than the diffusion length to avoid particle escape
- the radiative time-scale must be larger than the acceleration time-scale

**Acceleration time limit**

Using the acceleration rate derived for first order Fermi acceleration as a function of the compression ratio \( r \):
\[ E_{\text{max}} = \int_0^t \frac{dE}{dt} \, dt = \int_0^t \frac{r - 1}{r(r + 1)} Z e B V_s^2 \, dt \]  

(10.8)

For typical parameters of a supernova remnant and a strong shock, one obtains:

\[ E_{\text{max}} \sim 12 \text{ TeV} \left( \frac{B}{1 \text{ nT}} \right) \left( \frac{V_s}{5 \times 10^3 \text{ km s}^{-1}} \right)^2 \left( \frac{t}{100 \text{ yrs}} \right) \]  

(10.9)

**Radiation limit**

The radiative limit results from the comparison of acceleration and cooling time scales:

\[ \left( \frac{dE}{dt} \right)_{\text{acceleration}} + \left( \frac{dE}{dt} \right)_{\text{cooling}} = 0 \]  

(10.10)

For synchrotron cooling,

\[ \left( \frac{dE}{dt} \right)_{\text{cooling}} = -b U E^2 \text{ with } U = \frac{B^2}{2\mu_0} \]  

(10.11)

Thus:

\[ E_{\text{max}}^2 = \frac{1}{b U B} \frac{r - 1}{r(r + 1)} Z e B V_s^2 \]  

(10.12)

For typical SNRs, one finds:

\[ E_{\text{max}} \sim 240 \text{ TeV} \left( \frac{B}{1 \text{ nT}} \right)^{-1/2} \left( \frac{V_s}{10^4 \text{ km s}^{-1}} \right)^2 \]  

(10.13)

**Particle escape**

The particle escape limit results from the requirement that the diffusion length of the particles must be less than the thickness of the shell. The typical diffusion distance of the particles is \( L_d = D/v \) where \( D \) is the diffusion coefficient. In the Bohm diffusion regime (which corresponds to a minimum of the diffusion coefficient):

\[ D = \frac{p c}{3 Z e B} = \frac{E}{3 Z e B} \]  

(10.14)

Particles can escape if \( L_d \geq \eta R_S \), where \( \eta \leq 1 \) is a factor controlling the escape of particles. This results into a maximum energy:

\[ E_{\text{max}} = 3\eta Z e B V_s R_S \]  

(10.15)

In the adiabatic evolution phase:

\[ R_S = R_{ST} t^{2/5} \quad \text{and} \quad V_S = V_{ST} t^{-3/5} \]  

(10.16)

Thus,

\[ E_{\text{max}} = 3\eta Z e B V_{ST} R_{ST} t^{-1/5} \]  

(10.17)

For typical SNRs:
\[ E_{\text{max}} \sim 125 \text{ TeV} \left( \frac{\eta}{0.2} \right) \left( \frac{E}{10^{51} \text{ erg}} \right)^{2/5} \left( \frac{n_0}{1 \text{ cm}^{-3}} \right)^{-2/5} \left( \frac{t}{1000 \text{ yrs}} \right)^{-1} \] (10.18)

This is illustrated in Fig. [10.3] for a supernova remnant. The maximum energy is first limited by the age of the remnant (compared to the acceleration time-scale) in young supernova remnants. As the remnant becomes older, the shock slows down, the acceleration time increases and the maximum energy becomes limited by radiative losses.

![Figure 10.3: Maximum energy as a function of age for a Supernova Remnant.](image)

10.2 Magnetic field in SNRs

In recent years, a growing body of evidence for magnetic field amplification in the shocks of SNRs has emerged. Measurements of thin filaments in SN 1006 with Chandra [154] [155] strongly suggest rapid electron cooling in intensified magnetic fields of the order of 0.1 mG. However, alternate models [e.g. [156]] attribute the observed thin filaments to a damping of the magnetic field by downstream instabilities.

Magnetic field amplification has been observed in six different young supernova remnants [157]. Quantitative measurements of field amplitude versus shock speed tend to indicate the field amplification is larger for large velocities and thus for younger remnants. This is fully consistent with DSA theory which predicts that field amplification in SNRs should decrease linearly with the shock velocity.
More recently, the discovery of the brightening and decay of X-ray hot spots in the shell of the SNR RXJ1713.7-3946 on a one-year timescale [158] (Fig. 10.4) indicates magnetic field amplification factors of the order of 100 compared to typical values in the interstellar medium. This rapid variability indicates that the X-rays are produced by ultra-relativistic electrons through a synchrotron process and that electron acceleration does take place in a strongly magnetized environment. This variability also allows to put constraints on the electron acceleration time scales.

Figure 10.4: Chandra X-ray images of the western shell of SNR RX J1713.723946. b): A sequence of X-ray observations in July 2000, July 2005 and May 2006 for a small box (labelled b) in the main map a. c): Hard-band (3.5–6-keV) images of a box (labelled c) in a. Yellow arrows indicate bright X-ray hot spots that appear and decay on a one-year timescale. From [158]

Magnetic field amplification is a major result of DSA theory [e.g. 159 160 161], where it arises from non-linear retro-action of hadronic cosmic rays. It can increase the rate of acceleration and facilitate acceleration to higher energies, well above $10^{15}$ eV and possibly up to $10^{17}$ eV. Observation of high magnetic fields in SNRs therefore strongly support acceleration of hadrons up to PeV energies.

Recent measurements using thermal Doppler broadening of the Hα line on RCW 86 [102] indicate that the post-shock temperature is significantly lower ($2.3 \pm 0.3$ keV) than previously expected from the measured shock velocity ($42 - 70$ keV). This was attributed to a large acceleration efficiency, resulting in a cosmic-ray induced pressure that can exceed the thermal pressure behind the shock, thus giving further support to the hypothesis of magnetic field amplification in SNRs.

The amplified magnetic field also leads to a efficient cooling of electrons by synchrotron emission, thus leading to a relative depression of the inverse-Compton and non-thermal Bremsstrahlung γ ray emission, whereas the hadronic emission remains unaffected. This also implies that the highest-energy γ rays emitted by SNRs should only be observable
during the early evolution phases, roughly at the beginning of the Sedov-Taylor phase. Subsequent escape of high energy particles then strongly limits the maximum energy.

10.3 A particular SNR expanding in a relatively uniform environment: the case of SN 1006

Among the aforementioned SNRs, SN 1006 is, due to its position 500 pc above the Galactic Plane, an ideal case to study particle acceleration mechanisms. It indeed expands into a relatively uniform, low density \( n \sim 0.085 \text{ cm}^{-3} \) medium [163] and a uniform magnetic field. Moreover, SN 1006 is one of the best-observed SNRs with a rich data-set of radio, X-ray and optical measurements. The progenitor of SN 1006 is believed to be a type Ia supernova [163], probably the brightest supernova in recorded history. SN 1006 was detected by H.E.S.S. after a deep observation (130 h of data) and clearly exhibits a bipolar morphology (Fig. 10.5), strongly correlated with the non-thermal emission measured by XMM-Newton [160].

![Figure 10.5: H.E.S.S. correlated γ ray excess map of SN 1006. The linear colour scale is in units of excess counts per \( \pi \times (0.05^\circ)^2 \). The white contours correspond to a constant X-ray intensity as derived from the XMM-Newton flux map and smoothed to the H.E.S.S. point spread function, enclosing respectively 80%, 60%, 40% and 20% of the X-ray emission. Reproduced from [160].](image)

The close correlation between X-ray and VHE-emission demonstrated by the radial
profile (Fig. 10.6) points toward particle acceleration in the strong shocks revealed by the Chandra observation of X-ray filaments [155]. Moreover, the bipolar morphology of the VHE emission in the NE and SW regions of the remnant supports a major result of diffusive shock acceleration theory, according to which efficient downstream injection of supra-thermal charged nuclear ions is only possible for sufficiently small angles between the ambient magnetic field and the normal to the shock. Assuming a relatively uniform magnetic field oriented in the NE-SW direction, a higher density of accelerated nuclei at the poles is predicted [167] [168] [169].

Figure 10.6: Radial profile around the centre of the SNR obtained from H.E.S.S. data and XMM-Newton data in the 2 - 4.5 keV energy band smoothed to H.E.S.S.. PSF. Reproduced from [146].

Three different models were investigated to account for the spectral energy distribution (SED), and are compared in Fig. 10.7. In a purely leptonic model (dotted blue line), TeV emission results from inverse-Compton scattering of multi-TeV electrons. The resulting magnetic field then needs to be higher than 30 μG so that the IC emission does not exceed the measured VHE flux. In a second, dominantly hadronic model (dashed red line), TeV emission results from proton-proton interactions with π⁰-production and subsequent decay, whereas the X-ray emission is still produced by leptonic interactions. The resulting magnetic field needs to be higher (≈ 120 μG), which is consistent with magnetic field amplification at the shock, as indicated by the measurements of thin X-ray filaments mentioned above, but requires that a very high fraction (about 20%) of the supernova energy was converted into high-energy protons to account for the TeV emission. A mixed model (solid black line), in which hadronic and leptonic processes contribute almost equally to the very high-energy emission, gives a good description of the data, with a more reasonable overall acceleration efficiency of ≈ 12%. Discrimination between the different models might be possible from observations in the MeV domain, however no instrument with the required sensitivity is currently in operation.
Figure 10.7: Spectral energy distribution of the supernova remnant SN 1006, compared with various models.

Recent, high precision observation of the synchrotron emission covering the entire rim of the shell [170] led to the clear observation that the synchrotron cut-off frequency $\nu_{\text{break}}$ varies significantly with azimuthal angle (in a way that is consistent from earlier observations [166]), the cut-off frequency increasing in the non-thermal limbs where efficient acceleration is supposed to occur.

This allows the dependence of the connection between acceleration efficiency and particle injection in the shock on the orientation of the magnetic field to be tested.

### 10.4 Are the supernovae remnant hadronic accelerators?

Recent multi-wavelength observations of supernova remnants are progressively leading to the emergence of a consistent picture of particle acceleration up to very high energies in these systems. Young supernova remnants have been proven to accelerate particles (electrons and/or protons) up to 100 TeV at least.

Observation of the bi-polar morphology in SN 1006, aligned with the direction of the interstellar magnetic field, strongly supports the current understanding of particle injection in the shock.

There is growing evidence, consistent to the prediction of DSA, that the acceleration efficiency can be as large as 50% due to the retro-action of cosmic rays on the shock. Magnetic field amplification, predicted in the framework of DSA, is now convincingly observed in several objects, thus further supporting the hypothesis of hadron acceleration in SNRs. $\gamma$ ray emission at 100 TeV is difficult to achieve with inverse-Compton scattering due to Klein-Nishina suppression of the cross section at high energies.

The debate remains inconclusive however with some authors, using comprehensive
time-dependent models based on non-linear diffusive shock acceleration [e.g. [171]] arguing strongly in favor of hadronic acceleration, whereas others [172], using a consistent calculation of thermal X-ray emission, predict very intense X-ray thermal Bremsstrahlung emission in the hadronic scenario, in contradiction with observations.

Recent observation of RX J1713.7-3946 with the Fermi-LAT telescope [173], shown in Fig. 10.8 still suggest that the dominant emission mechanism might be leptonic, although a significant contribution from hadronic cosmic rays cannot be excluded.

![Figure 10.8: Fermi observation of the young supernova remnant RX J1713.7-3946 compared to leptonic (left) and hadronic (right) emission models. From [173]](image)

One hundred years after their discovery, cosmic rays still cannot be firmly associated with supernova remnants though there is a growing body of evidence that there are. Further progress might arise from the observation of cosmic rays escaping from old supernova remnants and interacting with interstellar molecular clouds.

