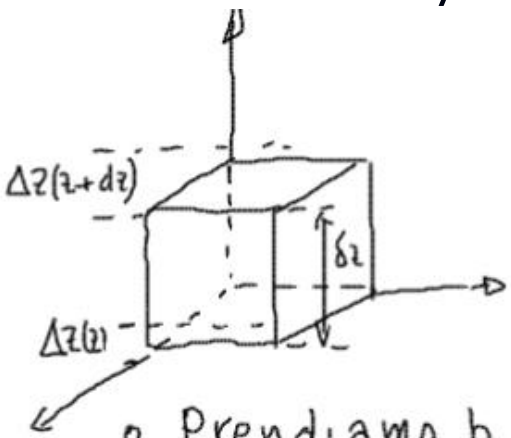
Lecture 11 131119

- Il pdf delle lezioni puo' essere scaricato da
- http://www.fisgeo.unipg.it/~fiandrin/didattica_fisica/cosmic_rays1920/

$dV/dt = V_r \bullet v$



- In un tempo dt, il volume si deforma a causa del moto di espansione

- Prendiamo p. es. il moto lungo z
  Lo spost. dipende dalla velocità locale

In dt si ha $\Delta z(z) = v(z)\,dt$

$$\Delta z(z+\delta z) = v(z+\delta z)\,dt$$

$\Rightarrow$ La variazione complessiva e'

$$\Delta z_t = \Delta z(z+\delta z) - \Delta z(z) = [v(z+\delta z) - v(z)]\,dt \Rightarrow$$

$$\Delta z_t = \frac{\partial v}{\partial z}\delta z\,dt \Rightarrow \frac{d\Delta z}{dt} = \frac{\partial v}{\partial z}\,dz$$

$$\Rightarrow \left(\frac{dV}{dt}\right)_z = \frac{\partial}{\partial z}\,\delta x\,\delta y\,\delta z = \frac{\partial v}{\partial z}\cdot\delta V$$

$$\Rightarrow \frac{dV}{dt} = \sum_{i=1}^{3}\left(\frac{dV}{dt}\right)_{x_i} = (\vec{\nabla}\cdot\vec{v})\,V$$

- Quindi $\frac{dE}{dt} = -\frac{z}{3}(\vec{\nabla}\cdot\vec{v})E$

- Esempio: sfera in espansione uniforme

$v_r = v_0\left(\frac{r}{R}\right)$ con $v_0 = v(R)$

$\vec{\nabla}\cdot\vec{v} = \frac{1}{r^2}\frac{\partial}{\partial r}\left(\frac{v_0 r^3}{R}\right) = 3\left(\frac{v_0}{R}\right) \Rightarrow -\frac{dE}{dt} = z\left(\frac{v_0}{R}\right)E$

$\Rightarrow -\frac{dE}{dt} = z\left(\frac{1}{R}\frac{dR}{dt}\right)E$

- Le perdite adiabatiche sono importanti quando il tempo scala dell'espansione $T_{esp} = \left(\frac{1}{R}\frac{dR}{dt}\right)^{-1}$ è dell'ordine del tempo in cui le part. sono rimaste nel volume T_{conf}

- Importante in:
  - fuga di particelle da SNR
  - quando le part. esercitano pressione, le perdite adiab sono importanti nell'ISM in generale

-Modulazione solare dei RC

Tempi scala di perdita (1)

o Il tempo scala per dimezzare l'energia e'

$$\tau_{1/2} = -\int_{E_i}^{E_f} \frac{dE}{(dE/dt)} \qquad E_f = E_i/2$$

- Ionizz. $\frac{dE}{dt} = 7.64 \times 10^{-15} N \, (3 \ln \gamma + 19.8) \; eV \, s^{-1}$

- Trascurando la dipendenza logaritmica dall'energia, una buona approx e'

$$\frac{dE}{dt} \approx 10^{-8} N \quad \text{eV/anno}$$

Quindi $\tau = -\int_{E}^{E/2} \frac{dE}{10^{-5} N} = \frac{E\,(eV)}{2 \cdot 10^{-8} N}$ anni

Per $E \sim 1$ GeV $\quad \tau \approx \frac{5 \times 10^{16}}{N}$ anni

Con $N \sim 10^{6}$ pm^{-3} $\quad \tau \sim 5 \times 10^{10}$ anni

Tempi scala di perdita

$1\text{x}10^{23}\text{N s}^{-1}$ in caso no screening

- Brems (H neutro) $\frac{1}{E}\frac{dE}{dt} = 3.66\times10^{-22}\,N\ s^{-1}$

$$\tau = -\int_{E}^{E/2}\frac{dE}{3.66\cdot10^{-22}EN} = \frac{\ln 2}{3.66\cdot10^{-22}N}\ s = \frac{1.9\times10^{21}}{N}\ s$$

$$\simeq \frac{6\times10^{14}}{N}\ \text{anni, indipendente da } E$$

- $N \approx 10^{6}\ \text{pm}^{-3}$ ➔ $\tau \approx 6\text{x}\ 10^{8}$ anni Screening completo

$1\text{x}10^{9}$ anni no screening

Tempi scala di perdita (3): IC+Sinc

$$< (\frac{dE}{dt}) >= \frac{4}{3}\sigma_T \; c \; U_B \; \beta^2 \; \gamma^2 \qquad < (\frac{dE}{dt})_{tot} >= \frac{4}{3} c \sigma_T U_{rad} \gamma^2 \beta^2$$

- $\frac{dE}{dt} = \frac{4}{3}\sigma_T c \, (U_R + U_B)\gamma^2 \qquad \tau = \int \frac{dE}{\frac{4}{3}\sigma_T c \, (U_R + U_B)\gamma^2}$

$E = mc^2(\gamma - 1) \Rightarrow dE = mc^2 \gamma \Rightarrow \tau = -\int_{\gamma}^{\gamma/2} \frac{mc^2}{\frac{4}{3}\sigma_T c} \cdot \frac{1}{(U_R + U_B)} \cdot \frac{d\gamma}{\gamma^2}$

Perciò $\tau = \frac{mc^2}{\frac{4}{3}\sigma_T c} \cdot \frac{1}{(U_R + U_B)} \cdot \frac{1}{\gamma}$

- $\frac{4}{3}\sigma_T c = 2.66 \times 10^{-20} \; m^3 s^{-1} \qquad mc^2 = 5.11 \times 10^5 \; eV$

$\tau = \frac{1.92 \times 10^{+25}}{(U_R + U_B)\gamma} \; s$

$U_B \approx 2.2 \times 10^5 \; eV m^{-3}$ @ $B = 3 \times 10^{-10} \; T$

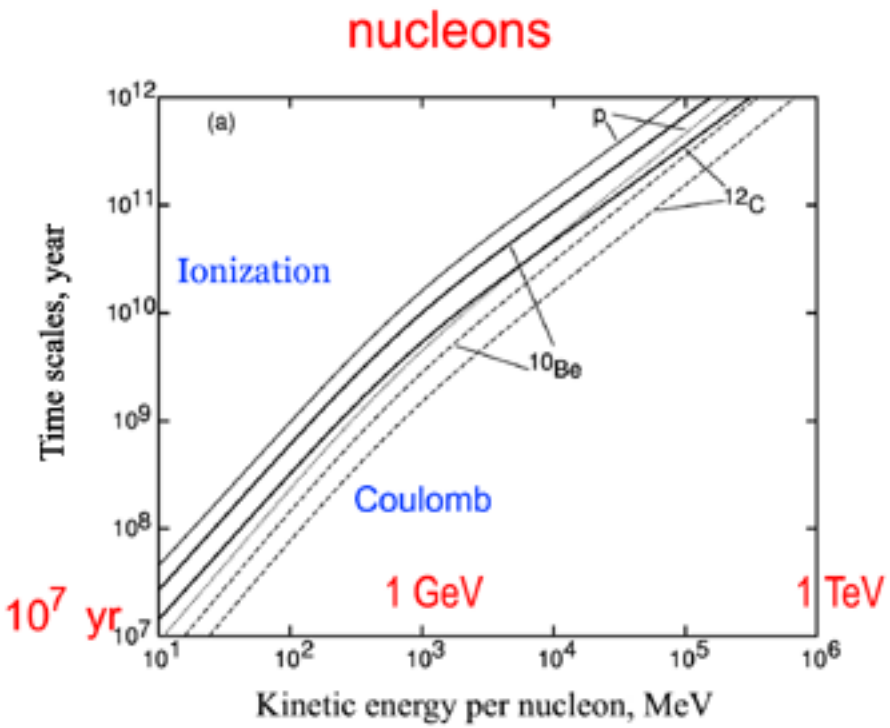
$U_{CMB} \approx 2.62 \times 10^5 \; eV m^{-3}$

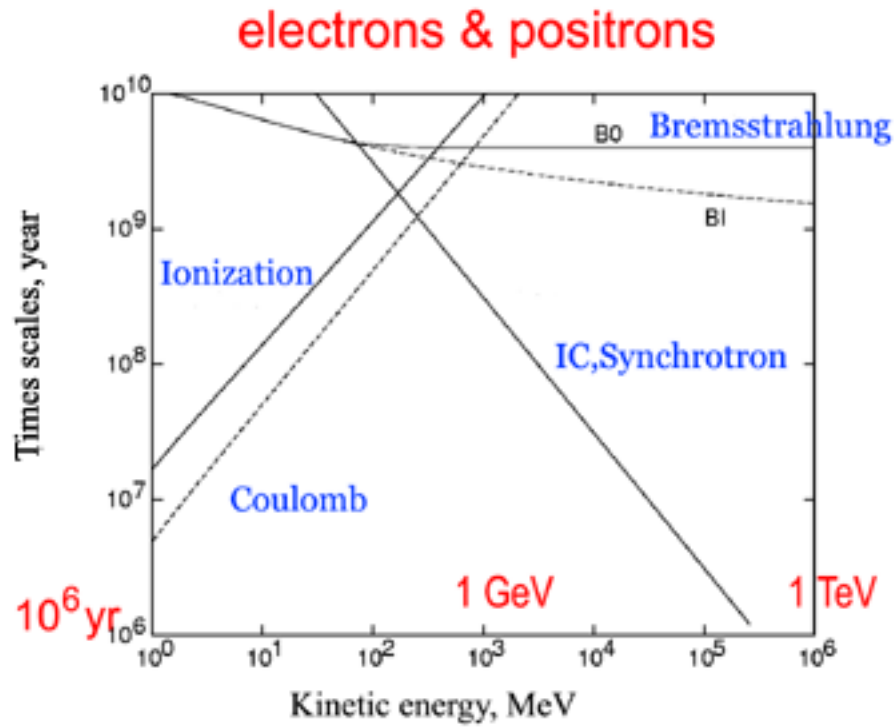
$U_{starr} \approx 6 \times 10^5 \; eV m^{-3}$

× CMB presente ovunque $\Rightarrow$ Agisce sempre

$\tau \leq \tau_{CMB} = \frac{7.328 \times 10^{19}}{\gamma} \; s = \frac{2.32 \times 10^{12}}{\gamma}$ anni $\left(= \frac{4.54 \times 10^9}{E \; (GeV)} \text{ anni} \right)$

- Per e^- a $E = 100 \; GeV$, $\tau \lesssim 10^7$ anni





Differenti regimi di perdite, dipendenti da E

Il tempo scala piu' breve domina

IC e' un lim superiore: nella galassia non c'e' solo CMB ma anche starlight (ottica, UV, X,...) e campi magnetici → $U \approx 10^6$ eVcm^{-3}

- Nel caso di diffusione, le particelle si spostano in media di $\lambda \approx (2D\tau)^{1/2}$
- In generale $D = D(E) \Rightarrow \lambda(E, E_0) = \left(\int_0^{\tau(E)} D(E')\, d\tau' \right)^{1/2} = \left(\int_{E_0}^{E} \frac{D(E')\, dE'}{b(E')} \right)^{1/2}$

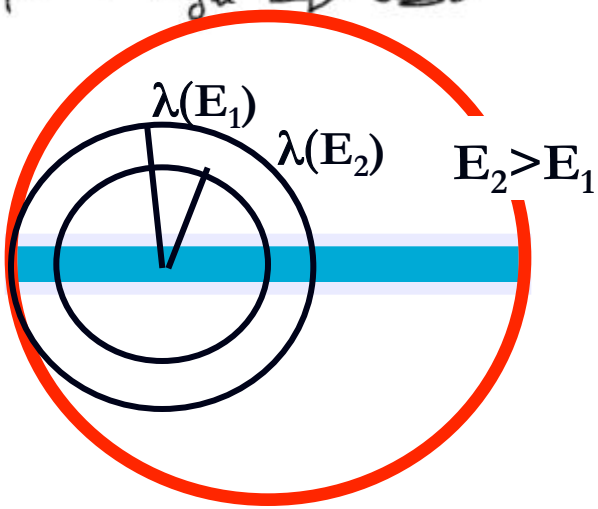
× $D \approx 10^{29}\ cm^2 s^{-1}$ (val. medio galattico) $\Rightarrow \lambda \approx 10$ kpc @ 1 GeV
(solo IC + Sincr. i.e. $\tau = \frac{1}{kE_0}$)

