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CHAPTER

1

COSMIC RAYS AND THE POSITRON EXCESS

Cosmic rays are charged particles traveling across the Universe. As observed in a series of experiment carried out from the beginning of the last century, they are made up of protons and fully ionized atoms for the 98% and of electrons and positrons for the remaining 2%. The 98% of protons and nuclei is made up of protons for the 87%, alpha particles for the 12% and nuclei with higher atomic number for the remaining 1%.

Their study begun in the first years of the XX century, with the purpose of finding a reason for the spontaneous discharge of electroscopes, even if they were kept distant from sources of natural radioactivity.

There is a first clue in the understanding of the nature of cosmic rays in 1910, when the Italian physicist Domenico Pacini performed a series of measurements to study the effect of ionization in water. Pacini finds that the intensity of radiation under sea level decreases with depth. In the interpretation of this phenomenon he proposed that air's ionization was caused by something in the atmosphere and not by radioactive substances on ground, as was suggested by others [1].

The big breakthrough came with the firsts balloon experiments of Vicktor Hess (1912) and Werner Kolhörster (1914), with the aim to measure the variation of atmosphere ionization with increasing altitude. The result of

their experiments was the finding of an increase of ionization starting from an altitude of 1.5 km above sea level. This allowed to exclude the hypothesis of a terrestrial nature of the observed ionizing radiation and to strengthen the idea of a cosmic origin [2].

However it was not clear if the observed radiation was made up of charged particles or, as hypothesized by Millikan, photons that produce electrons in their interaction with the atmosphere by Compton effect.

A second important discovery was made in 1927, when Jacob Clay demonstrates the dependence of cosmic rays intensity from magnetic latitude. Cosmic rays that penetrates in the atmosphere interact with Earth magnetic field: at the poles they arrive in parallel with magnetic field's lines, so the Lorentz force is zero; on the contrary, moving close to the equator we have a contribution to the Lorentz force due to the normal component of the velocity vector to the magnetic field and so a deflection causing a lower flux of cosmic rays. Clay's observation allowed to confirm the nature of charged particle of cosmic rays [3].

A crucial step for the understanding of the nature of cosmic rays arrives at the beginning of the '30s. Bruno Rossi suggested the presence of an east-west effect: if cosmic rays are made up of charged particles we should observe an asymmetry in their direction of arrival in the case that they were made up mainly of charges of the same sign. Because of a series of delay, dealing with Rossi's academic responsibilities, this effect was detected by T.H. Johnson and by L. Alvarez and A.H. Compton in two separate experiments carried out in Mexico. The results showed that the intensity of cosmic rays is higher towards west rather than east, proving that they were made up mainly of positive charges.

A few years later, in 1938, using a coincidence set-up of Geiger-Müller counters Pierre Auger and Ronald Maze observed that counters, even if placed at 300 m from one another, detect within a coincidence time of 1 μ s the arrival of cosmic rays. This showed the existence of the so called Extended Air Shower (EAS), that is cascades of secondary particles produced in the interaction of cosmic rays with atmosphere [1].

1.1 Chemical composition

To compare the chemical abundances of the elements in our Solar System with the ones present in the cosmic rays, we use data coming from meteoritic samples because they have the advantage of not being influenced by geophysical phenomena, as the terrestrial ones, and are believed to have the same chemical abundances present at the formation of the primitive solar nebula.

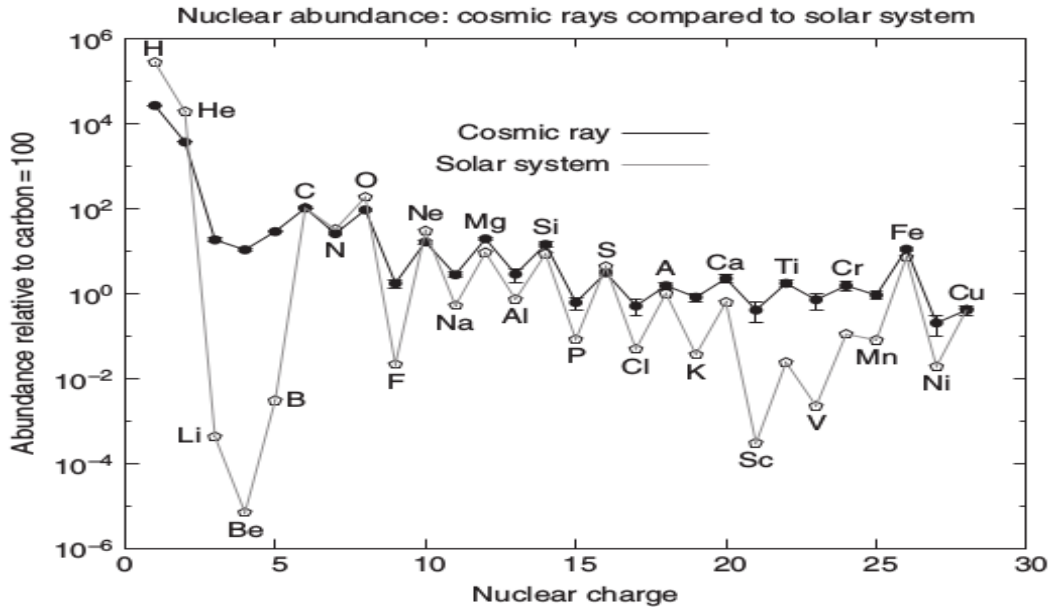


Figure 1.1: Chemical abundances, up to copper, present in the cosmic rays (black line) and in the Solar System (gray line), all relative to carbon $\times 100$ [4]

Comparing the chemical abundances of the elements present in the cosmic rays and in our Solar System, as shown in figure 1.1, we can notice some features:

- the odd-even effect, that is nuclei that have even atomic number are more stable and so are more abundant than the odd ones, is present in both cases
- light elements, such as lithium, beryllium and boron, are overabundant in the cosmic rays relative to the Solar System's abundances
- there is an overabundance of chemical elements also in the case of elements with atomic number just less than that of iron, in particular between calcium and iron

The overabundance of some elements present in cosmic rays relative to the Solar System abundances is due to the phenomenon called “spallation”: once accelerated from their sources, cosmic rays can interact with the interstellar medium causing their fragmentation in two or more nuclei with lower mass number [2].

1.2 Spectrum

The particles flux arriving in the energy interval $E, E + dE$ coming from the solid angle $d\Omega$ is defined as [4]

$$\Phi = \frac{dN}{dEdAdtd\Omega} \quad (1.1)$$

Frequently the flux of cosmic rays is called “differential energy spectra”, or simply “spectrum”, that is the number of particles arriving per unit area, time, solid angle and energy. An example is shown in figure 1.2 where can be seen the energy spectra of cosmic rays obtained from various experiments. The energy spectra of cosmic rays can be described in a first approximation by a power law

$$N(E) = KE^{-\gamma} \quad (1.2)$$

where γ is called “spectral index” and K is the normalization. By the way there are other models that can be used to describe the spectrum of cosmic rays:

a) A broken power law

$$N(E) = K\left(\frac{E}{E_0}\right)^{-\gamma_i} \quad (1.3)$$

where $i = 1, 2$. In this case the phenomenon is described by a power law with spectral index γ_1 up to the energy E_0 and with spectral index γ_2 for $E > E_0$.

b) A power law with an exponential cutoff at high energy

$$N(E) = KE^{-\gamma} e^{-\frac{E}{E_c}} \quad (1.4)$$

or at low energy

$$N(E) = KE^{-\gamma} e^{-\frac{E_c}{E}} \quad (1.5)$$

where E_c is the cutoff energy.

Often it can be used the so called S.E.D. (spectral energy distribution) representation multiplying the spectrum by E^x , where $x = 2$ for the proton spectrum or $x = 3$ for the electron spectrum. This representation is useful to get a better distinction between energy spectra with similar spectral index and permit to highlight the features of the spectrum.

Coming back to the cosmic rays spectrum in figure 1.2, we can underline some features:

- A cutoff can be seen at low energy, that is about 10^9 eV, phenomenon known as “solar modulation”. Following a periodicity of 11 years, the

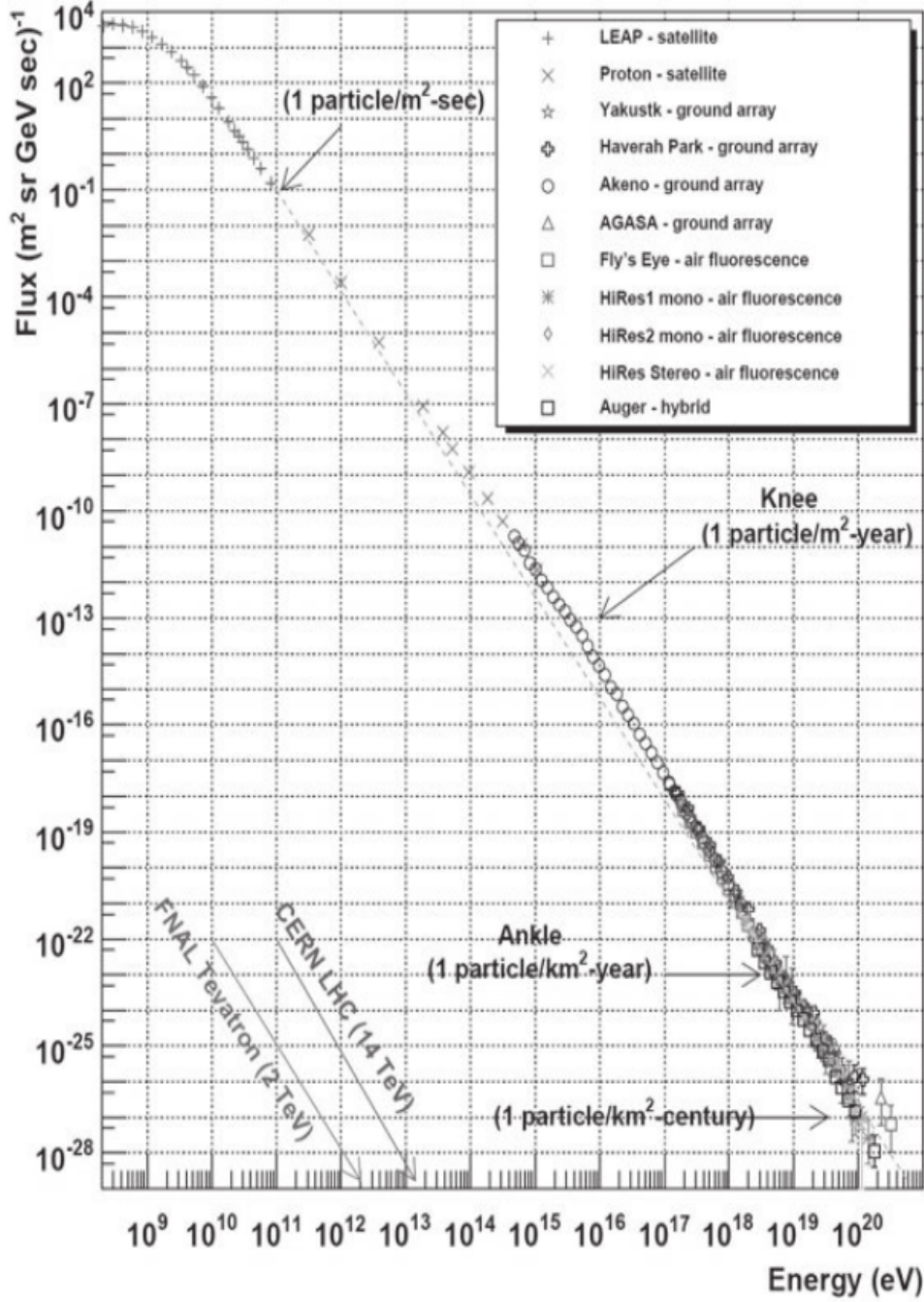


Figure 1.2: Differential energy spectra of cosmic rays from various experiments. The energy interval is taken from 10^9 up to 10^{20} eV. It should be noted two important features of the cosmic rays spectra: the Knee at 10^{15} eV and the Ankle at 10^{19} eV [2].

number of charged particles emitted by the Sun, called Solar wind, varies with the effect of shielding the Earth from the arriving cosmic rays with energy below $10^9 \text{ eV nucleon}^{-1}$.

- Up to 10^{15} eV the spectrum is described by a power law with spectral index included in the interval $2.5 < \gamma < 2.7$.
- Between 10^{15} and 10^{19} eV the spectrum is described by a power law with spectral index $\gamma = 3.1$.
- Beyond 10^{19} eV the spectral index came back to a value included in the interval $2.5 < \gamma < 2.7$ and a cutoff is observed.

A possible interpretation to explain the change of index in the cosmic rays energy spectra is related to the gyroradius of charged particles traveling through the Galaxy: particles with energy higher than 10^{15} eV have a gyroradius, for an average magnetic field in the Galaxy of order $3 \mu\text{G}$, such that decreases the containment efficiency of cosmic rays inside the Galaxy. Beyond 10^{18} eV the gyroradius became greater than the thickness of the Galactic disk; this is the reason why it is believed that cosmic rays of such energy are of extragalactic origin.

The GZK (Greisen-Zatsepin-Kuzmin) effect, permits to explain the observed cutoff at high energies as due to energy losses of cosmic rays with energy beyond $6 \times 10^{19} \text{ eV}$, because of their interaction with photons of the Cosmic Microwave Background (CMB) producing charged and neutral pions [3] [2].

1.3 Electrons and positrons

The electrons and positrons constitutes about 1-2% of the cosmic-rays. Quite the opposite of protons and ions, these particles are subjected to a series of energy loss processes which cause a distortion of the injection spectra as they propagate through the Galaxy. The use of cosmic rays experiments involving a magnetic spectrometer allow the distinction of the sign of the electric charges. Doing so is possible to distinguish the positrons from the electrons in cosmic rays. One of the most recent experiment using a magnetic spectrometer is AMS-02, launched on 2011 it operates onboard of the I.S.S. (International Space Station). The main background component evaluating the positron flux is given by cosmic-ray protons. The AMS-02 experiments has an overall proton-to-positron rejection power of order 10^6 for most of the energy range under study, which goes from 0.5 GeV up to 350 GeV .

The data from AMS-02 experiment [5] can be seen in figure 1.3, compared

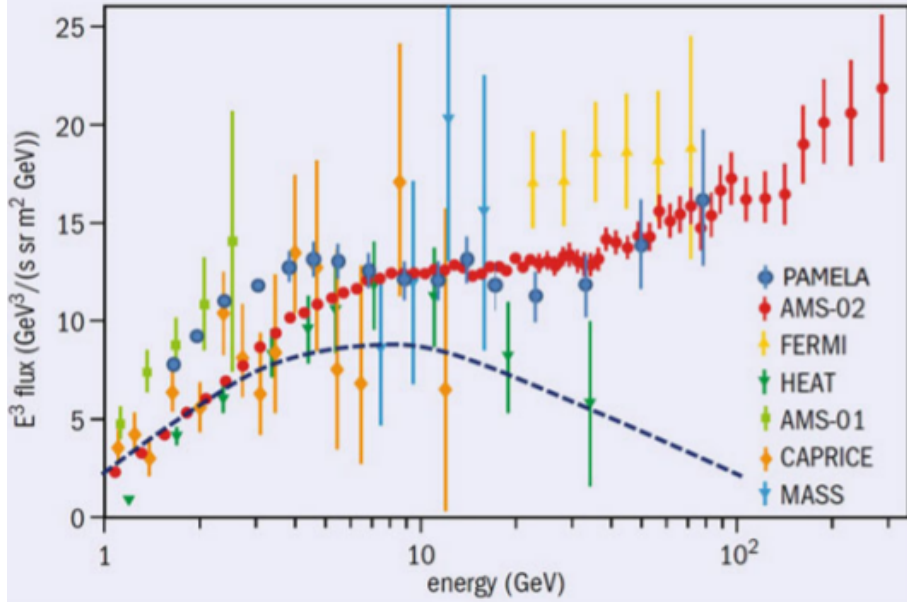


Figure 1.3: Positrons energy spectrum multiplied by E^3 by different experiments (from balloon and space based). The dashed lines represents the positron spectrum of secondary origin [5].

with data on positron energy spectrum from various balloon based and space based experiments. As can be seen, respect to the positron spectra obtained from a secondary production mechanism hypothesis, an excess starts above a few GeV. The increase is still more evident in figure 1.4, where the ratio of the positron flux to the total flux of electrons and positrons (positron fraction) rises above 10 GeV respect to a model based on secondary production of positron.

The physics of secondary production is known: cosmic-rays interact with interstellar matter producing secondary particles, with electrons and positrons produced via the initial decay of charged pions. Uncertainties on the secondary production model come from primary fluxes, cross sections and propagation parameters.

Quite the opposite, the physics at the origin of the positron excess is still uncertain. Some proposed that the positron excess can be an indirect signature of dark matter. If dark matter is composed by weakly interacting massive particles (WIMPs), their annihilation or decay can produce GeV-TeV cosmic-rays equally abundant in particles and antiparticles [6]. However some authors suggested that is possible to explain the positron excess as an effect of nearby pulsars and their wind of relativistic electrons and positrons. Anyway, since these particles are subjected to a series of energy

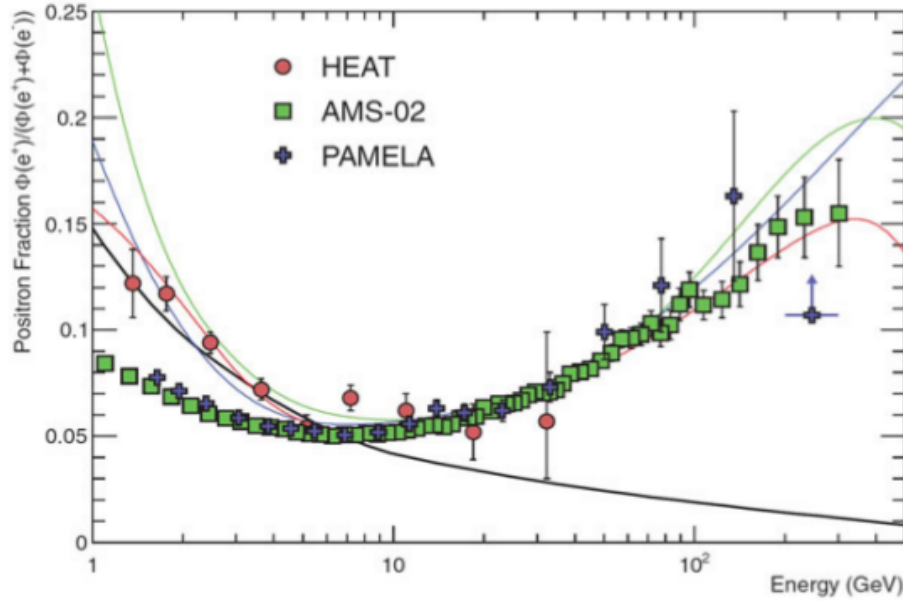


Figure 1.4: The positron fraction (ratio of the flux of e^+ to the total flux of $e^+ + e^-$) as a function of the energy measured HEAT (balloon based), PAMELA and AMS-02 (space based). The heavy black line is a model of pure secondary production. The three thin lines show three representative attempts to model the positron excess with different phenomena: dark matter decay (green); propagation physics (blue); production in pulsars (red). [5]

losses interacting with magnetic and photon fields, the maximum distance within sources like pulsars can contribute to the positron flux is found to be of a few hundreds parsec [7].

1.4 Gamma rays production mechanisms

Traveling through the Galaxy electrons and positrons interact with interstellar matter, the Galactic magnetic field and photon fields such as the CMB (Cosmic Microwave Background). All these processes generate high energy photons.

Bremsstrahlung

Bremsstrahlung, from the German word for “braking radiation”, is a process that consists in the emission of electromagnetic radiation from charged particles passing through matter, due to their interaction with atomic electromagnetic fields. Since the Bremsstrahlung cross section is inversely pro-

portional to the square of the particle mass, electrons and positrons are the only particles in which Bremsstrahlung losses contribute substantially to the energy loss of the particle [8].

The energy loss rate of Bremsstrahlung from the full relativistic quantum treatment is [2]

$$-(\frac{dE}{dt})_{Brems} = \frac{Z(Z+1.3)e^6 N}{16\pi^3 \epsilon_0^3 m_e^2 c^4 \hbar} E (\ln(\frac{183}{Z^{1/3}}) + \frac{1}{8}) \quad (1.6)$$

with Z atomic number of the material and N space density of nuclei. The term $(Z + 1.3)$ takes account of electron-electron Bremsstrahlung due to the interaction between electrons and the atomic electrons.

Synchrotron radiation

Synchrotron radiation are photons emitted by charged particles moving in magnetic fields, if their trajectories are not parallel to the magnetic field's lines. It is responsible of most of the radiation from radio to X-ray band from galactic and extragalactic sources.

In relativistic regime ($\beta \approx 1$) an electron moving in a magnetic field B loses energy to synchrotron radiation at a rate

$$-\frac{dE}{dt} = 2\sigma_T c \gamma^2 u_B \sin^2 \theta \quad (1.7)$$

with σ_T Thomson cross section, θ angle of electron trajectory relative to the magnetic field direction and $u_B = B^2/8\pi$ energy density of the magnetic field. Since the energy loss rate is proportional to the square of the Lorentz factor of the charged particle, it is inversely proportional to the square of the particle mass. Because of the strong dependence from the mass synchrotron radiation is not important for protons being $m_e/m_p \approx 10^{-3}$ [9].