### 10.5 Old SNRs in interaction with molecular clouds

Further insight into the acceleration mechanisms in SNRs may come from the observation of middle-aged supernova remnants in the vicinity of dense molecular clouds [174]. Indeed, production of $\gamma$ rays is expected both during the acceleration phase of cosmic rays in the SNR shock and during their subsequent propagation. Molecular clouds can act as cosmic-ray barometers, with an enhanced $\gamma$ ray emission proportional to the cloud mass, whereas inverse-Compton emission from accelerated electrons would not be enhanced. Furthermore, the $\gamma$ ray emission from a cloud depends on the propagation time and can therefore last much longer than the emission from the SNR, thus making the detection of clouds more probable [175]. A specific spectral signature, in the form of a concave energy spectrum, has been predicted by [174] but not observed so far.

Four old SNRs in the vicinity of molecular clouds have been detected by H.E.S.S.: W 28 [176], an old ($\sim 35 - 150$ kyr) mixed-morphology SNR, HESS J1745-303 [177], HESS J1714-385 [178] and more recently HESS J1923+141 [179]. In W 28, H.E.S.S. data show VHE emission from four different spots coincident with HII regions and dense molecular clouds revealed by NANTEN $^{12}$CO($J = 1 - 0$) data [180]. Interaction with
a molecular cloud along its north and north-eastern boundaries is further confirmed by
the high concentration of 1720 MHz OH masers [181]. Under the assumption of a cloud
distance between 2 and 4 kpc, a cosmic-ray overdensity in the range 13 to 32 is derived,
which is consistent with expectations.

Figure 10.9: Image (1.5° × 1.5°) of the VHE γ-ray excess counts (events), corrected for
exposure and smoothed with a Gaussian of radius 4.2′ (standard deviation). Overlaid
are solid green contours of VHE excess (pre-trial) significance levels of 4, 5, and 6σ, after
integrating events within an oversampling radius θ=0.1° appropriate for point-like sources.
From [176].

Similar conclusions are derived from observations of HESS J1745-303, HESS J1714-
385 and HESS J1923+141, thus confirming that shell-type SNRs are efficient accelerators
of hadronic cosmic rays. Three of these old SNRs in the vicinity of clouds are firmly
associated with bright Fermi sources [182]: W 28 is associated with 1FGL J1801.3-2322c
and 1FGL J1800.5-2359c; HESS J1714-385 is associated with 1FGL J1714.5-3830 and
HESS J1923+141 with 0FGL J1923.0+1411. A possible, less convincing, association also
exists for HESS J1804-216. Spectral continuity between GeV and TeV energies, large
cosmic-ray densities derived from the VHE flux and coincidence with the shocked clouds
strongly favour a hadronic origin of the GeV and TeV γ rays.

Further multi-wavelength studies with ACTs and Fermi will help to understand the
mechanisms at the origin of Galactic CRs.
Chapter 11

Diffuse Galactic Emission

The high-energy (GeV scale) emission of the Galaxy is fully dominated by the Galactic diffuse emission resulting from the interaction of charged cosmic rays (CR) with the interstellar molecular and atomic gas material filling the Galactic Plane. This is beautifully illustrated by the Fermi view of the Galactic Plane above 100 MeV (Fig. 11.1) [183].

Figure 11.1: Fermi LAT counts above 100 MeV, having in total more than $3 \times 10^7$.

Analysis of high energy sources in the Galactic Plane therefore relies on the existence and quality of a diffuse emission model built from several elements: emission produced by the nuclear interaction of cosmic rays with the interstellar medium, inverse-Compton (IC) scattering of electrons, Bremsstrahlung, and contribution of strong sources. Numerical models of cosmic-ray propagation and interactions inside the Galaxy such as GALPROP [184] [185] have been successfully developed for more than ten years. Although they have produced quite precise predictions and are used on a daily basis to compare predictions with actual observation, these methods suffer from large systematic uncertainties due to the incomplete knowledge of the interstellar medium (radiation field, matter
density, magnetic field). Analysis in the GeV domain is also considerably complicated by the confusion limit, thus requiring a superposition model adjusted on the data using log-likelihood approaches.

In the VHE domain, two prominent emission mechanism might contribute to the diffuse emission, as discussed below:

11.1 Diffuse hadronic component

A diffuse hadronic component is expected to result from the interaction of accelerated cosmic rays with the interstellar medium. Due to the small energy losses, nuclei and protons are characterized by a long range propagation dominated by the diffusion on the magnetic field inhomogeneities. Due to the increase of the diffusion coefficient with energy, the distribution of high-energy protons and nuclei becomes more and more homogeneous at the Galactic scale as the energy increases. The hadronic component therefore acts as a tracer of matter (target) density in the Galaxy. Observation of VHE diffuse emission could therefore bring a new measurement of the matter distribution in the Galaxy that could complement the existing spectroscopic surveys (CO, CS, ...). The interpretation of the latter depends in particular on the dynamic model of the Galaxy and are therefore affected by uncertainties that an independent measurement could help to reduce.

Due to a relatively flat high-energy proton-proton cross-section, the energy spectrum of gamma-rays resulting from the decay of neutral pions produced in cosmic-ray interaction almost reflects the original cosmic-ray spectrum. VHE gamma-ray observations give a handle on the cosmic-ray spectrum throughout the Galaxy. This view can complement the view from the Earth that we have from direct balloon and space-born experiments. Detailed analysis of the cosmic-ray spectrum as a function of Galactic latitude and longitude can provide precise tests for propagation models. In particular, due to the increase of the diffusion coefficient with energy and due to re-acceleration of cosmic rays in the interstellar medium, the cosmic-ray spectrum at the Earth is expected to be significantly steeper than the spectrum at the source. Although the high-energy diffuse emission is observed by Fermi in the GeV region (where the hadronic component most likely dominates), recent studies [e.g. 186] indicate that the very-high-energy spectrum (and composition) of cosmic rays throughout the Galaxy remains quite uncertain, making independent measurements toward the source highly desirable. Propagation effects that affect the cosmic-ray spectrum at the Earth are a composite of local (near sources, illuminated molecular clouds) and global effects (CR anisotropies, expanding bubbles, stochastic effects). Last but not least, very-high-energy measurements are also affected by the uncertainties on nuclear cross sections above 1 TeV. Uncertainties on the gamma-ray flux that arise from a particular choice of the photon production cross-section and the primary CR protons and helium flux a reach a factor of 2 at 1 TeV.

Although the hadronic component is expected to mainly reflect the large extended structures in galactic matter distribution, young sources could contribute to localized excess. Comparison of VHE diffuse emission with spectroscopic surveys could help identify new classes of cosmic-ray accelerators, thus adding new clues to the cosmic-ray puzzle.

Nearby galaxies, as they are viewed from outside, do not suffer problems with line of sight integration and are therefore good targets for investigating the propagation of cosmic rays. In that sense, the Large Magellanic Cloud (LMC) is an excellent target for studying the link between CR acceleration and gamma-ray emission since the galaxy is nearby (≈ 50 kpc), and is seen almost face-on. The LMC was detected with Fermi-LAT
Figure 11.2: Comparison of local effective emissivity of the ISM per hydrogen atom as measured by Fermi (black points) compared to various models of photon production. From [186].

(> 33σ) and for the first time the emission is spatially well resolved in gamma rays [187].

Surprisingly, the observations reveal little correlation of the gamma-ray emission with gas density, as expected if CRs propagate throughout the entire galaxy. The gamma-ray emission correlates more with tracers of massive star forming regions, supporting the idea that CRs are accelerated in these regions by stellar winds and supernova explosions of massive stars.

11.2 Inverse-Compton diffuse emission

In contrast to protons and nuclei, electrons are affected by strong radiative losses throughout their propagation in the Galaxy: ionization and Bremsstrahlung on the interstellar medium (below GeV energies), synchrotron radiation on the Galactic magnetic fields and inverse-Compton scattering on the cosmological, diffuse infrared and optical backgrounds which result into energy losses that quickly increase with energy \( (dE/dt \propto E^2) \) for synchrotron and IC scattering in the Thomson regime). As IC scattering depends on the electron and the interstellar radiation field densities, its intensity can trace the popula-
tion of freshly accelerated electrons in the neighbourhood of the sources and can therefore bring constraints on the particle propagation at very high energy in a domain that is still poorly known. It can also bring strong constraints on the distribution of sources [e.g. [158].

Outside the Galactic Plane, the diffuse target background for IC scattering is also dominated by the CMB and is therefore homogeneous. Diffuse VHE gamma-ray emission would therefore trace the population of accelerated electrons. In the Galactic Plane, radiation fields are dominated by infrared and optical emission that are strongly correlated with matter content, but in the VHE energy regime, Klein-Nishina effects will result into a strong suppression of the IC scattering on the visible and infra-red components. In that case as well, CMB will be the main target for IC photons and diffuse emission will also act as a tracer of the accelerated electrons population.

It has to be noted that, although the electron energy spectrum at the source is often much steeper than for hadronic cosmic rays, the $E^2$ dependence of the energy losses naturally produces a hard IC spectrum. In the very high energy range, the IC component might then dominate over the hadronic component. The possible observation of a break in the IC diffuse emission would bring strong constraints in the acceleration limit of leptonic sources.

11.3 Relative contributions

The relative contribution of hadronic and leptonic diffuse emission is not very easy to estimate accurately. However, at a first glance, supernova remnants mainly accelerate protons and nuclei. The fraction of accelerated electrons, estimated from the cosmic ray composition at the Earth, might be of the order of 1% above 1 GeV. This might result from a less efficient escape of electrons that are trapped by the large magnetic fields inside supernova remnants. In contrast, pulsar wind nebula mainly accelerate electrons and positrons created in the pulsar magnetosphere. In these systems, lower magnetic fields compared to those in SNRs result into a more efficient release of accelerated particles into the interstellar medium. Given the fact that pulsars mainly originate from core-collapse SNR, and taking into account a reasonable acceleration efficiency, the hadronic and leptonic diffuse contributions at TeV energies might be of the same order of magnitude.

11.4 Diffuse emission close to accelerators

VHE instruments such as H.E.S.S. have revealed a large population of accelerators belonging to the classes of SNRs and PWNs, among others. But the way accelerated particles are released into the interstellar medium remains largely unknown. Indeed, magnetic field amplification in supernova remnants leads to a highly amplified synchrotron emission inside the remnant. In contrast, particles are released into a low-magnetic-field interstellar medium ($B \sim 5 \mu G$) and become therefore almost invisible as soon as they leave the remnant. In a similar way, emission from neutral pions is intensified by the larger density at the shock and decreases behind the shock. IC emission of accelerated particles on the CMB (due to Klein-Nishina effects, IC emission on CMB might dominate over infra-red and visible) would, in contrast, not be extinguished outside the source and might therefore bring a handle on particles that escape the source [174] [189]. Detailed predictions of energy spectra from run away cosmic-rays and electrons are difficult to establish precisely.
due to competing phenomena: high energy particles tend to escape first from the source, but tend to get more diluted due to diffusion. Depending on the relative time scales, the spectrum of runaway cosmic-rays can become softer or harder than that of trapped CRs. In addition, electrons can suffer strong energy losses due to IC scattering and synchrotron cooling leading to a spectral steepening away from the source [15].