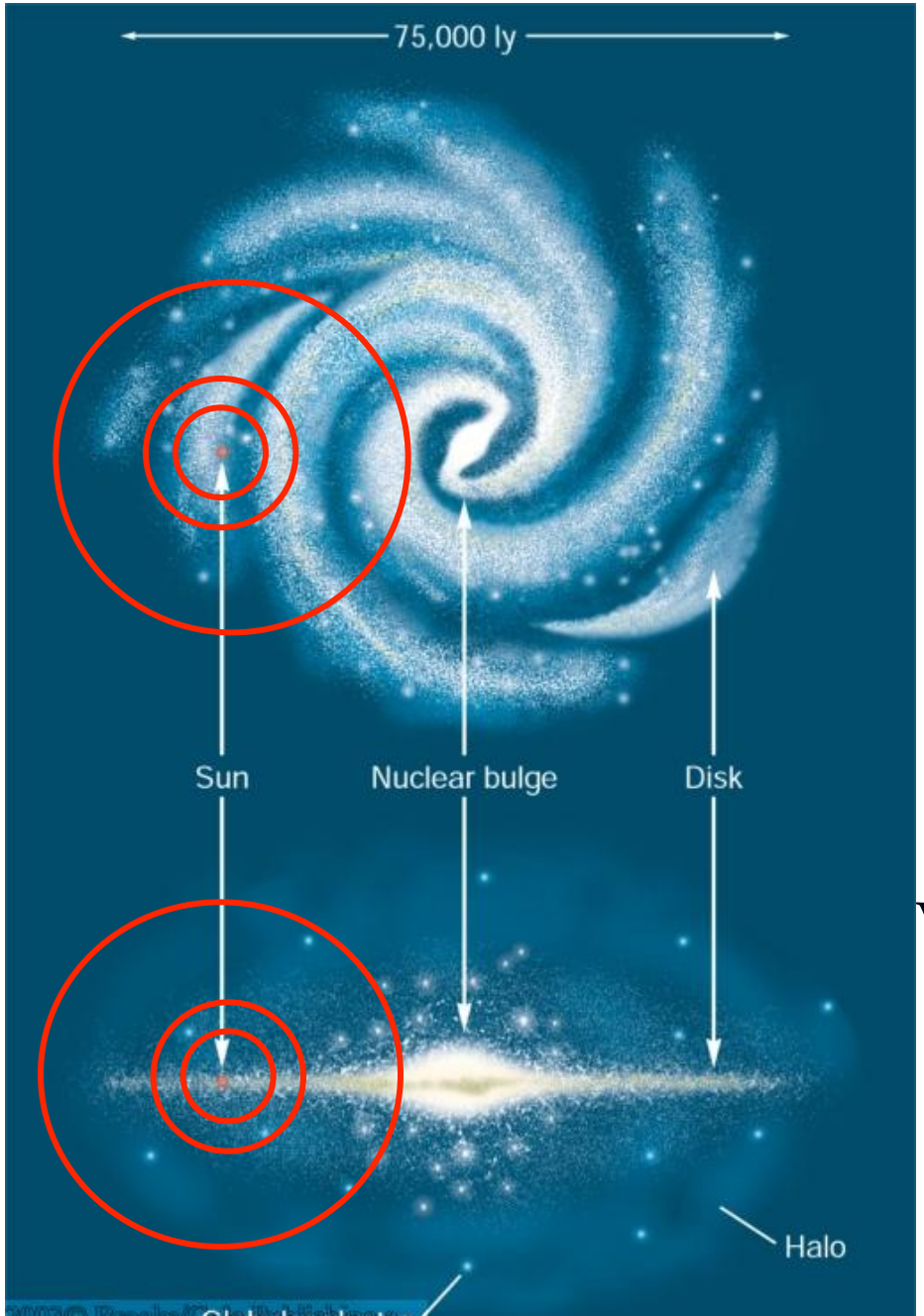
× Trascurando la dipend. di D da E $\lambda \propto E^{-1/2} \Rightarrow$
Lo spettro degli $e^{\pm}$ nei differenti intervalli di energia dipende alle condizioni medie di propagazione nell' ISM in una regione di raggio $R \sim \lambda(E)$

× Esempio: per e^- con $E > 1$ GeV, $\lambda(E) \lesssim 10$ kpc $= R_{gal} \Rightarrow$ ~~sono~~

$\Rightarrow$ Gli e^- sono galattici

- Quindi lo spettro osservato cambia parecchio durante la propagazione





$\approx 10^2$ ly @ 1 TeV

$\approx 3.16 \times 10^3$ ly @ 100 GeV

$\approx 10^4$ ly @ 10 GeV

$\approx 3 \times 10^4$ ly @ 1 GeV

Gli elettroni vengono da vicino!

Ancora propagazione

$$\frac{\partial N_i}{\partial t} - \vec{\nabla}\cdot(\hat{D}_i \vec{\nabla} N_i) + \frac{\partial}{\partial E}(b_i N_i) + n v \sigma_i N_i + \cancel{\frac{N_i}{\tau_i}} = q_i + \cancel{\sum_{j<i} n v \sigma_{ij} N_j + \sum_j \frac{N_j}{\tau_{ij}}}$$

σ_i e' la sez d'urto per il processo $p_{CR}p_{ISM} \rightarrow \pi^{\pm} \rightarrow \mu^{\pm} \rightarrow e^{\pm}$ ed n e' la densita' di p nell ISM, N_i e' la densita' di elettroni

Da dove vengono gli elettroni?

- Da interazione dei RC con i protoni del mezzo interstellare
- Da sorgenti primarie, cioe' siti di accelerazione
- Da sorgenti esotiche(?)

Per gli elettroni, a differenza della componente nucleare, sono importanti i processi di perdita di energia durante la propagazione.Il termine che descrive le perdite di energia continue non puo' essere trascurato ma anzi diventa dominante

Toy Model

- Caso semplice di: – soluz. stazionarie $dN/dt = 0$

× C'e' una distrib infinita e uniforme di sorgenti che iniettano e^- con spettro di iniezione $Q(E) = k\,E^{-p}$

⟹ La diffusione non e' importante ⟹ L'equ. diventa

$$-\frac{d}{dE}\left[b(E)\,N(E)\right] = Q(E) \quad \Longrightarrow \quad \int d\left[b(E)\,N(E)\right] = -\int Q(E)\,dE$$

× Assumiamo $N(E) \to 0$ per $E \to \infty$ ⟹

$$N(E) = \frac{kE^{-(p-1)}}{(p-1)\,b(E)} \quad {}^{1)} \qquad \text{dove} \quad b(E) = A_1\left(\ln\frac{E}{mc^2} + 19.8\right) + A_2 E + A_3 E^2$$

⟹ Se domina ionizz. $N(E) \propto E^{-(p-1)}$. spettro di E più piatto

Se " Brems $N(E) \propto E^{-p}$ (o adiab. loss). lo spettro non cambia

Se IC o sincr dominano $N(E) \propto E^{-(p+1)}$ spettro più ripido

Spettro a bassa E

Ad E<10 GeV, lo spettro locale di e- e' distorto dal vento solare (densita' soppressa e decel adiabatica)→ il flusso misurato sulla Terra non e' rappresentativo del flussso nell ISM

Deducibile da misure indirette, come lo spettro radio nell' alone della galassia dovuto ad emissione di sincrotrone nel campo magnetico galattico

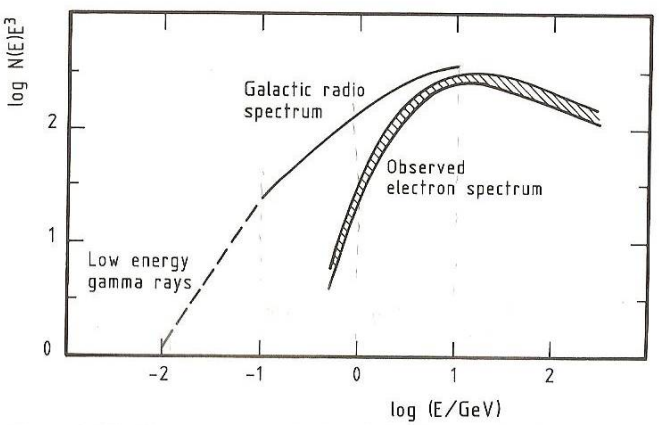
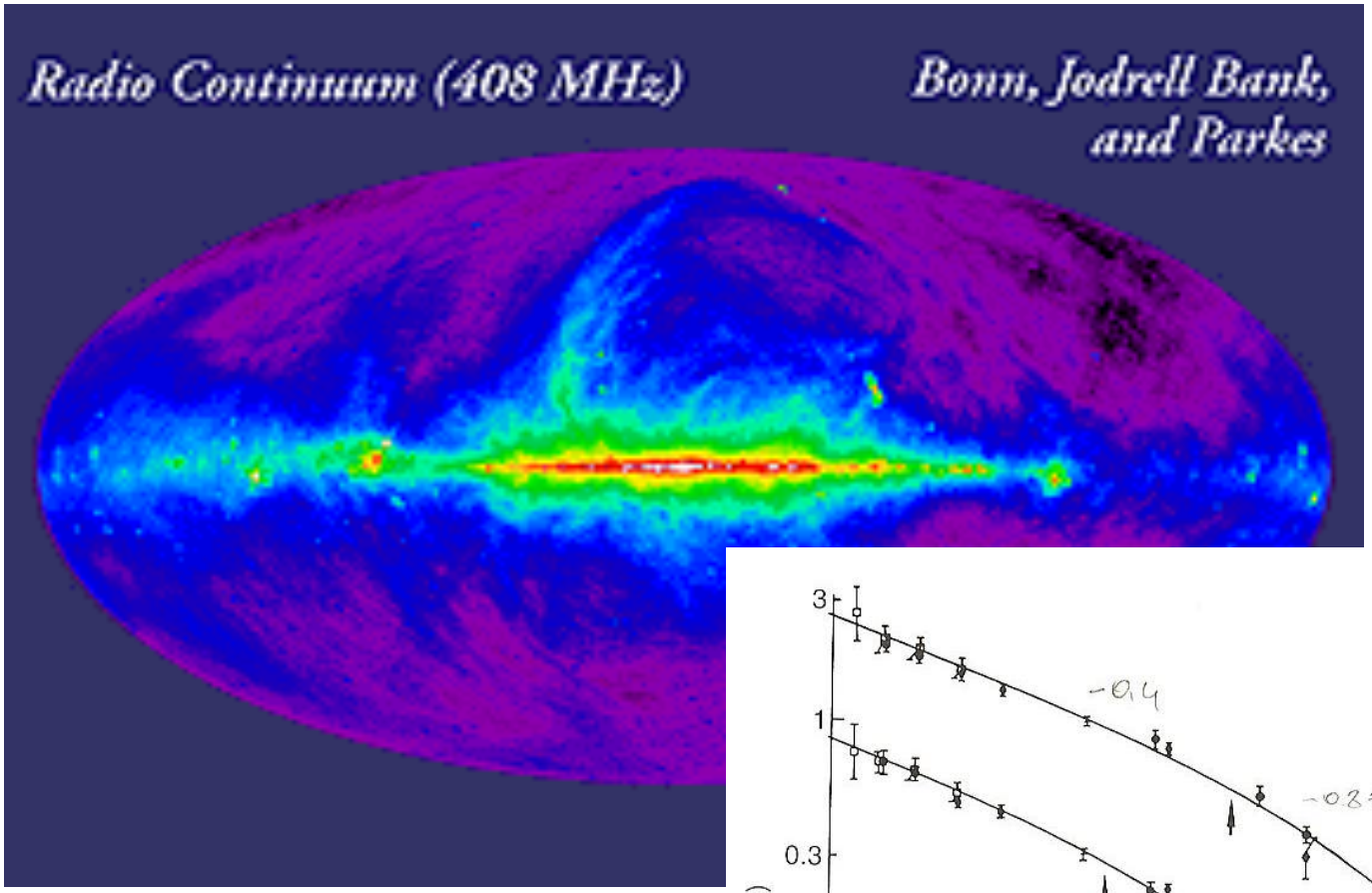


Figure 18.14. The spectrum of relativistic electrons in the local interstellar medium (see also Fig. 9.3). The observed electron spectrum is indicated by the hatched area. The spectrum deduced from the spectrum of the Galactic radio emission is indicated by a solid line. The low energy spectrum deduced from observations of low energy γ-rays is shown as a dashed line. The difference between the observed electron spectrum and that inferred from the Galactic radio emission provides a measure of the solar modulation of the flux of cosmic ray electrons. The units of $N(E)E^3$ are particles m^{-2} s^{-1} GeV^2. (From W. Webber (1983). In *Composition and origin of cosmic rays*, ed. M.M. Shapiro, p. 83. Dordrecht: D. Reidel Publishing Co.)