If the angle distribution is isotropic, like it is believed to be due to irregularities in the magnetic field distribution, an expression for the average energy loss rate is obtained [2]

$$-(\frac{dE}{dt})_{Sync} = \frac{4}{3} \sigma_T c u_B \gamma^2 \quad (1.8)$$

In figure 1.4 is shown an example of synchrotron spectra obtained from an electron distribution modeled with an exponential cut-off power law with index $\alpha = 2.1$ and cut-off energy at 13 TeV, at a distance of 1.5 kpc and with a magnetic field of $100 \mu G$.

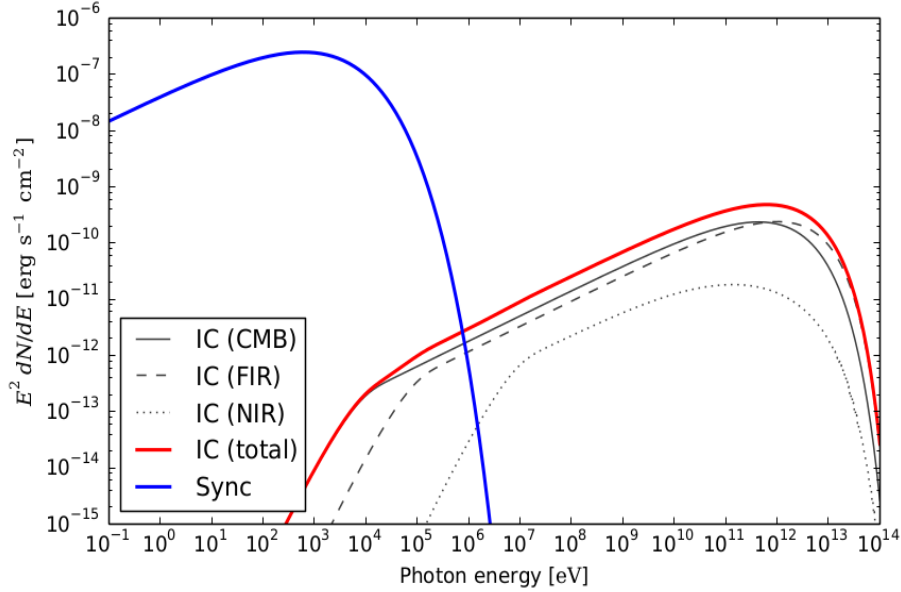


Figure 1.5: Example of Synchrotron and inverse Compton spectra. Both spectra are obtained from the same electron distribution modeled as an exponential cut-off power law with index $\alpha = 2.1$ and cut-off energy at 13 TeV at a distance of 1.5 kpc. The Synchrotron spectra is obtained with a magnetic field of 100 μG , while the IC target photon fields are: Cosmic Microwave Background, Far IR and Near IR [10]

Inverse Compton scattering

In the Compton scattering an high energy photon, colliding with a stationary electron, transfers some of it energy to the electron; thus the scattered photon results to have a lower frequency. If the photon energy is much less than the electron rest mass energy $h\nu \ll m_e c^2$ in the center of momentum frame of reference, the scattering may be treated as Thomson scattering. Anyway if the photon energy is equal or higher than the electron rest mass energy $h\nu \geq m_e c^2$ the quantum relativistic cross-section has to be used (figure 1.5)

$$\sigma_{K-N} = \pi r_e^2 \frac{1}{\gamma} \left[1 - \frac{2(\gamma + 1)}{\gamma^2} \right] \ln(2\gamma + 1) + \frac{1}{2} + \frac{4}{\gamma} - \frac{1}{2(2\gamma + 1)^2} \quad (1.9)$$

called Klein-Nishina cross-section, with $\gamma = h\nu/m_e c^2$ and $r_e = e^2/4\pi\epsilon_0 m_e c^2$ is the classical electron radius. It can be seen that in the low energy limit $\gamma \ll 1$

$$\sigma_{K-N} = \frac{8\pi}{3} r_e^2 (1 - 2\gamma) = \sigma_T (1 - 2\gamma) \simeq \sigma_T \quad (1.10)$$

the Klein-Nishina cross section reduces to the Thomson cross section, instead in the ultra-relativistic limit $\gamma \gg 1$

$$\sigma_{K-N} = \pi r_e^2 \frac{1}{\gamma} \left(\ln(2\gamma) + \frac{1}{2} \right) \quad (1.11)$$

the Klein-Nishina cross section decreases as E^{-1} .

As suggested by its name inverse Compton scattering is a process in which

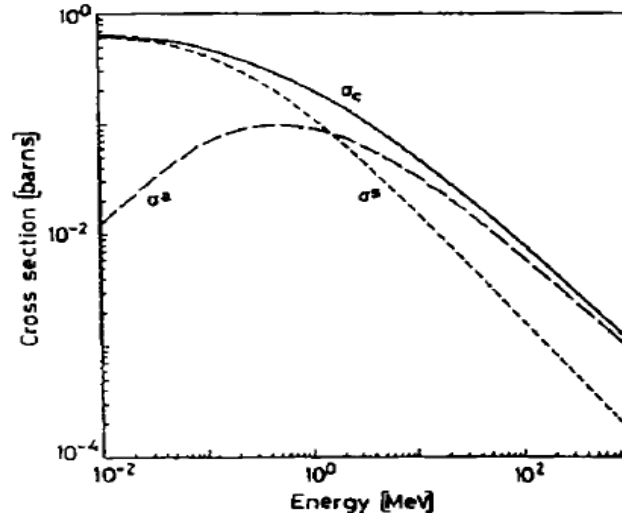


Figure 1.6: Total Compton scattering cross section. The dashed lines are the Compton scattering cross section, σ^s , defined as the average fraction of the total energy contained in the scattered photon, and the absorption cross section, σ^a , defined as the average energy transferred to the recoil electron [8].

a low energy photon gain energy once scattered by relativistic electrons. If photons have energy lower than the electron rest mass energy $h\nu \ll m_e c^2$ and the electrons have $\gamma \gg 1$, the centre of momentum frame is very closely that of the relativistic electrons and in this frame the photons have energy $\gamma h\nu$. Thus if $h\nu < m_e c^2 / \gamma$ the Thomson scattering cross section can be used to describe the probability of scattering. In the frame of reference in which the electron is at rest, the energy loss rate of the electron is the rate at which energy is re-radiated by the electron

$$-\left(\frac{dE}{dt}\right)' = \sigma_T c u'_{rad} \quad (1.12)$$

with u'_{rad} energy density of radiation in the rest frame of the electron. Since dE/dt is a Lorentz invariant between inertial frames of reference and since

the energy density of radiation in the electron frame is related to the energy density of radiation in the laboratory frame of reference u_{rad} by the relation $u'_{rad} = \frac{4}{3}u_{rad}(\gamma^2 - \frac{1}{4})$, we have

$$-\frac{dE}{dt} = \frac{4}{3}\sigma_T c u_{rad}(\gamma^2 - \frac{1}{4}) \quad (1.13)$$

The energy gain of the photon field is the energy gained by the photon field due to scattering of low energy photons by relativistic electrons minus the initial energy of the low energy photons. Since the loss rate of energy of the low energy photon field is $\sigma_T c u_{rad}$ we have that the inverse Compton loss rate is

$$-(\frac{dE}{dt})_{IC} = \frac{4}{3}\sigma_T c u_{rad}(\gamma^2 - 1) \quad (1.14)$$

This expression is incredibly similar to the energy loss rate due to synchrotron radiation. The reason is that the energy loss rate depends upon the electric field causing the acceleration of electrons and it doesn't matter which is the origin of the electric field. It can be seen that the average energy of the scattered photon is

$$\overline{h\nu} \propto \gamma^2 h\nu \quad (1.15)$$

proportional to the square of the electron Lorentz factor. This is why inverse Compton scattering is an effective process to create very high energy photons. In both cases, synchrotron and inverse Compton, the spectrum of the emitted radiation will have a slope $\alpha' = (\alpha + 1)/2$, if α is the index of a power law distribution of electrons. In the case of inverse Compton scattering when passing to the Klein-Nishina regime the spectrum will be steeper with a slope $\alpha' = \alpha + 1$; therefore even with a power law distribution of electrons a break will be produced in the gamma ray spectrum due to the Klein-Nishina effect [11].

CHAPTER

2

GAMMA RAYS DETECTORS

The gamma-ray energy domain covers an extended range of energies, above about 10^5 eV, more than all the rest of the known electromagnetic spectrum. This enormous energy band can be divided in 6 categories: low energies below 30 MeV, high energies between 30 MeV and 30 GeV, very high energies between 30 GeV and 30 TeV, ultra high energies between 30 TeV and 30 PeV and extremely high energies above 30 PeV.

Only a limited fraction of the Universe can be observed in the gamma domain. As can be seen in figure 2.1 the gamma mean free path is strongly energy dependent and is determined by interaction of gamma-rays with the diffuse extragalactic photon fields. For example gamma rays of 1 PeV have mean free path of 8.5 kpc, thus all objects beyond our Galaxy are not visible in PeV gamma rays; for TeV gamma rays the visibility range increases to about 100 Mpc [12].

In order to cover the range of the gamma ray band described before, different detection methods have to be applied, since different interaction phenomena concern gamma-rays of such different energies. The categories cited are themselves defined by the techniques employed in their detection, rather than the interaction phenomena implied. For example below 30 MeV Compton telescopes are used to detect gamma rays since the Compton process is the dominant interaction process at such energies. The high and very high energies band observations are based on the pair-production interaction, even if

in different ways: high energy telescopes identify the pair in satellite detectors, whereas very high energy telescopes detect the resulting electromagnetic cascade produced by gamma rays in their interaction with the upper layers of the atmosphere [13].

In this chapter I give an overview of the main physical processes regarding the interaction of charged particles with matter, photons pair production and cascade, and the space and ground based telescopes used to detect high and very high energy gamma rays.

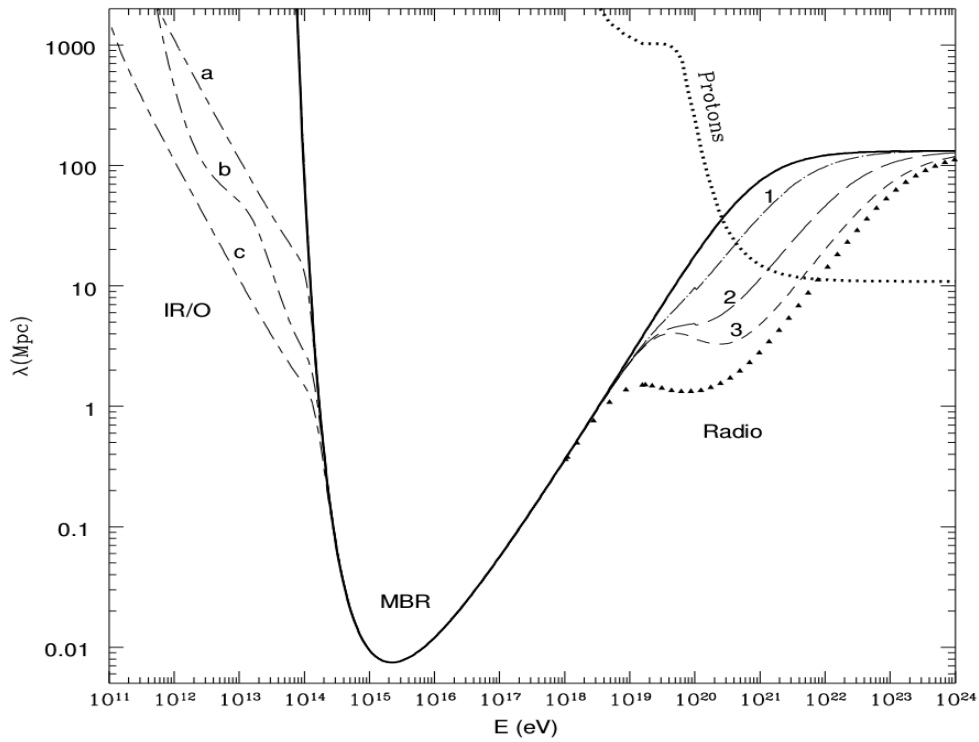


Figure 2.1: Mean free paths of gamma-rays in the intergalactic medium at redshifts $z \ll 1$ (non cosmological epochs). Below 10^{14} eV γ -rays interact with infrared and optical photons, above 10^{19} eV with low frequency radio emission. Large uncertainties in predicted mean free paths are the result of poorly known fluxes of the extragalactic diffuse background radiation at these wavelengths. Between 10^{14} eV and 10^{19} eV, γ -rays interact mainly with the 2.7 K CMB radiation, therefore the mean free paths can be predicted with very high accuracy. [12]

2.1 Charged particles interaction with matter

There are two main processes that characterize the passage of charged particles through matter: inelastic collisions with the atomic electrons of the material, which are responsible of a loss of energy of the particles, and elastic scattering from nuclei, which are responsible of a deflection of the particles from their incident direction. Other processes are: Bremsstrahlung, emission of Cherenkov light and nuclear reactions [8].

2.1.1 Ionization energy loss

Inelastic collisions of charged particles with matter result in a transfer of energy to atomic electrons, causing the excitement or the ionization of the atoms. Inelastic collisions are statistical in nature but, since their number is generally large, the fluctuations are small and one can work with the average energy loss per unit path length, also called stopping power dE/dx . The correct quantum-mechanical treatment leads to

$$-\frac{dE}{dx} = z^2 \frac{Z}{A} \rho f(\beta, I) \quad (2.1)$$

called Bethe-Bloch formula. Here z is the charge of the incoming particle, Z and A are the atomic and mass number of the material with density ρ and $f(\beta, I)$ is a function of the particle velocity and the so-called mean excitation potential I , which is essentially the average orbital frequency of the atomic electron times Planck's constant. Anyway some corrections have to be made to the Bethe-Bloch formula:

- A density effect, due to the polarization of the atoms of the material caused by the electric field of the incoming charged particle. This causes the shielding of the atomic electrons far from the path of the particle, which results in a lesser contribution to energy losses.
- A shell correction that accounts for effects which rise when the velocity of the incident particle is comparable or smaller than the atomic electron's orbital velocity. In such a case the assumption that the atomic electron is stationary falls off.
- In case of electrons corrections must be made for two reasons: because of their small mass the assumption that the particle is undeflected during the collisions is no more valid; the second is the indistinguishability of incoming electrons and the atomic electrons.

2.1.2 Cherenkov radiation

A charged particle passing through a dielectric medium induce the polarization of the molecules in the vicinity of particle's trajectory. If the particle moves slow, that is $\beta = v/c \ll 1$, the polarization is symmetrical around and along the trajectory resulting in no radiation. The situation changes if the particle is relativistic and pass through the medium with velocity

$$v_p > \frac{c}{n} \quad (2.2)$$

greater than the speed of light in the medium with refractive index n . In this case the polarization along the trajectory is no more symmetrical and a resulting radiation is emitted with an angle

$$\cos\theta = \frac{1}{\beta n} \quad (2.3)$$

respect to the particle direction of motion [13].

2.2 Pair production

Pair production involves the transformation of a photon in a electron-positron pair. It can take place if the photon energy is grater than $2m_e c^2 \approx 1.022$ MeV and only in the presence of a third body. The reason why pair production cannot take place in free space is that energy and momentum cannot be both conserved, thus a third body is needed. The mean free path of a gamma-ray for pair production is found to be

$$\lambda_{pair} = \frac{1}{N\sigma_{pair}} \approx \frac{9}{7}L_{rad} \quad (2.4)$$

similar to radiation length for Bremsstrahlung, with N density of atoms [2] [8].

2.3 Space based detectors

Since the atmosphere is not transparent to gamma rays, their direct detection requires detectors placed in space. At energies above a few tens of MeV, the detection of gamma-rays is based on their conversion into an electron-positron pair thanks to foils of dense metal. A successive tracking system and a calorimeter measure the arrival direction and the energy of the incoming gamma ray. To reduce the the background events caused by the higher

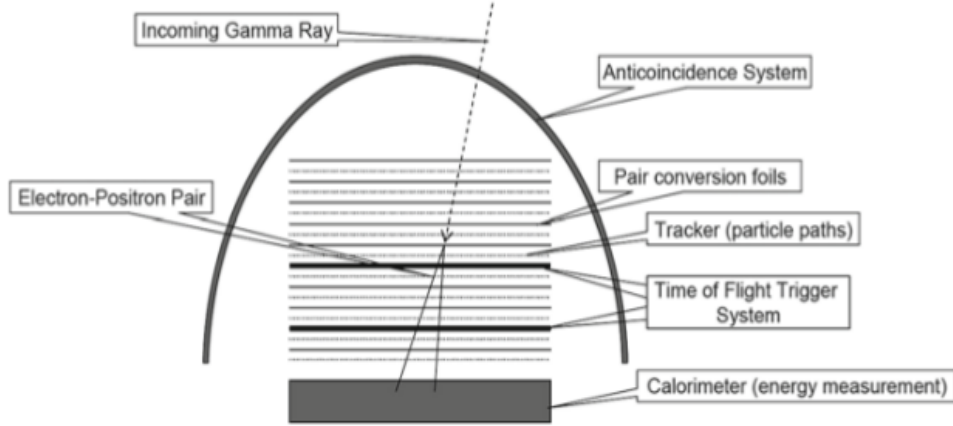


Figure 2.2: Schematic diagram of a telescope which use the conversion of a gamma-ray into an electron-positron pair. [5]

cosmic-ray flux, usually an anticoincidence shield covers the instrument [14].

The effective detection area of satellite experiments is limited to a few m^2 , because of the limitations on size and weight for the payloads of current launchers [5]. The steep spectra of gamma-ray sources limits the operative energy range of these detectors to some hundreds of GeV. Above these energies only very few photons can be detected during the lifetime of the instrument [14].

2.3.1 Fermi

Fermi is a space-based gamma-ray experiment launched by NASA on 2008. It consists of two instruments: the Gamma-ray Burst Monitor (GBM) and the Large Area Telescope (LAT). The GBM is sensitive to X-rays and gamma-rays from 8 keV to 40 MeV and it permits to complements the LAT observations of transients sources. The LAT is a telescope for gamma-rays in the energy range between 20 MeV and 300 GeV, that allows measures of their arrival time, direction and energy [5]. It consists of a tracker composed of 36 layers of silicon strip detectors interleaved with 16 layers of tungsten foil, 12 thin and 4 thick layers adding up to a total radiation length of $1X_0$, in order to facilitate the pair creation. The calorimeter is composed of 8 layers of CsI crystal summing up to a total radiation length of $8.5X_0$. The anticoincidence system surrounding the telescope, consisting of plastic scintillators read out by photomultiplier tubes, allows the cuts of the cosmic-ray background. It has effective area of $0.65 m^2$ at 1 GeV with a wide field of view of about 2

sr and an unprecedented angular resolution that goes from 0.8° at 1 GeV to 0.2° above 10 GeV [11].

2.3.2 AGILE

AGILE (Astro-rivelatore Gamma a Immagini LEggero) is a small telescope, projected by the Italian Space Agency (ASI) launched in 2007. AGILE's gamma-ray imager (GRID) is similar to the Fermi-LAT but on a smaller scale. It consists of a Tungsten-Silicon tracker composed of 12 planes of area of about $40\text{cm} \times 40\text{cm}$ and a CsI(Tl) calorimeter [11]. It is devoted to gamma-ray observations in the energy range 30 MeV - 50 GeV with a peak effective area of about 600 cm^2 up to about 10 GeV. Moreover it can perform hard X-ray imaging in the 18-60 keV energy range [15].

2.4 Electromagnetic cascade

When a relativistic electron or positron or a high energy photon interacts with matter, because of pair production and Bremsstrahlung effects, it leads to the formation of the so-called electron-photon shower. What happens is that an high energy photon interacting with matter converts in a electron-positron pair. Those particles, in turn, will produce photons by Bremsstrahlung. If these photons are enough energetics, they lead both to the creation of a electron-positron pair and so on [2].

The development of the cascade is a statistical process. We can describe it through a simple model using the notion of radiation length, as can be seen in figure 2.2. In the relativistic limit we can approximately consider $\lambda_{pair} \simeq L_{rad}$. Starting with a photon of energy E_0 , after a mean free path λ_{pair} it will convert in an electron-positron pair interacting with matter. These particles, each with mean energy $E_0/2$, will interact with matter producing Bremsstrahlung photons with mean energy $E_0/4$. If the process continues for n mean free path, the total number of particles (photons, electrons and positrons) will be

$$N = 2^n \quad (2.5)$$

each with average energy

$$E(n) = \frac{E_0}{2^n} \quad (2.6)$$

The shower development reaches a maximum at a penetration depth $R_{max} = n_{max}\lambda_{pair}$ and than stops abruptly at a critical energy

$$E_c = \frac{E_0}{2^{n_{max}}} \quad (2.7)$$

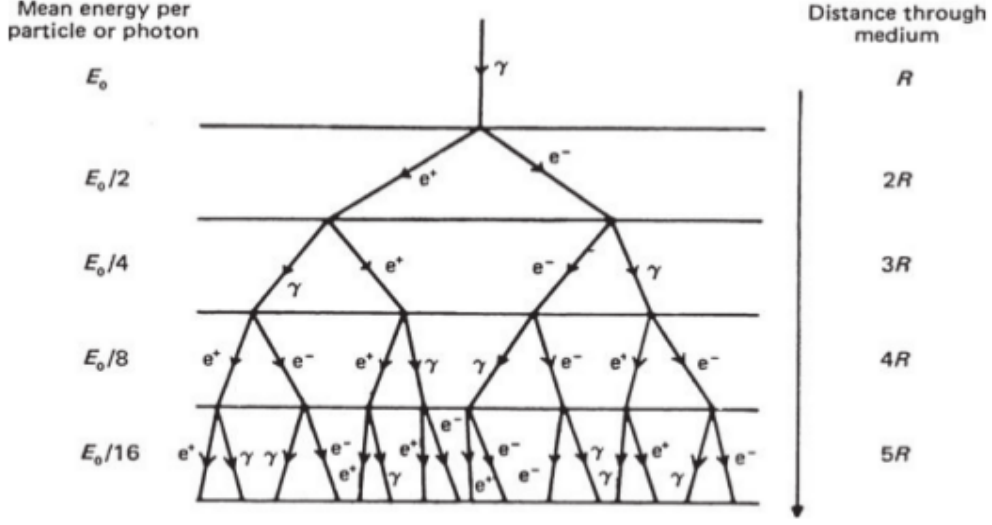


Figure 2.3: Schematic picture of a simple model for electromagnetic shower development [2]

that represent the energy level under which electrons and positrons start losing energy through ionization rather than Bremsstrahlung and photons interact through Compton scattering and photoelectric effect instead of pair production [8].