11.5 Observational evidence for TeV diffuse emission

Due to the steep spectrum of cosmic rays, the large effective area and the good angular resolution (better than 0.1°) of Imaging Cherenkov Telescopes the assessment of diffuse emission at relatively small scales (a few degrees) becomes more favourable at higher energies. The H.E.S.S. view towards the Galactic Center (Fig. 11.3 [26]) revealed diffuse emission associated with molecular clouds illuminated by cosmic rays, however at considerably smaller scale than the diffuse emission seen in GeV. The hard spectrum observed in TeV γ rays and the high cosmic-ray flux (a factor of 3-9 higher than in the Earth neighbourhood) have been interpreted as a strong support to Galactic Diffusion models: clouds illuminated by cosmic rays produced in the neighbourhood reflect the energy spectrum of cosmic rays at the source, whereas cosmic rays at the Earth are affected by propagation through the Galaxy. As the Larmor radius of charged cosmic rays in the Galactic magnetic field increases with energy, the diffusion coefficient increases and high energy particles are less confined inside the Galaxy. This leads to a steepening of the cosmic ray spectrum away from the sources.

At energies around 10 TeV, Milagro revealed a extended gamma-ray emission localized in the Cygnus region [190] near Galactic longitude 80°. This region is in a direction tangential to our spiral arm located approximately at the same distance from the Galactic centre as the solar system. It provides the opportunity for the most accurate determination of the gas distribution based on velocity information and the Galactic rotation curve. It also contains a large column density of interstellar gas and is also the home of potential

![Figure 11.3: TeV Diffuse emission around the Galactic Centre. Reproduced from [26].](image)
CR accelerators (Wolf-Rayet stars, OB associations and SNRs), thus making it a natural laboratory for the study of cosmic-ray origin.

Observations from Milagro showed that TeV emission is correlated with matter density (shown as black contours on Fig. 11.4) and is a factor of ~7 more intense compared to the conventional GALPROP model. Solutions to this excess invoking an underestimation of the diffuse emission by GALPROP have been proposed [191]. To disentangle possible solutions to this excess (incorrect diffuse emission prediction from GALPROP at these energies, contribution of unresolved TeV sources, population of high energy electrons or population of dark cosmic-ray accelerators), deeper observations with higher sensitivity instruments (such as ACTs) are required.

At lower energies, around 200 GeV, deep observations to explore the low energy limits of VHE astronomy revealed that, close to the threshold of the instrument, diffuse emission seems to appear at many places in the Galactic Plane, as illustrated in Fig. 11.5.

This emission, which seems to be well concentrated around the Galactic Plane, is however difficult to establish firmly due to the large range of possible systematic effects occurring at low energies. In particular, ACTs have a rather limited field of view (of the order of a few degrees), and are background dominated. Moreover, the background, mainly consisting of cosmic-ray-induced showers (and electron-induced showers) is subject to large fluctuations related in particular to atmospheric variations. Background subtraction techniques, such as those described in section 7.2, rely on the estimation of the background level in the same field of view, either using different particle populations (template model) or using different locations around the source (ring model). In both cases, any region possibly containing γ-ray events (including diffuse emission) must be excluded from the background estimation, thus making the analysis of structures larger than or similar in size to the field of view impossible at present.

Figure 11.4: Cygnus region of the Galaxy as seen in TeV gamma rays by Milagro. Reproduced from [190]
11.6 Towards diffuse emission at VHE energies

Recent progress in the understanding of the response of Atmospheric Cherenkov Telescopes, in particular through the work presented in this thesis, should hopefully lead to the firm identification and characterization of diffuse emission at VHE energies along the Galactic plane. Although this appears to be possible on relatively short time scales, careful and precise work on several issues needs to be done with special care and precision:

- Determination of acceptance: detailed treatment of systematic effects, such as shower distortion due to magnetic field effect and variation of acceptance related to fluctuations of night sky background have to be incorporated, if possible, at the reconstruction phase. Effects on the expected level of background need to be assessed quantitatively.

- Model of instrument response: identification of large scale diffuse emission would require that the current need for a normalization region in every observation be overcome. This would require significant work in the analysis of the relation between atmospheric conditions and background event rate in cuts. Detailed monitoring of the atmosphere would also have to be used at the analysis stage, not only to select good quality data.

- Assessment of unresolved source contribution: In order to disentangle true diffuse emission from the contribution of unresolved sources, the contribution of the latter needs to be determined as accurately as possible. From the deep H.E.S.S. observation in the Galactic Plane, it is already possible to derive a luminosity function (Log-N-log-S) for several source classes (Pulsar Wind Nebulae, Supernova Remnants). Incorporating information from other wavelengths (population of young pulsars, catalogue of supernovae remnants, ...), an assessment of the contribution of unresolved sources could be obtained. Also, latitudinal and longitudinal projections of the observed gamma-ray fluxes will provide another handle on the precision of the unresolved component estimation. After subtraction of the homogeneous diffuse
component, such as that of cosmic electrons, we would then be able to derive the true diffuse emission from the Galactic Plane, and to compare it with propagation models.

Although this is very challenging, we believe that the firm identification of Galactic diffuse emission along the Galactic plane will be one of the major challenges of the Atmospheric Cherenkov Telescopes during the next decade.
Chapter 12

Active Galactic Nuclei

Introduction

Active Galactic Nuclei are a major class of VHE sources with nearly 50 objects detected to date, and they constitute the vast majority of extra-galactic VHE sources. AGN are comprised of a super-massive black hole with a mass from millions to billions of solar masses surrounded by a large accretion disk. Recent estimates indicate that nearly 1% of all galaxies are categorized as active and have an emission from the core that significantly outshines all the stars in the galaxy.

In about 10% of the cases, they additionally exhibit two large scale jets of collimated, highly relativistic outflows that can reach sizes of several Mpc, far larger than the extension of the Galaxy.

Most (∼ 90%) of active galaxies are radio-quiet Seyfert Galaxies with a bolometric luminosity \( L_{bol} \sim 10^{36-38} \text{W} \). The remainder are radio-loud, seen face-on, and are significantly brighter (bolometric luminosity \( L_{bol} \sim 10^{38-41} \text{W} \)), with a contribution from the jet \( L_{jet} \sim 10^{48} \text{W} \).

Active Galactic Nuclei have been classified according to spectral properties (presence or absence of emission and absorption lines, Doppler broadening of the lines . . .), but this classification is mainly the result of the angle with which the system is seen. Fig. [27] shows the unified model of active galactic nuclei, basically divided in two intrinsic classes: active galactic nuclei with jets (upper part), showing intense radio emission, and without (lower part).

Seyfert galaxies are in particular categorized in two classes depending on the viewing angle:

- Seyfert 1 galaxies exhibit both broad and narrow emission lines. Broad emission lines are caused by Doppler broadening \( (V \sim 1000 \text{ km/s}) \) from high velocity clouds above the accretion disk.

- Seyfert 2 galaxies only exhibit narrow lines, the broad-line regions being hidden by an obscuring torus.

AGNs that have their jets approximately pointing toward the Earth are called blazars. This class comprises, among others, BL Lac (Lacertae) objects, that are characterized by almost featureless spectra showing very few emission lines originating from very weak emission from the accretion disk. The luminosity of these objects is completely dominated
by the featureless non-thermal emission from the jets. Optically Violent Variable (OVV) and other radio loud quasars belong to the class of blazars.

Blazars are characterized by strong and rapidly variable γ ray emission (variable on time scales of minutes or even less, as shown in Fig. [12.2] originating from the jets, and thus beamed toward the earth.

With the exception of two radio galaxies (M 87 [193] and Centaurus A [194]) all AGNs observed at VHE energies belong to the class of blazars. High-frequency-peaked BL Lac objects (HBLs) dominate the sample.

12.1 Very High Energy emission from AGN

12.1.1 Variability of AGNs in Very High Energy

The variability time scale of the VHE emission from AGNs can be used to set limits on the size of the emission region. For a emission region of size $R$, moving at relativistic speed with a Lorentz factor $\Gamma$ describing bulk motion of the jet toward the earth, the variability time scale cannot exceed, from causal arguments, $\Delta t = R/(\Gamma c)$. Short-time variability such as that observed in PKS 2155-304 (Fig. [12.2]) can constitute an important probe for cosmological and fundamental physics studies such as Lorentz invariance.

A too small size of the emitting region, given the observed luminosity, would however lead to unavoidable absorption of VHE photons by pair creation on the CMB or on the stellar light [e.g. 196]. The compactness (opacity) parameter at the energy $\epsilon_2$ for a source of luminosity $L_0(\epsilon_1)$ and size $l$ is given as a function of the Thomson cross section by:
12.1. VERY HIGH ENERGY EMISSION FROM AGN

Figure 12.2: Exceptional Very High Energy γ ray Flare of PKS 2155-304, in which the emission is found to be variable at time scales of a minute. From [195].

\[
C = \frac{L_0(\epsilon_1)\sigma_T}{16\pi\epsilon_1lc}
\]

where \(\epsilon_1 = m_e^2c^4/\epsilon_2\) is the energy of the photons that are able to pair create on the high energy photons. At 100 GeV, absorption is predominantly produced by visible photons of energy of the order of \(\sim 1\) eV. Compactness factors greater than 1 result in highly opaque sources, thus giving a lower limit on the source size:

\[
l_{\text{min}} = \frac{L_0(\epsilon_1)\sigma_T}{16\pi\epsilon_1c}
\]

For a typical Blazar in the MeV domain (e.g. 3C 279), the energy in the jet frame (before boosting) is \(\epsilon_1 \sim 511\) keV and \(L_0 \sim 10^{41}\) W. The minimum source size is then found to be \(l_{\text{min}} \sim 5.4 \times 10^{15}\) m = 0.17 pc and would then correspond to a variability time scale of the order of one year. Very large Doppler factors, of the order of 100 or more, are therefore required to solve this opacity paradox and to reconcile it with the observed variability, for emission regions with size comparable to the Schwarzschild radius of the super-massive massive black hole (\(\sim 10^{-5}\) pc). The emission in the direction of motion is then amplified by a factor \(\delta^4\), with \(\delta\) denoting the Doppler factor \(\delta = 1/(\Gamma(1 - \beta\cos\theta))\), thus drastically reducing the power requirement. The opening angle of the emission is then of the order of \(1/\Gamma\) and the compactness parameter is scaled as \(\delta^{4+\alpha}\) where \(\alpha \sim 1\) is the photon index of the source.

Unlike Galactic VHE sources, AGN all appear point-like given the \(\sim 5'$ resolution of IACTs, with position consistent with the nominal location of the AGN. As a consequence, Blazars cannot be used to pin-point the localization of the VHE emission. In these blazars, VHE γ ray emission is thought to arise inside the jets, from “blobs” of high-energy particles moving along the jet axis and leading to internal shocks. Flares would be created when high-energy electrons are injected into the blobs [e.g. 197].
12.1.2 Radio-Galaxies

Radio-Galaxies are active galactic nuclei that are not seen face-on, thus allowing, for those that are relatively nearby, the emission region to be identified. In contrast to blazars that are visible from cosmological distances due to strong Doppler boosting ($\delta^4$) of their high energy luminosity, Radio-Galaxies are however expected to be visible only at small distances (of the order of 20 Mpc).

M 87, the central galaxy of the Virgo Cluster, was the first radio galaxy detected in the TeV regime \[193\]. M 87 has a jet structure that is highly inclined from the line of sight, and is spatially resolved in X-ray (by Chandra), in optical and in radio observations. However the TeV emission had not been very well localized (Fig. 12.3) due to a insufficient angular resolution.

Figure 12.3: The 90 cm radio data (32) measured with the Very Large Array, together with the TeV position from H.E.S.S. with statistical and 20" pointing uncertainty errors (white cross) and again the 99.9% confidence level extension upper limit (red circle). The size of the emission region deduced from the short-term variability is smaller by a factor of $\sim 10^6$. The black cross marks the position and statistical error of the $\gamma$ ray source reported by HEGRA. The green ellipse indicates the host galaxy seen in the optical wavelengths with an extension of $8.3 \times 6.6$ arcmin in diameter. From [193].