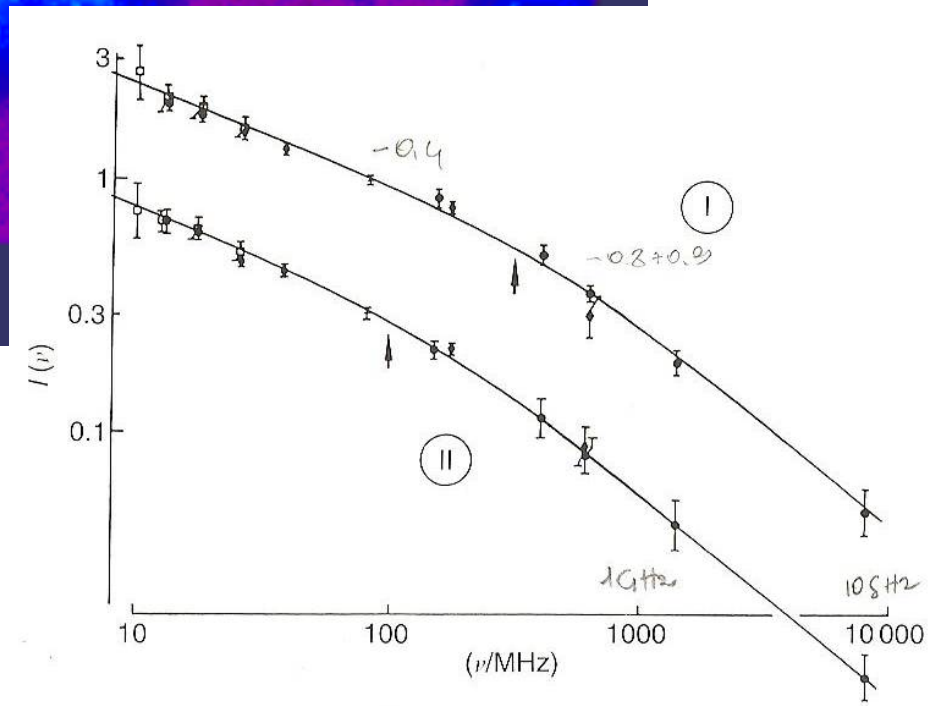
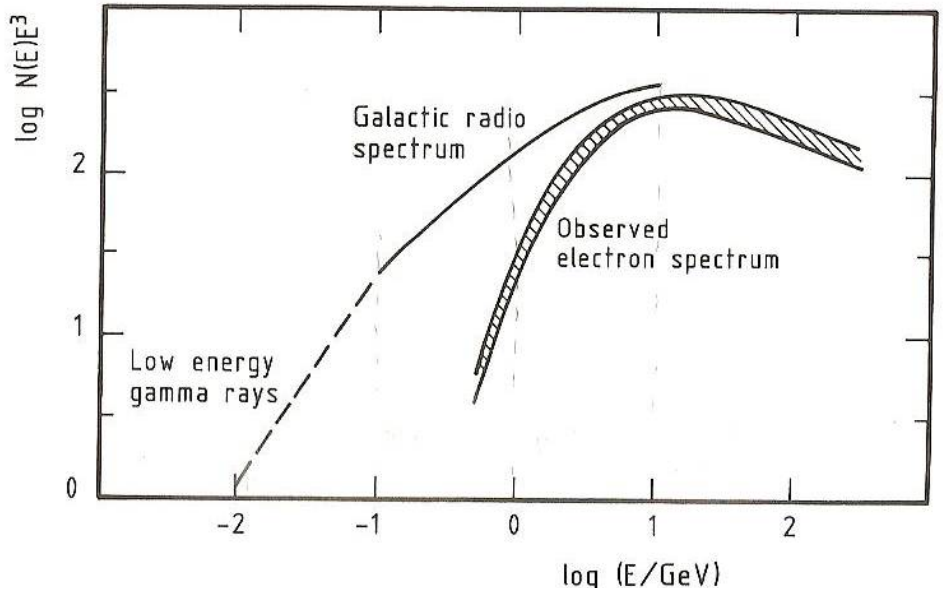


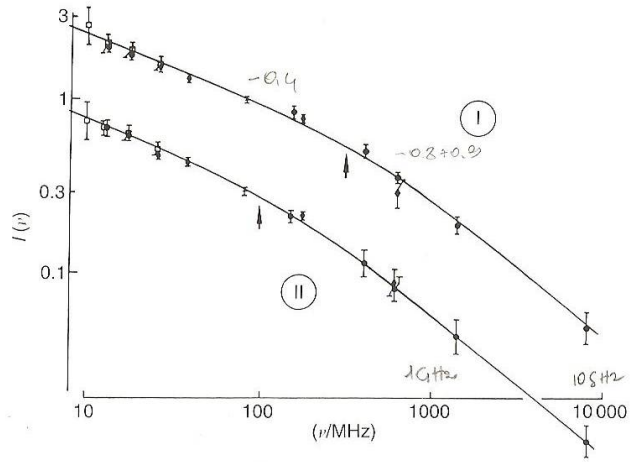
Figure 18.13. The spectrum of the Galactic radio emission. Region I corresponds to the anticentre direction at high galactic latitudes, and region II corresponds to the interarm region. (From A.S. Webster (1971). *Cosmic ray electrons, and Galactic radio emission.* Ph.D. dissertation; and A.S. Webster (1974). *Mon. Not. R. Astron. Soc.*, **166**, 355.)

Lo spettro di radiazione misurato e' una legge di potenza $I(\nu) \approx \nu^{-\alpha}$ con $\alpha = 0.4$ @$\nu < 200$ MHz e $=0.8$-0.9 a $\nu > 200$ MHz



Per emissione di sincrotrone $I(\nu) \approx \nu^{-(p-1)/2}$ dove p e' l'indice spettrale della popolazione di e- che mette

Un e- irraggia quasi tutta l'energia ad una frequenza $\nu = 28\gamma^2 B$ GHz ($B \approx 1.5$-6×10^{-10} T, val medio 3×10^{-10}) → $\nu \approx 100$ Mhz corrispondono ad e- di $E \approx 1$ GeV, $\nu \approx 1$ GhZ ad e- di ≈ 4 GeV



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- Da $(p-1)/2 = 0.4$ → $p=1.8$ @ $E \approx 1$ GeV
- $(p-1)/2 = 0.9$ → $p = 2.8$ @ $E \approx 10$ GeV

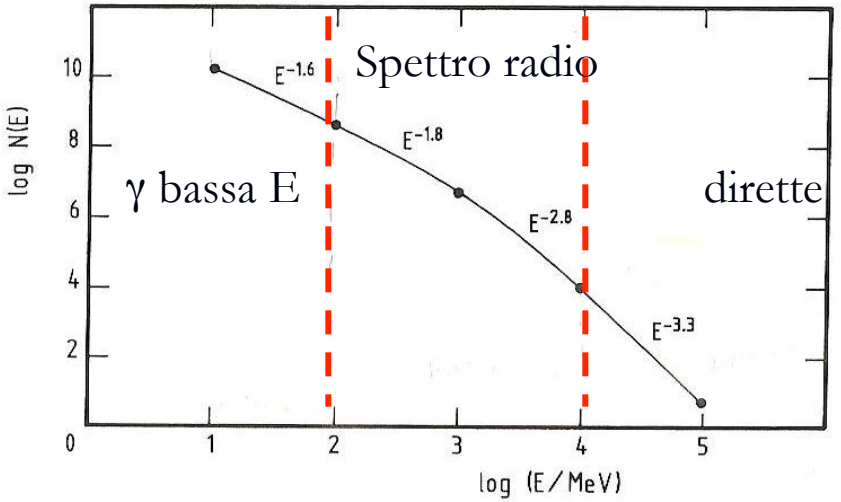


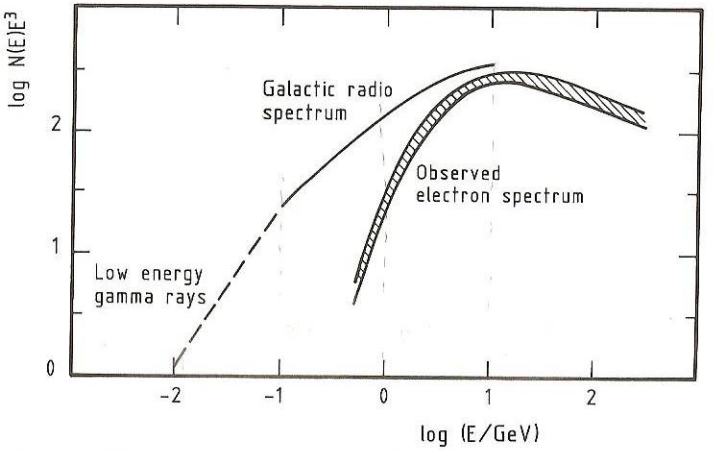
Figure 19.3. A schematic representation of the electron energy spectrum in the local interstellar medium from the data discussed in Sections 18.2 and 18.3. This spectrum has been subject to energy losses at high and low energies during propagation of the electrons through the interstellar medium. The units on the ordinate are relative units (see Fig. 18.14 for physical units).

Remember:

$$N(E) = (4\pi/c)\phi(E)$$

Lo spettro diventa piu' ripido all'aumentare di E

Al di sotto di 10 MHz il mezzo interstellare diventa opaco alla radiazione a causa dell'auto-assorbimento (cfr. nota black body) → si usa l'emissione gamma a bassa energia (10-100 MeV) per stimare lo spettro degli elettroni tra 30- 100 MeV



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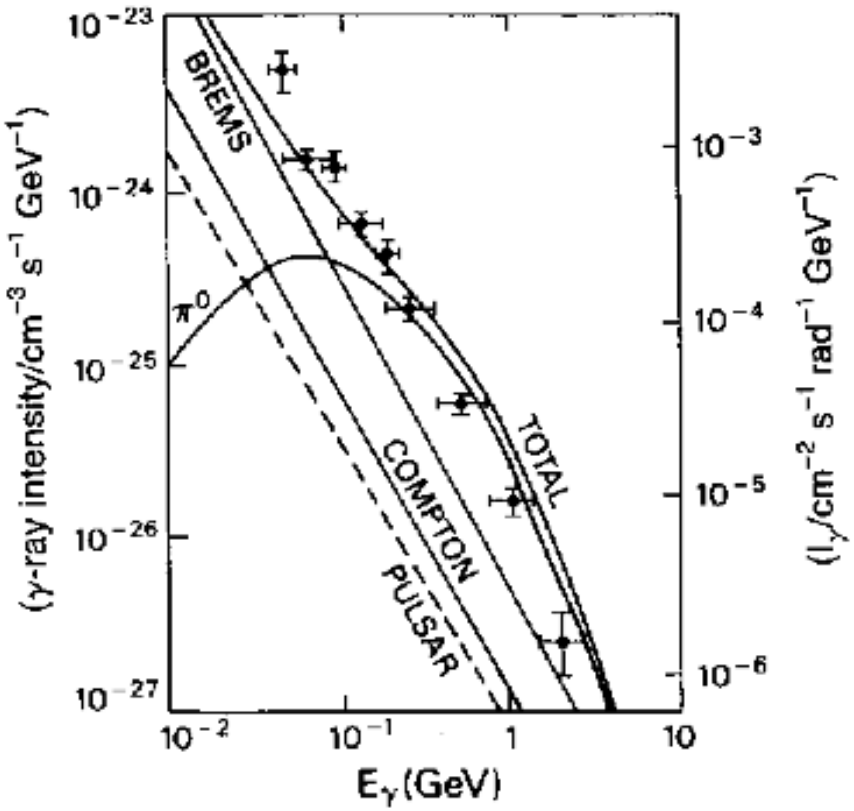
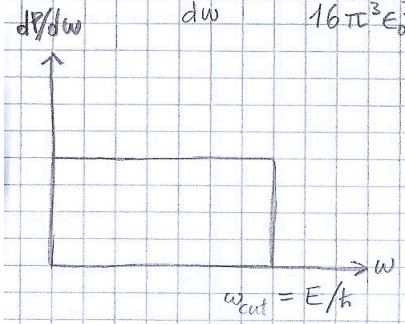


- Lo spettro γ dal disco galattico ad alta E e' dominato dall'emissione dovuta alla produzione di $\pi 0$
- A bassa E (30-100 MeV) il contributo dominante e' quello di brems di e- sui nuclei di H nell'ISM: l'emissivita' γ del disco dovuta a brems puo' essere valutata sottraendo il contributo dei $\pi 0$
- Lo spettro risultante e' $\kappa_\gamma(\varepsilon) \approx 10^{-27} \times \varepsilon^{-1.6}$ fotoni/cm^{3}secGeV

Spettro a bassa E (≈30-100 MeV)

$$\frac{dP}{d\omega} = \frac{Z(Z+1.3)e^6 N}{16\pi^3\epsilon_0^3 m_e^2 c^4} \ln\left(\frac{183}{Z^{1/3}}\right) \qquad \hbar\omega < E \quad [W\,Hz^{-1}]$$

Spettro di potenza di singola part. di energia E



Il # di fotoni con ϵ fra $\epsilon \div \epsilon + d\epsilon$ è

$$\frac{dN_\gamma(\epsilon)}{dt\,d\epsilon} = \frac{1}{\epsilon}\frac{dP}{d\omega}\cdot\frac{d\omega}{d\epsilon} = \frac{Z(Z+1.3)e^6 N}{16\pi^3\epsilon_0^3 m_e^2 c^4 \hbar} \ln\left(\frac{183}{Z^{1/3}}\right)\cdot\frac{1}{\epsilon} \qquad \#\text{fotoni}/\text{sec MeV}$$

$$= \frac{\alpha N}{\epsilon} \qquad \text{dove} \quad \alpha = 3.62\times10^{-22} Z(Z+1.3)\ln\left(\frac{183}{Z^{1/3}}\right) \quad m^3 s^{-1}$$

Assumendo 90% H e 10% He $\quad \alpha \approx 10^{-21}\ m^3 s^{-1}$

L'emissività è quindi (integro sulla popolazione di e^-)

$$K_\gamma(\epsilon) = \int_\epsilon^\infty \frac{dN_\gamma(\epsilon)}{dt\,d\epsilon}\cdot \underbrace{n_e(E)dE}_{dn_e} \qquad \text{con} \quad n_e(E) = K E^{-p} \quad \Rightarrow \quad K_\gamma(\epsilon) = \frac{\alpha N K \epsilon^{-p}}{(p-1)} \qquad \#\text{fotoni}/m^3\,\text{sec MeV}$$

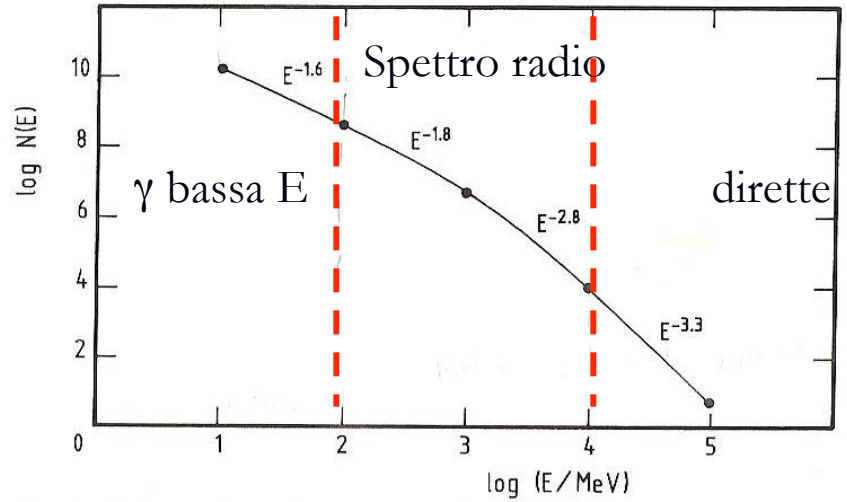
Sol e^- con $E \geq \epsilon$ contribuiscono

$$= \frac{\alpha N}{(p-1)} n_e(\epsilon) \quad \Rightarrow \quad \frac{n_e(\epsilon)}{p-1} = \frac{K_\gamma(\epsilon)}{\alpha N} \qquad cm^{-3}\,MeV^{-1}$$

$$\Rightarrow \quad \phi_e(\epsilon) = \frac{c}{4\pi} n_e(\epsilon) \qquad cm^{-2} s^{-1} MeV^{-1} sr^{-1}$$