2.5 Ground based detectors

When the photons energy exceed 100 GeV the gamma-ray fluxes are very faint. Space based experiments are thus limited by their detection area that cannot exceed few m^2 . Gamma-rays cannot be directly detect since the atmosphere is opaque to them. Fortunately above this energy an alternatively method can be used to detect, indirectly, gamma-rays: observation of electromagnetic showers caused by their interaction with the upper layers of the atmosphere. These cascade of relativistic particles can be observed through Cherenkov radiation by ground based detectors in two ways: detecting the Cherenkov light of the shower particles in the atmosphere or through water tanks detectors that detect Cherenkov light of the shower particles in water. In both cases the instrumented area A should be such that

$$N_{ev} = \Phi(> E) \times A \times \Delta\Omega \times T \quad (2.8)$$

gives a reasonable number of events [5] in a time T , with $\Delta\Omega$ covered solid angle and $\Phi(> E)$ has units of $photons\ s^{-1}\ cm^{-2}\ sr^{-1}$.

The atmosphere presents a density profile that decreases with altitude. Typically the photon initiated shower starts at an altitude of about 20 km above sea level (a.s.l.) and reaches a maximum that depends on the photon energy. For 1 TeV gamma-rays this maximum corresponds to an altitude of 10 km a.s.l. [11]. As pressure even the refraction index n decreases with altitude. As consequence the angle of Cherenkov light emitted respect to the particle direction in electromagnetic cascade increases with depth from the top of the atmosphere, since $\cos\theta_c \simeq 1/n$. Typically the Cherenkov angle increases from 0.66° at an height of 10 km to 0.74° at 8 km [5] .

2.5.1 IACT

The Imaging Air Cherenkov Technique (IACT) is based on wide-field optical telescopes. They consist in a large reflector of about 10 m radius that reflect the Cherenkov light into high speed multi-PMT camera positioned in the focal plane. This technique permits the full reconstruction of the arrival direction of photons, discrimination of photon events against the cosmic-ray background and calorimetric measurement of the shower, since the Cherenkov light intensity is proportional to the energy of the primary photon [11].

The Cherenkov light emitted by charged particles in shower arrives with a spectrum that peaks at 300-350 nm (blue light). The flux of this light is very faint: for a 1 TeV gamma-ray Cherenkov light has density in the range from 100 to hundreds photons m^{-2} above 2000m and this number drops to 10 photons m^{-2} at sea level, thus large collection area are needed [12].

Another problem arising from the faintness of Cherenkov light is related to the background optical photons. The night sky background, even at the darkest sites, is of order 10^{12} photons $m^{-2} s^{-1} sr^{-1}$. For camera with field of view of diameter 0.2° and effective mirror area of $60m^2$ the background can be reduced to 10^9 photons s^{-1} [11]. For this reason, in order to detect the faint Cherenkov light against the optical background, IACTs have to operate with exposure time of a few nanoseconds. The need of dark nights affect the duty cycle of the imaging technique reducing it to a about 10-15%, since even moonlight make it impossible to detect the radiation.

Once dealt with these problems, the main issue to overcome is related to the background of cosmic rays, also producing air showers, that is up to 10^4 times larger than gamma rays rate [11]. Two image criteria for single imaging telescope cosmic-ray background rejection are: the intersection of the shower track with the source location and the shower shape. The first is valid only for point sources, the second is based on the larger momentum transfered to secondary particles produced in hadronic interactions [12], as can be seen in figure 2.4.

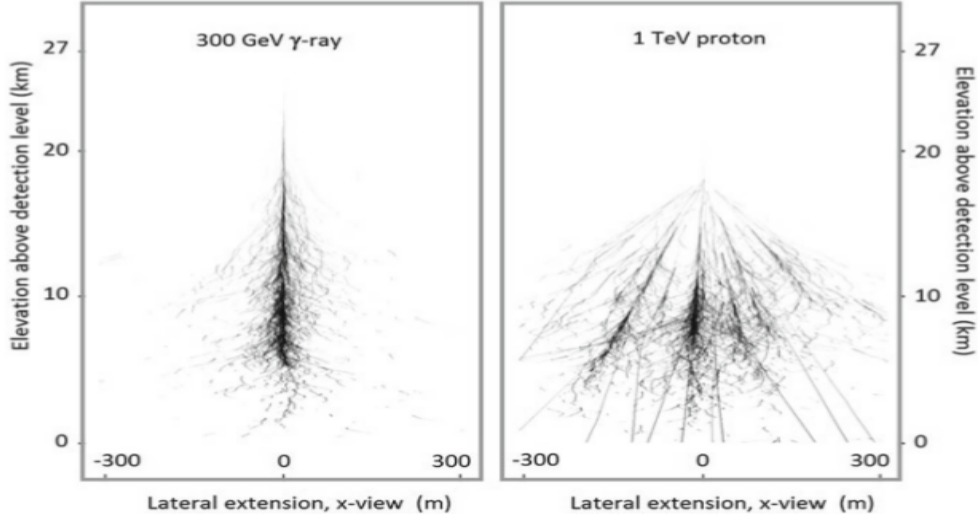


Figure 2.4: Comparison of a pure electromagnetic shower (from a 300 GeV gamma-ray) and a shower initiated by a 1 TeV proton. The plot shows the projection of secondary particle trajectories onto a plane in which the ordinate corresponds to the elevation [5].

A technique that permits to reduce the cosmic-ray background by a further factor of 100, consists in the stereoscopic observation of the shower. Basically, the concept of stereo imaging is based on simultaneous detection of a single shower by, at least, two telescopes [12] separated by a proper distance. Since the Cherenkov angle at the shower maximum is 0.66° , this results in a ring-like region at the ground of radius $R \simeq 10\text{km} \times 0.012\text{rad} = 120\text{ m}$ [5]. Thus, the positioning of imaging telescopes at a distance of about 120 m using field of view (FoV) of IACT of about 4° , allows to obtain different images of particles shower. Moreover, quite the opposite of single telescope observations that measure only the shower inclination in the focal plane, the stereoscopic technique allows full reconstruction of the arrival direction of the primary photon with an accuracy of about 0.1° [12], as can be seen in figure 2.5. However this technique implies a non-negligible loss in the detection rate because of the overlap of shower detection areas of individual telescopes [12]. Currently there are three IACT stereoscopic systems operating around the world: H.E.S.S. (High Energy Stereoscopic System) in Namibia, MAGIC (Major Atmospheric Gamma-ray Imaging Cherenkov Telescopes) in La Palma on the Canary Island and VERITAS (Very Energetic Radiation Imaging Telescope Array System) in southern Arizona.

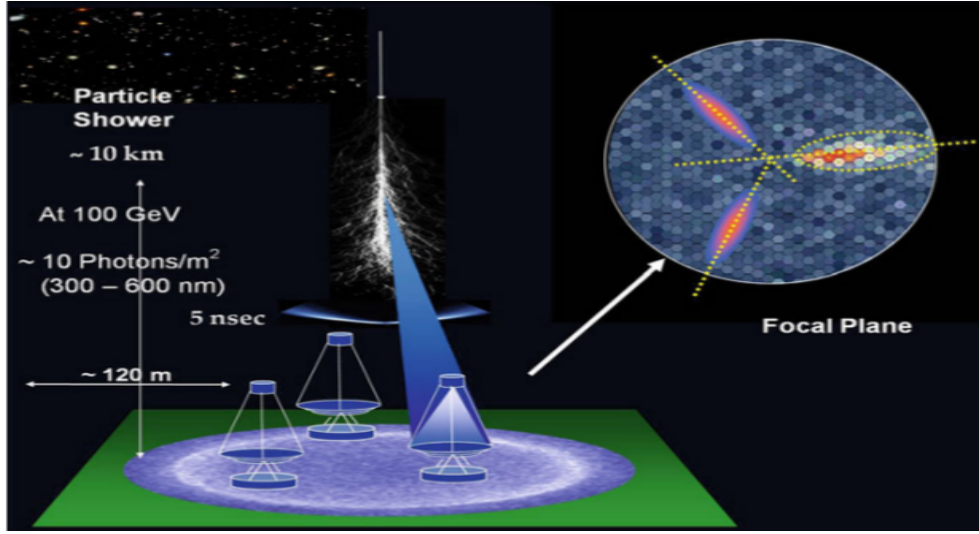


Figure 2.5: Sketch of the imaging atmospheric Cherenkov telescopes technique. A shower initiated by a gamma-ray of energy ≥ 100 GeV is observed through the Cherenkov radiation emitted by charged particles with about 10 m diameter reflectors positioned about 100 m away from each other. The images at the focal plane of different telescopes (upper right inset) allow the determination of the shower direction and energy [5].

2.5.2 CTA

The Cherenkov Telescope Array (CTA) is the next generation IACT that will be finished in the next few years and that will make it possible to observe the gamma sky for the next 30 years. CTA is a worldwide project composed of more than 1000 members working in 210 institutes from 32 countries. Present Cherenkov telescope arrays are sensitive in the energy range between 80 GeV and 50 TeV, with field of view between 3° and 5° . CTA will improve all these characteristics and the energy sensitivity at 1 TeV will increase of about one order of magnitude, better energy and angular resolution will be reached.

The use of different types of telescope permits to optimize costs and to respect the scientific requirements covering the full energy range in a cost effective way [16]. The three types of telescopes are distinguished by the acronyms: LST, MST and SST. The LSTs (Large Size Telescopes) will cover the energy range between 20 GeV and 100 GeV with a field of view of 4° . With a mirror diameter of 23 m and a focal length of 28 m will permit to collect as many photons as possible from the low energy showers. The MSTs (Medium Size Telescopes) will cover the medium energy range between 100 GeV and 10 TeV, having a mirror diameter of 12 m and a focal length of 16 m with a

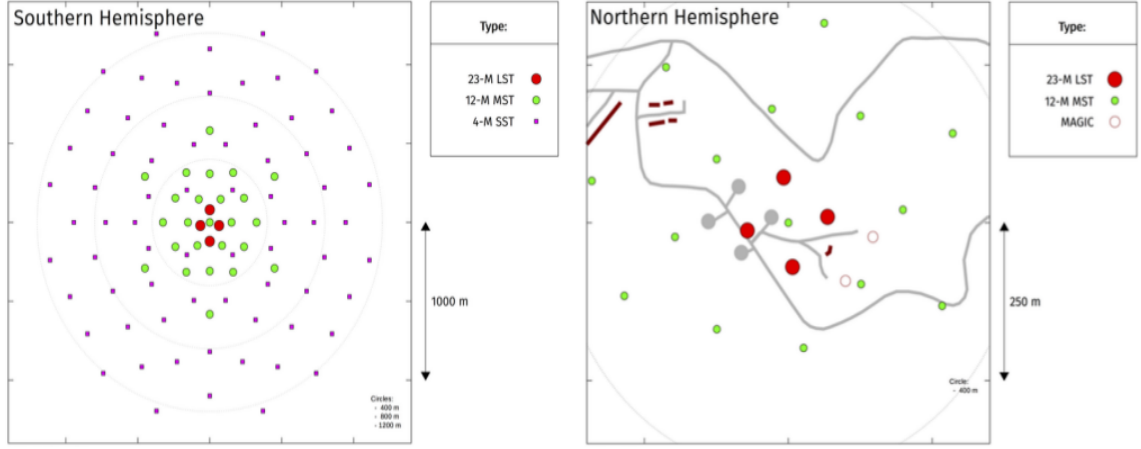


Figure 2.6: Possible layout of telescope array in the Southern (left) and Northern (right) hemisphere [16].

field of view of 8° . The SSTs (Small Size Telescopes) will cover the energy range between 10 TeV and 200 TeV, with mirror diameter of 4 m and field of view of 9° . As can be seen in figure 2.6 more than 100 telescopes will be located in two sites in both hemispheres: in the Atacama Desert in Chile and in La Palma in the Canary Islands. This choice will permit to reach a full sky coverage. The Northern site using only LSTs and MSTs will be optimized in the energy range from about 20 GeV to about 20 TeV. The Southern site will be equipped with about 100 telescopes of all sizes in order to observe the sky in the entire energy range of CTA.

2.5.3 EAS arrays

The Extensive Air Shower (EAS) technique is based on the direct detection of particles through scintillators or water Cherenkov detectors, spread over large areas. Since the particles have to reach the ground for a direct detection, these detectors are typically located at high altitude. This technique was first developed in order to detect cosmic-rays at PeV (10^{15} eV) and EaV (10^{18} eV) energies. Anyway it can be used in the gamma-ray astronomy field if the energy threshold is reduced; this can be achieved using dense particle arrays [17].

Water Cherenkov arrays are based on the detection of Cherenkov light generated by the passage of relativistic secondary particles in EAS. This is pursued using photomultiplier tubes installed deep in tanks filled with pure water. Typically these detectors have lower sensitivity to point sources, due to their high energy threshold and worse angular resolution, respect to IACTs. Any-

way they presents the unique advantage of continuously monitor the sky, having a duty cycle of about 100%. This is possible even thanks to the large solid angles achievable with this technique. If l is the latitude of the detector and Θ_{max} is the maximum zenith angle of the observation, the solid angle covered by a water Cherenkov detector is [11]

$$\Omega = 4\pi \cos(l) \sin(\Theta_{max}) \quad (2.9)$$

typically of order of 1 sr or more.

HAWC

The HAWC [18] (High Altitude Water Cherenkov) observatory is water Cherenkov detector located in Mexico at an elevation of 4 km a.s.l.. It consists of 300 water Cherenkov detectors 5 m high and with 7.32 m diameter. Each tank contains 190,000 liters of purified water and four photomultiplier tubes at the bottom. The PMTs are covered by 4 m of water, corresponding to 10 radiation length. This guarantees the full absorption of electrons, positrons and gamma-rays of the electromagnetic shower within the detector. The layout of HAWC detector can be seen in figure 2.7.

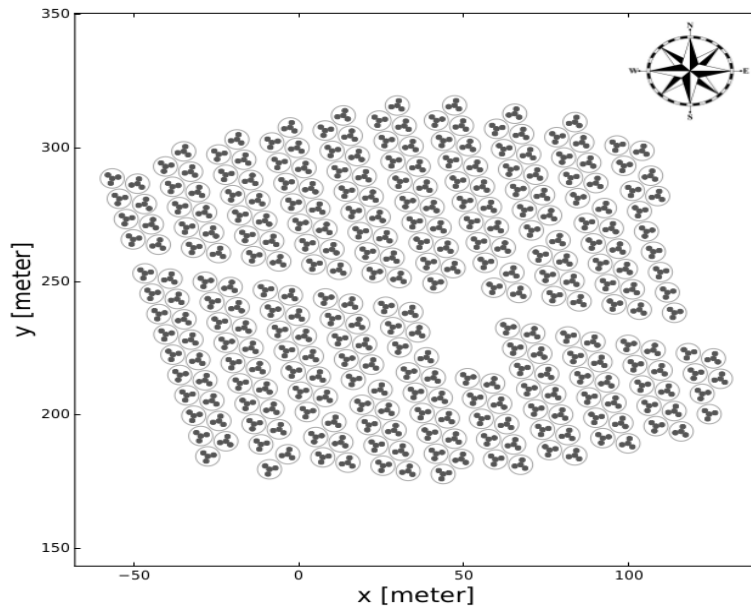


Figure 2.7: Layout of HAWC water Cherenkov detectors and position of the photomultiplier tubes. The conspicuous gap indicates the location of the counting house, which is centrally located to minimize the cable length. [18].

The relativistic particles will produce Cherenkov radiation in water. This

light generates a trigger when a sufficient number of PMTs record a hit within a window of 150 ns; for example in the HAWC catalog paper, that will be discussed in the next chapter, 28 hits were required. The angular resolution depends of the event size: from about 0.2° for events that hit almost all the PMTs to about 1° for the events near the trigger threshold.

The rejection of cosmic-rays background can be done on the basis of the morphology. The HAWC detector can reach a 99% level of rejection of the hadronic background in case of largest event and at least 85% near the trigger threshold.

Milagro

Milagro [5] was a water Cherenkov detector that operated between 2000 and 2008 in New Mexico. It consisted of a large water reservoir of $60 \times 80m^2$ located at an altitude of 2.6 km a.s.l.. In 2004 an array of 175 small tanks were added, spread over an area of $200 \times 200m^2$ around the central reservoir.

CHAPTER

3

A MULTI-ZONE DIFFUSION MODEL FOR THE GEMINGA TEV HALO

Pulsar Wind Nebulae are nebulae generated by the relativistic wind of electrons and positrons accelerated by a pulsar. These particles interact with various photon fields and with the Galactic magnetic field, generating wide regions visible in the gamma domain that could extend for tens of parsec. The recent HAWC observation of an extended gamma emission, coincident with the position of Geminga, constrain the diffusion coefficient around these regions to be about 2 order of magnitude lower than the standard diffusion coefficient in the interstellar medium. This could have important implications both for the resulting electrons and positrons flux measured at Earth and for the search of extended gamma regions around other pulsars.

3.1 Pulsars and Pulsar Wind Nebulae

Stars sustain their mass through the radiation pressure generated from the nuclear fusion processes that occur in their core. When a massive star runs out of its hydrogen it collapse under its own gravity causing an increase of

temperature in its central zone. This increase in temperature allows the beginning of helium fusion processes. When the stars run out of helium too, starts another contraction which cause another increase in temperature allowing the beginning of the nuclear fusion of other elements and so on [3]. The successive fusion processes lead to the formation of ^{56}Ni which undergoes successive electron-capture and β^+ decays to form ^{56}Co and ^{56}Fe . From this point the fusion processes starts being endothermic, the star's radiation pressure can no more sustain its gravitational collapse and it implodes releasing an enormous amount of energy of about 10^{53} erg.

If the core mass is less than $3M_{\odot}$ the newborn dead star is a neutron star.

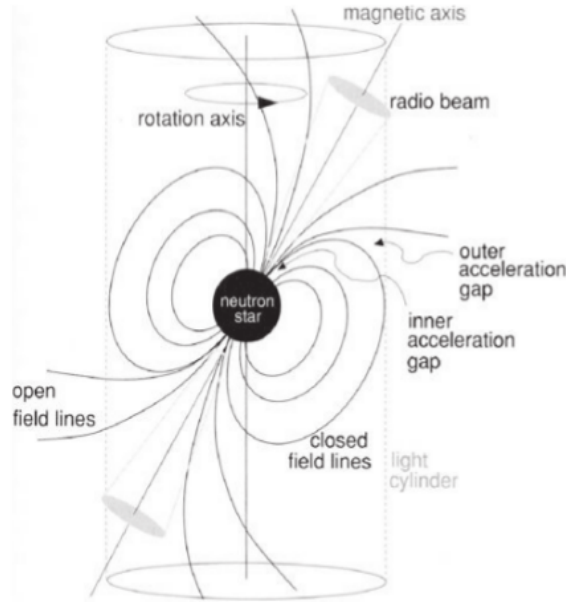


Figure 3.1: A schematic model of a pulsar as a magnetised rotating neutron star in which the magnetic and rotation axes are misaligned. The radio pulses are assumed to be due to beams of radio emission from the poles of the magnetic field distribution and are associated with the passage of the beam across the line of sight to the observer [2].

These are objects extremely compact, with radii of 10-20 km, and with rotational period of a few milliseconds, as seen in chapter 1. Since there is no internal heat source, the mechanism that sustain these stars is the degeneracy pressure of neutrons [2]. If a pulsed radio emission is observed these objects are also referred to as pulsar. The radio pulses are emitted along the magnetic axis of the neutron star, which does not coincide with its rotational axis as can be seen in figure 3.1.

These objects are also responsible of high and very high energy gamma-ray

emission. Pulsed gamma rays are generated in the pulsar magnetosphere and can be observed up to a few GeV. The very high energy component is due to a wind of relativistic particles, mainly electrons and positrons, produced in the magnetosphere of the pulsar and accelerated by the rotationally induced electric field. These particles can eventually escape the magnetosphere and are further accelerated by the surrounding shock with the interstellar medium. This relativistic wind may interact with the Galactic magnetic field or with various photon field, producing a spectrum of photons from radio wavelengths to gamma rays via synchrotron and inverse Compton processes, creating structures surrounding the star called Pulsar Wind Nebulae (PWNe) [12].

3.2 Geminga TeV halo

Geminga is a radio quiet neutron star, since its radio beam does not point towards the Earth. Because of this it was one of the first example of unidentified gamma-ray source.