A joint observation campaign from the three major ACTs took place in 2008 [198], leading to a total of 120 hours of data accumulated within 5 months. The multi-wavelength picture was provided by an intensive VLBA monitoring of the core of M 87 at 43 GHz
every 5 days. In February 2008, high TeV activity with rapid flares was detected and M 87 was contemporaneously monitored with high spatial resolution instruments in the X-ray band (Chandra). Simultaneous increased X-ray and radio activity were detected from the nucleus of M 87 whereas the central knot HST-1 was found to remain quiet. (Fig. 12.4)

This consistent increase of radio, X-ray, and TeV emission suggested that the emission originates from the central core and allowed to exclude HST-1 as the main source of emission, leading, for the first time, to the identification of the emission region in an active galactic nucleus.

Figure 12.4: Combined M 87 light curves from 2007 to 2008, allowing to pin-point the source of emission. From [198].
12.2 Possible extended VHE emission from Centaurus A

This section contains confidential results that are not yet published, and still under investigation. Given their potential high impact, they will be probably removed from the published manuscript.

12.2.1 Centaurus A

Centaurus A is the closest radio-galaxy \((d = 3.8 \text{ Mpc}, \text{ nearly ten times closer than M 87})\) and is classified as a type I Fanaroff-Riley radio galaxy which would correspond, according to the unified classification of AGN, to a mis-aligned BL Lac. The angle of the jet axis to the line of sight is estimated to be \(15 - 80^\circ\).

At radio wavelengths rich jet structures are visible, extending from the core and the inner pc and kpc jet to giant outer lobes with an angular extension of \(8^\circ \times 4^\circ\). Inner, kpc-scale jets are resolved in X-rays \([199]\), revealing a complex structure of bright knots and diffuse emission that are within the angular resolution capabilities of recent IACTs. For a review, see \([e.g. 200]\).

12.2.2 Indications for extension

The analysis of recent data with H.E.S.S., using the Model Analysis, seems to indicate the presence of an extended emission feature aligned with the X-ray and radio jets. Moreover, an enhanced emission is seen at the location of the radio lobes.

![Figure 12.5: Smoothed Excess Map of Centaurus A. Unpublished, confidential result.](image)
Deep exposure on Centaurus A was acquired by the H.E.S.S. experiment, between 2004 and 2010, for a total of 192 hours (corrected for dead time) passing quality selection cuts. Using the Model Analysis, 516 excess events were obtained, corresponding to a statistical significance of 11 \sigma. No significant variability was found at any time scale. The relatively large signal over background rate ($S/B \sim 0.3$) allowed the hypothesis of a point like emission to be tested quantitatively, using the tools described in section 7.6.

The excess map of Centaurus A obtained with H.E.S.S. and smoothed using the actual PSF of the instrument, is shown in Fig. 12.5. A comparison with the PSF shown as an inset suggests an extension of the emission in the NE-SW direction.

![Figure 12.6: Direction for projections](image)

Projections in the two diagonal directions shown in Fig. 12.6 are displayed in Fig. 12.7 with uncorrelated binning of the event counts. The projection along the NE-SW axis, with a Gaussian width of $0.04^\circ = 2.4'$ is fully compatible with the angular resolution of the array, whereas the extension along the NW-SE axis, with a Gaussian width of $0.072^\circ = 4.3'$ is not, thus confirming in a more quantitative manner the observed extension. The H.E.S.S. PSF for Model Analysis is $R_{68} = 0.06^\circ$ (68\% containment radius) corresponding to an RMS of 0.03 degrees, compatible with the width measured along the NW-SE direction.

Quantitative studies of the $\theta^2$ distribution exclude a point like emission at a statistical significance of 6 \sigma. Fit of the uncorrelated excess distribution by an elliptical shape yielded the following results:

\[
\begin{align*}
\text{RA} & = 13h25m33.15s \pm 2s \\
\text{Decl} & = -43^\circ 0'40.7'' \pm 22''
\end{align*}
\]
Figure 12.7: Projection of the TeV emission from Centaurus A projected along the NW-SE (minor) and NE-SW (major) axis. Unpublished, confidential result.
\[ \sigma_{\text{minor}} = 0'5'' \pm 30'' \]
\[ \sigma_{\text{major}} = 3'42.2'' \pm 30'' \]
\[ \phi = 37^\circ \pm 4^\circ \]

The extension in one direction is compatible with zero, whereas a significant extension of 3' is found in the other direction, with a statistical significance of 7 \( \sigma \). The best-fit position, slightly shifted by 1' with respect to the nucleus, is compatible within statistical uncertainties with the nucleus position. The resulting best-fit model (elliptical emission region) is shown in Fig. 12.8. The obtained azimuthal angles lines up perfectly in the direction of the radio jets at 21 cm, thus strengthening the case for extended emission along the jet.

![Fit Result Sky Map](image)

Figure 12.8: Result of the fit of a elliptical shape on the uncorrelated excess map of Centaurus A, and superimposed with the VLA emission at 21 cm. Unpublished, confidential result.

The confidence regions for the parameters \( \sigma_{\text{minor}} \), \( \sigma_{\text{major}} \) and \( \phi \) are shown in Fig. 12.9 with 1 and 3 standard deviation intervals indicated by green lines. Thus further confirms that observations are compatible with a null extension along the minor axis and a significant extension along the major axis. Fig. 12.10 shows the residuals after subtraction of the best fit model presented in Fig. 12.8. The residuals are compatible with purely statistical fluctuations, thus further strengthening the case for an elongated emission.

In order to investigate the possible superposition of two contributions, one originating from the core and the other one from the jets, a subtraction of a point source at the nucleus position has been performed. The residuals of this analysis are shown in Fig. 12.11 and
Figure 12.9: Confidence regions for the fit parameters: width of the minor axis (left), major axis (middle) and azimuthal angle (right). Unpublished, confidential result.

Figure 12.10: Residual emission from Centaurus A, after subtraction of an elliptical source shape, and superimposed with the VLA emission at 21 cm without any smoothing. Unpublished, confidential result.
Figure 12.11: Residual emission from Centaurus A, after subtraction of a point source, and superimposed with the VLA emission at 21 cm without any smoothing (upper) and with the VLA emission smoothed to the H.E.S.S. PSF (lower). Unpublished, confidential result.
are compatible with the VLA emission smoothed to match the H.E.S.S. PSF (Fig. [12.11] Bottom). The alignment of the H.E.S.S. structure with the jets is obvious and the bright lobes seem to correspond to enhanced TeV emission as well. The black dot indicates the best fit position for a point source.

12.2.3 Cross-checks

Various cross-checks have been performed to check the robustness of these results. These include in particular:

- Other analysis techniques, such as Hillas parameter based analysis. These do not have a sufficiently good angular resolution to reproduce these results, but do not however, exclude the analysis presented here.

- The data set has been divided in two parts with equal excess count. The extension appears in both data sets, although, as expected, with lower significance.

- The data set has been divided into high and low energy bands at the median energy of \( \sim 450 \text{ GeV} \). Similar shapes have been obtained in both bands with a possibly more pronounced effect at low energies.

- Other sets of cuts have been employed, such as the High Resolution Cuts, which keep only the events with the best angular resolution at the expense of statistics. Once again, very similar results are obtained.

- To investigate a possible problem arising from background subtraction, the ring background has been split in four quadrants corresponding to different regions on the sky. Results remained consistent.

- A full simulation of a point source has been performed to quantitatively assess the chance probability of finding such an extension. The results of this simulation is shown in Fig. [12.12] for 500 independent tries. The ingredients of the simulation are the following:

  - Poisson ON and OFF background derived from the actual exposure maps
  - Amount of gamma-ray derived according to Poisson statistics, on a run-by-run basis, with mean value set to the observed excess in each run
  - Actual PSF taken for each run to simulate event position. The PSF is obtained from gamma-ray Monte-Carlo using the actual parameters of the run (zenith angle, off-axis angle and optical efficiency)
  - Same binning and configuration as for real data

The width of the \( \lambda \) and \( \beta \) (position) distributions are very similar to the uncertainty derived on the best fit position, thus confirming that the statistical treatment is correct. The obtained extension (right plot) is \( 0.34 \pm 0.0053^\circ \) whereas the distribution obtained on simulations gives an average extension of \( 0.0023 \pm 0.03^\circ \), thus indicating that the measured extension is at 10 standard deviations away from expectation for point-like sources.

- Detailed analysis of the Blazar 1ES 1312-423, visible in the same field of view, and using the same data set, does not show any sign of elongation
Figure 12.12: Results of the 500 simulations of a point like source at the nominal position of Cen A, reconstructed with the same procedure. Unpublished, confidential result.

Conclusion

Significant extension of the emission from Centaurus A, aligned with the direction of the kpc-scale jets, has been obtained, thus allowing for the first time to investigate the localization of the VHE emission in blazars. Preliminary results indicate an emission at the base of the jets and close to the central nucleus. The statistical significance of the results is, depending on the tests, between 6 and 7 σ and extensive systematic cross-checks have been performed. In the absence of conclusive cross check from another analysis, further checks are still on-going. Further data will be accumulated in 2012, which will hopefully confirm this result.
Chapter 13
Starburst galaxies

13.1 Introduction

Starburst galaxies are essentially ordinary galaxies in which episodes of massive star formation occur, usually in a very localized region (referred as start-burst region) of a few thousand light years in diameter. During these star-burst events, the rate of star formation is increased by a factor of 10 to 100 times over that in normal galaxies.

Figure 13.1: Chandra X-ray image (inset) of the central region of the star-burst galaxy NGC 253 in comparison to the optical view. Chandra detects a proportionally high number of suspected intermediate-size black holes. Credit: X-ray: NASA/SAO/CXC, Optical: ESO
This star-burst activity is believed to be either triggered by external events, such as a galaxy merger, or induced by galaxy-bar instabilities that could distort the dynamical equilibrium of the interstellar gas. The archetypal star-burst galaxy is M 82, which interacted with the more massive galaxy M 81.

Starburst galaxies are therefore characterized by large concentrations of gas and dust and by a significant population of young and very massive stars. Winds from those stars lead to the formation of many HII regions through ionization of the surrounding gas in OB associations. Formation of massive, short-live stars leads to a large rate of supernova explosions which produce more shock waves and trigger more star formation. A chain reaction of star formation and supernova can sweep through the central region of the galaxy, where most of the gas is located. Episodes of intense stellar formation are usually relatively short, and do not last more than 10 million years, after which the interstellar gas in the formation region becomes exhausted or blown away by the explosions. Due to their population of very massive stars, star-burst galaxies are among the most luminous of all galaxies.

Some star-burst galaxies also show activity from their nucleus. It is possible that both types of phenomenon – burst of star formation and nucleus activity – may be triggered by galactic collisions and mergers.

New-born stars usually remain surrounded by dust and gas for million of years, which absorbs their light and radiates the heat generated as infrared radiation, making star-burst galaxies very luminous infrared objects. The infrared satellite IRAS discovered thousands of star-burst galaxies. Starburst galaxies are rare among nearby galaxies, but they were common many billions of years ago, when episodes of galaxy collisions or close encounters were more frequent, causing more star bursts.