- **L'indice spettrale degli elettroni e' lo stesso dei fotoni ≈1.6**

Propagazione degli e-



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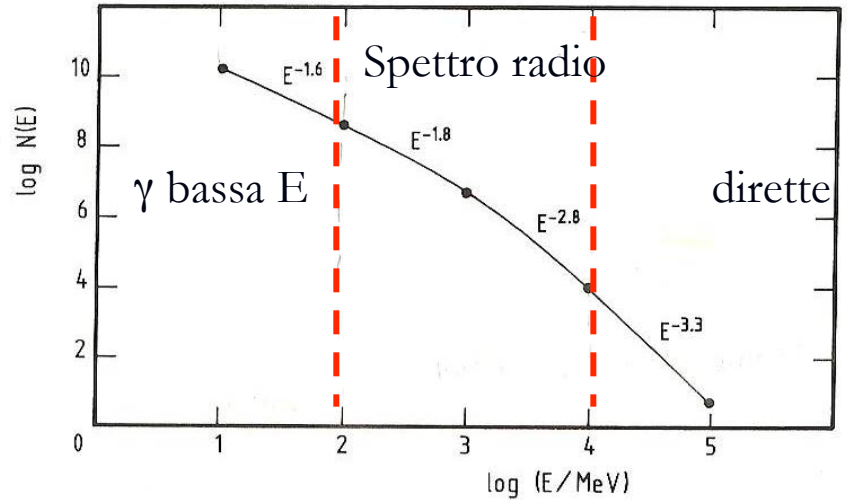
Potremmo risolvere l'equ di diffusione completa...ma sarebbe piuttosto lungo...

Usiamo invece il modello Leaky Box in cui il termine di diffusione e' rimpiazzato da un term N(E)/T(E), T tempo di fuga

- Distr gaussiana di path length→ simula la diffusione
- Distr exp per il caso in cui non c'e' diffusione ma solo prop nel volume

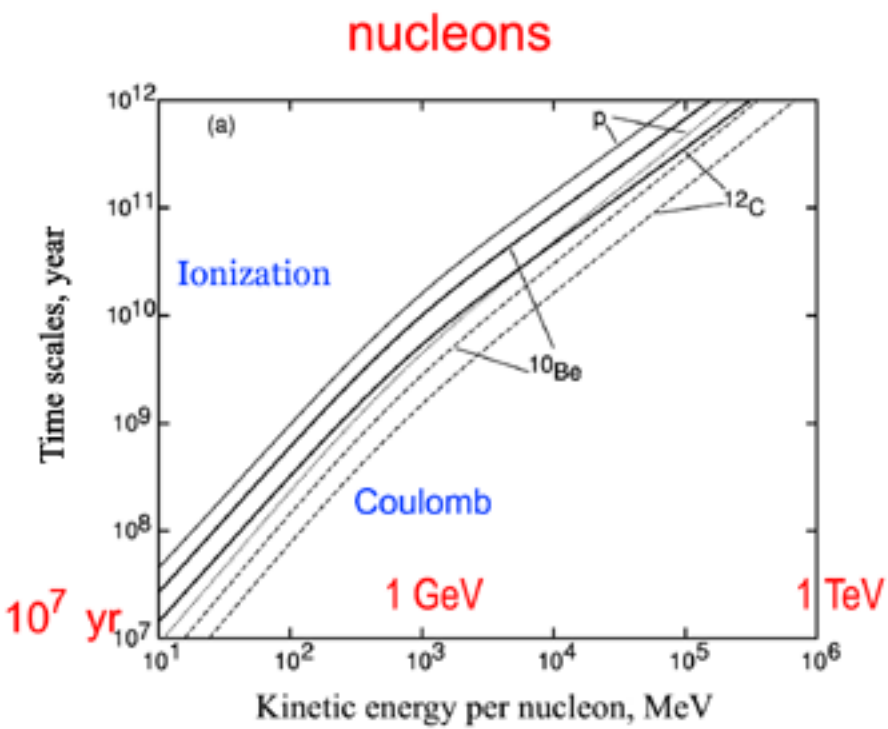


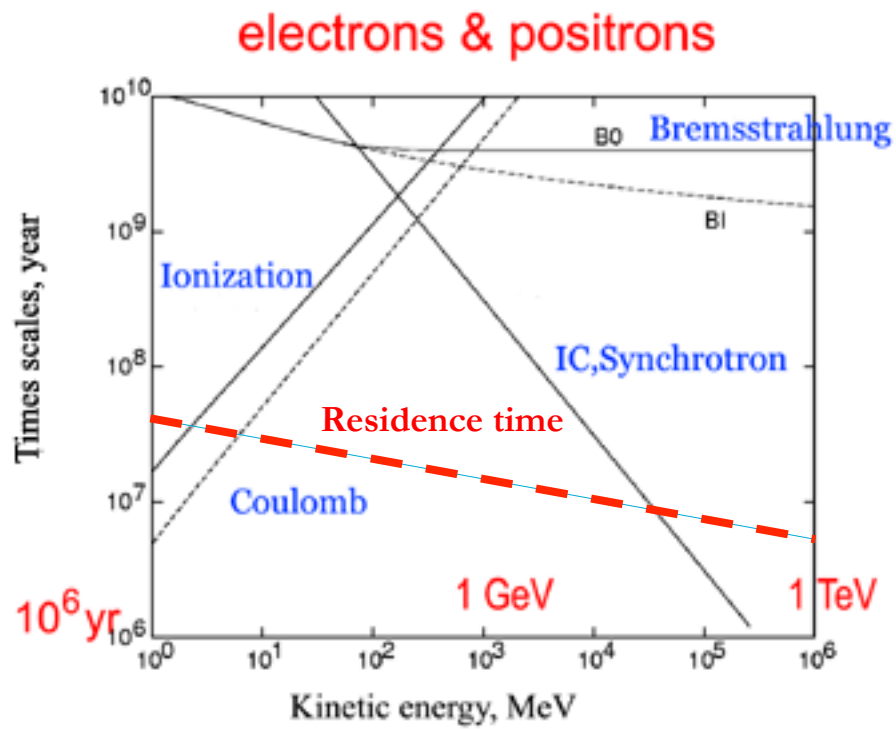
Abbiamo visto che i tempi di confinamento dei RC sono dell'ordine di $T_{esc}(E) \approx 1\text{-}3 \times 10^7$ anni

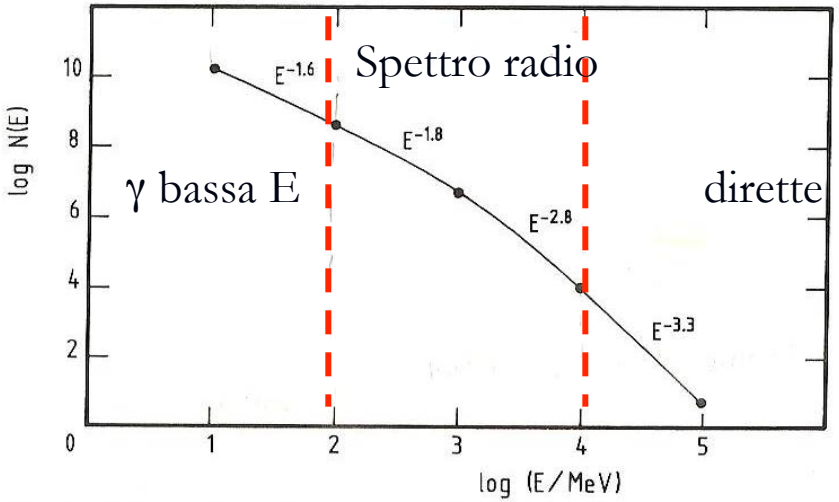
Se gli e- hanno tempi simili, il tempo scala per le perdite di energia τ e' < T_{esc} (E) per differenti ragioni in differenti intervalli di energia:

Ad E< 1Gev dominano le perdite per ionizz. $\tau(E) \approx 10^7$ anni per e- ad E=300 MeV

Ad E>10 GeV, IC e sincr dominano ed $\tau(E) < 10^7$ anni per e-


nucleons
(a)
p
12C
10Be
Ionization
Coulomb
1 GeV
1 TeV
10^7 yr
Time scales, year
Kinetic energy per nucleon, MeV


electrons & positrons
B0
Bremsstrahlung
BI
Ionization
IC,Synchrotron
Residence time
Coulomb
1 GeV
1 TeV
10^6 yr
Times scales, year
Kinetic energy, MeV



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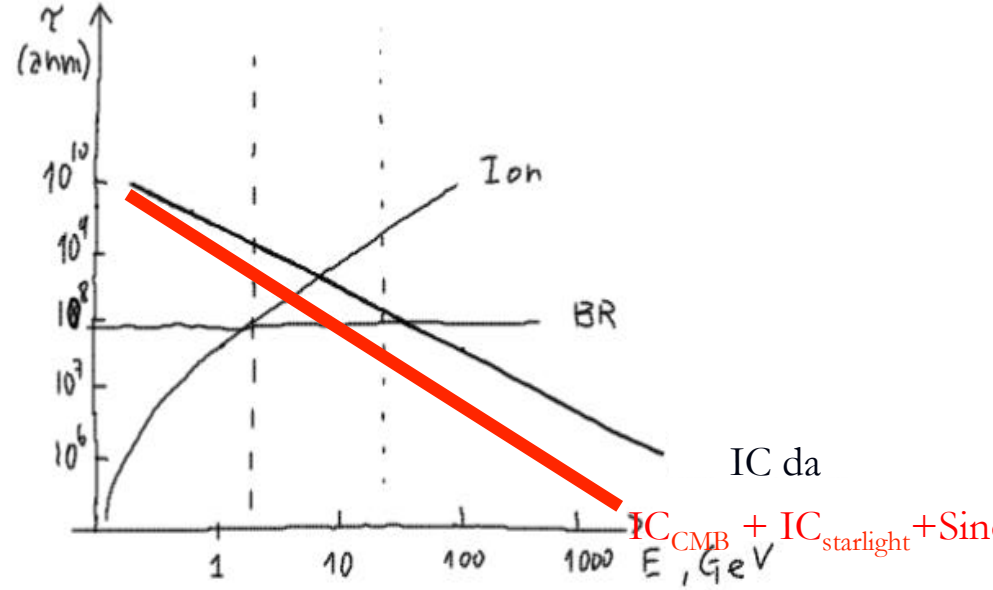


Quindi, se gli e- sono iniettati continuamente nell'ISM, dovrebbero raggiungere una situazione stazionaria di equil fra iniezione e perdite, dN/dt = 0

A bassa E, dove la ionizz domina quindi $N(E) \approx E^{-p-1} \approx Q(E)E^{-1}$ → (p-1) = 1.6 → $Q(E) \approx E^{-2.6}$

Ad alta E, IC e Sincr dominano quindi $N(E) \approx E^{-(p+1)}$ → p+1 = 3.3 → $Q(E) \approx E^{-2.3}$

I due indici spettrali sono piuttosto simili: cio' suggerisce che lo spettro alla sorgente degli e- sia una legge di potenza con indice spettrale p≈ 2.5 fra 100 MeV e 100 GeV

Possibili Break del flusso

- Un processo crea una distr. di e^- alla sorgente di tipo $Q(E) = kE^{-p}$ $[m^{-3}\,GeV^{-1}]$

- Gli e^- possono rimanere nella sorgente per un tempo t_0 prima di sfuggire e propagarsi fino a noi. Per esempio SNR

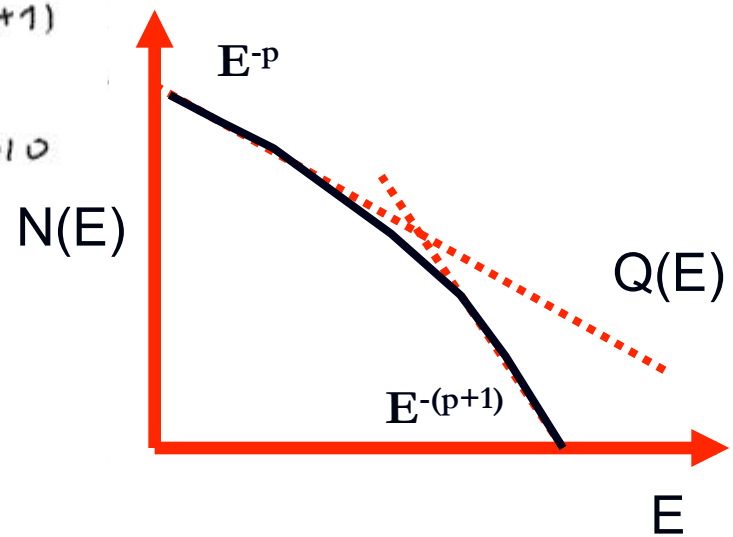
- Durante la permanenza nella sorgente possono perdere energia, per es. per sincrotrone

$\Rightarrow$ e^- con $E \geq E_0 = \frac{1}{k t_0} = \frac{0.125}{B^2 t}$ eV $[T]\,[anni]$ hanno perso una fraz. significativa di E <u>prima</u> lasciare la sorgente

- per $E < E_0$, gli e^- non hanno perso frazioni importanti di E
  $\Rightarrow$ lo spettro degli e^- che lasciano la sorgente dopo t_0 è lo stesso di $Q(E) \Rightarrow N(E) = k E^{-p}$

- per $E > E_0$, osserviamo solo gli elettroni prodotti nell'intervallo τ precedente t_0, cioè $\tau = 1/E$
  $\Rightarrow N(E) = Q(E) \cdot \tau(E) \propto E^{-(p+1)}$

- All'energia E_0 c'è un cambio di pendenza nello spettro (un "break")



Electron mass small compared to protons and heavy nuclei, ➔ lose energy more rapidly

Lifetimes are short, ➔ electron sources are **Galactic**.