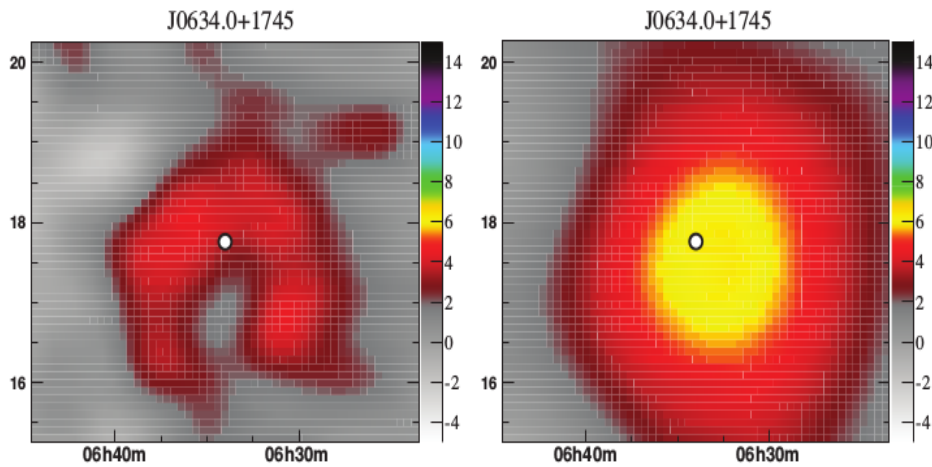


Figure 3.2: Significance map of the Milagro data in a $5^\circ \times 5^\circ$ region around Fermi source J0634.0+1745 the Geminga pulsar. The location of the Fermi source is identified by a white dot. The figure on the left shows the significance map after smoothing by the Milagro point-spread function. The figure on the right shows the same region smoothed by an additional 1° Gaussian in order for an extended emission region. The color scale shows the statistical significance. [19]

The first observation of TeV emission from Geminga date back to 2009 from

the Milagro observatory. Emission over 1 TeV was initially reported but with a too low significance to be classified as a detection. However, since Milagro observed an emission from several degrees, if the map is convolved with a 1° Gaussian to account for an extended emission region (figure 3.2), the significance increases to 6.3σ . This allowed the Milagro collaboration to regard it as a detection of extended emission from Geminga [19].

In 2017 a couple of articles signed by the HAWC collaboration opened the debate about Geminga's multi-TeV emission. In the first one was presented the second HAWC catalog based on 507 days of observations by the complete HAWC detector. Once again the HAWC gamma source associated with Geminga presented a detection with 11.22σ when fitted with a uniform disk model of 2° in radius [18]. However at a later time the same collaboration published a second and more discussed paper based on a detailed study of the extended emission around the Geminga region.

The HAWC collaboration tested three different morphological models, as can be seen in figure 3.3, in order to explain the extended gamma emission observed: an extended disk with a constant average flux within a distance from the source; a symmetrical two dimensional Gaussian with a flux depending on distance; a diffusion model where electrons and positrons diffuse from the central source through the interstellar medium [20]. Of the three models only

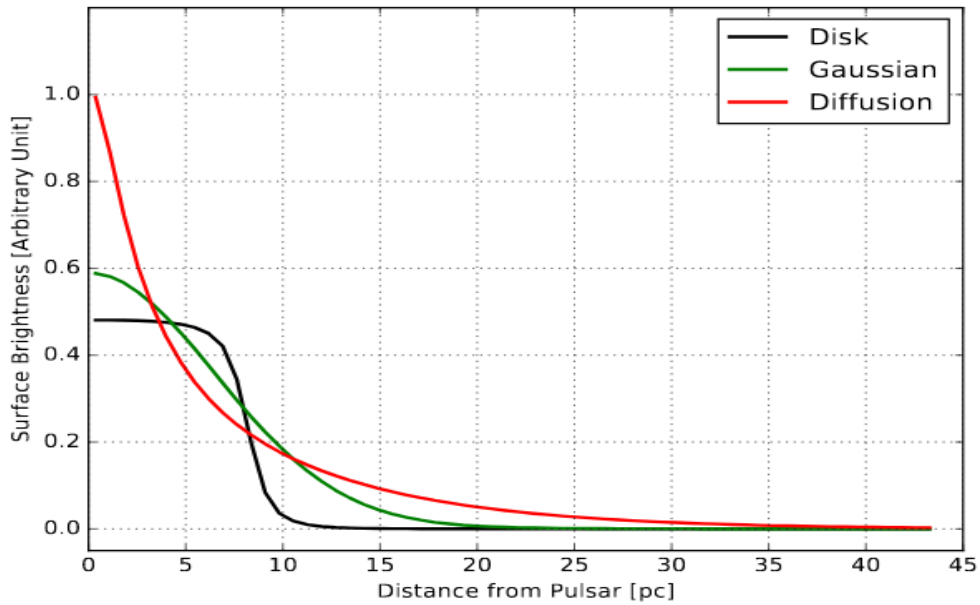


Figure 3.3: Surface Brightness profile of Geminga for different morphological models as a function of distance from the pulsar [20]

the last depends on physical assumptions, since it is related to the particles' diffusion rather than to a flux distribution.

As can be seen in figure 3.4 the observed gamma emission extends for about 10° from the source. The gamma emission origin is ascribed to electrons and positrons diffusing away from the pulsar, interacting with various photon fields (cosmic microwave background, infrared and optical) through inverse Compton scattering. The fit is performed with a spatial diffusion model with

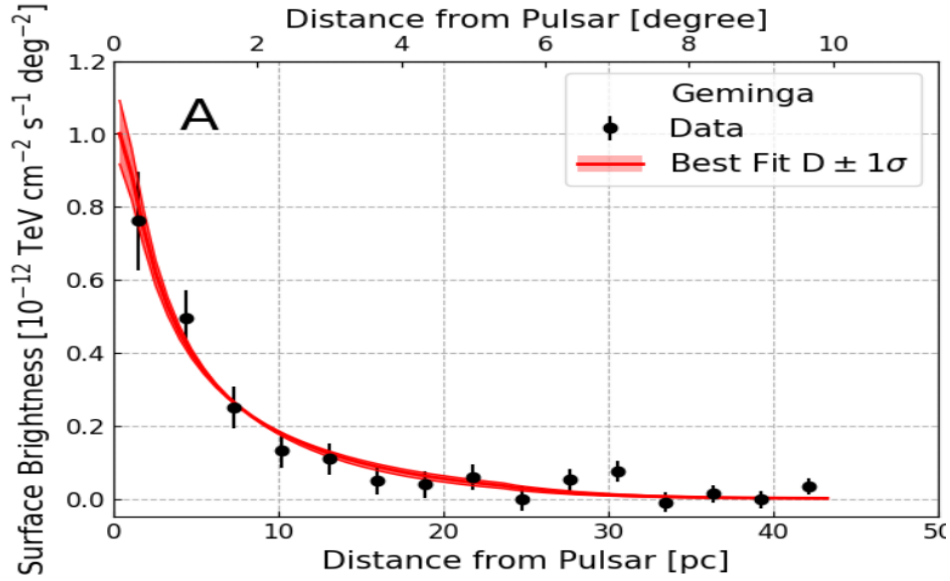


Figure 3.4: Surface brightness profile of TeV gamma emission from Geminga as a function of the distance from the source. [21]

the gamma ray flux N as a function of the angle from the source θ and the gamma ray energy E

$$\frac{d^2 N}{dE d\Omega} = N_0 \left(\frac{E}{20 \text{ TeV}} \right)^{-\alpha} \frac{1.22}{\pi^{3/2} \theta_d(E) (\theta + 0.06 \theta_d(E))} e^{-(\theta/\theta_d(E))^2} \quad (3.1)$$

where N_0 is the flux normalization at 20 TeV, Ω is the solid angle and $\theta_d = \frac{\pi}{180} \frac{r_d}{d_{src}}$ is the diffusion angle with d_{src} distance of the source from the Earth and $r_d = 2\sqrt{D(E)t}$ diffusion radius [21]. The model parameters obtained from the HAWC fit are presented in figure 3.5. The value of the diffusion coefficient obtained from the fit is $D_{HAWC} = 3.2 \times 10^{27} \text{ cm}^2 \text{ s}^{-1}$ at 100 TeV, smaller by a factor of about 500 with respect to the diffusion coefficient in the local interstellar medium found from Boron-to-Carbon ratio measurements. Assuming this diffusion coefficient the HAWC collaboration calculated the

Pulsar Parameters		Geminga
(Right ascension, declination) (J2000 source location)	[degrees]	(98.48, 17.77)
τ_c (characteristic age)	[years]	342,000
T (spin period)	[seconds]	0.237
d (distance)	[parsecs]	250^{+120}_{-62}
dE/dt (energy loss rate due to pulsar's spin slowing)	[$\times 10^{34}$ ergs/sec]	3.26
Model Values		
θ_0 (θ_a for 20 TeV γ -ray)	[degrees]	5.5 ± 0.7
N_0	[$\times 10^{-15}$ photons/TeV/cm ² /sec]	$13.6^{+2.0}_{-1.7}$
α		2.34 ± 0.07
D_{100} (Diffusion coefficient of 100TeV electrons from individual fit of PWN)	[$\times 10^{27}$ cm ² /sec]	$3.2^{+1.4}_{-1.0}$
Energy Range	[TeV]	8 to 40
Luminosity in gamma-rays over this energy range	[$\times 10^{31}$ erg/sec]	$11 \times (d/250 \text{ parsec})^2$
Assumed Parameters		
L_0 (initial spin down power)	[$\times 10^{36}$ ergs/sec]	27.8
W_e (total energy released since pulsar's birth)	[$\times 10^{48}$ ergs]	11.0

Figure 3.5: Pulsar parameters, values of parameters from the model fitting to the observed extended gamma-ray emission, and assumed parameters of HAWC model [21]

positron flux at the Earth with their EDGE code [22] and concluded that Geminga cannot produce a measurable contribution to the positron flux at the Earth measured by AMS-02.

However H.E.S.S. has detected electrons and positrons up to about 20 TeV and these particles cannot travel more than 10-20 pc accounting for their cooling time if such a low diffusion coefficient is assumed. However what is suggested by other authors [23] is that, since HAWC data probe a region extended up to a few tens of pc around Geminga, it is more likely that the slow diffusion scenario is limited around the source. Beyond this slow diffusion region, particles diffuse faster as typical Galactic cosmic rays

$$D(r, E) = \begin{cases} D_1(E), & r < r_b \\ D_2(E), & r \geq r_b \end{cases} \quad (3.2)$$

with r_b breaking radius between the slow and the fast diffusion zone. Within this picture, also called two-zone diffusion model, we can explain both the HAWC gamma observations and the diffusion coefficient in interstellar medium suggested by B/C measurements.

3.3 Diffusion-loss equation

Calculations to explain the observed abundances of elements in cosmic rays show that they should pass through a column density of 5 g cm^{-2} before reaching the Earth. The column density of the Galaxy is of about $10^{-3} \text{ g cm}^{-2}$, so cosmic rays propagation cannot occur along straight lines and should be present some mechanism to confine them within the Galaxy [24]. Once injected by their source, charged particles experience random scattering by irregularities in the Galactic magnetic field due to fluctuations in the field or to instabilities caused by the motion of the same cosmic rays, hence they diffuse from their sources through the interstellar medium. Since once injected the number of particles should be conserved, from the continuity equation

$$\frac{\partial N(r, t)}{\partial t} + \nabla \cdot J(r, t) = 0 \quad (3.3)$$

where N is the particles density function and J is the particles flux. Given that the particles flux is related to the gradient of density through the diffusion coefficient

$$J(r, t) = -D(r)\nabla N(r, t) \quad (3.4)$$

we have

$$\frac{\partial N(r, t)}{\partial t} = \nabla \cdot [D(r)\nabla N(r, t)] = D\nabla^2 N(r, t) \quad (3.5)$$

if a constant diffusion coefficient is assumed [4].

Throughout their diffusion through the Galaxy cosmic rays are subject to various energy gains and losses

$$-\frac{dE}{dt} = b(E) \quad (3.6)$$

if $b(E)$ is positive particles lose energy, if negative particles gain energy. In order to take into account for the energy losses, consider an elementary volume dV in which particles are injected. $N(E)\Delta E$ is the number of particles with energies in the interval $E, E + \Delta E$ at time t , at a later time $t + \Delta t$ these particles are replaced by those in the energy range $E', E' + \Delta E'$ where

$E' = E + b(E)\Delta t$ and $E' + \Delta E' = (E + \Delta E) + b(E + \Delta E)\Delta t$. For $\Delta E \rightarrow 0$ and subtracting E'

$$\Delta E' = \Delta E + \frac{db(E)}{dE} \Delta E \Delta t \quad (3.7)$$

So the change in the number of particles $N(E)\Delta E$ in the time interval Δt is

$$\Delta N(E)\Delta E = -N(E, t)\Delta E + N(E + b(E)\Delta t, t)\Delta E' \quad (3.8)$$

Thus for small $b(E)\Delta t$ and substituting $\Delta E'$

$$\Delta N(E)\Delta E = \frac{dN(E)}{dE} b(E)\Delta E \Delta t + N(E) \frac{db(E)}{dE} \Delta E \Delta t \quad (3.9)$$

that is, for $\Delta t \rightarrow 0$

$$\frac{dN(E)}{dt} = \frac{d}{dE} [b(E)N(E)] \quad (3.10)$$

Including a source term $Q(E, t)$ to account for the injection of particles we obtain

$$\frac{dN(E)}{dt} = D\nabla^2 N(E) + \frac{d}{dE} [b(E)N(E)] + Q(E, t) \quad (3.11)$$

called diffusion-loss equation, that describes the time evolution of the energy spectrum of particles subject to diffusion and energy losses [2].

3.4 Numerical solution of the diffusion-loss equation

The aim of this work is to find models that give results compatible with the HAWC observations. As a first step I solved the diffusion-loss equation for electrons and positrons emitted by Geminga.

I assumed that particles are emitted by the source in all direction with a spherical symmetry. In this picture the diffusion equation to solve is not the one seen in the last equation, but reduces to the form

$$\frac{\partial f}{\partial t} = \frac{1}{r^2} \frac{\partial}{\partial r} (r^2 D(r) \frac{\partial f}{\partial r}) + \frac{\partial}{\partial E} (b(E)f) + Q(E) \quad (3.12)$$

with $f=f(r, E, t)$ energy distribution function of particles at instant t and distance r from the source. Here I neglected the convection term, which will be taken into account as an indirect effect as will be soon clear [25]. In order to solve the diffusion equation I approximated the derivatives in time and energy as follow

$$\frac{\partial f}{\partial t} \simeq \frac{f_j^{n+1} - f_j^n}{\Delta t} \quad (3.13)$$

$$\frac{\partial}{\partial E}(b(E)f) \simeq \frac{f_{i+1}^n b_{i+1} - f_i^n b_i}{\Delta E} \quad (3.14)$$

where n index stands for the time step, i for the energy step and j for the distance step. For the diffusion term I used the discretization proposed in [26]

$$\frac{1}{r^2} \frac{\partial}{\partial r} (r^2 D(r) \frac{\partial f}{\partial r}) \simeq \frac{1}{r_j^2} \frac{1}{\Delta r^2} (r_{j+\frac{1}{2}}^2 D_{j+\frac{1}{2}} (f_{j+1}^n - f_j^n) - r_{j-\frac{1}{2}}^2 D_{j-\frac{1}{2}} (f_j^n - f_{j-1}^n)) \quad (3.15)$$

this approximation gives the chance to take into account a diffusion coefficient that varies with the distance from the source and reduces to

$$\frac{D}{r_j^2} \frac{1}{\Delta r^2} (r_{j+\frac{1}{2}}^2 (f_{j+1}^n - f_j^n) - r_{j-\frac{1}{2}}^2 (f_j^n - f_{j-1}^n)) \quad (3.16)$$

for a diffusion coefficient that does not depend on distance. Otherwise one could evaluate the derivative in the diffusion term and approximate them one by one.

The source term is a power law

$$Q(E) = N_0 E^{-\alpha} \quad (3.17)$$

that describes the spectra of particles accelerated by Geminga from $E_{min} = 1GeV$ to $E_{max} = 500TeV$. Moreover I performed all the simulation under the condition that the energy integral of the spectrum, that gives the total energy released by the pulsar, is

$$\int_{E_{min}}^{E_{max}} EQ(E)dE = 1.1 \times 10^{49} erg \quad (3.18)$$

as assumed by the HAWC collaboration [21].

The diffusion coefficient has a power law energy dependence

$$D(E) = D_0 E^\delta \quad (3.19)$$

with δ called diffusion index, usually in the range $0.3 < \delta < 0.6$. In this work I used $\delta = 1/3$, which is compatible with recent results from AMS-02 for the spectrum of hadronic cosmic rays [21]. However, traveling through the Galaxy, cosmic rays particles will be subjected to convection too, which has been previously neglected. The effect of convection affects only low energy particles because their diffusion is slower. In a pure diffusion scenario the only way to simulate convection is that the diffusion coefficient must become

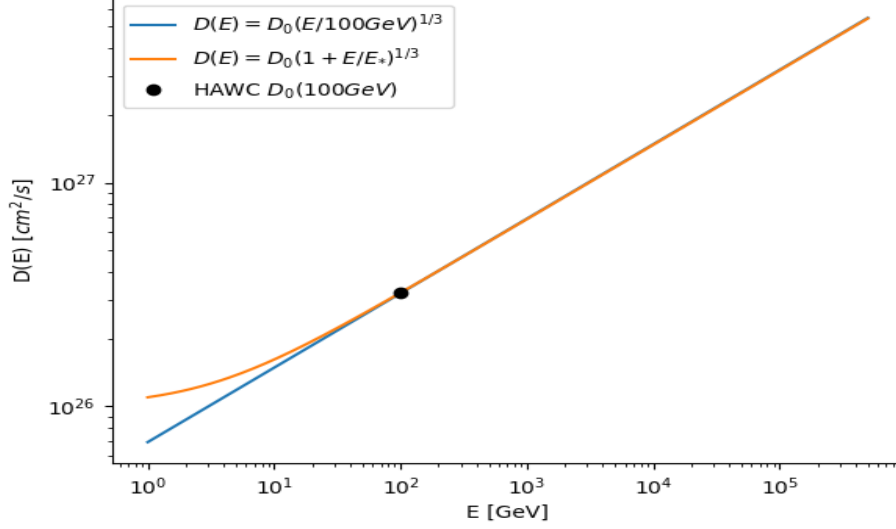


Figure 3.6: Effect of convection on diffusion coefficient (orange line) and diffusion coefficient without convection correction (blue line). The black dot represent the value of diffusion coefficient at 100 GeV from HAWC best fit.

constant at low energy [25], as can be seen in figure 3.6. This is taken into account if a diffusion coefficient of the form

$$D(E) = D_0 \left(1 + \frac{E}{E_*}\right)^\delta \quad (3.20)$$

is used, with $E_* = 3$ GeV.

The energy losses term has the form

$$b(E) = -\left(\frac{dE}{dt}\right) = A_1 \left(3 \ln\left(\frac{E}{m_e c^2}\right) + 19.8\right) + A_2 E + A_3 E^2 \quad (3.21)$$

where A_1 describes the ionization energy losses and depends weakly on energy, A_2 describes the bremsstrahlung losses and A_3 the inverse Compton and synchrotron losses. To make explicit the losses terms, I used from [2]

$$-\left(\frac{dE}{dt}\right)_{ion} = 7.64 \times 10^{-18} \left(\frac{n_H}{cm^{-3}}\right) \left(3 \ln\left(\frac{E}{m_e c^2}\right) + 19.8\right) \quad \frac{GeV}{s} \quad (3.22)$$

$$-\left(\frac{dE}{dt}\right)_{brems} = 3.66 \times 10^{-16} \left(\frac{n_H}{cm^{-3}}\right) \times E \quad \frac{GeV}{s} \quad (3.23)$$

for ionisation and bremsstrahlung for the case of atomic hydrogen. Here n_H stands for interstellar density that typically has value of order $1 cm^{-3}$.

For the third term I used the expression described in the first chapter

$$-\left(\frac{dE}{dt}\right)_{sync} = \frac{4}{3}c\sigma_T u_B \frac{E^2}{(m_e c^2)^2} \quad \frac{GeV}{s} \quad (3.24)$$

for the synchrotron energy loss, with the variation in the magnetic energy density given by [22]

$$u_B = \frac{B^2}{4\pi} = 0.22 \left(\frac{B}{3\mu G} \right) \quad \frac{eV}{cm^3} \quad (3.25)$$

I used $B = 3 \mu G$ as typical value of Galactic magnetic field.

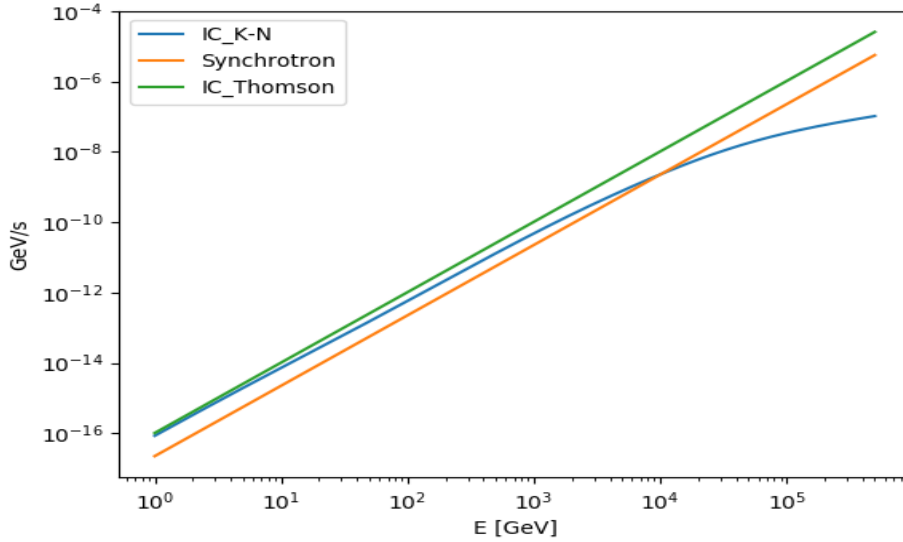


Figure 3.7: Synchrotron and Inverse Compton energy loss terms. Inverse Compton term is presented for both cases: Thomson approximation and accounting for Klein-Nishina effect.