The rapid rate of supernova explosions in star-burst galaxies leads to the formation of hot and energetic super-bubbles of rapidly expanding winds (super-winds), that are

![Figure 13.2: Left: Simulation of super-wind evolution. From [201]. Right: Schematic of the Hα and X-ray emission arising in a star-burst wind and their spatial relationship. From [201].](image)
observed through their intense X-ray emission. Fig. [13.1] shows for instance the X-ray emission from the central region of the star-burst galaxy NGC 253 as seen by the Chandra satellite. One of the most striking results of these observations is the close spatial relationship between the X-ray emission (of thermal origin) and the He-emitting gas [202]. Super-winds are thought [e.g. 203] to be driven by the thermal pressure of hot ($\approx 10^8$ K) ejecta from supernova remnants and to be therefore comprised largely of carbon, nitrogen, oxygen, iron and other heavy elements. Shocked disk or halo gas that has been swept up into the wind, in the form of dense clouds or shells, could also account for the correlation between these two emissions. Adiabatic expansion cools the gas, converting the thermal energy into bulk motion at velocities up to $v \approx 3000$ km s$^{-1}$. Thermal conduction and/or turbulent mixing between hot gas regions and cool dense gas, swept up into the wind, can also lead to filamentary X-ray emission at the boundaries.

Simulations of the formation of such winds are shown in Fig. [13.2] [204]. The outflow resembles a super-bubble in the "snowplow" phase, sweeping a substantial amount of the ambient hot ISM. Five zones can be identified:

- The central, hot, injection zone ($T > 5 \times 10^7$ K).
- A supersonic free wind ($T \gtrsim 10^6$ K, $v \approx 1000$–2000 km/s).
- A region of hot, shocked, turbulent gas ($T \gtrsim 10^7$ K, $v \approx 500$–1200 km/s).
- A cooler, dense outer shell ($T \approx 7 \times 10^6$ K, $v \approx 300$–400 km/s).
- The undisturbed ambient gas ($T = 5 \times 10^6$ K).

Due to their characteristically high star-formation rates leading to strong winds, Starburst galaxies are the best laboratories for studying the development of these winds. When seen edge-on, they also allow detailed studies of the transport mechanisms of cosmic rays in galaxies. Together with their more extreme cousins the Ultraluminous Infrared Galaxies, they could provide new tests concerning the acceleration of cosmic rays.

### 13.2 NGC 253

#### 13.2.1 Introduction

NGC 253 is an archetypal star-burst galaxy, and is the closest to the Earth in the southern hemisphere, at a distance of 2.6 – 3.9 Mpc [204, 205, 206], the more recent value being 3.5 Mpc [207]. NGC 253 is seen edge-on and as such it represents the best local laboratory to study the effects of stellar feedback from strong star formation on the gaseous interstellar medium (ISM), processes which are intimately connected to galaxy formation and galaxy evolution. NGC 253 hosts a galactic wind indicated by various ISM phases in its halo.

NGC 253 was claimed as a TeV emitter by the Cangaroo collaboration [208] but this detection was later withdrawn [209]. Early observation from H.E.S.S. with 28 hours of data resulted in a constraining upper limit [210].

Compared to the Milky Way, NGC 253 exhibits an increased overall star formation rate (SFR), with the SFR in the star-burst nucleus comparable to that in the entire remaining disk of the galaxy. For NGC 253 as a whole the SN rate is estimated to be $\approx 0.08$ yr$^{-1}$. Estimates for the SN rate in the star-burst nucleus range from 0.03 yr$^{-1}$ [211] up to 70%
of the total supernova activity [212] of the galaxy. Therefore, the SFR in the star-burst nucleus alone could amount to \(5 \, M_\odot \, \text{yr}^{-1}\). The star-burst region itself (Fig. 13.3) has a cylindrical shape with a radius of \( \approx 150 \, \text{pc} \) and a height of \( \approx 60 \, \text{pc} \) perpendicular to the disk of the galaxy and is symmetric with respect to its mid-plane with a volume \(\text{Vol} \approx 1.2 \times 10^{62} (d/2.6\, \text{Mpc}) \, \text{cm}^3\) [e.g. 213].

![Figure 13.3: Cartoon of the central region of NGC 253. From [213].](image)

### 13.2.2 TeV observation

NGC 253 was observed with the full four-telescope H.E.S.S. array in 2005 and from 2007 – 2009 for a total of 241 hours. After usual data quality selection and dead-time correction, the total live time amounts to 177 hours of 3,- and 4-telescope observations which were used for the generation of sky images and the reconstruction of energy spectra. Observations were carried out at zenith angles of \(1^\circ\) to \(42^\circ\), with a mean value of \(12^\circ\). This makes the NGC 253 field one of the deepest fields observed with H.E.S.S. so far.

A VHE \(\gamma\) ray excess map for a \(1.5^\circ \times 1.5^\circ\) Field-of-View centred on the optical position of NGC 253 and produced with Model Analysis, standard cuts, is shown in Fig. 13.4. The map has been smoothed with a 2D Gaussian kernel of \(3'\) RMS to reduce the effect of statistical fluctuations. A total of \(329 \pm 49\) excess events corresponding to a significance level of \(7.1 \sigma\) are found at the nominal position of NGC 253. The best fit position of the source is RA \(00^h47^m34.3^s\pm1.4^s\), Dec \(-25^\circ17'22''6\pm0'3\) (J2000), compatible at the < \(1 \sigma\) level with the optical centre of NGC 253 at RA \(00^h47^m33^s.1\) and Dec \(-25^\circ17'18''\) (J2000).

The detected signal is consistent with point-like emission and is, in any case, much less extended than the size of the galaxy itself (shown as white contours in Fig. 13.4). No variability on any time-scaled has been detected, which is consistent with the standard picture of \(\gamma\) ray emission originating from the interaction of diffuse CRs accelerated in the star-burst region.
Figure 13.4: Map of NGC 253 emission (excess smoothed by the PSF, shown as an inset). From [24].

13.2.3 Multi-wavelength picture

Based on eleven months of Fermi-LAT data, NGC 253 has also been detected in the HE \( \gamma \) ray regime [214] together with the other archetypal star-burst galaxy M 82. It also appears on the Fermi Source Catalogs [215, 216]. After proper modelling of the diffuse Galactic and extra-galactic background components with the files `gll_iem_v02.fit` and `isotropic_iem_v02.txt` [217], the energy spectrum is best described by a power law in energy with \( \Gamma = 2.24 \pm 0.14 \text{stat} \pm 0.03 \text{sys} \) and integral flux between \((0.2 - 200)\) GeV of \( F(> 200 \text{ MeV}) = (4.91 \pm 1.03 \text{stat} \pm 0.30 \text{sys}) \times 10^{-9} \text{ cm}^{-2} \text{ s}^{-1} \).

The H.E.S.S. and LAT spectra are shown in Fig. 13.5 and are, within the 1 \( \sigma \) statistical errors, compatible, in spectral index as well as extrapolated flux level. A simultaneous, single power law fit to the H.E.S.S. and LAT spectral points results in a differential spectral index of \( \Gamma_{\text{sim}} = 2.34 \pm 0.03 \) and an energy flux at 1 GeV of \( 1.51 \pm 0.2 \times 10^{-12} \text{ erg cm}^{-2} \text{ s}^{-1} \). This corresponds to an integral flux above \( 200 \text{ MeV} \) of \( F_{\gamma}(> 200\text{MeV}) \approx 5.3 \times 10^{-12} \text{ erg cm}^{-2} \text{ s}^{-1} \) and to a \( \gamma \) ray luminosity \( L_{\gamma}^E(> 200\text{MeV}) \approx 7.8 \times 10^{39} \text{ ergs}^{-1} \), for a distance of 3.5 Mpc. This is about a factor of 10 larger than the \( \gamma \) ray luminosity \( L_{\gamma}^\text{MW}(> 200\text{MeV}) \simeq (7 - 10) \times 10^{38} \text{ erg s}^{-1} \) for the Milky Way galaxy [218]. The fit has a \( \chi^2 \) of 8.27 for 7 d.o.f.

\[\text{http://fermi.gsfc.nasa.gov/ssc/data/access/lat/BackgroundModels.html}\]
and a probability of 30%. Note that the somewhat steeper index for the simultaneous fit compared to the individual measurements could originate from a systematic difference in energy scale between the two experiments. This result implies, that in order to explain the available data, a spectral model which includes a break or turnover is not needed.

13.2.4 Interpretation

In a simplified scenario, $\gamma$ ray production results from the interaction of CR accelerated in the shocks from the SN explosions inside the star-burst region. In particular, production due to particle acceleration in the remaining disk of NGC 253 is neglected because of the low average gas density and radiation field intensity of the average interstellar medium. Indeed, the HE $\gamma$ ray luminosity of the Milky Way galaxy is an order of magnitude lower than the $\gamma$ ray luminosity of the star-burst region of NGC 253 [218].

The stellar winds from the early-type stars and the subsequent core collapse SN explosions heat the lower-density parts of the surrounding material, causing them to expand rapidly from the star-burst region in the form of a supersonic wind. The pressure induced by the shocks from the SN explosions adds to the excess thermal gas pressure. The dense and massive material in the star-burst region outside the SN remnants remains essentially neutral in this process and will not participate in the flow. The CRs from the percolating wind flow are nevertheless likely to penetrate also the dense gas in the star-burst region which therefore provides a massive target for $\gamma$ ray production, giving rise to significant HE emission, consistent with LAT observation.

Due to the very high gas density $n_g$ in the star-burst region ($n_g \approx 580$ cm$^{-3}$) [211], production of VHE $\gamma$ rays is expected to arise principally from interaction of accelerated CR with the interstellar medium, at least up to several TeV. Adopting this $n_g$ corresponds to an average loss time due to inelastic p-p collisions of $\langle \tau_{pp} \rangle \approx 1.1 \times 10^5$ yr. Contribution of discrete $\gamma$ ray sources is not expected to drastically change this overall picture.

Figure 13.5: Spectral Energy Distribution of NGC 253.
In a picture where a quasi-steady equilibrium exists between production and losses, the population of CRs accelerated in NGC 253's star-burst nucleus is removed from the star-burst region predominantly via three different processes:

- advective removal of particles in the star-burst super-wind, with speeds of the order of several hundred of km/s.
- diffusion of particles from the source region,
- quasi-catastrophic inelastic (“p-p”) interactions.

Energetic electrons/positrons suffer, in addition, radiative losses. The contributions of these components and the resulting γ ray spectra have been extensively discussed by several groups [e.g. 219, 220, 221, 222, 223].

The lack of structure in the HE-VHE spectrum strongly supports the idea that advective removal of particles dominates. Indeed, diffusion-dominant mechanisms would result in an escape time that would depend on the particle energy. Empirically, for the Milky Way galaxy, the measurements of the CR secondary-to-primary ratio [224] suggest a softening of the CR source spectrum by a factor $\propto E^{-0.6}$ above a few GeV per nucleon as a result of energy-dependent diffusion, which is not observed here. The observed spectrum ($\Gamma \sim 2.3$) would, under this assumption, reflect the primary spectrum of the accelerators.

Assuming a lifetime of the star-burst of about $(2 - 3) \times 10^7$ yr [211], and a convective loss time of about $10^5$ yr a steady state approximation similar to the Leaky Box can be safely used. This allows the fraction on non-thermal energy that is channelled into production of $\pi^0$ to be estimated from the supernova rate (and the total CR input) and the convective loss time. Depending on the chosen distance and the assumptions made, a fraction of the order of 20 – 40% is derived. This constitutes a measure of the extent to which the star-burst region is calorimetric with respect to its hadronic interactions, dissipating its own non-thermal output. Similar conclusions can be drawn for M 82, observed jointly by VERITAS and Fermi-LAT.