Observed energy density ~4×10^3 eV m^{-3} (total for cosmic rays ~10^6 eV m^{-3})

Assuming Crab pulsar-like sources…

can Galactic pulsars source CR electrons?

Need first to calculate how many electrons produced by the Crab nebula.

Observed synchrotron X-rays from SNR, $\omega_c = eB\gamma^2/m$

$$\nu = 4 \times 10^{36}\, E^2 B \text{ Hz}$$

@ $\nu \sim 10^{18}$ Hz (X-ray) assume $B_{SNR} = 10^{-8}$ Tesla

$$\rightarrow E_e \sim 5 \times 10^{-6}\, J = \boxed{3 \times 10^{13} \text{ eV}}$$

$P = 2.4 \times 10^{12} E^2 B^2$ J/s

$= 2.4 \times 10^{12} \times 2.5 \times 10^{-11} \times 10^{-16}$ J/s

$= 6 \times 10^{-15}$ J/s

Observed flux = 1.6×10^{-11} J $m^{-2}s^{-1}$

Distance = 1kpc = 3×10^{19} m

Total luminosity, L = $1.6 \times 10^{-11} \times 4\pi d^2$ J/s = $1.6 \times 10^{-10} \times 10^{2} \times 10^{38}$ J/s = 1.6×10^{30} J/s

Number of electrons = luminosity/power per e^- = 1.6×10^{30}/ 6×10^{-15} = 2.6×10^{44}

Synchrotron lifetime, τ_{sinc}=$5 \times 10^{-13} B^{-2} E^{-1}$ s = 30 years

Thus in 900 yrs since SN explosion, must be 30 replenishments of electrons and these must be produced by the pulsar.

Total no. electrons = $2.6 \times 10^{44} \times 30 \sim 8 \times 10^{45}$

each with E = 5×10^{-6} J

Total energy is thus 4×10^{40} J

Assume 1 SN every 100 years for 10^{10} years ➔ total energy due to pulsars $4 \times 10^{40} \times 10^{8}$ J = 4×10^{48} J in a volume of $\sim 10^{63}$ m^{-3} (ie. the Galaxy)

➔ energy density of electrons produced by pulsars :

$= 4 \times 10^{48} / 10^{63}$ Jm^{-3} $= 4 \times 10^{-15}$ Jm^{-3}

$= 4 \times 10^{-15} / 1.6 \times 10^{-19}$ eVm^{-3} $= 2.5 \times 10^{4}$ eVm^{-3}

Remember that the observed electron energy density is 4×10^{3} eV/m^{3}, so electron emissions from pulsars are sufficient to produce observed levels, given the assumptions made.

SNRs vicine

Vicino al noi ci sono numerose potenziali sorgenti di elettroni

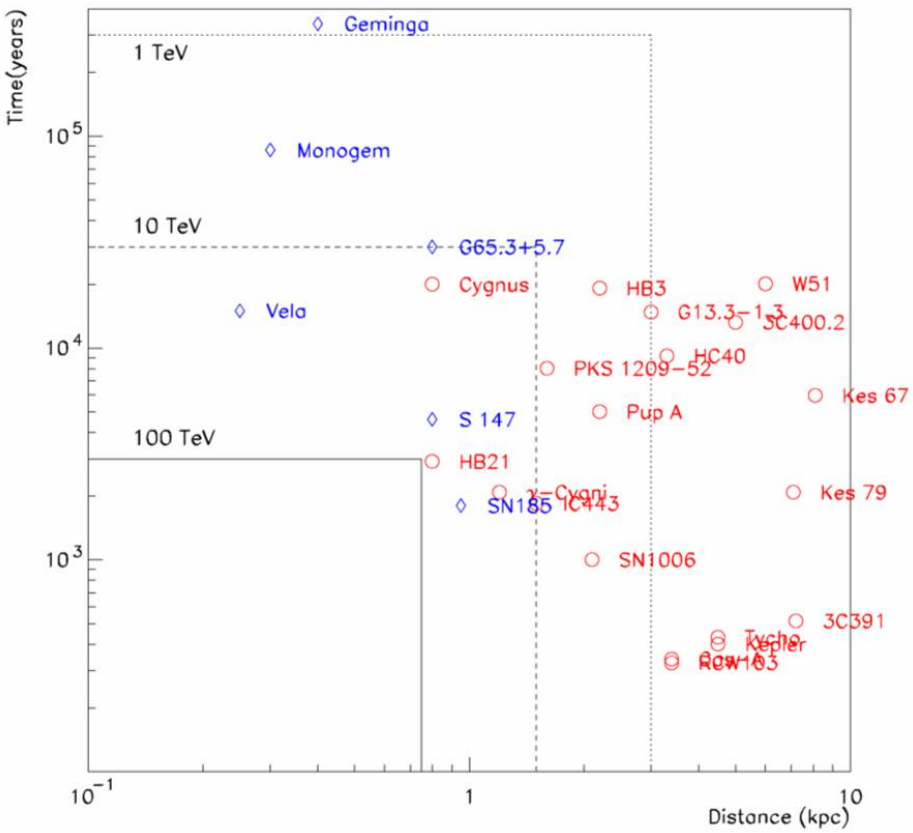


TABLE 1

LIST OF NEARBY SNRs

SNR	Distance (kpc)	Age (yr)	E_{max} [a] (TeV)	Reference
SN 185	0.95	1.8×10^3	1.7×10^2	1
S147	0.80	4.6×10^3	63	2
HB 21	0.80	1.9×10^4	14	3, 4
G65.3+5.7	0.80	2.0×10^4	13	5
Cygnus Loop	0.44	2.0×10^4	13	6, 7
Vela	0.30	1.1×10^4	25	8
Monogem	0.30	8.6×10^4	2.8	9
Loop1	0.17	2.0×10^5	1.2	10
Geminga	0.4	3.4×10^5	0.67	11

[a] Maximum energy limited by the propagation of electrons in the case of the prompt release after the explosion. The delay of the release time gives the larger value.

REFERENCES.—(1) Strom 1994; (2) Braun et al. 1989; (3) Tatematsu et al. 1990; (4) Leahy & Aschenbach 1996; (5) Green 1988; (6) Miyata et al. 1994; (7) Blair et al. 1999; (8) Caraveo et al. 2001; (9) Plucinsky et al. 1996; (10) Egger & Aschenbach 1995; (11) Caraveo et al. 1996.

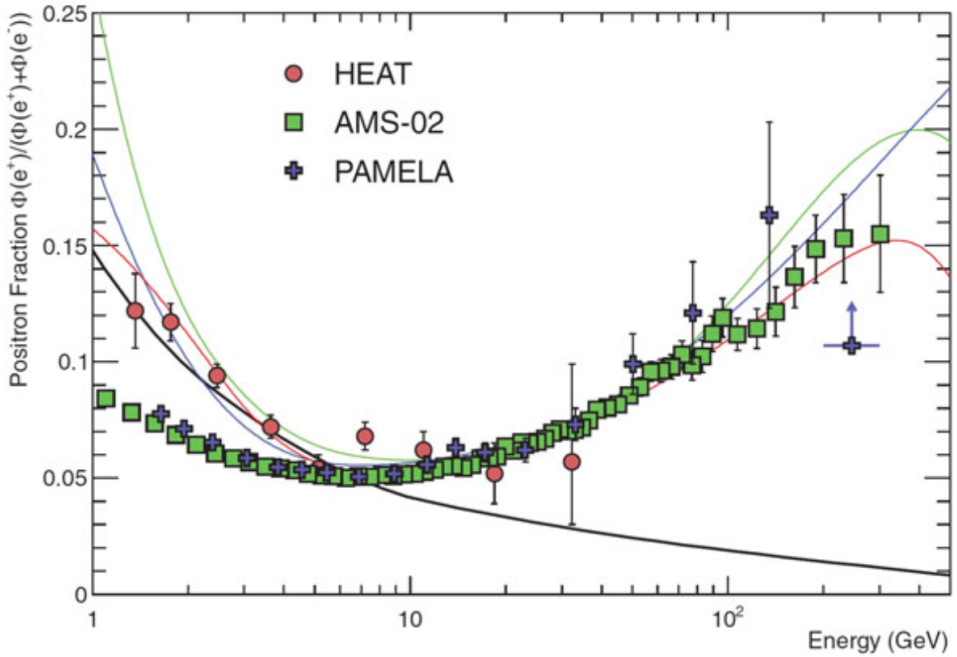


Fig. 3.14 The positron fraction (ratio of the flux of e^+ to the total flux of $e^+ + e^-$) as a function of the energy measured HEAT, PAMELA, and AMS-02. The heavy *black line* is a model of pure secondary production using a detailed propagation model of CRs (Sect. 5.4). The *three thin lines* show three representative attempts to model the positron excess with different phenomena discussed in Sect. 13.9.3: dark matter decay (*green*); propagation physics (*blue*); production in pulsars (*red*). The ratio below 10 GeV is dependent on the polarity of the solar magnetic field. Figure from the Sect. 27. Cosmic Rays of Beringer et al. (2012)

- The positron-fraction spectrum does not exhibit fine structures and steadily increases in the region between 10 and 250 GeV. In the high statistics AMS-02 sample the ratio is of the order of ~10 % above a few tens of GeV. As a consequence, since positrons are always created in pair with an electron *about 90 % of the observed electrons must be of primary origin*.

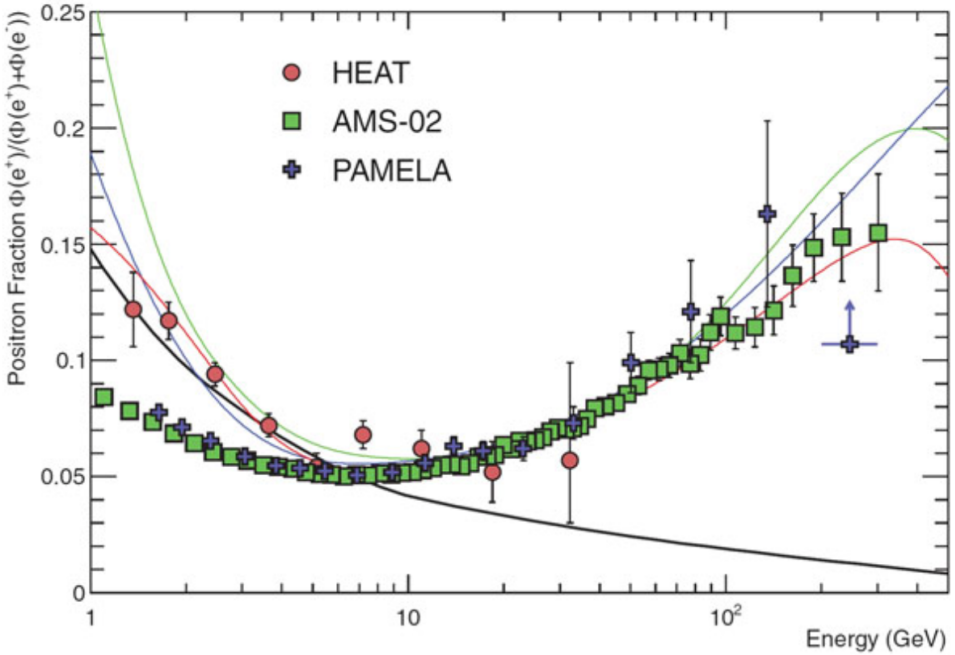
E. Fian

Eccesso rispetto a cosa?