The inverse Compton term, as described in the first chapter, is

$$-\left(\frac{dE}{dt}\right)_{IC} = \frac{4}{3}c\sigma_T u_p \frac{E^2}{(m_e c^2)^2} \quad \frac{GeV}{s} \quad (3.26)$$

with u_p energy density of the photon field. Anyway this expression is valid only in the Thomson limit. If $\gamma_e h\nu \simeq m_e c^2$ the inverse Compton loss term should include the Klein-Nishina effect as [27]

$$-\left(\frac{dE}{dt}\right)_{IC} = \frac{4}{3}c\sigma_T \left[\sum_i \frac{u_i}{(1 + 4\pi\gamma_e\epsilon_i)^{3/2}} \right] \frac{E^2}{(m_e c^2)^2} \quad \frac{GeV}{s} \quad (3.27)$$

with u_i energy density of the photon field for different energy channel and $\epsilon_i = 2.8k_B T_i/m_e c^2$ normalized photon energy for a Black Body radiation with temperature T_i , with k_B Boltzmann constant. The energy densities and Black Body temperatures used are the ones presented in table 3.1.

As can be seen in figure 3.7, if the Thomson approximation is used electrons will be subjected to higher energy losses resulting in an electron density underestimation at higher energies. Anyway even if one takes into account the Klein-Nishina effect, energy loss at higher energy will still be subjected to synchrotron loss which dominate above about 10 TeV. In figure 3.8 can be seen each energy loss contribution and the overall loss term.

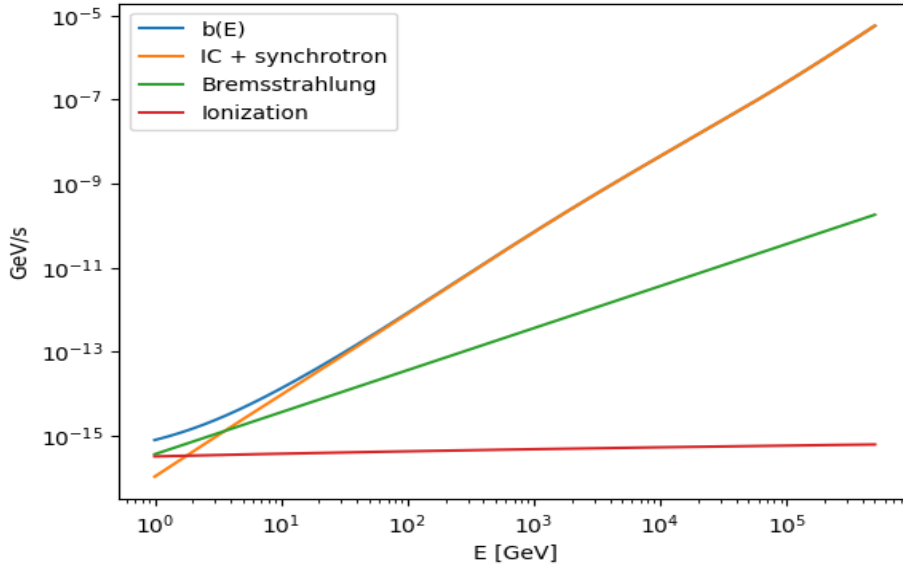


Figure 3.8: Ionization (red line), bremsstrahlung (green line) and radiative losses (orange line) terms. The overall energy loss term is also present (blue line).

Photon field	T(K)	$u(\text{eV cm}^{-3})$
CMB	2.7	0.26
IR	20	0.3
Optical	5000	0.3

Table 3.1: Physical parameters for the photon fields [21].

3.5 Validation of the algorithm

There are two models for which there are analytical solutions of the diffusion-loss equation: a burst like injection, in which all the particles are emitted at the same time $t_0 = 0$ and then start diffusing and losing energy; a continuous injection, in which the particles are injected for all the time of the simulation, meanwhile diffusing and losing energy. I used these cases as tests for the algorithm implemented to solve numerically the diffusion equation. Since the continuous injection model is the one that best describes the emission of particles from a pulsar, here I'll show only this one.

I started neglecting the diffusion of particles setting $D(E) = 0$. If the energy loss term is $b(E) = A_3 E^2$, that is only energy losses due to synchrotron and inverse Compton scattering, and it is assumed that there is continuous injection of electrons, the energy spectrum after time t_f is [2]

$$N(E, t_f) = \begin{cases} \frac{N_0 E^{-(\gamma-1)}}{A_3 t_f (\gamma-1)} [1 - (1 - A_3 E t_f)^{\gamma-1}] & \text{if } A_3 E t_f \leq 1 \\ \frac{N_0 E^{-(\gamma-1)}}{A_3 t_f (\gamma-1)} & \text{if } A_3 E t_f > 1 \end{cases} \quad (3.28)$$

if a power law injection spectrum with index γ is assumed.

I performed this simulation setting $t_f = 1.5 \times 10^4$ years, $\Delta r = 1$ pc and $\Delta t = 4$ years. As can be seen in figure 3.9 the numerical solution follows the analytical one, with a break in the spectrum at an energy of 20 TeV as expected since at this energy $A_3 E t_f > 1$.

The second check concerns the diffusion term. This time I neglected energy losses setting $b(E) = 0$ considering a pure diffusion of electrons continuously emitted by a point source. In this case the energy spectrum of electrons is [12]

$$\begin{cases} f(R, T, E) = \frac{N_0 E^{-\gamma}}{4\pi D(E) R} \text{erfc}\left(\frac{R}{2\sqrt{D(E)t_f}}\right) \\ \text{erfc}(z) = \frac{2}{\sqrt{\pi}} \int_z^\infty e^{-x^2} dx \end{cases} \quad (3.29)$$

I performed this simulation setting $t_f = 1.5 \times 10^5$ years, $\Delta r = 1$ pc, $\Delta t = 4$ years and with a diffusion coefficient as in equation 3.20 with $D_0(100\text{GeV}) = 3.2 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$. In figure 3.10 can be seen the numerical and analytical solutions at a distance of $R = 100$ pc from the source.

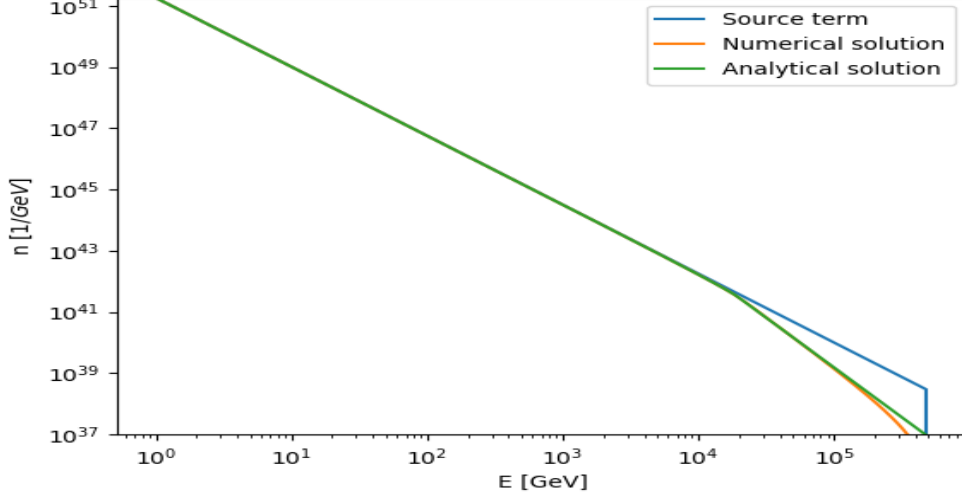


Figure 3.9: Number of particles per unit energy in GeV as a function of particles energy. It can be seen the effect on spectrum of energy losses of the form $b(E) = A_3 E^2$ after a time $t_f = 1.5 \times 10^4$ years. As expected a break in the spectrum occurs at the energy of about 20 TeV since at this energy $A_3 E t_f > 1$.

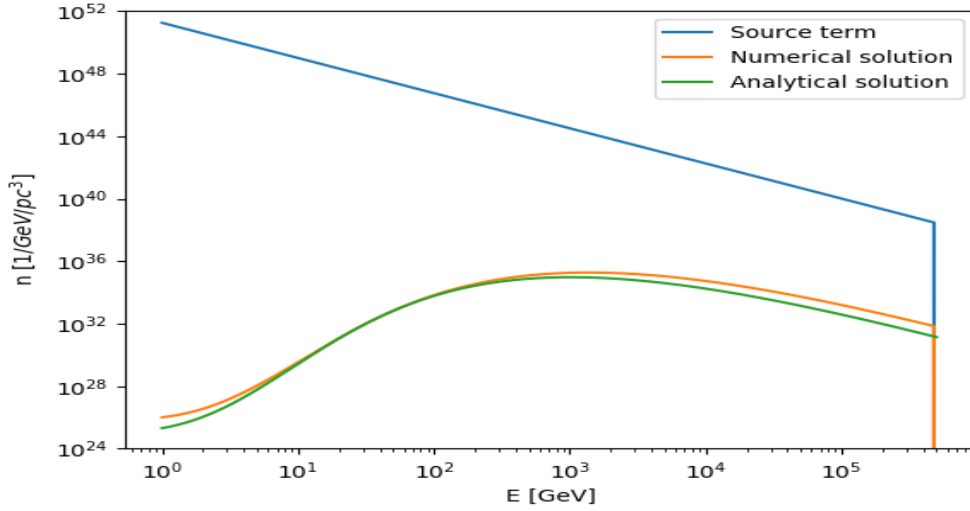


Figure 3.10: Number of particles per unit energy in GeV per unit volume in pc^3 as a function of particles energy. The figure show the effect of diffusion on spectrum at a distance $R = 100$ pc from the source after a time $t_f = 1.5 \times 10^5$ years.

3.6 Predicted positron flux at Earth

Once validated the code to solve the diffusion-loss equation I performed many simulations in order to get the energy distribution function of electrons to calculate the corresponding gamma flux as will be seen in the next sections. Here I present the results I obtained using the HAWC physical parameters in the 1-zone and the 2-zones models framework. If $f(R_{gem}, t_{gem}, E)$ is the electrons and positrons energy function distribution at distance $R_{gem} = 250$ pc and at time $t_{gem} = 3.42 \times 10^5$ years, the total flux is [21]

$$\Phi(E) = \frac{c}{4\pi} f(R_{gem}, t_{gem}, E) \quad (3.30)$$

Since electrons and positrons are considered to be produced in the same proportions, the positron flux at Earth is $\Phi(E)/2$.

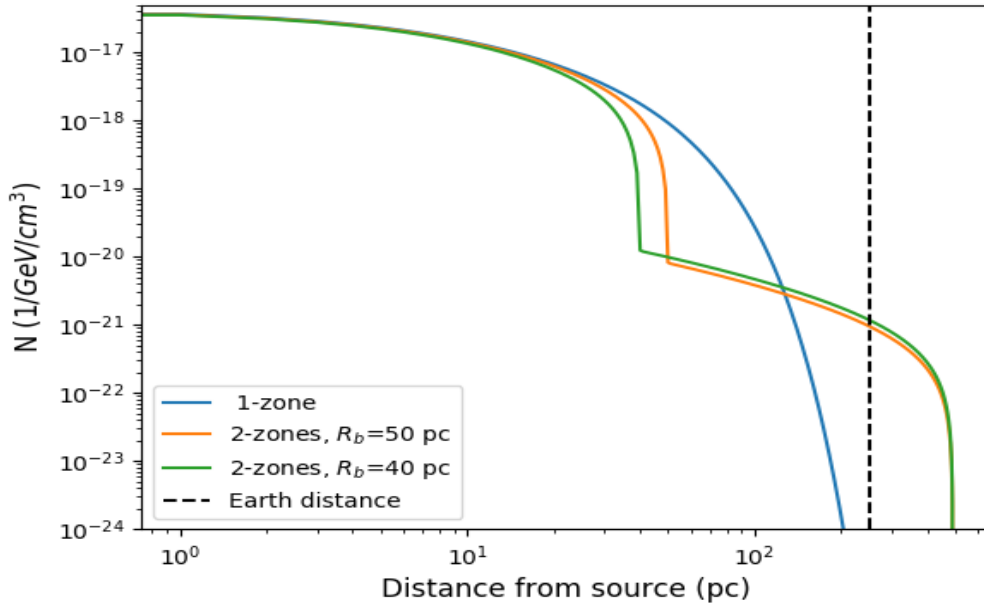


Figure 3.11: Density of positrons in the energy bin corresponding to 1 TeV as a function of the distance from Geminga.

In figure 3.11 I show the radial distribution of 1 TeV positrons. In figure 3.12 can be seen the positron flux at Earth from Geminga compared to the positron flux data from AMS-02. If a 1-zone model is used the positron flux at Earth results too low to produce a contribution to the positron flux measured by AMS-02, as claimed by the HAWC collaboration. Anyway this scenario

is quite unlikely. As explained before it is more likely that particles diffuse slower in a confined region of tens pc from Geminga and then diffuse faster in the interstellar medium. This way the positron flux can be significantly enhanced. In the hypothesis of a breaking radius of 50 pc Geminga could contribute to the positron flux above 100 GeV from about 30% to about 85%. For a breaking radius lower than 50 pc the Geminga contribution can be further enhanced, as can be seen in figure 3.12 for $R_b = 40$ pc.

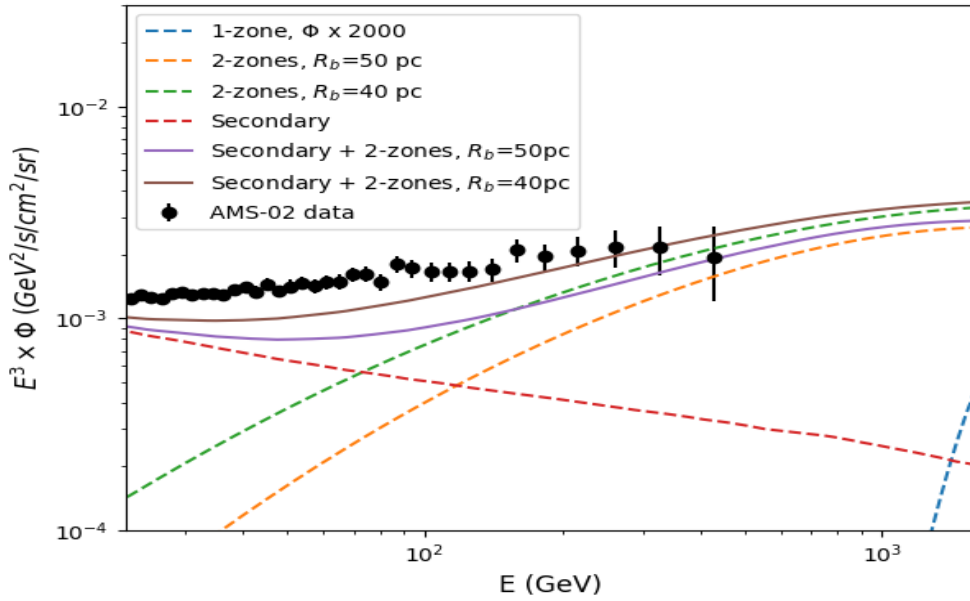


Figure 3.12: Positron flux at Earth from Geminga in the 1-zone and 2-zones models framework, compared to AMS-02 data from [28]. The positron spectrum of secondary origin is taken from [23].

3.7 A gamma-ray sky model for the Geminga TeV halo

The aim of this section is to find a way to calculate the diffuse gamma emission around Geminga due to electrons and positrons diffusion. A way to solve the diffusion-loss equation was found in section 3.4. Since I solved the equation in spherical symmetry, the solution $f(r, t_{gem}, E)$ is a matrix containing the energy distribution function at the radial distance r from the source after

a time equal to the age of Geminga. Once solved the diffusion-loss equation I saved the solution in a fits file. The fits (“Flexible Image Transport System”) format is the standard archival data format for astronomical data sets and images [29].

3.7.1 Spatial and spectral informations

The position of an astrophysical object should be specified in three dimensions: two angular coordinates and the distance from the observer. In this work I decided to use the galactic coordinate system. In this system the galactic equator on the celestial sphere is chosen to be the plane of the Galaxy. The angles that define the position of an object on the celestial sphere are: the galactic longitude l measured in degrees between 0° and 360° eastward from the direction of the galactic center; the galactic latitude b measured in degrees between 0° and $\pm 90^\circ$ from the galactic equator [30].

To compute the gamma-ray emission from the electron and positron populations I used Naima [10], which is an algorithm for computation of non-thermal radiation from relativistic particle populations. In order to use the radiative models one have to give them a particle distribution function as first argument. To do this, particle distribution models can be found in one of the naima’s modules as: power law, power law with exponential cutoff, log parabola, broken power law and exponential cutoff broken power law. Otherwise one can use the TableModel naima class to define a particle distribution function from an array of energies and an array of particles per unit energy. Since I previously resolved the diffusion-loss equation obtaining the particle density distribution function I used the TableModel class.

Since the gamma emission around Geminga is ascribed to inverse Compton scattering of relativistic electrons and positrons I used only the inverse Compton naima model. In addition to the particle distribution function, calling the inverse Compton function one has to declare the photon fields to use for the computation. Naima has three default photon fields: cosmic microwave background, near infrared and far infrared. These should be called with the strings ‘CMB’, ‘NIR’ and ‘FIR’ for which radiation fields with temperatures of 2.72 K, 30 K, and 3000 K, and energy densities of 0.26, 0.5 and 1 eV cm^{-3} will be used. Otherwise, as I’ve done, the photon field parameter can be provided through a list of three items, for an isotropic source: a name for the photon field, its temperature in K and its energy density in eV cm^{-3} units [31].

3.7.2 Map creation and estimation of the surface brightness profile

Here I want to find models' parameters that could explain the surface brightness profile, that is the flux emitted per solid angle unit, observed by HAWC. To do so in addition to the *naima* package I used the *numpy* and *astropy* packages. *Numpy* is a Python package for scientific computing. It is a Python library that provides multidimensional array objects, routines for fast operations on arrays, basic linear algebra, random generators and so on [32]. The *astropy* package contains various classes and tools for astronomy computation.

Before starting the computation I have to define the space around the source. To do so I used the *astropy* class "SkyCoord" to define the position of Geminga in galactic coordinates, with $l = 195.13^\circ$ and $b = 4.26^\circ$. At this point I created a map cube of coordinates centered at the source. Each map has an extension of $20^\circ \times 20^\circ$ in longitude and latitude. The third dimension is the distance from Earth. Since the source is located at 250 pc from Earth and I am interested in the gamma emission within a certain distance from the source, due to computational time reasons I considered only the points between $r_{inf} = 200$ pc and $r_{sup} = 300$ pc. Once defined the space around the source I used the *astropy* method `l.separation_3d(2)`, that computes the three dimensional separation between the point 1 and the point 2. So, at this point, I have a map cube that contains for each point in the sky its distance from the source.

Thanks to the solution of the diffusion-loss equation I know the electrons and positrons populations at a certain distance from the source after a time t_{gem} . At this point I can use the *naima* `TableModel` class to select the particles energy function distribution for each distance and the `InverseCompton` class to compute the photons emissivity corresponding to the particles distribution. What I will create is a series of map in galactic coordinates each of those represent the photon emission at a certain energy. To do so I have first to compute, for each pixel of the map, the contribution to the photons emissivity given by those points that lays within a maximum distance $\rho_{max} = 50$ pc from the source. So in the beginning I started selecting from the map cube of distances the indexes of the points that satisfy the condition

$$r_{point} < \rho_{max} \quad (3.31)$$

Since I know the particles population with a spatial step $\Delta r = 1$ pc, I approximated the distances corresponding to the selected indexes with the nearest integer.

If the input particles distribution (figure 3.13) is in particles per unit energy

per unit volume in pc^3 , the naima's output (figure 3.14) will be a photons emissivity in units of $photons\ eV^{-1}\ s^{-1}\ pc^{-3}$.

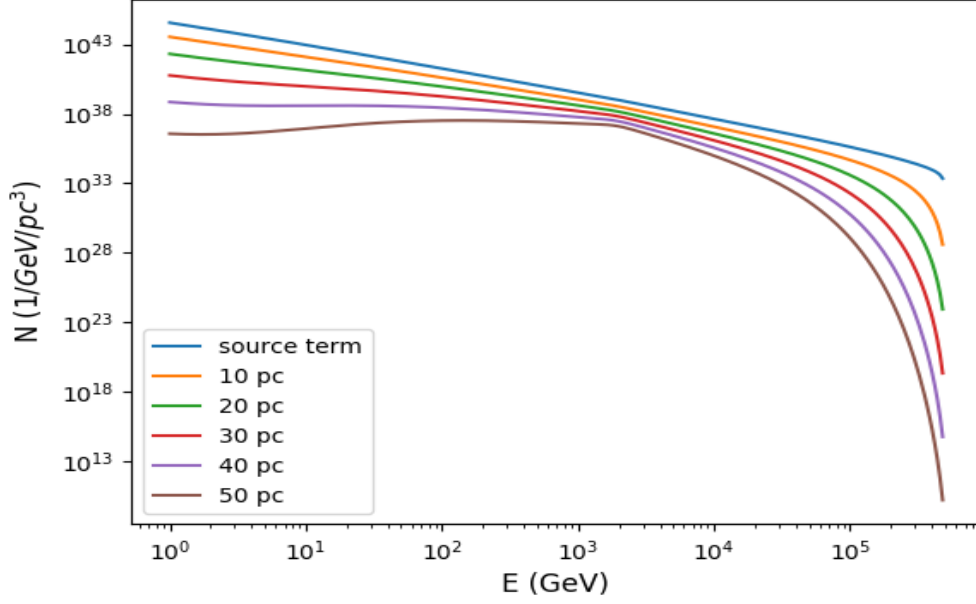


Figure 3.13: Naima's input particles distribution of the source term and at distances of 10 pc, 20 pc, 30 pc, 40 pc and 50 pc, in particles per GeV per unit volume in pc^3 .