13.3 Conclusions

Star-burst galaxies are recent additions to the list of VHE γ ray sources. Smooth alignment of the HE and VHE spectra points towards an energy-independent particle removal mechanism from the star-burst region, most likely under the form of advection by the super-winds. They allow, due to the partially calorimetric nature of this measurement, the mechanisms of cosmic ray acceleration at the sources to be investigated. Detection of star-burst galaxies by IACTs was achieved using very long exposures, of the order of 200 h. The quality of the results illustrates the level of precision of the background estimation and the great improvements made in understanding of the systematic effects, thus allowing for further steps towards precision VHE γ ray astronomy.
Part III
Perspectives
Chapter 14

Towards Low Energy with H.E.S.S.-II

Towards low energy

Below the energy range covered by the existing ACTs, the $10 - 100$ GeV domain has been explored almost exclusively by Fermi with very limited statistics. As the energy decreases, the shower images contain fewer and fewer photons, and get diluted by the night sky background. Moreover, the standard reconstruction techniques require that well-defined images remain after cleaning of the raw data, and perform badly near the instrument threshold. Most of the recent VHE results were obtained with an analysis threshold well above the instrument capabilities (typically above 200 GeV).

Figure 14.1: Photo-montage of H.E.S.S. Phase II (courtesy W. Hofmann). The 28 m telescope in the center of the original H.E.S.S. array should increase the sensitivity of the system by a factor $\sim 2$, at a threshold energy of $\sim 70$ GeV in coincidence mode.

This energy domain is of fundamental interest not only for diffuse emission, but also for several classes of sources, among which are the pulsars. In particular, H.E.S.S. demonstrated [26] that most energetic pulsars with large spin-down luminosity are associated with VHE $\gamma$ ray sources. This implies that these pulsars emit on the order of 1% of their spin-down energy as TeV gamma-rays. These young pulsars also consist of a large fraction (at least 19 out of 46) of the Fermi pulsar population\footnote{The second population consisting of millisecond, re-accelerated, old pulsars}} in which the GeV emission is dominantly pulsed [141], thus stressing the GeV-TeV connection.
Moreover, pulsar emission models predict a cut-off in the pulsar high energy emission in the range of 10 to 50 GeV, depending on the emission location. Detailed observation in this energy range would bring strong constraints on the emission mechanism. The detection of pulsed signals from the Crab pulsar (in the northern sky) with MAGIC \[225\] \[226\] and more recently by VERITAS \[227\] above 100 GeV strengthen the need for deeper coverage of this intermediate energy domain.

Due to the absorption of VHE $\gamma$ rays through pair creation on the infra-red background, the observable VHE Universe also shrinks as the energy increases. Lowering the energy threshold of the instrument should therefore lead to a large increase of the number of detected Active Galactic Nuclei, thus leading to a better understanding of their properties. Micro-quasars, consisting of stellar-mass accreting black holes, are often depicted as scaled-down version of the extra-galactic AGNs. They could exhibit similar VHE behaviour, but on much faster time scales (smaller by the ratio of masses) and probably at lower energies. Detection of micro-quasars at TeV energies could improve our understanding of accretion-ejection phenomena.

Last but not least, H.E.S.S. excluded a massive ($E \sim 10$ TeV) neutralino or Kaluza-Klein particle as the agent responsible for the bulk of the VHE emission from the Galactic Centre. Searches for dark matter from dwarf galaxies and other potential dark matter wells are so far inconclusive but only explored dark matter mass ranges well above the instrument threshold. Lowering the energy threshold would increase the search range, and therefore probe dark matter particles with lower masses, but detailed understanding of the Galactic diffuse emission is a prerequisite.

Several projects, such as H.E.S.S.-II and MAGIC-II, aim at lowering the energy threshold of ACTs down to 20 GeV through the use of larger dishes (28 m diameter for H.E.S.S.-II), in order to address these physics goals. H.E.S.S.-II is currently under construction, with commissioning scheduled for mid 2012, and MAGIC-II is now in operation. The expected sensitivity of H.E.S.S.-II is shown in Fig. \[142\] compared to the sensitivity of the current experiment and to the flux of the Crab Nebula.

Although simulations show that current analysis techniques should perform adequately down to 80 GeV and perhaps 50 GeV, the lower energy range remains much more uncertain. Advanced reconstruction techniques are most probably required to improve both the reconstruction precision and the hadronic-shower rejection in this range. The Model Analysis \[22\], based on the log-likelihood comparison between the actual shower images and a pre-calculated shower model, uses, in contrast to previous techniques, all the information available in the camera. It takes into account the actual night sky background in each pixel and does not require any image cleaning. For all these reasons the Model Analysis performs significantly better at low energies, with an analysis threshold close to the trigger threshold, an energy resolution better than 10% in most of the energy range and an improved angular resolution. It also provides an accurate measurement of the altitude of the development of the shower in the atmosphere, information that can help to distinguish (on a statistical basis) the gamma-ray induced showers from those induced by electrons, which become the primary source of irreducible background at low energies. However, some systematic effects such as the Earth’s magnetic field, which bends the path of charged particles and introduces distortions in the low energy shower development, are not yet taken into account. A detailed treatment of these effects might be required to achieve good performance at low energies and to improve the understanding of the background.

H.E.S.S.-II will be able to record not only the integrated signal, but also its time
development in each pixel. As shown in particular by the MAGIC experiment [228], usage of timing information can minimize the influence of the background from the light of the night sky background and reduce, through an improvement of the reconstruction of the shower impact point, the background from unwanted hadronic shower by a factor of $\sim 2$. Implementation of a time-dependent Model Analysis is expected to significantly improve the sensitivity of the H.E.S.S.-II experiment at low energies and might constitute an essential element toward the detection of diffuse emission.

Last but not least, detailed comparison of existing reconstruction techniques shows that their discriminating parameters are only poorly correlated, thus demonstrating that none of them makes use of all available information. The ultimate analysis for ACTs still remains to be invented.

### 14.1 Timing information

#### 14.1.1 Time structure in showers

The time structure in showers results from the competition between two different delays, illustrated in Fig. 2.7.
• Due to the refractive index of the air, Cherenkov photons travel slightly slower than the charged particles comprising the shower. As a result, Cherenkov photons emitted high in the atmosphere tend to arrive comparatively late.

• For a given impact parameter, photons emitted high in the atmosphere are emitted at a lower angle and thus have a smaller geometrical path length than those emitted at low altitude.

To first order, the time difference can be written, as a function of altitude of emission \(z_e\) and angle \(\alpha = r/z_e\):

\[
\Delta t = \frac{\Delta z}{c} \times \left( \frac{n}{\cos \alpha} - 1 \right)
\]  
(14.1)

For a dry atmosphere, the refractive index reads

\[(n - 1) \approx 2.92 \times 10^{-4} \times \frac{P}{P_0} \times \frac{288.15 \, \text{K}}{T}\]

(14.2)

And in a simple isothermal atmosphere:

\[P = P_0 \exp(-z/z_0) \quad \text{with} \quad z_0 = \frac{RT}{gM} = 8.4 \, \text{km}\]

(14.3)

Thus the time difference is approximately:

\[
\Delta t = \frac{\Delta z}{c} \times \left( \frac{1}{\cos \alpha} - 1 + \frac{2.92 \times 10^{-4} \exp(-z_e/z_0)}{\cos \alpha} \right)
\]

(14.4)

the leading order approximation gives:

\[
\Delta t \approx \frac{r^2}{2cz_e} + 2.92 \times 10^{-4} \frac{z_0}{c} (1 - \exp(-z_e/z_0))
\]

(14.5)

Figure 14.3: Duration of showers as a function of impact distance.

The result of this competition is illustrated in Fig. 14.3 for small impact distances, the first delay dominates over the second, and the photons emitted at low altitude reach the ground first. The wave front typically lasts a few nanoseconds. For large impact distance, the opposite applies. At a specific distance, around 125 m the two delays almost compensate each other and the wave front is very short.\footnote{This distance depends on the altitude of the observatory}
14.1.2 Timing information in analysis

The timing information can be used in single telescope mode to break the degeneracy between distant, high-energy showers and nearby, lower-energy showers. This was pioneered by the HEGRA experiment which first measured a time gradient along the major axis in the Cherenkov images. The MAGIC telescope more recently demonstrated that, in single telescope mode, the use of timing lead to an improvement of the background rejection by a factor of $\sim 2$ due to the combination of two different effects:

- By integrating the signal only around the time of its maximum, the effect of the night sky background can be minimized.
- The time gradient across the field of view can be used as an additional discriminating parameter.

Since February 2007 the data acquisition of the MAGIC telescope uses FADCs digitizing the signal at the speed of 2 GSamples/s. In order to minimize the contamination from the NSB, the signal in each FADC track is fitted using a simple cubic spline; and its integral round the highest peak is used as a measure of the charge recorded by the pixel. The image cleaning procedure has been improved to use the timing information. Requiring a time coincidence between the pixel arrival time and the mean arrival time in the shower helps to disentangle the NSB signal from the actual shower tails.

![TIME GRADIENT vs IP](image)

Figure 14.4: Time gradient as a function of shower impact, from simulations. From [228].

The Time gradient measures how fast the arrival time changes along the major image axis and permits to distinguish showers falling far from the telescope from those falling in the neighbourhood. This is confirmed by simulations performed by the MAGIC collaboration (Fig. 14.4) showing the time gradient as a function of impact parameters; nearby showers have a negative gradient and distant showers a positive one. The time gradient is also correlated with the DIST parameter (Fig. 14.3), which is the angular distance from the image center of gravity to the source location on the camera, and which is sensitive to
the impact distance. Hadron images are distributed more isotropically and do not exhibit correlation between time gradient and DIST parameter (Fig. 14.5, right), thus allowing the gradient to be used as a discriminating parameter.

Similar studies performed by the VERITAS collaboration in stereoscopic mode [230] did not, however, lead to any improvement.

![Figure 14.5: Time gradient as measured by the MAGIC experiment as a function of DIST parameter, for simulated γ rays (left), ON-OFF events (middle) and background events (right). From [228].](image)

## 14.2 H.E.S.S.-II Telescope

The H.E.S.S.-II telescope is the biggest Atmospheric Cherenkov Telescope built to date, with 28 m diameter and 35 m focal length. Its design is based on the success of the H.E.S.S.-I telescopes, with however some significant improvement to the electronics.

### 14.2.1 Mount

Similarly to the current H.E.S.S. telescopes, the H.E.S.S.-II telescope (Fig. 14.6) is based on a steel structure on an alt-azimuthal mount, weighing approximately 560 tons. Azimuthal motion is performed with a system of six bogies, each comprising four wheels, rolling on a circular rail of 36 m diameter. The maximum movement speed is 200° mm⁻¹ for a total range of 540° allowing continuous tracking of a source, whatever its position on the sky. The elevation axis is made of two circular arcs equipped with rack-and-pinion rails centered on a horizontal axis installed 24 m above the ground and allowing orientation of the dish between −35° from horizontal (parking position) up to 180° (with the possibility of reverse pointing for sources culminating around zenith). The accuracy of the tracking system is designed to be within 1′ of the nominal position. A view of the telescope structure is shown in Fig. 14.6.

The very large focal length of the H.E.S.S.-II telescope (36 m) and the weight of the camera and supporting structure introduce, due to the large level-arm, strong mechanical constraint leading to elastic deformation of the structure. A CCD system installed at the center of the dish and imaging the camera, coupled with a second CCD looking at the star field, is used to correct offline for the deformations of the structure. As for H.E.S.S.-
I, absolute pointing precision of the order of 5 to 10" is expected using this correction procedure.

As the showers imaged by ACTs develop in the atmosphere at a distance of \( \sim 10 \) km, focus should not be tuned at infinity \([232]\). The average shower distance also increases with zenith angle (due to larger slant depth). An auto-focus system installed at the top of the main quadrupod automatically adjusts the camera position by as much as a few cm to correct for the variation of the average shower distance between zenith and elevations of about 20°.