Produzione secondaria nell'ISM di e+ ed e-

Sorgenti primarie di e-: $n_-(E)=K_{ep} N_{cr}(E) + \sigma_{pe-} n_{ISM} c\, n_{cr}(E)$, $K_{ep} \approx 10^{-3}$

No sorgenti primarie di e+: $n_+(E) = \sigma_{pe+} n_{ISM} c\, n_{cr}(E)$



Leaky box equilibrium density $n_{cr}(E) = N_{cr}(E)\tau_{esc}(E)$, $\tau_{esc}(E) = \tau_0 E^{-\delta}$

$$n_-(E)/n_+(E) = (K_{ep} N_{cr}(E) + \sigma_{pe-} n_{ISM} c\, N_{cr}(E)\tau_{esc}(E))/\sigma_{pe+} n_{ISM} c\, N_{cr}(E)\tau_{esc}(E) =$$
$$= \sigma_{pe-}/\sigma_{pe+} + K_{ep}/\tau_{esc}(E)$$

$$n_+(E)/(n_+(E) + n_-(E)) = 1/(1 + n_-(E)/n_+(E)) = 1/(1 + \sigma_{pe-}/\sigma_{pe+} + K_{ep}/\tau_{esc}(E))$$

$\sigma_{pe-}/\sigma_{pe+} \approx 0.5 - 0.3$ ➔ $n_+(E)/(n_+(E) + n_-(E)) \approx 1/(1.5 + (K_{ep}/\tau_0)E^{\delta})$

➔ **Senza produzione primaria di e+ alle sorgenti la frazione decresce con l'energia**

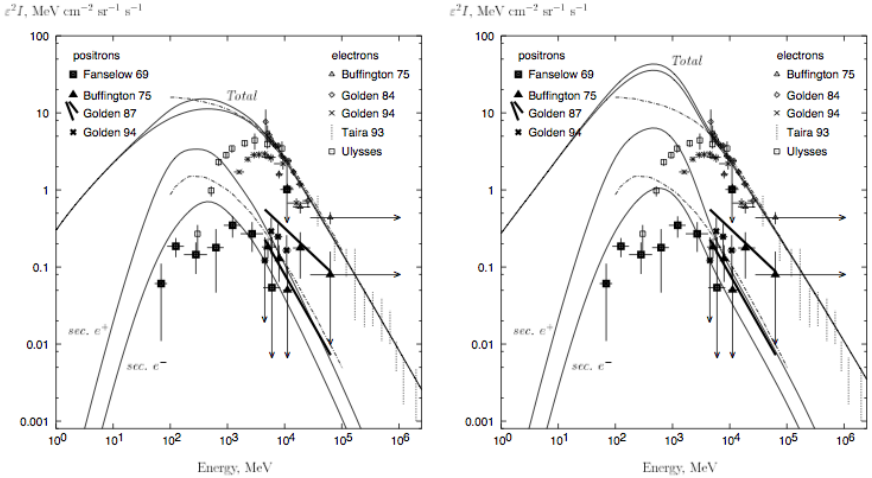


Fig. 5.— Spectra of secondary positrons and electrons, and of primary electrons. *Left panel*: model with no reacceleration (08–005). Electron injection spectrum index 2.1, 2.4 above and below 10 GeV respectively. Model: upper curves: primary electrons and primary+secondary electrons and positrons. Lower curves: secondary positrons, electrons. Lower dashed-dot line: Protheroe (1982) leaky-box prediction. Data: electrons: Buffington, Orth, & Smoot (1975), Golden et al. (1984), Golden et al. (1994), Taira et al. (1993), Ulysses (Ferrando et al. 1996), upper dashed-dot line: Protheroe (1982); positrons: Fanselow et al. (1969), Buffington, Orth, & Smoot (1975), Golden et al. (1987), Golden et al. (1994). *Right panel*: same, model with reacceleration (08–006).

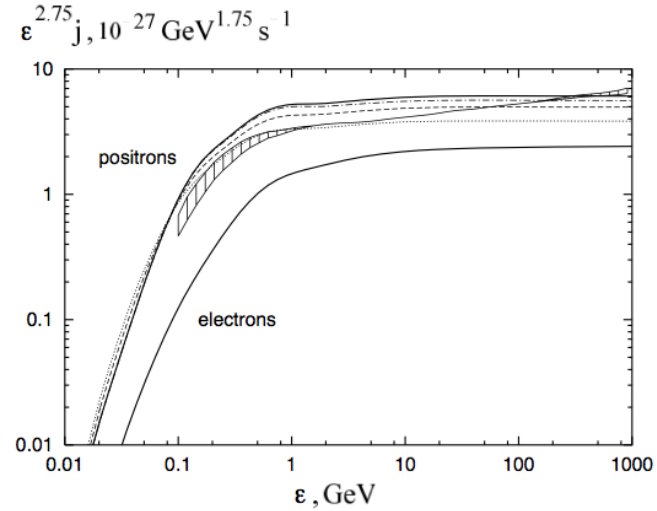


Fig. 4.— Illustration of the effect of the different positron distributions in the muon rest system. The thick solid lines show the production spectrum of secondary positrons and electrons per hydrogen atom for the cosmic-ray proton spectrum given by Mori (1997). The positron spectrum with no kaon contribution is shown by the dash-dotted line. The dashed line shows the spectrum calculated assuming an isotropic distribution of e^+ in the muon rest system ($\xi = 0$), the dotted line shows the spectrum calculated for $\xi = -1$. The thin solid line with hatched regions shows
osmic rays in interstellar medium (including contribution of He

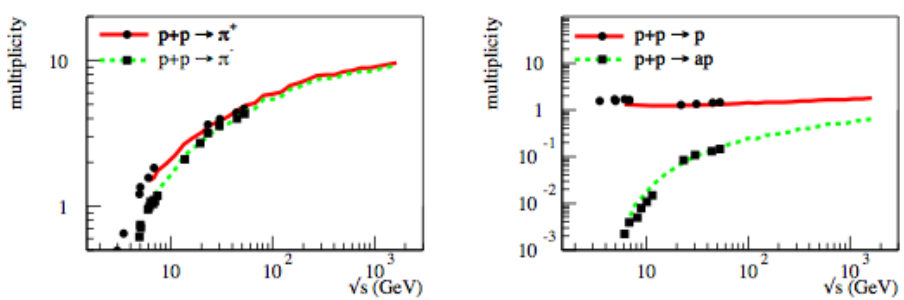


Figure 1: The 4π multiplicities of π^+, π^+, proton, antiproton produced in a pp collision as a function of $\sqrt{s}$. The full and dashed lines show the result of the χ^2 fit. The points are data from [15].

Comparing our data with a minimal model, as an example.

In this model the e^+ and e^- fluxes, Φ_{e+} and Φ_{e-}, are parametrized as the sum of individual diffuse power law spectra and the contribution of a single common source of $e^{\pm}$:

$$\Phi_{e^+} = C_{e^+} E^{-\gamma_{e^+}} + C_s E^{-\gamma_s} e^{-E/E_s}$$ ▪ pulsar source **Eq(1)**

$$\Phi_{e^-} = C_{e^-} E^{-\gamma_{e^-}} + C_s E^{-\gamma_s} e^{-E/E_s}$$ (E in GeV) **Eq(2)**

Coefficients C_{e+} and C_{e-} correspond to relative weights of diffuse spectra for positrons and electrons.

C_s is the weight of the source spectrum.

γ_{e+}, γ_{e-} and γ_s are the corresponding spectral indexes.

E_s is a characteristic cutoff energy for the source spectrum.

With this parametrization the positron fraction depends on 5 parameters.

A fit to the data in the energy range 1 to 350 GeV yields a $\chi^2/d.f.$ = **28.5/57** and:

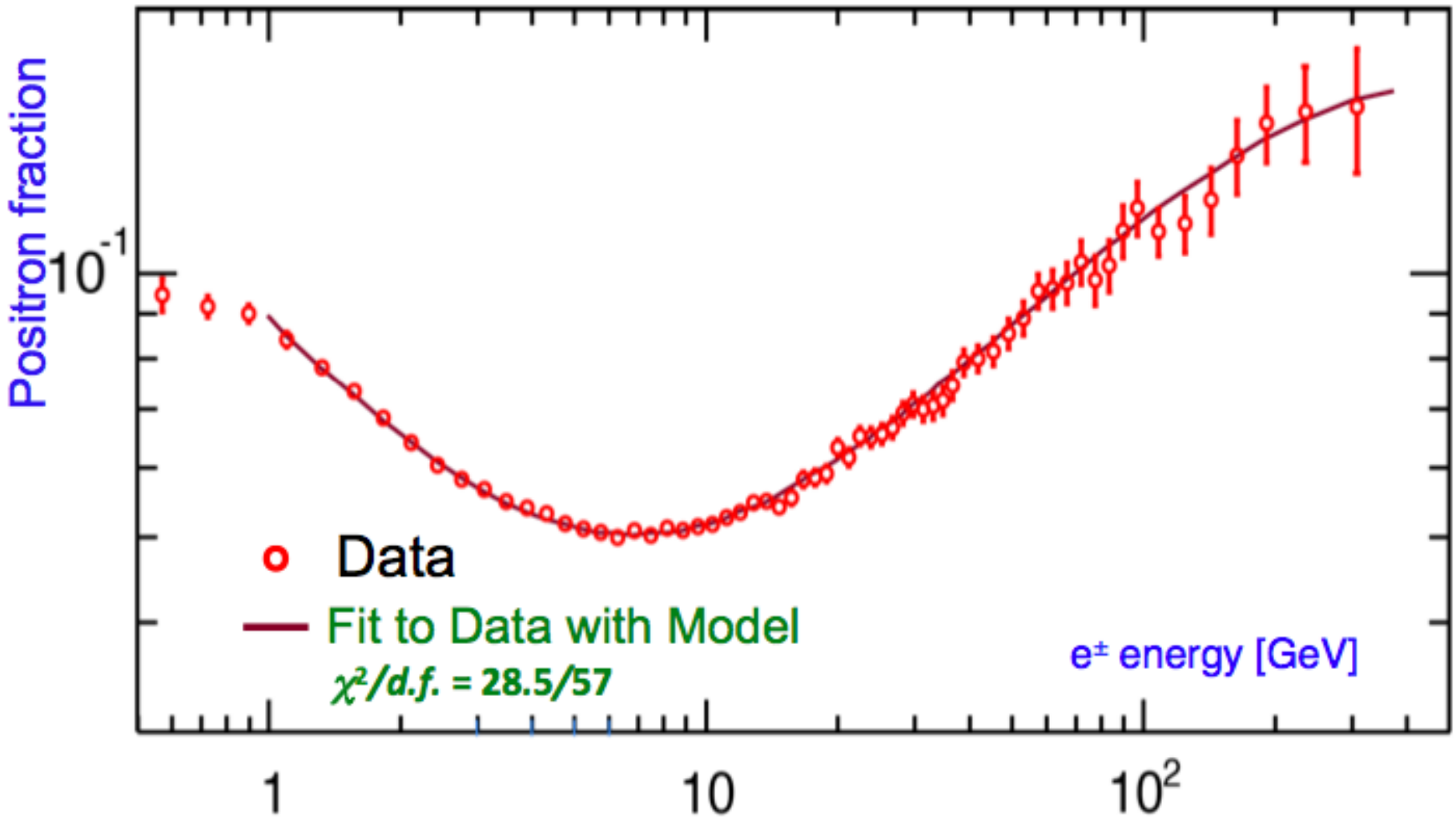
$\gamma_{e-} - \gamma_{e+}$ = **−0.63 ± 0.03**, *i.e.*, the diffuse positron spectrum is less energetic than the diffuse electron spectrum;

$\gamma_{e-} - \gamma_{s}$ = **0.66±0.05**, *i.e.*, the source spectrum is more energetic than the diffuse electron spectrum;

C_{e+}/C_{e-} = **0.091 ± 0.001**, *i.e.*, the weight of the diffuse positron flux amounts to ~10% of that of the diffuse electron flux;

C_s/C_{e-} = **0.0078 ± 0.0012**, *i.e.*, the weight of the common source constitutes only ~1% of that of the diffuse electron flux;

$1/E_s$ = **0.0013 ± 0.0007 GeV^{-1}**,
corresponding to a cutoff energy of 760^{+1000}_{-280} **GeV**.



The agreement between the data and the model shows that the positron fraction spectrum is consistent with $e^{\pm}$ fluxes each of which is the sum of its diffuse spectrum and a single common power law source.

The origin of the excess …

…is there any privileged arrival direction?

Analysis of possible deviation of the measured ratio as a function of the arrival direction in galactic coordinates (b,l)

$$\frac{r_e(b,l)}{\langle r_e \rangle} - 1 = \sum_{\ell=0}^{\infty} \sum_{m=-\ell}^{\ell} a_{\ell m} Y_{\ell m}(\pi/2 - b, l)$$

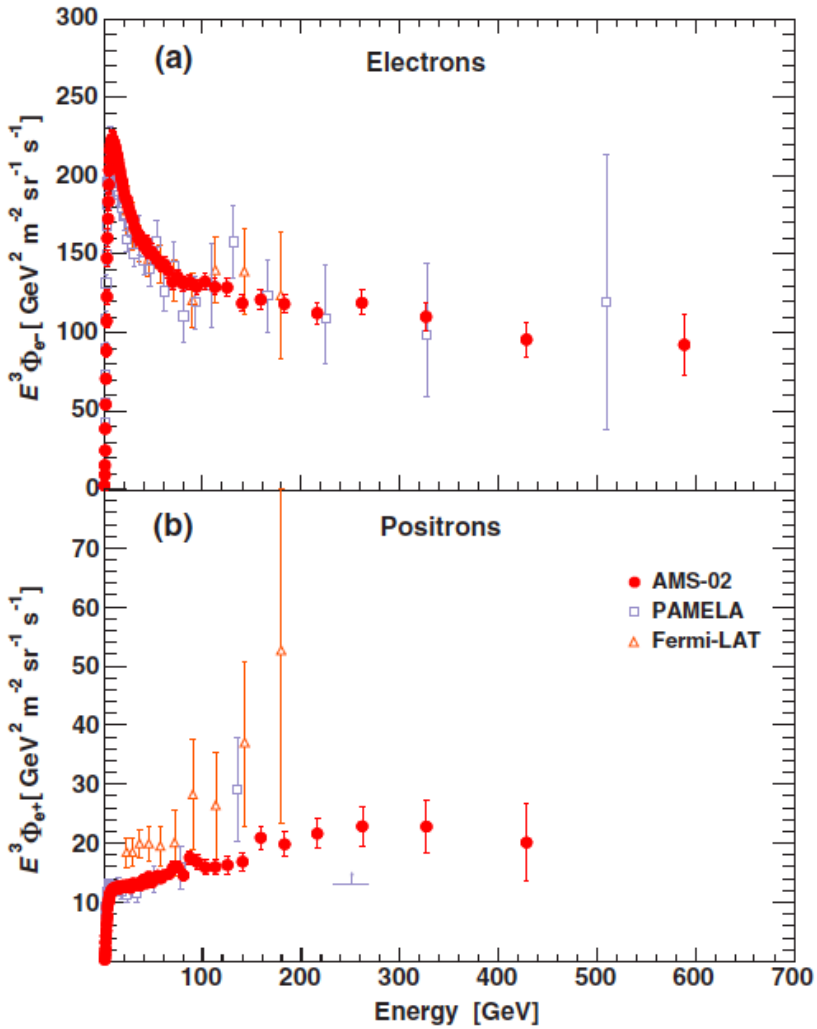
Power spectrum from the coefficient of the spherical armonics:

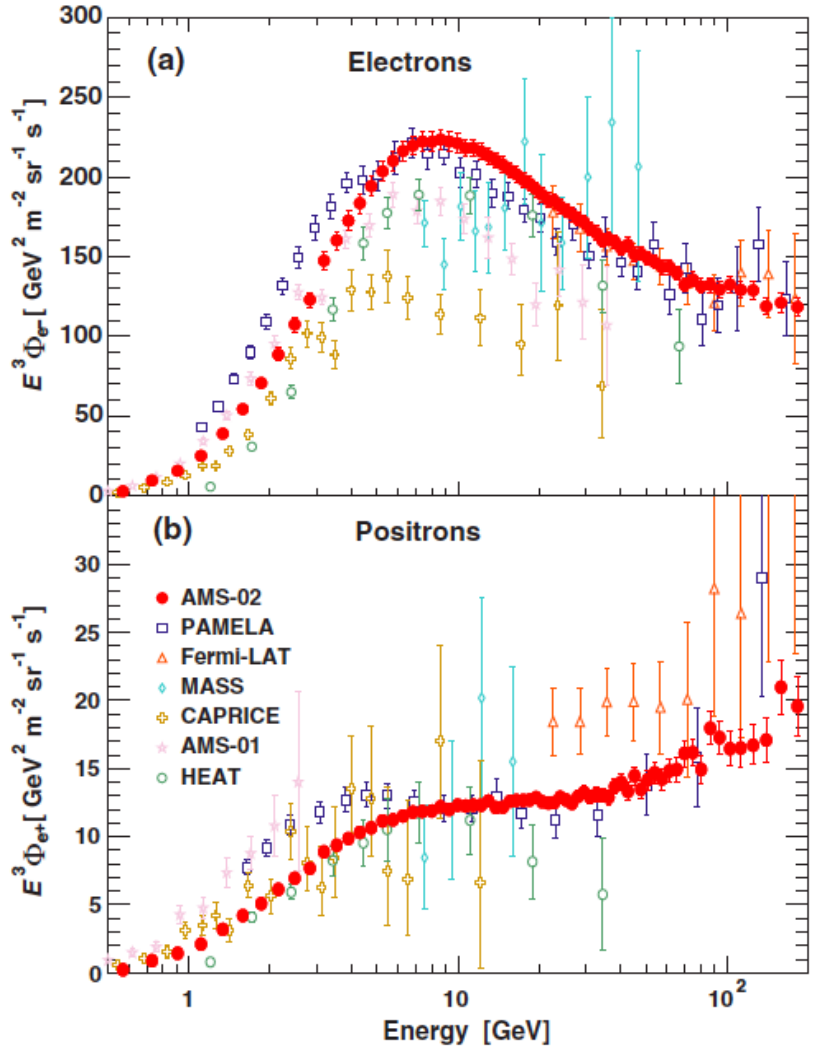
$$C_\ell = \frac{1}{2\ell+1} \sum_{m=-\ell}^{\ell} |a_{\ell m}|^2.$$

Compatible with isotropy in the dipolar mode $\delta = 3\sqrt{C_1/4\pi}.$

δ ≤ 0.036 at the 95% confidence level

Oggi:


(a)
Electrons
(b)
Positrons
AMS-02
PAMELA
Fermi-LAT
$E^3\Phi_{e^-}$[GeV^2 m^{-2} sr^{-1} s^{-1}]
$E^3\Phi_{e^+}$[GeV^2 m^{-2} sr^{-1} s^{-1}]
Energy [GeV]


(a)
Electrons
(b)
Positrons
AMS-02
PAMELA
Fermi-LAT
MASS
CAPRICE
AMS-01
HEAT
$E^3\Phi_{e^-}$[GeV^2 m^{-2} sr^{-1} s^{-1}]
$E^3\Phi_{e^+}$[GeV^2 m^{-2} sr^{-1} s^{-1}]
Energy [GeV]

Possiamo dare un quadro generale degli spettri alle sorgenti:

- Per la componente nucleare $p \approx 2$
- Per la componente e^{+-} $p \approx 2.5$

Fino ad almeno qualche centinaio di GeV

Gli spettri alla sorgente sono differenti per e- e protoni (cum grano salis)

Per entrambe le componenti c'e' evidenza sperimentale di hardening dello spettro a qualche O(100 GeV)

Per e+ ed e- l'evidenza e' netta: c'e' una sorgente che immette e+ ed e- di alta energia nell'ISM

Per i nuclei la situazione e' piu' incerta: nuova classe di sorgenti? Effetto di propagazione? Servono piu' dati.

Le pulsar accelerano e+ ed – ma non protoni (per esempio)

Occorre quindi trovare uno o piu' processi astrofisici che siano in grado di accelerare le particelle fino ad energie ultrarelativistiche con la distribuzione osservata

Pulsars

Gamma-ray pulsars are predicted to produce energetic electron-positron pairs with a harder spectrum than that from secondary cosmic-ray induced origin, leading to the possibility that such sources may dominate the cosmic ray positron spectrum at high energies.

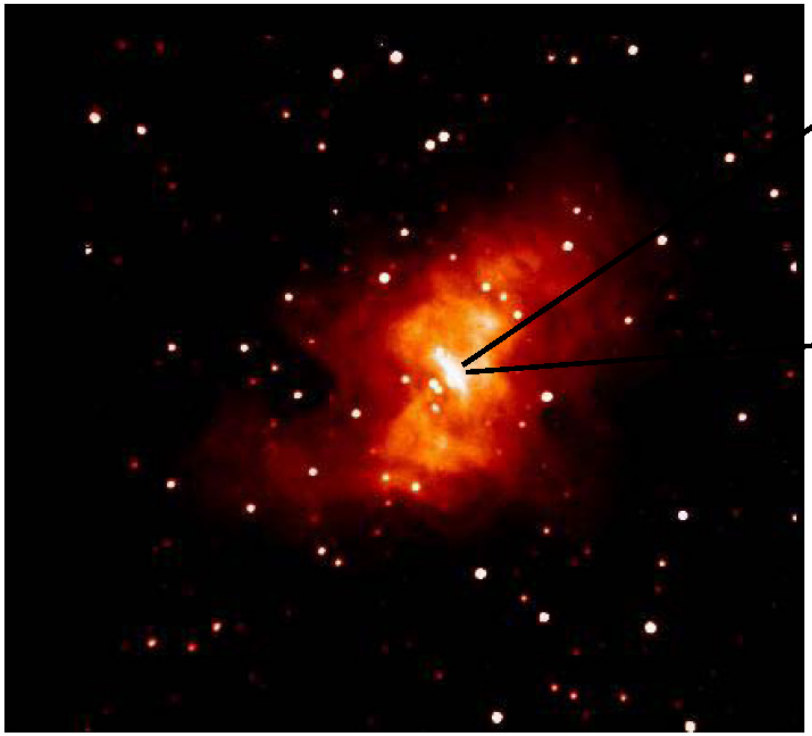
Una Pulsar è una giovane *stella di neutroni (NS)* rapidamente ruotante rispetto ad un asse.

Una NS ha un'altissima densità (quella dei nuclei), massa pari a ~1.4 $M_¤$ (massa sole), e raggio R_{NS}~10 km

Supponendo un campo magnetico della stella prima del collasso pari a B ~ $B_¤$ ~ 10^{-2} T, per la legge di Gauss:

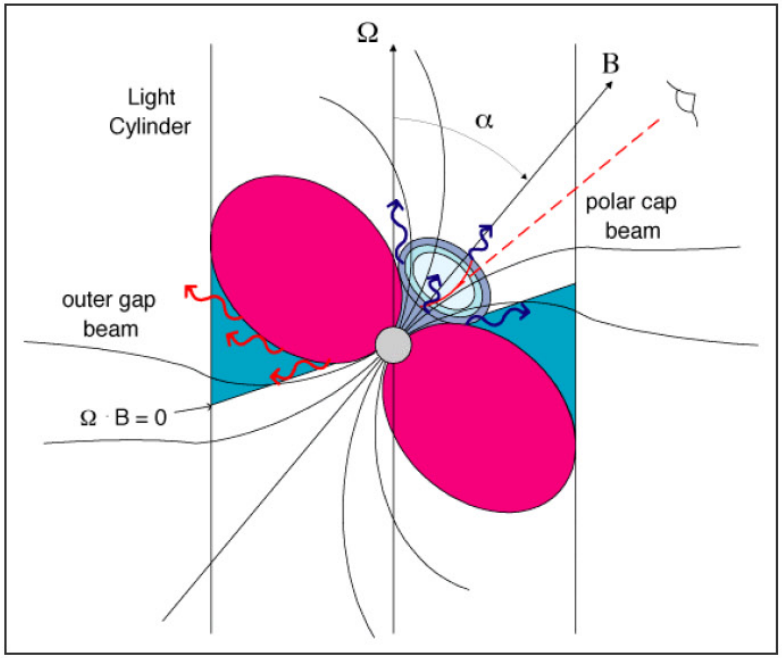
$$B_{NS} = B\left(\frac{R}{R_{NS}}\right)^2 \cong 10^{-2}T \cdot 10^{10} \cong 10^8 T$$