For each pixel of the map, the total emissivity is given by the integral along the line of sight as

$$\epsilon_{l.o.s.}(l, b, E) = \int_{r_{inf}}^{r_{sup}} \epsilon(l, b, r, E) dr \quad (3.32)$$

It should be noted that another approach consists in computing the photons emissivity with naima and then integrating along the line of sight only those points that satisfy the condition 3.31. Anyway this approach shouldn't be recommended since, because of naima, the CPU time for the computation could be significantly higher.

If dl and db denote the longitude and latitude step, the solid angle covered by each pixel is

$$\Omega = 4arcsin(sin(\frac{dl}{2})sin(\frac{db}{2})) \simeq dl^2 \quad (3.33)$$

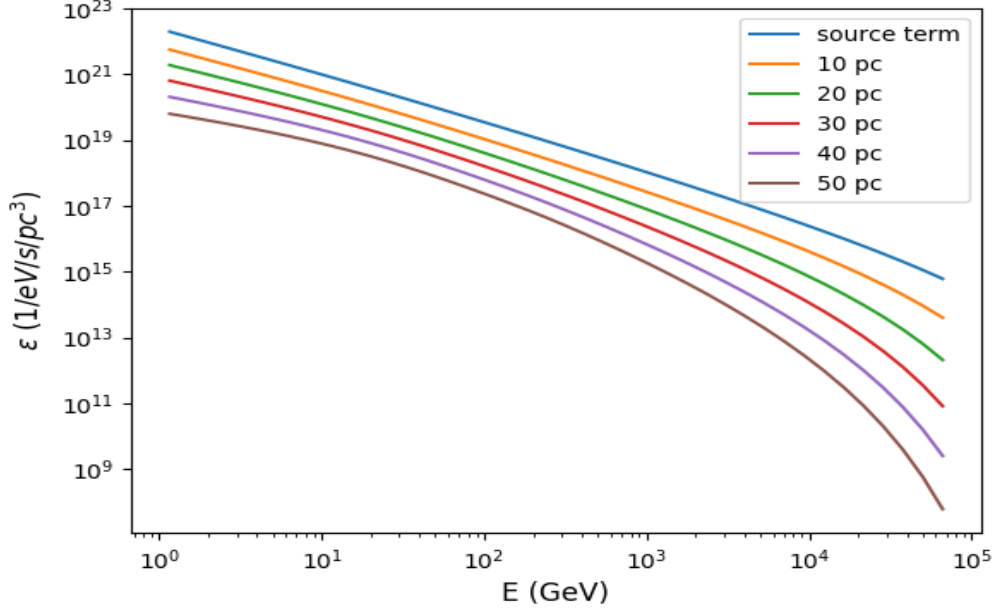


Figure 3.14: Naima's output emissivity of the source term and at distances of 10 pc, 20 pc, 30 pc, 40 pc and 50 pc, in photons per eV per second per unit volume in pc^3 .

if $dl = db$ and in case of small angles. Once integrated the emissivity along the line of sight and knowing the solid angle covered by each pixel, the photons luminosity is given by

$$L(l, b, E) = \frac{\Omega_{sr}}{4\pi} \epsilon_{l.o.s.}(l, b, E) \quad (3.34)$$

with $\Omega_{sr} = \Omega \times (\pi/180)^2$ solid angle in steradian. This is an intrinsic property of the source, which does not depend on the distance of the observer.

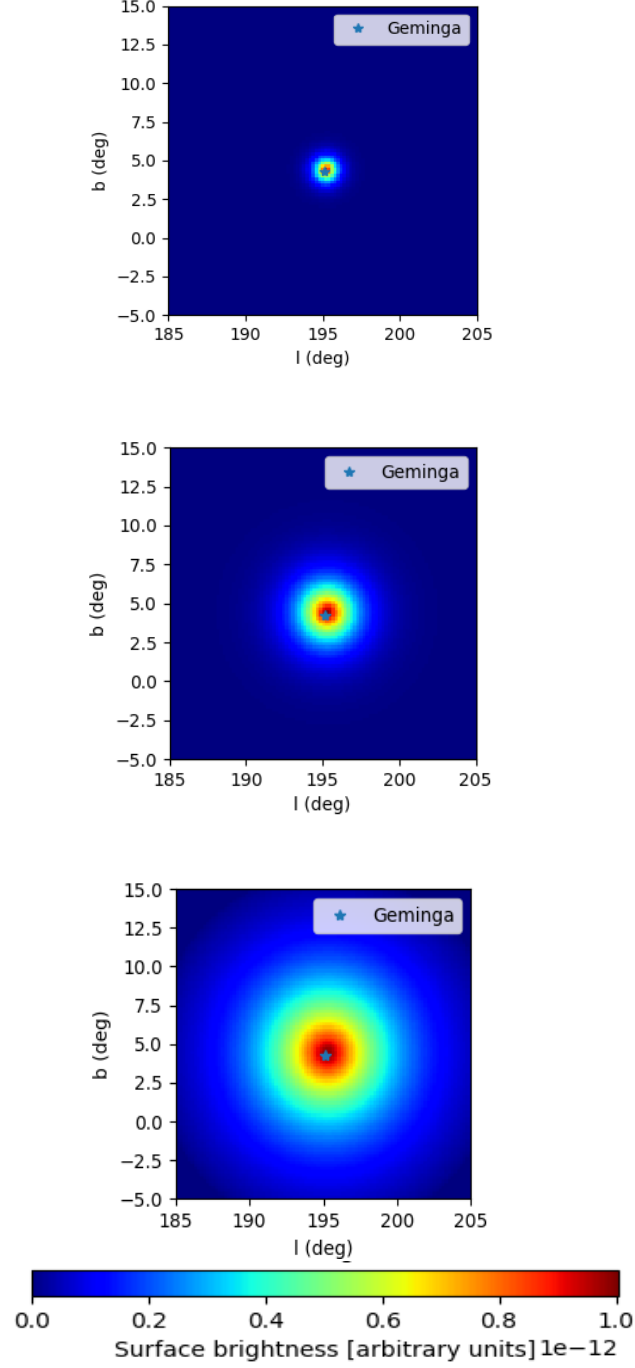


Figure 3.15: Maps of the surface brightness in the energy interval 5-50 TeV around the Geminga region. Maps are presented for three different values of the diffusion coefficient at 100 GeV: $D_{0,1} = 1 \times 10^{25} \text{ cm}^2 \text{ s}^{-1}$, $D_{0,2} = 1 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$, $D_{0,3} = 1 \times 10^{27} \text{ cm}^2 \text{ s}^{-1}$

Doing so I created a map cube that contains for different energy values the map in galactic coordinates of the gamma-ray emission around the source, as can be seen in figure 3.15. At this point, using the map cube I can evaluate

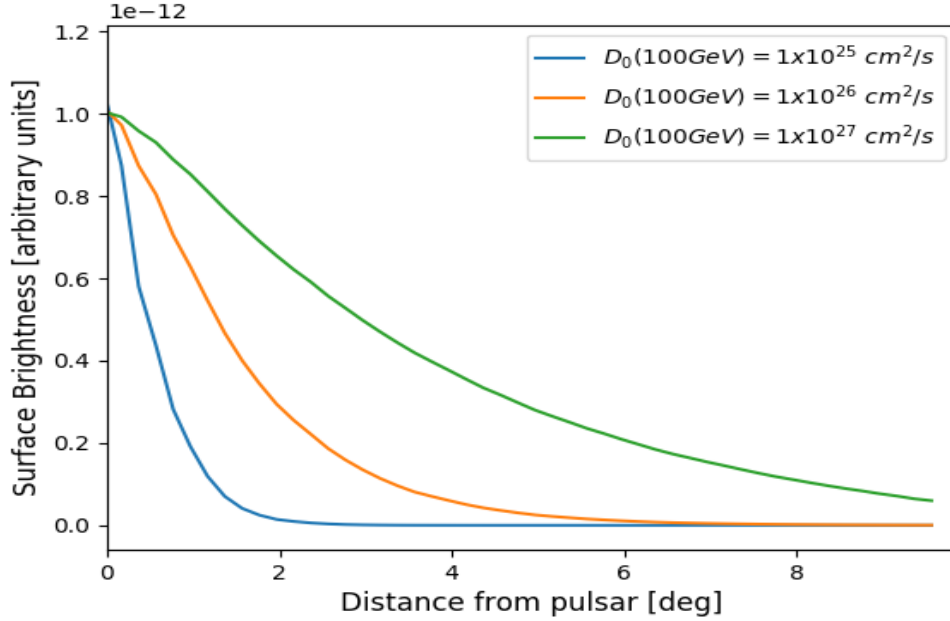


Figure 3.16: Surface brightness profile in the energy interval 5-50 TeV for three different values of the diffusion coefficient at 100 GeV: $D_{0,1} = 1 \times 10^{25} \text{ cm}^2 \text{ s}^{-1}$, $D_{0,2} = 1 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$, $D_{0,3} = 1 \times 10^{27} \text{ cm}^2 \text{ s}^{-1}$.

the surface brightness profile of the source. If $A = dl \times db$ denotes the area of each pixel, the surface brightness between E_{min} and E_{max} is given by

$$SB(l, b, E) = \frac{1}{A} \int_{E_{min}}^{E_{max}} E \times L(l, b, E) dE \quad (3.35)$$

Once centered at the source location, the surface brightness profile can be estimated as the surface brightness of each pixel as a function of the radial distance of the pixel from the central pixel of the map, that defines the source position. In figure 3.16 can be seen an example of the surface brightness profile obtained with this method. In this figure are presented the profiles for three different values of the diffusion coefficient at 100 GeV: $D_{0,1} = 1 \times 10^{25} \text{ cm}^2 \text{ s}^{-1}$, $D_{0,2} = 1 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$, $D_{0,3} = 1 \times 10^{27} \text{ cm}^2 \text{ s}^{-1}$. As can be seen, for small values of the diffusion coefficient the gamma-ray profile is confined in a region close to the source. Quite the opposite, for higher values of the

diffusion coefficient the gamma-ray profile become broader because the high energy particles can travel faster within the cooling time.

3.8 Predicted gamma-rays flux

As pointed out previously a 2-zones model is the one that best describes both the surface brightness of Geminga and the contribution of Geminga to the positron flux at Earth. Anyway as can be seen in figure 3.11 as long as we consider distances from the source within the breaking radius R_b between the two zones, the electrons and positron flux is the same in both models. This is why from now on I decided to use a 1-zone model searching for the parameters that best fit the surface brightness profile with the HAWC data. Here I tested a power-law model for the injection of electrons with different values of the spectral index. Since the particles are supposed to be accelerated with energies in the interval 1 GeV - 500 TeV for a total energy released since the Geminga birth of 1.1×10^{49} , the hardest the spectral index the highest is the number of particles at higher energies. This is why a lower fraction of the total energy released by Geminga into electrons and positrons is required to fit the HAWC data, as the spectral index goes from 2.25 to 1.5. The photon fields involved in inverse Compton scattering used for this computation are the ones in table 3.1. From figure 3.17 to 3.20 are presented the maps of the surface brightness and the surface brightness profile in the energy interval 5-50 TeV (that is the energy interval considered in the HAWC fit), obtained with the method described in the previous section, for the four spectral index values. As can be seen in these figures all the models can fit the data under the assumption of a lower fraction of energy released by Geminga into electrons and positrons (table 3.2).

Spectral index	1.5	1.7	2	2.25
E (erg)	4.23×10^{46}	5.24×10^{46}	1.34×10^{47}	6.87×10^{47}

Table 3.2: Spectral index of the injection spectrum and corresponding energy released in electrons and positrons from the fits of the surface brightness profile with the HAWC data.

The value of the diffusion coefficient from the fit is

$$D(100\text{GeV}) = 1.4 \times 10^{26} \quad \text{cm}^2\text{s}^{-1} \quad (3.36)$$

In order to obtain the spectral energy distribution of the source within a certain distance from the source, for each energy layer of the cube map I

selected only the indexes of the map that satisfy the condition

$$(l_i - l_{Gem})^2 + (b_i - b_{Gem})^2 < \Theta \quad (3.37)$$

with l_i and b_i coordinate of each pixel, l_{Gem} and b_{Gem} galactic coordinate of Geminga and $\Theta = 1^\circ, 2^\circ, 3^\circ, \dots, 10^\circ$ angular distance from the source. This way I was able to select only the pixels that give a contribution to the total flux within the region of interest. Then I integrated the surface brightness to obtain the spectral energy distribution as

$$\Phi(E) = \int \int SB(l, b, E) dl db \times A \quad (3.38)$$

with $A = dl \times db$ pixel area. The total spectral energy distribution, obtained integrating the surface brightness within a radius of 10° from the source, is presented in figure 3.22 for the four spectral indexes tested. The integration over the solid angle of the HAWC model in equation (3.1) gives the spectral energy distribution in figure 3.22.

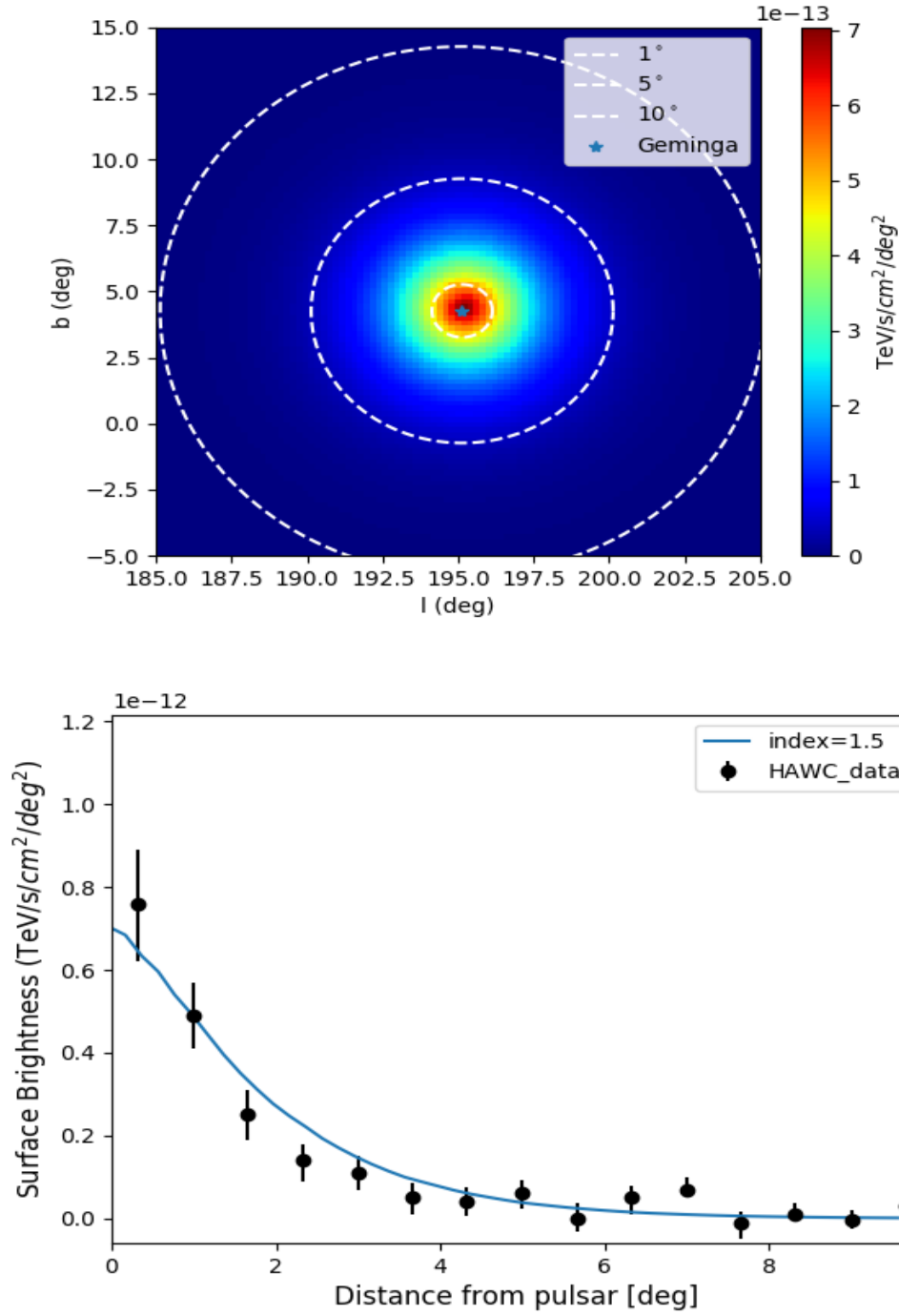


Figure 3.17: Map of the surface brightness around Geminga (up) and surface brightness profile (down) of Geminga for an injection power-law model with spectral index $\gamma = 1.5$, total energy $E = 4.23 \times 10^{46}$ and for a diffusion coefficient at 100 GeV $D_0 = 1.4 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$

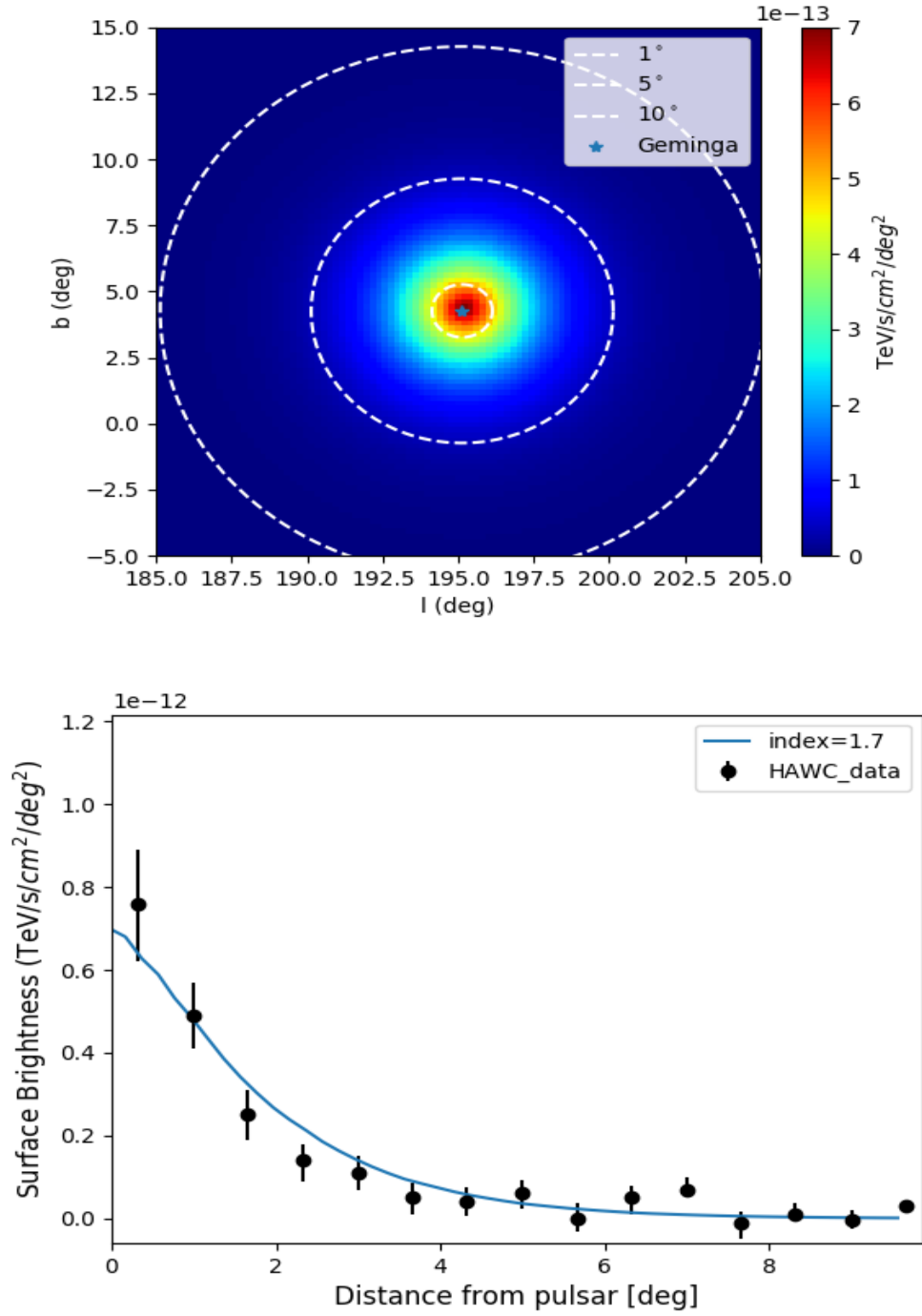


Figure 3.18: Map of the surface brightness around Geminga (up) and surface brightness profile (down) of Geminga for an injection power-law model with spectral index $\gamma = 1.7$, total energy $E = 5.24 \times 10^{46}$ and for a diffusion coefficient at 100 GeV $D_0 = 1.4 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$