### 14.2.2 Dish & Mirrors

The dish, of focal length 36 m and diameter 28 m, is built from 850 hexagonal facets of 90 cm diameter having the same focal length, arranged on a parabolic shape of 596 m\(^2\) area. A Davis-Cotton layout of this size would, according to equation \([3.2]\) have introduced anisochronisms of the order of 9 ns, thus resulting in a significant degradation of the signal over background ratio. The drawback is an increase of off-axis aberrations. Up to angles

![Figure 14.6: Structure of the HESS-II Telescope. From [231].](image-url)
of $\theta < 0.8^\circ$, the PSF is still smaller than the individual pixel size, but at larger angles, the smearing effect will have to be taken into account in shower image analysis. As for the H.E.S.S. telescopes, two actuators behind each facet are used to align the mirrors.

## 14.3 H.E.S.S.-II Camera

### 14.3.1 Structure

The camera, of 2 m diameter corresponding to 3.17$^\circ$ field of view diameter, is comprised of 2048 photo-multipliers of 29 mm (0.07$^\circ$ FOV) diameter each. The smaller field of view is imposed both by the aberration induced by the parabolic mount and the requirement of a reasonable camera size and number of pixels. It will somewhat limit the survey capabilities of H.E.S.S.-II. However, the core program of H.E.S.S.-II, due to its comparatively poorer angular resolution, is more centered on low energy point-like sources. Fig. 14.7 shows an exploded view of the camera structure.

![Structure of the HESS-II Camera](image)

**Figure 14.7:** Structure of the HESS-II Camera.

### 14.3.2 Front-End Electronics

The H.E.S.S.-II camera is based on similar concepts that proved to be successful in H.E.S.S.-I. In particular, the front-end electronics is contained in 128 drawers of 16 PMTs each. The main changes, compared to the H.E.S.S.-I camera, are an improved readout speed and reduced dead time to be able to handle about $\sim 3 - 4$ kHz of triggering rate.
The back plane of the camera contains, among others, power supplies and 3 cPCI crates dedicated respectively to the data acquisition and the instrument slow control, the trigger and the security system. In order to improve the readout speed, the acquisition crate contains 3 different CPUs, one dedicated to general slow control and the two others dedicated to camera readout and data transfer.

14.3.3 PMTs

The H.E.S.S.-II camera will be equipped with XP29600 PMTs from Photonis, similar but slightly shorter (by 1 cm) than the XP2960 used in H.E.S.S.-I in order to limit the oscillations of the signal. A fully automated test bench equipped with flashing LED and white light sources has been conceived and implemented to systematically test all 2500 PMTs. Performed tests included in particular:

- Calibration of the relation between the gain the and high voltage
- Measurement of the rate of after-pulses as a function of illumination
- Measurement of the peak-over-valley ratio
- For a sub-sample of the PMTs, measurement of the spatial homogeneity of the response of the photo-cathode.

The results of all tests has been archived into a database for historical bookkeeping.

14.3.4 Analogue Memories

Although H.E.S.S.-I proved that the strategy of digitization with analogue memories installed close to the photo-sensors was very successful, several problems were identified in the ARS analogue memories (unlocked ARS, . . .). A specific ASICs (SAM, for Swift Analogue Memory) has been designed by the SEDI-Saclay in cooperation with the LPNHE. The major differences with the ARS chips are:

- The use of two differential channels per SAM reduces the pick-up noise and the cross-talk between channels to less than 0.1%, thus allowing a single chip to handle both the high and low gains. 8 chips are therefore installed on each analogue card (one per PMT) versus 4 for the H.E.S.S.-I camera.
- The depth has been increased from 128 to 256 cells per channel.
- The sampling frequency can reach 3.2 GHz. However, the constraint of transit time to the central trigger and back limits the usable rate to \( \sim 2 \) GHz, and the normal operation will be done with a rate of 1 GHz.
- The input bandwidth reaches 300 MHz.
- The readout speed has been improved by a factor of 10.

The SAM analogue memories are based on a matrix of 16 independent lines of 16 cells each. Cells are written one after the other in each column, but the readout of a complete column can be done in parallel over the 16 lines within \( \sim 90 \) ns, thus improving the readout time. Read charges are converted by an external ADC at a rate of 11 MHz. This parallel
readout is largely responsible of the reduction of the dead-time from $\sim 64 \ \mu s$ to $\sim 1.5 \ \mu s$. The pedestal of each line can be adjusted by an external DAC, thus requiring a precise calibration. This calibration is part of the various tests that have been systematically applied to each chip, and can also be performed online within $\sim 2 \ \text{mn}$.

### 14.3.5 Systematic test of analogue memories

An automated test-bench has been designed in LPNHE to test all 6000 analogue memories. Tests were performed and exploited at LUMP laboratory [233]. The test bench was composed of several elements connected by a GPIB bus to a central controller. Controller elements included in particular:

- power supplies
- programmable attenuator to investigate the linearity of the response of the chips
- programmable pulse generator, to investigate the response with the shape of the signal (jitter, ...) as well as the response as a function of the position in the memory (through adjustable delays)
- splitter to investigate cross-talk between high and low gain channels

Thanks to the complete automatizing of the procedure, the complete test of a drawer from H.E.S.S.-II took only 20 minutes compared to several hours for H.E.S.S.-I. Nearly 90% of the SAMs were found to be suitable for use in H.E.S.S.-II, based on their properties of linearity, cross-talk, electronic noise and jitter. Results of the tests were systematically archived in a MySQL database to record the evolution of the parameters with time.

### 14.3.6 Timing information

In the H.E.S.S.-I camera, only the charge in the high and low gains, integrated in a window of 16 ns, are read out and transferred to the ground. In H.E.S.S.-II, the larger capabilities of the FPGA installed in the analogue cards allows to derive timing information from the accumulated samples. Two additional pieces of information will be transferred to the ground, thus allowing to exploit the timing information in the shower:

- **Time-Of-Maximum**: The sample with the largest charge within the sampling window of 16 ns will be identified, separately in high and low gains

- **Time-Above-Threshold**: The number of samples for which the charge is above a pre-defined threshold will be calculated inside the FPGA and transferred to the ground, independently in the high and low gain channels. This measurement, coupled with the integrated charge, allows to derive an estimation of the signal duration in each pixel.

One has, however, to take into account that the clock signal is generated independently in each drawer. Given that the relative precision of the clocks is $\sim 10^{-5}$, this results in a random jitter of $[0-1]$ ns (with flat distribution) for each drawer, thus limiting the timing resolution to about 300 ps. The fact that 16 PMTs share the same clock could possibly be used at the analysis stage, via a minimization procedure, to derive an estimation of the time offset of each drawer having a significant signal in several pixels.
One of the nice features of the design of the H.E.S.S.-II front-end electronics is that the algorithms used to derive the timing information are stored in FPGA chips and can therefore be easily upgraded if a new algorithm is proven to bring significant improvement. Examples of timing information obtained from simulated showers (with realistic simulation of electronics) are shown in Fig. 14.8 and confirm that the direction of the gradient can easily be obtained. For distant showers (top), information from the top of the shower (close to centre of camera) arrive first, whereas for nearby showers (bottom), information from the bottom of the shower arrive first.

![Image](image_url)

**Figure 14.8:** Simulated amplitude (left) and time of maximum (right) for a shower of 1 TeV falling at a distance of 175 m from the telescope (top) and for a shower of 200 GeV falling at a distance of ~30 m from the telescope (bottom)

### 14.3.7 Trigger

#### Level 1 Trigger

The philosophy of the L1 trigger is very similar to that of H.E.S.S.-I: the trigger is based on a minimum number of pixels in a sector with signal exceeding a given threshold. The L1 trigger decision is made within 80 ns, and the event readout starts immediately after a positive decision (in a similar way to H.E.S.S.-I).

#### Level 2 Trigger

With only the L1 trigger, the expected triggering rate of the H.E.S.S.-II camera is expected to be in the range 20 – 50 kHz, mainly due to isolated muons and night sky background.
With an event size of $\sim 8$ kB, this rate would result in an unsustainable data rate of $160 - 400$ MB/s. As in H.E.S.S.-I, a second level trigger is required to reduce the rate of events induced in particular by isolated muons to an acceptable rate of $3 - 4$ kHz. Two different strategies are implemented:

- In the *hybrid* mode, as in the case of H.E.S.S.-I, a coincidence between several telescopes is required, after correction for time delays induced by the pointing. This mode makes use of the existing *Central trigger* facility. This mode can lead to a reduction of the threshold to about 50 GeV and an improvement by a factor of $\sim 2$ of the effective area around 100 GeV.

- Below $\sim 50$ GeV, the H.E.S.S.-I telescopes will not trigger anymore and the decision has to be made from the information recorded with the H.E.S.S.-II telescope only. In this *single telescope* mode, a second level trigger based on a pattern analysis is used to reduce the trigger rate. Two binary images (corresponding to predefined *low* and *high* thresholds) are generated in the analogue cards and passed to a specific multiprocessor board installed on the trigger cPCI crate. Algorithms based on the image topology (and eventually simplified Hillas parametrization) are implemented in an FPGA and can be dynamically loaded. The L2 trigger decision varies between 20 $\mu$s and 100 $\mu$s depending on the complexity of the algorithms used. This rather long latency means that the events have to be buffered until the L2 decision is made. FIFO buffers with a depth of $\sim 10$ events are installed on each analogue card and are handled by an FPGA, thus reducing the dead-time (the camera is ready to restart acquisition as soon as the event is transferred into the FIFO). Events are transferred on the BoxBUS to the camera CPU only when the L2 trigger decision is made. At 3.5 kHz triggering rate, the overall dead-time of the telescope is less than 1% which is a considerable improvement over the H.E.S.S.-I electronics (dead time of $\sim 10\%$ for a single telescope rate of $\sim 1.4$ kHz and a system rate of $\sim 300$ Hz). The limited size of the FIFO imposes strong constraint on the average transfer rate to the camera CPU (that rate cannot exceed 10 kHz), but allows for short bursts of a limited number of events within much shorter time scales.

14.3.8 Overall synchronization

Since the information are distributed over several crates, a precise event stamping mechanism is required. A *Time-Stamp* card is used to perform the synchronization of the different crates, to read and distribute to the various crates the information from the central trigger (unique event number, bunch number, destination machine on the farm), and to trigger a local GPS card allowing to provide an unique camera time-stamp for each event. This time-stamping mechanism is used offline to build the events from the different pieces of information.

14.3.9 Camera DAQ

The overall scheme of the Data acquisition for H.E.S.S.-II is shown in Fig. [14.9] The main differences with respect to H.E.S.S.-I come from the fact that the acquisition is distributed over several CPUs installed on several crates:
• The slow control process (BIGD) is responsible for the overall control of the camera (power supplies, drawer configuration, high voltage, ...) as well as monitoring (temperature, high voltage, scalers, ...). During the acquisition, this process regularly sends monitoring information to the farm.

• The acquisition process (ZORA) is responsible for the event readout and transfer to the farm. In order to increase the overall speed, two instances (master & slave) are used on separate CPU boards. An interruption mechanism on the cPCI bus is used so that one instance sends its data to the ground (via its optical fiber) while the other is reading the camera via the box-bus.

• The trigger process (EMILIE) is responsible for configuring the overall trigger board (L1 and L2 trigger) and for sending the trigger information (patterns, ...) to the ground.

• In addition, a security process runs on a dedicated crate. This process is responsible for the overall health of the camera. It adjusts, in particular, the speed of the various fans in the camera to ensure optimal cooling, and it can decide to switch off the high voltage or even a power supply in order to prevent damage to the camera. Various sensors (cells for ambient luminosity, temperature sensors, ...) provide input for the security decisions.