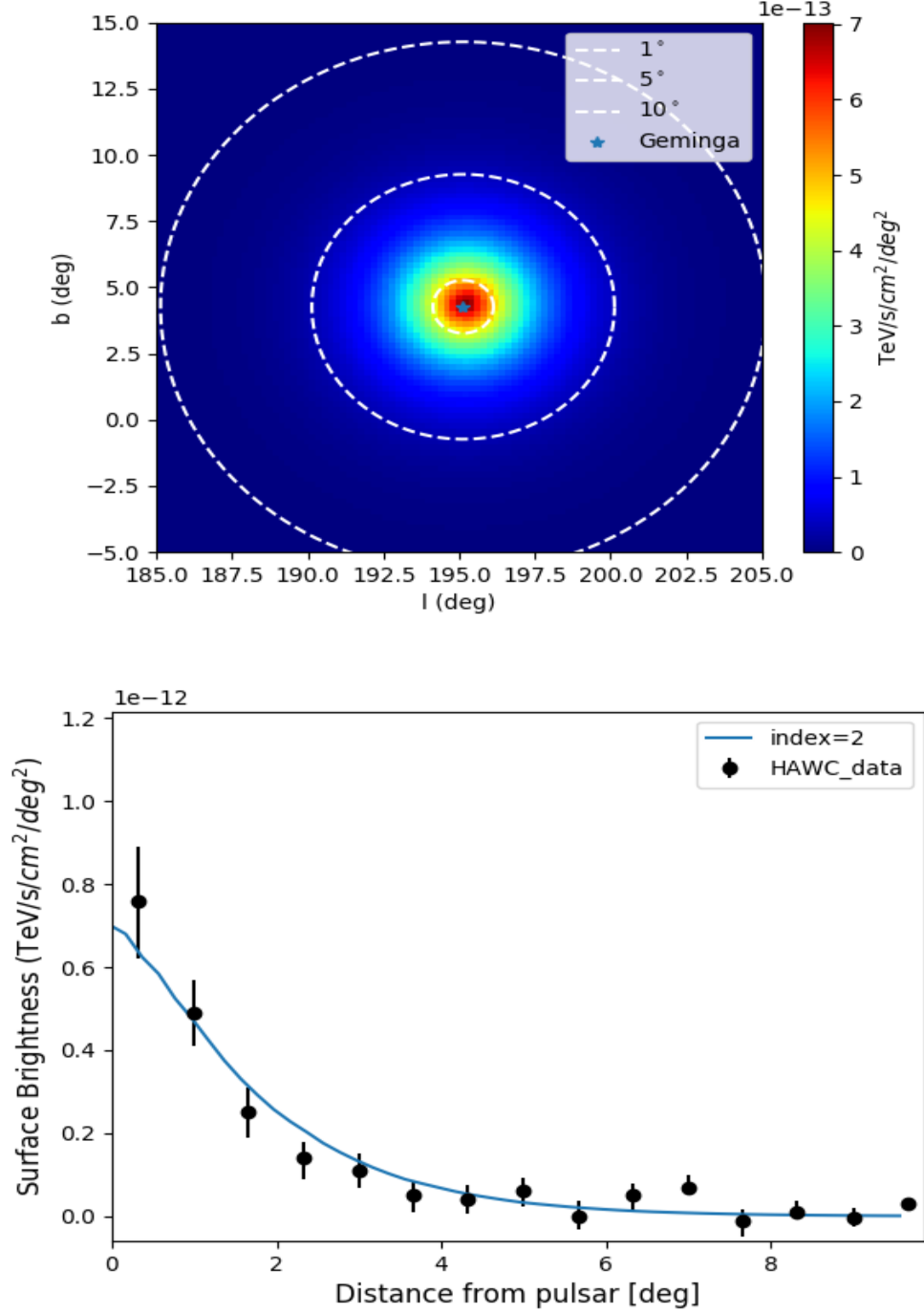


Figure 3.19: Map of the surface brightness around Geminga (up) and surface brightness profile (down) of Geminga for an injection power-law model with spectral index $\gamma = 2$, total energy $E = 1.34 \times 10^{47}$ and for a diffusion coefficient at 100 GeV $D_0 = 1.4 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$

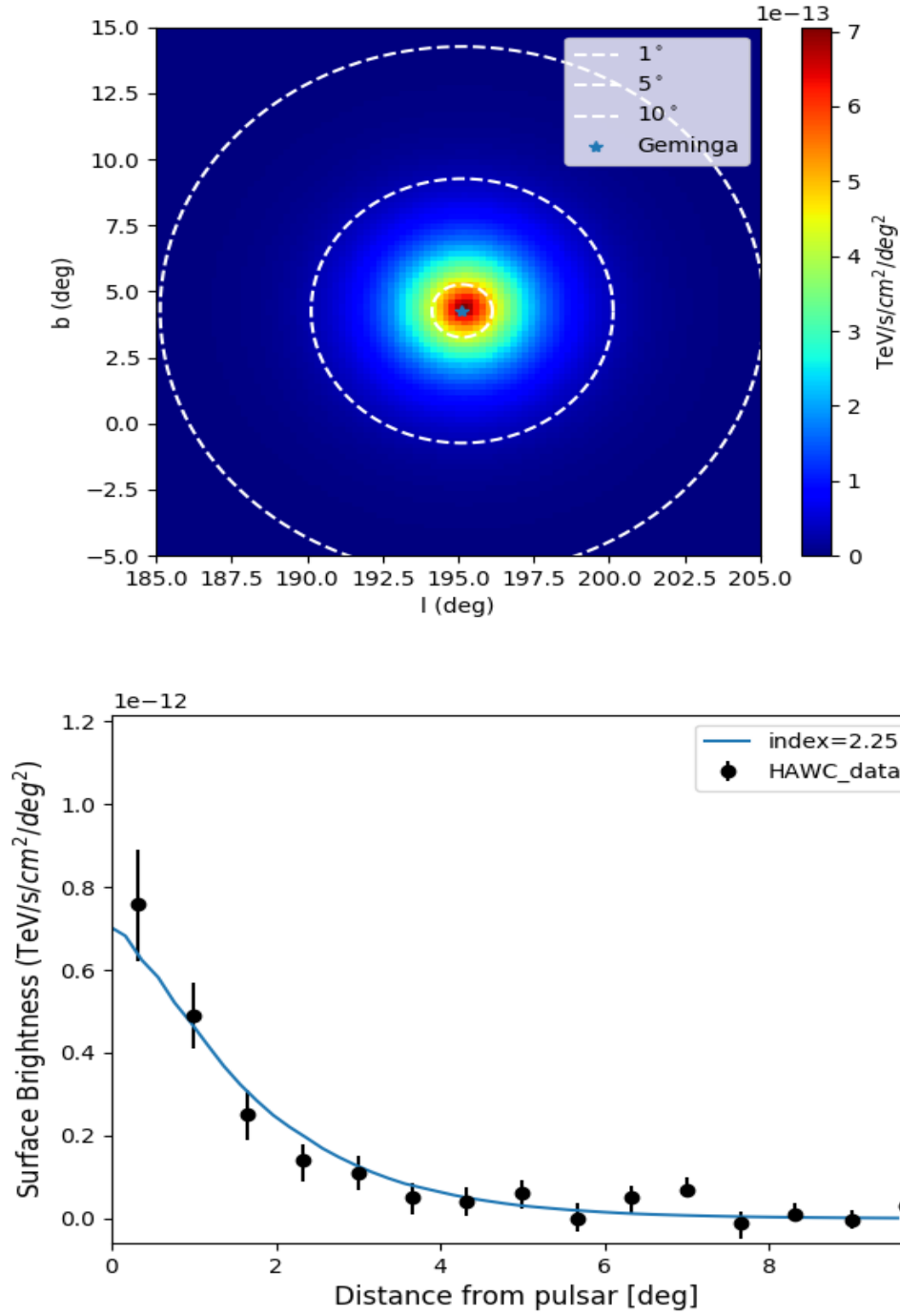


Figure 3.20: Map of the surface brightness around Geminga (up) and surface brightness profile (down) of Geminga for an injection power-law model with spectral index $\gamma = 2.25$, total energy $E = 6.87 \times 10^{47}$ and for a diffusion coefficient at 100 GeV $D_0 = 1.4 \times 10^{26} \text{ cm}^2 \text{ s}^{-1}$

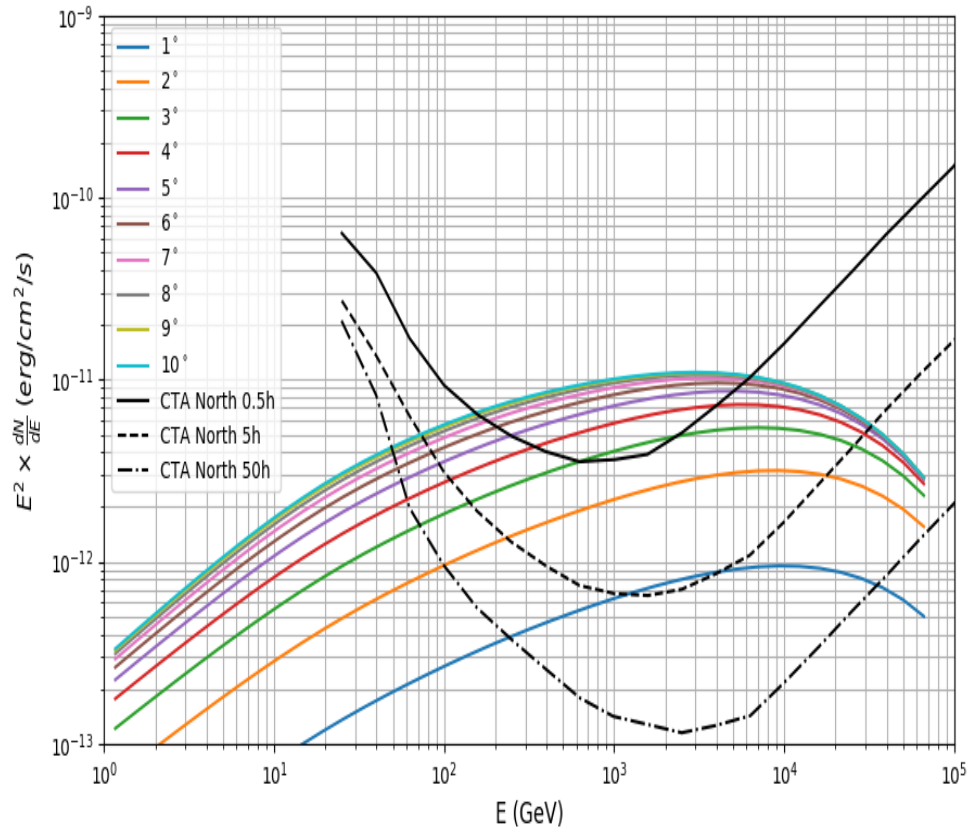


Figure 3.21: Inverse Compton spectral energy distribution as function of photon energy for different integration radii and sensitivities of the CTA north array for 0.5 h, 5 h and 50 h.

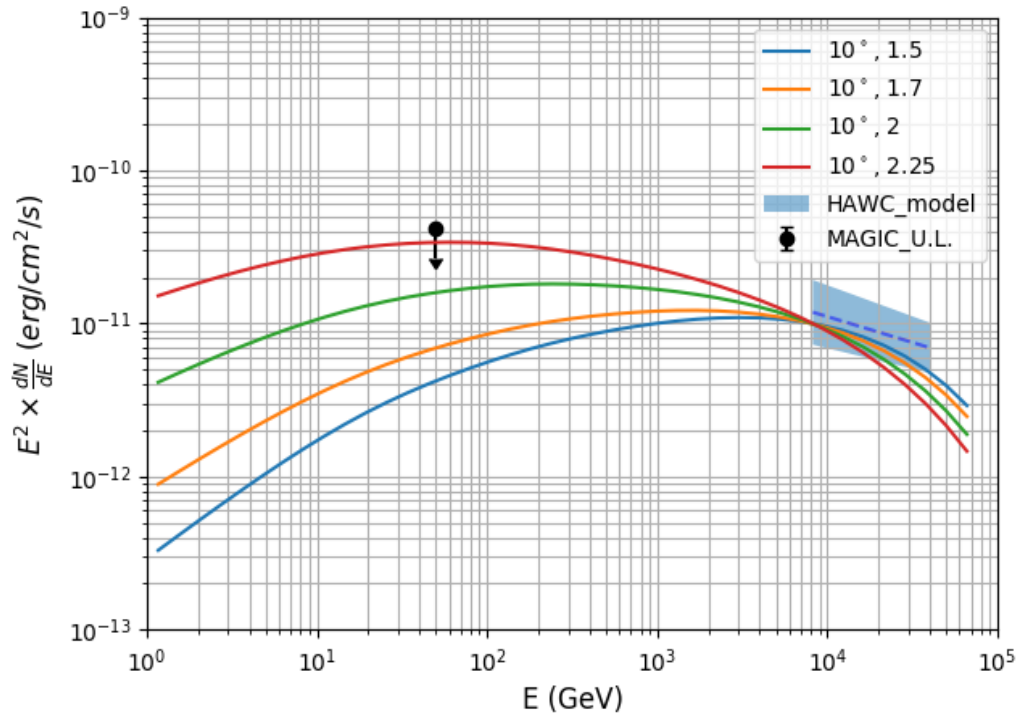


Figure 3.22: Inverse Compton spectral energy distribution as a function of photon energy, obtained with integration radius of 10° for the spectral indexes tested. The HAWC spectral energy distribution is given by integration over solid angle of equation (3.1). The MAGIC upper limit is taken from [27].

3.9 Conclusions

In this chapter I introduced two of the subjects of very high energy gamma ray astronomy: pulsars and their nebulae. Then I explained the Geminga science case arose a few years ago when the HAWC collaboration published the article about the Geminga TeV halo. In order to explain the HAWC data, in a scenario where the relativistic leptons emitted by the pulsar produce gamma ray through inverse Compton scattering with ambient photon fields, I first solved the diffusion-loss equation. I validated the code with analytical solution for pure diffusion and pure energy-loss cases. I then performed simulation using HAWC parameters in a 1-zone model and a 2-zone model with the common diffusion coefficient of $D_{ISM}(100\text{GeV}) = 1.8 \times 10^{29} \text{ cm}^2 \text{ s}^{-1}$ in interstellar medium. In the 2-zones scenario the positron flux can be significantly enhanced. Even if the only Geminga cannot account for the positron flux excess above 10 GeV measured by AMS-02, it can give an important contribution to the positron flux above 100 GeV ranging from about 30% to about 85%. If a lower breaking radius is assumed, the positron flux can be further enhanced. As can be seen in figure 3.21 the contribution to the total flux between 5 and 50 TeV above a distance of 7° , that corresponds to about 30 pc, is negligible. Thus, in order to reproduce the HAWC surface brightness profile the breaking radius shouldn't be lower than 30 pc.

At this point I used the 1-zone model to compute the gamma-ray emission around Geminga. Once created the map of the surface brightness between 5 and 50 TeV, I obtained the surface brightness profile for four different values of the spectral index. The HAWC data can be reproduced by all the spectral index values, if a lower fraction of the total energy released by Geminga from its birth for the hardest ones is taken into account. Then I evaluated the spectral energy distribution within a radius compatible with the extended region observed by HAWC. Future CTA observation in the multi-GeV and TeV region of the spectrum will permit to constrain the spectral index of the injection spectrum of electrons and positrons.

CHAPTER

4

SIMULATIONS AND ANALYSIS OF CTA OBSERVATIONS

In the previous chapter I derived a theoretical model for the electrons and positrons population in the interstellar space around the Geminga pulsar and their gamma-ray emissivity. In this chapter I show the work done in order to simulate and analyze an observation of CTA toward this source. To account for the instruments response two ways are possible: Monte Carlo simulation of the shower development and the instrument response or events calculation from sky model and simplified model of the instruments response. A series of tool based on this second choice have been developed for the CTA data analysis.

4.1 Instrument Response Function (IRF)

If $I(\mathbf{p})$ is the gamma ray intensity arriving at Earth as a function of the photon properties $\mathbf{p}$, usually true photon energy, true photon incident direction and true photon arrival time,

$$e(\mathbf{d}) = \int (R \times I(\mathbf{p})) d\mathbf{p} \quad (4.1)$$

is the expected event rate observed from the instrument with $\mathbf{d}$ event properties like measured photon energy, measured or reconstructed photon incident direction and measured photon arrival time. Here $R = R(\mathbf{d}|\mathbf{p})$ represents the instrumental response function (IRF) that describes the transformation from the physical properties of a photon to the measured characteristics of an event.

The instrument response function can be factorized as [34]

$$R(p', E', t' | p, E, t) = A_{eff}(p, E, t) \times PSF(p' | p, E, t) \times E_{disp}(E' | p, E, t) \quad (4.2)$$

where the primed symbols denote reconstructed or measured quantities (direction, energy and arrival time) and unprimed symbols true quantities. $A_{eff}(p, E, t)$ is the effective area in cm^2 , $PSF(p' | p, E, t)$ and $E_{disp}(E' | p, E, t)$ are the point spread function and the energy dispersion that satisfy

$$\begin{cases} \int PSF(p' | p, E, t) dp' = 1 \\ \int E_{disp}(E' | p, E, t) dE' = 1 \end{cases} \quad (4.3)$$

The effective detection area of a Cherenkov telescope is roughly the area of the Cherenkov light pool, thus of about $45000 m^2$ since the radius of the Cherenkov ring is of about 120 m. Since it depends on the light pool dimensions, it is a function of the energy of the incoming photon and of the arrival direction respect to the telescope axis. The dependence of A_{eff} on energy and offset angle can be seen in figure 4.1. The telescope has the largest effective area at the higher energies and for on-axis observations.

The point spread function represents the size of the image of a point source. It can be described in two variants [35], both depending on energy and off-axis angle: a superposition of three 2-D Gaussian functions each characterised by a width and a relative amplitude; or by

$$PSF(p' | p, E, t) = \frac{1}{2\pi\sigma^2} \left(1 - \frac{1}{\gamma}\right) \left(1 + \frac{1}{2\gamma} \frac{\delta^2}{\sigma^2}\right)^{-\gamma} \quad (4.4)$$

with δ angular separation between the true and measure photon directions, σ describes the width and γ the tail of the distribution.

The energy dispersion take account for the fact that the reconstructed event energy will differ from the true photon energy [34].

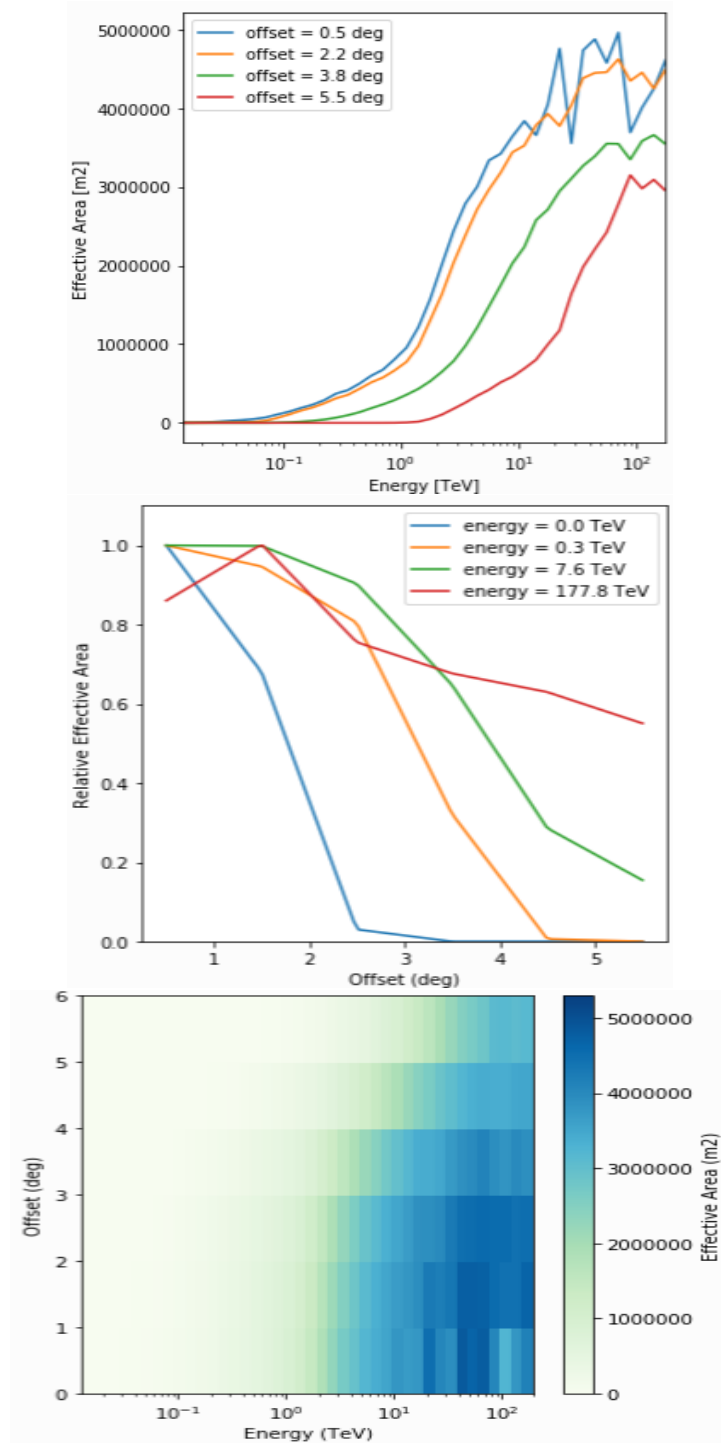


Figure 4.1: **Top:** Effective area as a function of energy for different offset angles. **Middle:** Effective area as a function of offset for different energies. **Bottom:** Effective area in the offset-energy space. [33]

4.2 CTA data analysis

4.2.1 Maximum Likelihood and detection significance

Methods based on the search of the best model that adapts to a set of data are called “best fit” methods. If $\mathbf{x} = (x_1, \dots, x_n)$ are the independent observed values of a random variable X , each with pdf (probability density function) $p(x_i; \boldsymbol{\theta})$ depending from an p -dimensional parameter $\boldsymbol{\theta}$, the function

$$L(\boldsymbol{\theta}; \mathbf{x}) = \prod_{i=1}^n p(x_i; \boldsymbol{\theta}) \quad (4.5)$$

is called “likelihood function”. This function represents the probability of obtaining the observed values $(x_1, \dots, x_n)$ for a set of parameters $(\theta_1, \dots, \theta_p)$. Since the properties of the likelihood function are the same of its logarithm, usually is used the log-likelihood function defined as

$$\mathcal{L} = -\ln(L(\boldsymbol{\theta}; \mathbf{x})) = -\sum_{i=1}^n \ln(p(x_i; \boldsymbol{\theta})) \quad (4.6)$$

because of the minus sign the search of a maximum of the likelihood corresponds to the search of a minimum of the log-likelihood.