The data receiver on the ground now receives information from many different sources:

• For the H.E.S.S.-I cameras, events are received in a unique block from each camera

• For the H.E.S.S.-II camera, events are received from two different instances of ZORA processes, imposing the usage of two independent buffers and a sorting mechanism

• Trigger information from the H.E.S.S.-II camera are received from independent data blocks from the Emilie process

Figure 14.9: Architecture of HESS-II DAQ system.
Central trigger information, concerning both the events from the H.E.S.S.-I and H.E.S.S.-II camera, are sent by the central trigger controller at the end of a bunch. This results in an increased complexity of the event building process, and strengthens the need for proper event stamping and redundancy.

### 14.3.10 First Events

First images of muons crossing the H.E.S.S.-II camera are shown in Fig. [14.10]. In each case, the upper panel shows the collected charge, the middle one the time of maximum and the lower one the number of triggered sectors.

The time gradient is clearly apparent in the middle panel, thus confirming the timing capabilities of H.E.S.S.-II. Extensive tests have been carried out to investigate the

Figure 14.10: First muons recorded by the HESS-II camera. **top:** Charge in High Gain. **Middle:** Time of maximum. **Bottom:** Number of triggered sectors.
behaviour of the camera.

14.4 Analysis for H.E.S.S.-II

14.4.1 Hillas based analysis

Fig. [14.11] and [14.12] show typical images that should be recorded by the H.E.S.S.-II telescope compared to the capabilities of H.E.S.S.-I, and illustrate the superior image definition.

Several studies of the performance of H.E.S.S.-II with elaborated analysis techniques based on Hillas parametrization have been performed [e.g. 234], and resulted in an energy resolution varying between 40% at 30 GeV to less than 15% above 1 TeV and an angular resolutions of the order of 0.25° at 30 GeV (R68) to above 0.15° at 300 GeV for an energy threshold around 30 GeV. Introduction of new discriminating variables also improved the discrimination by about 40% compared to standard techniques.

14.4.2 Model Analysis for H.E.S.S.-II

A further development of the Model Analysis is currently underway to adapt it to the specific constraints of H.E.S.S.-II. This includes in particular:

- The parametrization of the effect of the parabolic dish, leading to significant distortion of shower images away from the optical axis. This effect, which could be ignored for the H.E.S.S.-I telescopes, introduces some additional complexity as the effect of image broadening has to be calculated during the minimization procedure, taking into account the actual localization of the shower image in the camera.

- The single telescope fit for monoscopic events. In that case, the starting point is difficult to estimate and a quasi-degeneracy occurs between shower direction, impact and depth of primary interaction. It is not fully clear yet whether the amount of available information in the camera (and in particular the longitudinal development) is sufficient or whether some additional constraint have to be put in.

14.4.3 Time-dependant model analysis

In monoscopic mode, the timing information proved to be useful to break the degeneracy between nearby, low-energy showers and distant, higher-energy showers. In H.E.S.S.-II, the timing information could be relatively easy to implement in a specific flavour of the Model Analysis. Some preliminary work toward the incorporation of timing information has been done through the development of a time-resolved model. This is illustrated in Fig. [14.13] which shows the evolution of predicted image shape for times of 7.5, 10, 12.5, 15, 17.5 and 20 ns after impact of the primary particle on the ground.

In the near future, this information will be used to provide templates for the time gradients, that will be then used in the minimization procedure. The overall likelihood will be composed as a product of the likelihood on the charge by a likelihood on the time of maximum, on a pixel per pixel basis. This should hopefully greatly improve the Model analysis in monoscopic mode.
Figure 14.11: Simulated images for a $\gamma$ ray of energy 300 GeV at an impact distance of 120 m.
Figure 14.12: Simulated images for a $\gamma$ ray of energy 90 GeV at an impact distance of 90 m. Only two of the four H.E.S.S.-I telescope triggered.
Conclusion

H.E.S.S.-II, because of its size and complexity, represents a major challenge compared to previous instruments. The camera is in the commissioning phase and seems to full-fill the requirements. The telescope is currently being built and installation of the camera is foreseen within the next month for a first light around summer 2012. Promising analysis techniques already exists, and further ideas, using in particular the timing information,
should further improve the expected performances. H.E.S.S.-II will undoubtedly represent a major step towards the next major, worldwide, facility.
Chapter 15

CTA - the next generation facility

Introduction

The present generation of Atmospheric Cherenkov Telescopes, composed principally of H.E.S.S., MAGIC and VERITAS, opened a new window on the non-thermal universe and established the IACT technique as a robust, efficient and reliable tool for Very High Energy γ-ray astronomy. Despite the enormous amount of new discoveries, some questions are still unanswered and further steps are required.

We are in particular still lacking detection of Gamma-ray bursts as well as signatures of annihilation of dark matter particles. Galaxy clusters, which are supposed to act as storehouses of cosmic rays on cosmological time-scales, still resist any detection.

After having unveiled the VHE sky, precision measurements using large samples of sources are also required to fully understand the acceleration and escape mechanisms in SNR or binary systems. Fundamental studies such as Lorentz violation or distance Universe tomography also require much larger samples.

The IACT technique is well enough understood, and sufficiently scalable to propose now a fourth generation instrument which would be able to address these questions. This instrument would need to have a sensitivity improved by a factor of ~ 10 compared to the current instruments, resulting into a multiplication by about 1000 of the volume of the accessible Universe. An angular resolution improved by a factor ~ 5 is also required to address the question of confusion limit. Moreover, extension towards lower energies (for the physics of pulsars and micro-quasars in particular) and larger energies (to address, for instance, the question of acceleration limits of the SNRs) are needed to optimize the scientific return of the project.

15.1 CTA collaboration and road-map

The CTA collaboration was formed under the initiative of the H.E.S.S. and MAGIC collaboration which joined to propose a large network of Atmospheric Cherenkov Telescopes. With the joining of the US groups of the Advanced Gamma-ray Imaging System (AGIS) project, and of the Brazilian and Indian groups in Spring 2010, and with the strong Japanese participation, the CTA collaboration already comprises over 700 participants from 25 countries and more than 50 laboratories and will now represent a genuinely world-wide effort, extending well beyond its European roots. This project is well ranked on major road-maps such as those from ASPERA, ASTRONET and ESFRI.

The main scientific objectives of the CTA project are [235]:

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• decrease the energy threshold to few 10s of GeV,
• acquire sensitivity beyond 50 TeV,
• increase the sensitivity by a factor of 10 in the core energy range from 100 GeV to 50 TeV,
• improve the energy and angular resolution by a factor of \( \sim 5 \),
• build a large, opened observatory to serve a wide community of users,
• enhance the sky survey capability, monitoring capability and flexibility of operation.

Given our current knowledge of the technologies, the final sensitivity of CTA could look like the curve displayed in Fig. [15.1]

The result of a simulation of a galactic plane survey [52] is shown in Fig. [15.2] illustrating the richness of the field and the crucial importance of angular resolution to beat the confusion limit.

![Achievable sensitivity for CTA](image.png)

**Figure 15.1:** Achievable sensitivity for CTA. From [235].
Figure 15.2: Simulated image of a 240 hr long CTA Galactic Plane survey. From [52]

Improvement in sensitivity will not only result into a larger sample of sources, but also into a better precision for already known sources. A simulated phasogram of the binary system LS 5039, shown in Fig. [15.3], illustrates the fact that the uncertainties on individual measurement will be reduced by a factor of more than 2, thus resulting into much more precise constrains on the model.

Due to the very large size of the project, several phases are envisaged over the CTA’s lifetime, including in particular a design study phase, prototyping and preparatory phase, construction phase, and operation phase. In October 2010, the CTA consortium entered a preparatory phase, with a budget of ~ 10 million Euro, funded within the PF7 European program. The goals of the preparatory phase are:

- provide a governance scheme,
- optimize the performance and availability of the instrument,
- optimize the telescope structure, camera design, optics and mirrors,
- reduce the costs via deep involvement of industry,
- select the potential sites,
Figure 15.3: Expected light-curve for LS 5039, illustrating improvement in precision. From [230].

- prepare the infrastructure (software, procedures, data format, ... ) for observatory operations.

The construction phase is expected to start in 2015, and to last about 10 years. However, the array of telescopes will be able to start operation immediately after installation of the first telescopes, with sensitivity progressively increasing with deployment of the array.

## 15.2 Overall design concepts

In order to view the whole sky, at minimum two sites of operation are required. In each site, the goals in terms of sensitivity, resolution and energy coverage can only be achieved by combining many telescopes, of different sizes, and distributed over a large area (of at least 1 km$^2$).

The main site is foreseen in the southern hemisphere, focusing on the physics of galactic sources. A second complementary northern site will be primarily devoted to the study of Active Galactic Nuclei and cosmological galaxy and star formation and evolution.

The array (Fig. 15.4) will most probably comprise about a 100 telescopes of 2–3 different sizes: several small size telescopes (SST) of 6 m diameter distributed over a very large area to access the highest energies, many medium size telescopes (MST) of 12 m in a more packed array to improve the sensitivity in the core energy range, and few large size telescopes (LST) of 23 m diameter to push down the threshold to the $\sim 10$ GeV domain.

A possible layout, resulting from Monte-Carlo optimization, is shown in Fig. 15.5.
Figure 15.4: Artist view of the CTA concept. From [237].

Figure 15.5: A quadrant of possible array schemes promising excellent sensitivity over an extended energy range, as suggested by Monte Carlo studies. In the upper right part, clusters of telescopes of the 12 m class are shown at the perimeter, while in the lower left part an option with wide-angle telescopes of the 3-4 m class is shown. From [237].
The final number of the telescopes, their size, their configuration and the overall performance will be one of the outcome of the current preparatory phase. Several different designs are currently taken into account and the construction of the first prototypes is currently starting.

Conclusion

CTA will undoubtedly be one of the major astronomical instruments of the coming decades. The experiences gained in H.E.S.S. and H.E.S.S.-II will naturally be exploited in CTA. The model analysis, for instance, is currently being investigated by several groups within CTA and is expected to significantly improve the sensitivity of the instrument. Moving towards precision VHE astronomy will however not be an easy task; detailed monitoring of the atmosphere (and its fluctuations) will be required at each stage, from the data taking to the analysis. The experience gains in data acquisition technique and instrument control, calibration, and data analysis will definitely help moving toward precision measurement in this relatively new field.
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RÉSUMÉ

Les dix dernières années ont été marquées par l’arrivée à maturité de la technique d’imagerie Cherenkov atmosphérique, qui a permis, notamment grâce au réseau de télescopes HESS, l’ouverture d’une nouvelle fenêtre sur l’Univers non thermique. Ce mémoire d’habilitation à diriger des recherches retrace dix années de recherche en Astronomie Gamma de très haute énergie avec H.E.S.S.puis CTA. Les aspects techniques tels que la conception de l’instrument, son calibrage, la reconstruction des événements et l’analyse de données sont présentées dans une première partie, tandis que la seconde brosse un panorama des grandes découvertes dans ce domaine.

ABSTRACT

The last ten years saw the emergence of Imaging Atmospheric Cherenkov Telescopes as a mature and efficient technique for the study of the Very High Energy Universe, leading to the successful opening, mainly by the HESS experiment, of our understanding of the non-thermal Universe. This habilitation thesis summaries ten years of research in Very High Energy gamma-ray astronomy with HESS and CTA. In the first part, instrumental aspects such as the experiment design, its calibration, the reconstruction of the events and the data analysis are presented. The second part draws a panorama of the main discoveries in the domain.