The maximum likelihood method consists in the search of the parameter values that maximize the likelihood function: if $\mathbf{x} = (x_1, \dots, x_n)$ are the realizations of a random variable $\mathbf{X} = (X_1, \dots, X_n)$, with pdf $p(\mathbf{x}; \boldsymbol{\theta})$ depending from the parameter $\boldsymbol{\theta}$

$$L(\hat{\boldsymbol{\theta}}; \mathbf{x}) = \max(L(\boldsymbol{\theta}; \mathbf{x})) \quad (4.7)$$

where $\hat{\boldsymbol{\theta}}$ is the maximum likelihood estimates, if it exists, of the $\boldsymbol{\theta}$ parameter. Alternatively the same thing can be done minimizing the log-likelihood, solving the likelihood equations with respect to θ_k for $k=1,2,\dots,p$

$$\frac{\partial \mathcal{L}}{\partial \theta_k} = \sum_{i=1}^n \frac{1}{p(x_i; \boldsymbol{\theta})} \frac{\partial p(x_i; \boldsymbol{\theta})}{\partial \theta_k} = 0 \quad (4.8)$$

The maximum likelihood method can be used only if the pdf is a priori known and it is derivable. Under these assumptions this method is largely used because it gives estimators $\hat{\boldsymbol{\theta}}$ that are:

- Consistent, which means that the estimator of the parameter θ converges in probability to the parameter.

- Asymptotically unbiased, which means that $\langle T \rangle = \theta + b_n$ but the bias b_n goes to zero as the sample size n goes to infinity.
- Asymptotically normal
- Between the estimators of a parameter the maximum likelihood estimator is the more efficient, that is the one with minimum variance.

These and more informations about maximum likelihood techniques can be found on [36].

4.2.2 Hypothesis test and Test Statistic

Hypothesis tests can be used in order to decide if a starting model describing an event X is compatible with the realization x of that event. Let consider a pdf $p(x, \theta)$ of a random variable X depending on the parameter θ . If the starting model, hereafter null model, assumes the parameter to be $\theta = \theta_0$, to decide if the realization x is compatible with the null model the variability interval of X is divided into two regions: one favorable to the null hypothesis, one against it that is called “critical region”. Defining an a priori test level α , the quantile that defines the boundaries of the acceptance and critical regions are $x_{1-\alpha}$ in case of a one-tailed test and $x_{\alpha/2}$ and $x_{1-\alpha/2}$ in case of a two-tailed test. The observed significance level α_0 , also called p-value, is defined as $P(X \leq x)$ in a one-tailed test or $P(|X - \mu| > |x - \mu|)$ in a two-tailed test, with μ mean of the distribution. The null hypothesis is rejected if $\alpha_0 < \alpha$, otherwise it is accepted with a maximum probability α of being wrong.

Sometimes two or more model have to be tested respect to a set of data. Considering only two models, H_0 represents the null hypothesis and H_1 represents the alternative one. If $p(x|H_0)$ and $p(x|H_1)$ are the pdf corresponding to the two models

$$P(X < c|H_0) = \int_{-\infty}^c p(x|H_0)dx = \alpha \quad (4.9)$$

represents the probability of being wrong rejecting H_0 when it is true, called type I error, whereas

$$P(X > c|H_1) = \int_c^{\infty} p(x|H_1)dx = \beta \quad (4.10)$$

represents the probability of being wrong accepting H_0 when it is false, also called type II error.

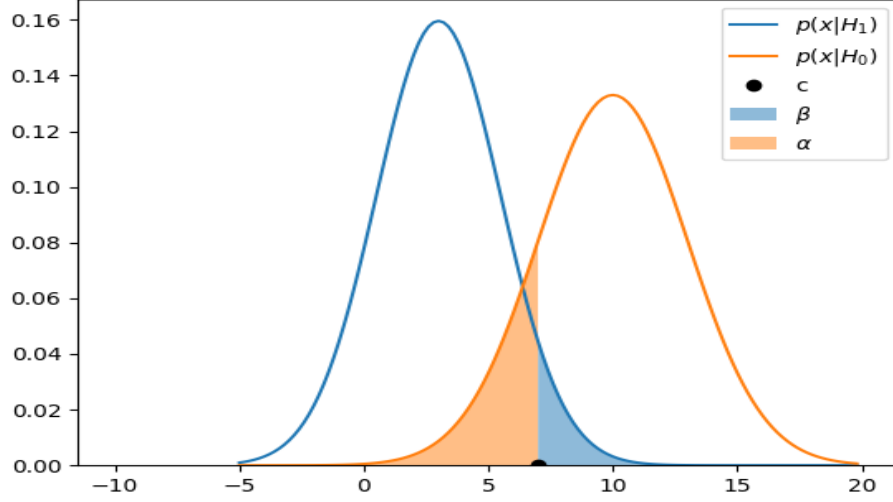


Figure 4.2: Schematic representation of the quantities in a two hypothesis test: the null model (orange line), the alternative model (blue line), the probability of being wrong rejecting the null model when it is true (orange shaded region) and the probability of being wrong accepting the null model when it is false (blue shaded region).

When one has to test two models with a set of data a good choice is to consider the likelihood ratio

$$\lambda = \frac{L(\boldsymbol{\theta}_0; \mathbf{x})}{L(\boldsymbol{\theta}_1; \mathbf{x})} \quad (4.11)$$

with $0 < \lambda < 1$, $\boldsymbol{\theta}_0$ parameter of the null model and $\boldsymbol{\theta}_1$ parameter of the alternative model. If the likelihood ratio λ is near 1 the null hypothesis is true, if it is near 0 the alternative hypothesis is true. Equivalently it is useful to consider the statistic

$$TS = -2\ln(\lambda) \quad (4.12)$$

as the basis of a statistical test. Now low values of TS imply that the null hypothesis is true, higher values imply an increasing incompatibility between data and the null model. Now the problem is to find the distribution of TS and thus the associated confidence levels. The answer is given by the Wilks' theorem which states that $-2\ln(\lambda)$ will have a probability distribution like χ^2 with q degrees of freedom, except for an additive term of order $1/\sqrt{n}$ with n sample size, q being the difference between the number of changed parameters in the null and alternative models [37].

4.2.3 CTA science tools

The ctools package is a software tools for the analysis of IACT event data. It is developed as science tools software for CTA, but can be used to perform analysis of data from existing IACT's (H.E.S.S., MAGIC, VERITAS) once that the data and the response functions are converted into the proper format. The ctools package is based on GammaLib, a versatile toolbox for the high-level analysis of astronomical gamma-ray data. The tools can be combined in various way to perform the needed analysis, since each tool perform a single defined analysis step. Moreover they are all available as Python classes through the call of the “ctools” Python module [35]. The tools I used for my work are ctobssim and ctlike.

Ctobssim simulates photon events from astrophysical source and background events that must be specified in an input model. The input model is defined as an XML file, where the spatial extension and the spectrum of the source can be specified. In addition to the input model must be specified as input parameters the pointing direction, the field of view, the duration of the observation, an energy interval and an instrumental response function (irf). As output ctobssim creates a fits file containing a list of photon events, their energy and arrival directions.

Ctlike performs a maximum likelihood fitting of an input model to event data simulated with ctobssim. All the models parameter in the XML input file that have the attribute “free=1” will be treated as free parameters and adjusted to the values that maximizes the probability that the observed data came from the input model [34]. Ctlike uses the Poisson statistics for likelihood computation and, if requested, it computes the Test statistics value defined as

$$TS = 2\ln(L(M)) - 2\ln(L(M_{-j})) \quad (4.13)$$

with $L(M)$ maximum likelihood value for the model M, $L(M_{-j})$ maximum likelihood value for the model M without the model component j. If the model M provides a satisfactory fit of the data, TS follows a χ_p^2 distribution with p degrees of freedom, that is the number of model parameters in the model component j. For p=1 the source significance is the square root of the TS [34]. As output ctlike generates an XML file with the best fit of the model parameters set as free, an error attribute providing their statistical error and the value of the test statistics [35].

4.3 The Geminga TeV halo with CTA

The spectral and spatial model in the XML file can be specified using the models that are available in ctools. Otherwise these models can be specified through a fits file, if they have been computed in other ways. Since up until now I have resolved the diffusion-loss equation and found a way to create maps of the sky around Geminga for different energies, I used the second way.

To create the map-cube I used Gammapy, an open-source Python package for gamma-ray astronomy [33]. Since I already have a map-cube containing the maps of the sky for different energies, all I have to do is to create the file. Of course, one more time, the ideal file type is a fits file. I created an empty map with the same pixel sizes, width, source location and energy bins through the Gammapy tools. Than I filled it with the photon flux computed in the previous chapter.

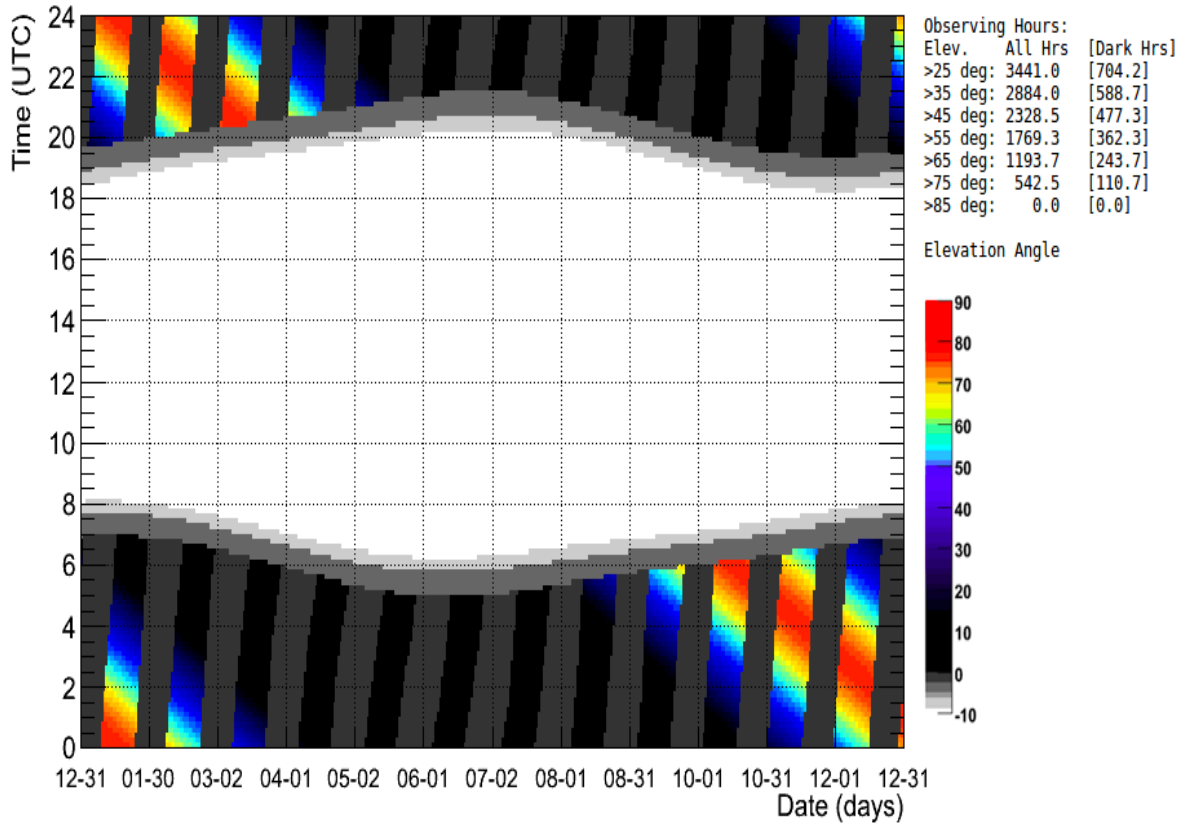


Figure 4.3: Observation time over the year 2026, observing hours and dark hours for Geminga observations [38].

One must pay attention to the physical units of the source intensity and energy, in order to make the map-cube compatible with ctools. Energy must be in MeV, while source intensity in *photons* $\text{MeV}^{-1} \text{ s}^{-1} \text{ cm}^{-2} \text{ sr}^{-1}$.

I performed simulation with ctobssim for 1, 3, 5, 10 and 15 hours, considering a region of interest of 10° around the source location in the energy interval between 50 GeV and 100 TeV. Since all the spectral model I've tested can reproduce the surface brightness profile observed by HAWC, I used the model with spectral index $\gamma = 1.5$ because is the one that best fit both the surface brightness profile and the HAWC spectral energy distribution.

To perform the simulation I have to specify the instrument response function. I decided to use the IRF of the CTA north array, since Geminga is observable from the north hemisphere, optimized for 50h and a zenith angle of 20° . Since for zenith angles higher than about 40° the shower pass through an increasing amount of atmosphere, elevation angles higher than 50° are needed. As can be seen in figure 4.3, above 55° about 360 dark hours will be available in 2026 for Geminga observations, one year after the end of the CTA array construction. In figure 4.4 are presented the density of events from Geminga and background for 1h, 3h, 5h, 10h and 15h of observation. As can be seen the hadronic background is about 2 order of magnitude higher than the signal from the source. Then, for each of the observations, I used ctlike to perform the test statistic and thus the significance of the source. The test statistic is computed as in equation (4.13), with the log-likelihood value obtained for a fit that includes the source (alternative model) and a fit that excludes the source (null model, only background). The significance versus observation time is plotted in figure 4.5. As can be seen, for an observation time higher than about 2 hours a detection of the Geminga TeV halo could be claimed.

It should be noted that different spectral model could lead to a different minimum observation time to claim for a detection.

Even if a few hours could be sufficient to claim for a source detection, CTA will need more hours in order to perform a detailed study of the source spectrum. In figure 4.6 I show the input model used to perform simulations throughout this chapter with the spectral points obtained with csspec, a script that computes the source spectrum by fitting a model in a given set of spectral bins, for an observation time of 15h.

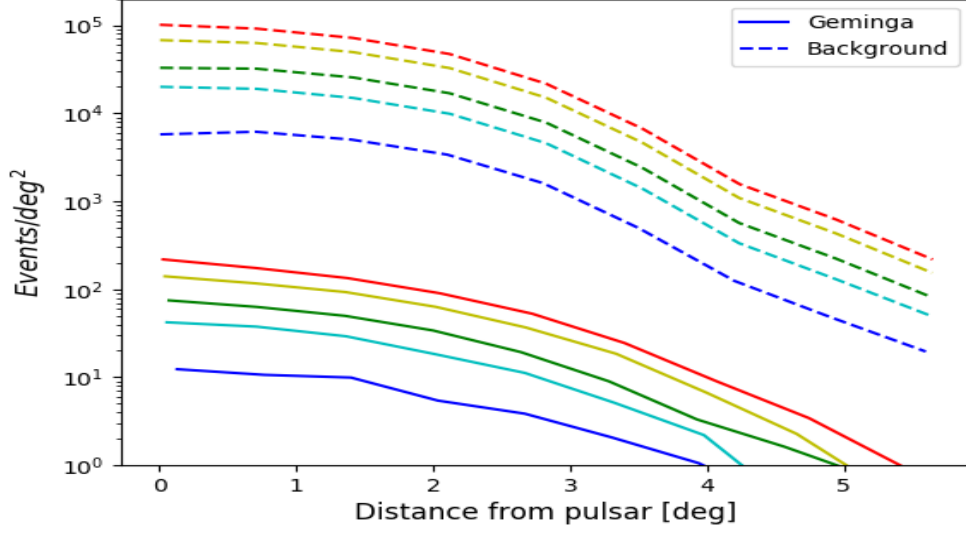


Figure 4.4: Events density as a function of angular distance from Geminga (solid lines) and background (dashed lines) for 1h (blue), 3h (cyan), 5h (green), 10h (yellow) and 15h (red) of observation time.

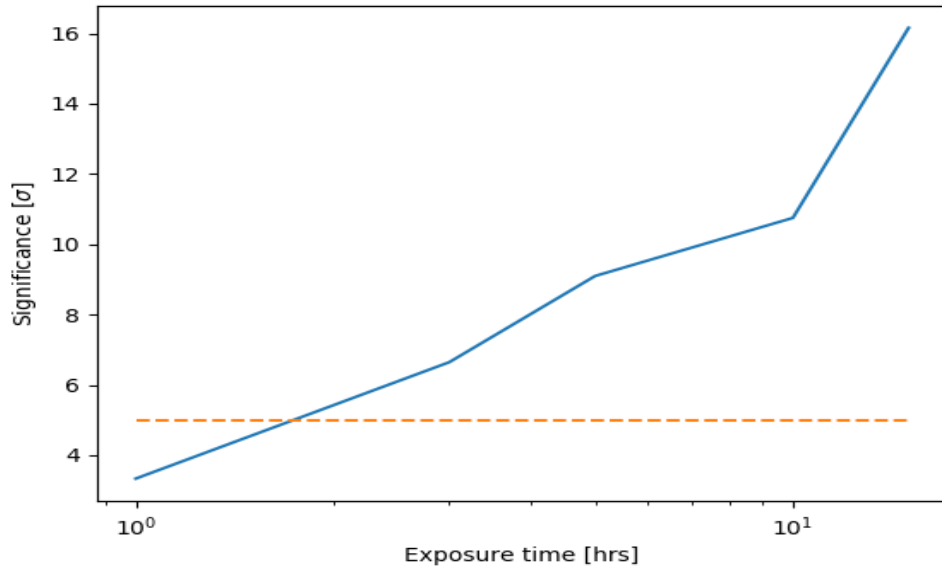


Figure 4.5: Significance of the source as a function of the observation time.

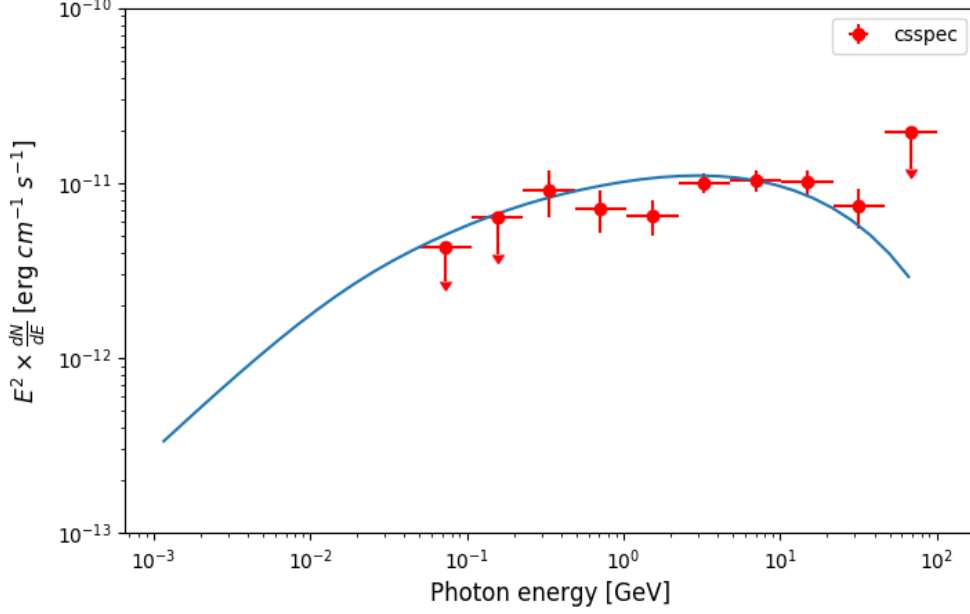


Figure 4.6: Spectrum of the Geminga TeV halo as observed by CTA with an exposure of 15 hours. The blue curve is the spectrum as obtained from the theoretical model used for the simulations.

4.4 Conclusions

In this chapter I introduced some statistical and instrumental concepts on which the CTA science tools are based. Once created a map cube compatible with the ctools I performed simulation for different observation times using the instrument response function of the CTA north array. Doing so I analyzed the signal from the source and the background and performed a likelihood analysis in order to obtain the minimum observation time for which the Geminga TeV halo will be detected by CTA. Using the spectral model that best fit both the surface brightness profile and the spectral energy distribution observed by HAWC, in about 2 hours CTA could claim the detection of the TeV halo.

Taking into account all these consideration Geminga will be an ideal target for future CTA observation. Analysis on this and other pulsars will be crucial not only to understand the physics of PWNe but also to shed light on the problem of the positron excess observed at Earth.

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