Power Source: Pulsar



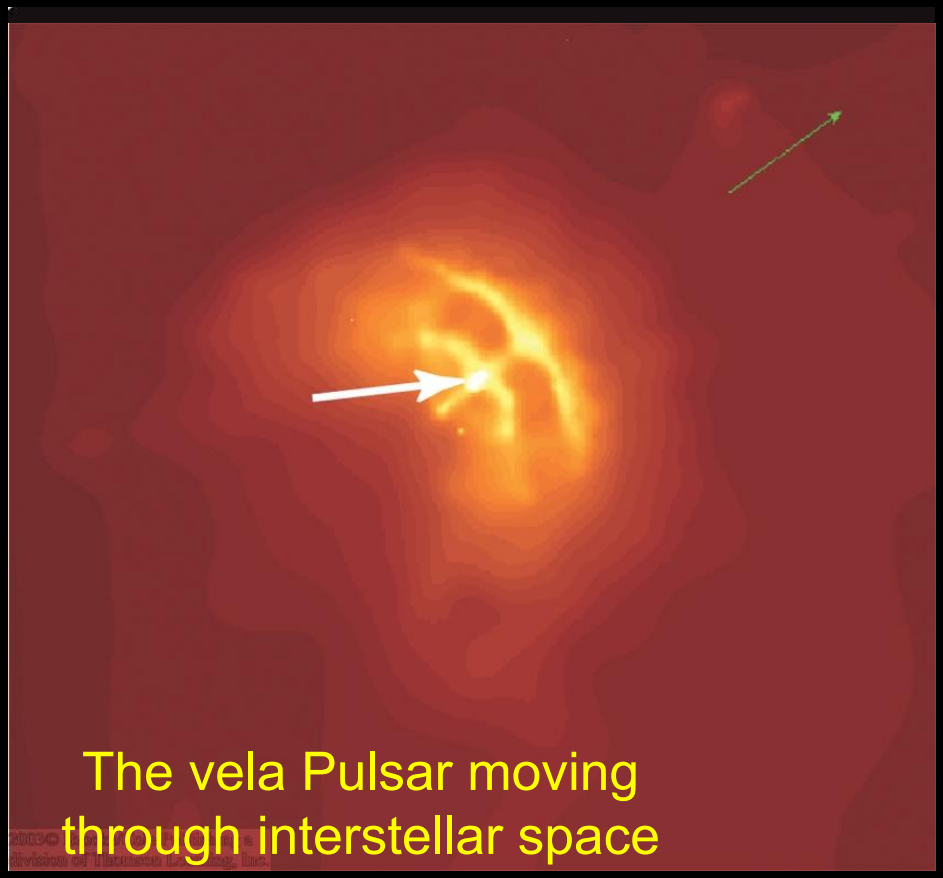
Crab Nebula

Pulsar

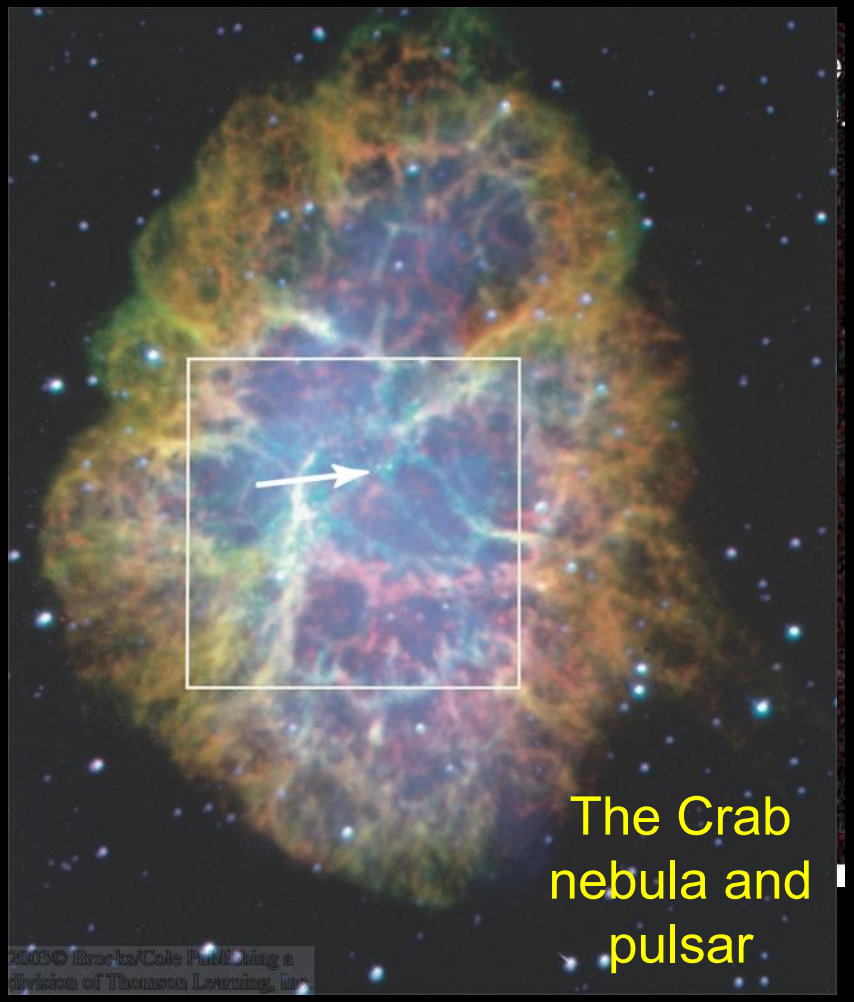


Emitted beams

Images of Pulsars and other Neutron Stars

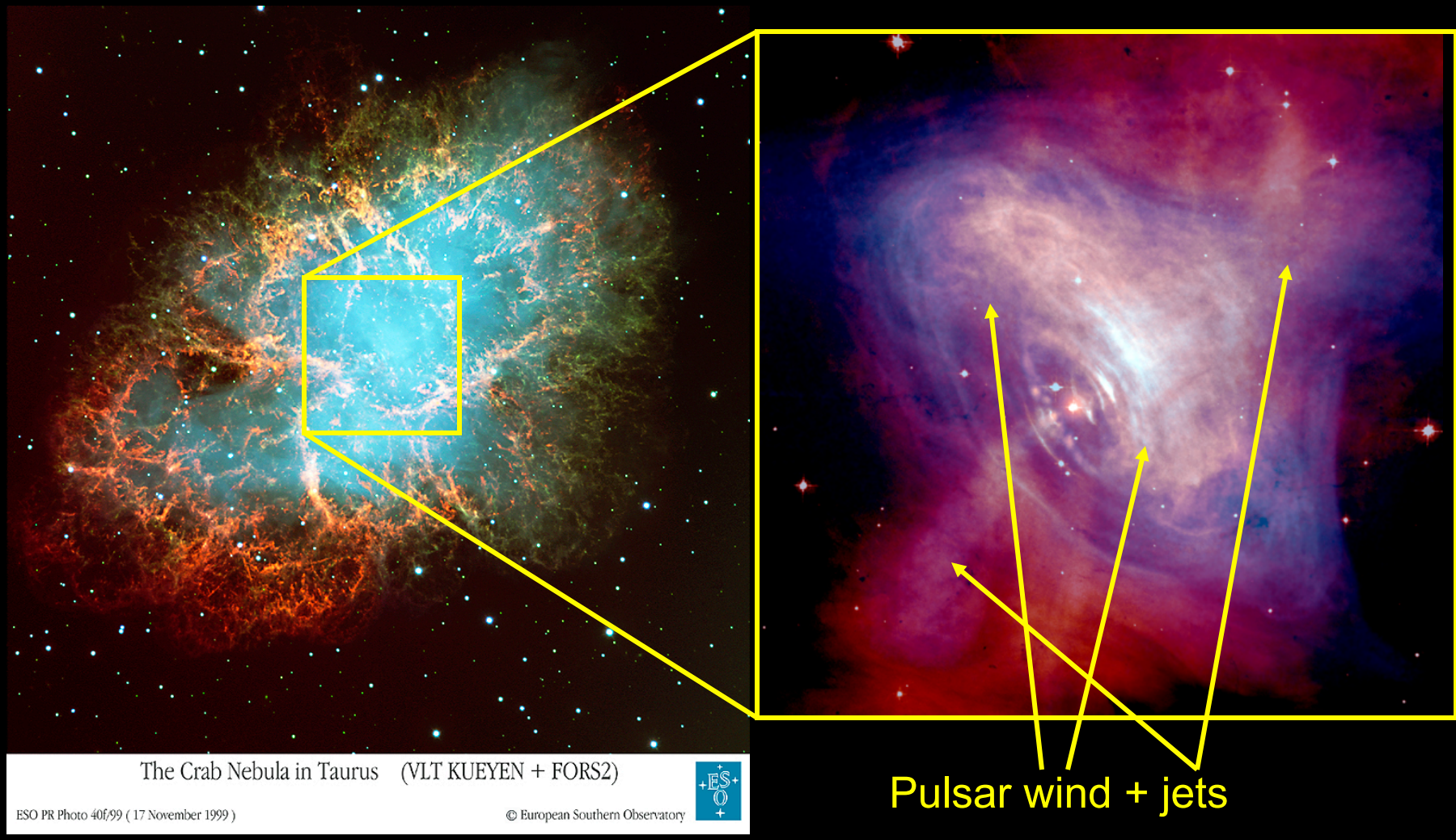


The vela Pulsar moving through interstellar space



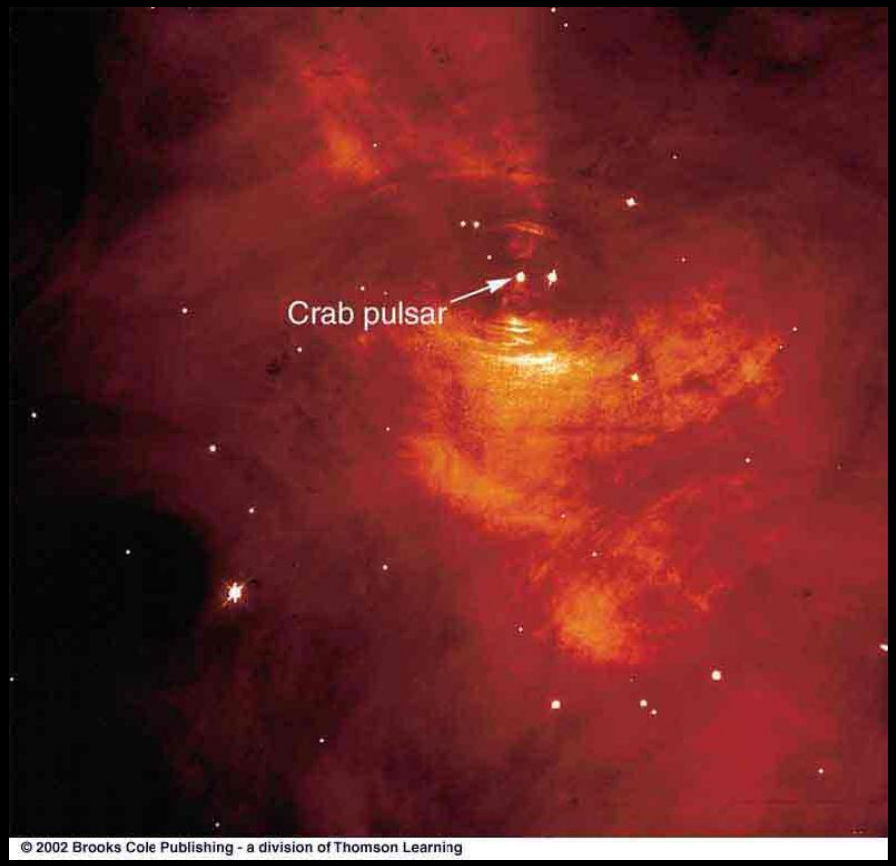
The Crab nebula and pulsar

The Crab Pulsar

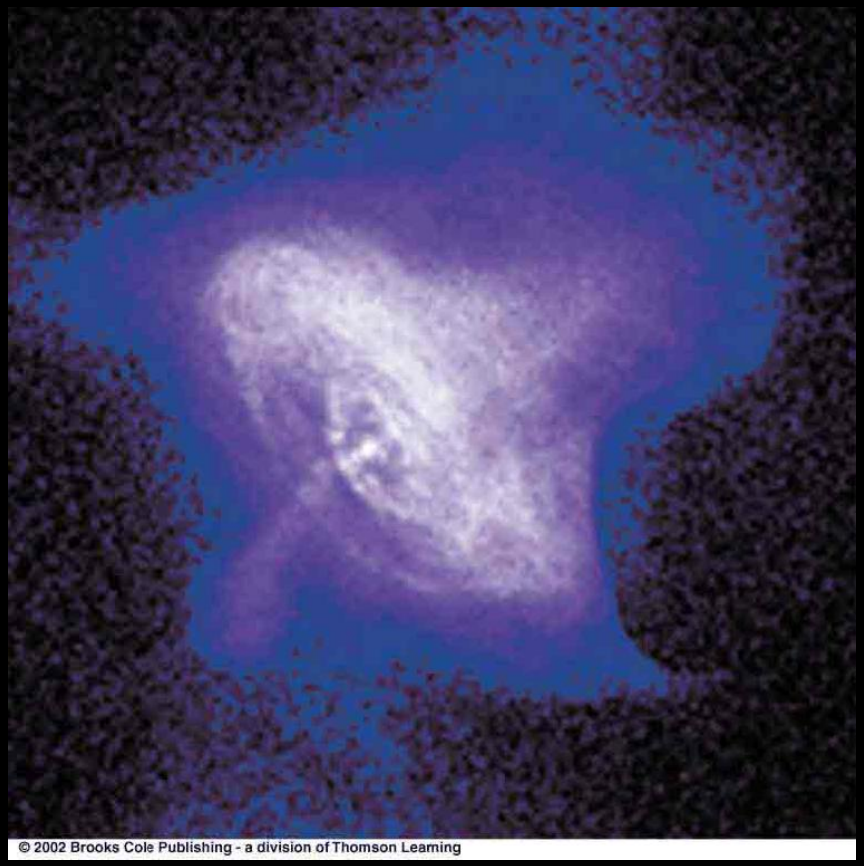


Remnant of a supernova observed in A.D. 1054

The Crab Pulsar



Visual image



X-ray image

Properties of Neutron Stars

Typical size: R ~ 10 km

Mass: M ~ 1.4 – 3 M_{sun}

Density: ρ ~ 4×10^{14} g/cm^3

→ 1 teaspoon full of NS matter has a mass of ~ 2 billion tons!!!

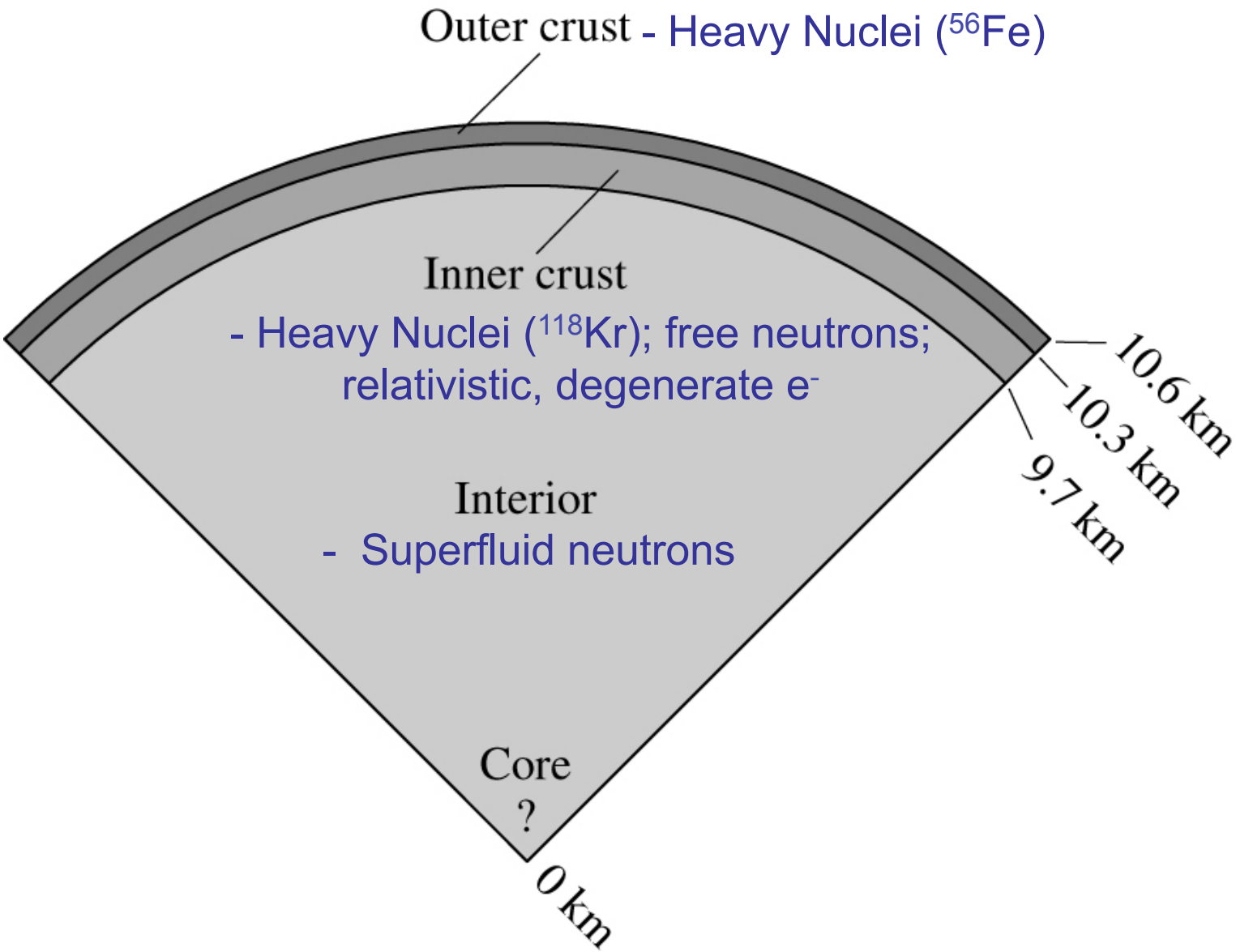
Rotation periods: ~ a few ms – a few s

Magnetic fields: B ~ 10^8 – 10^{15} G

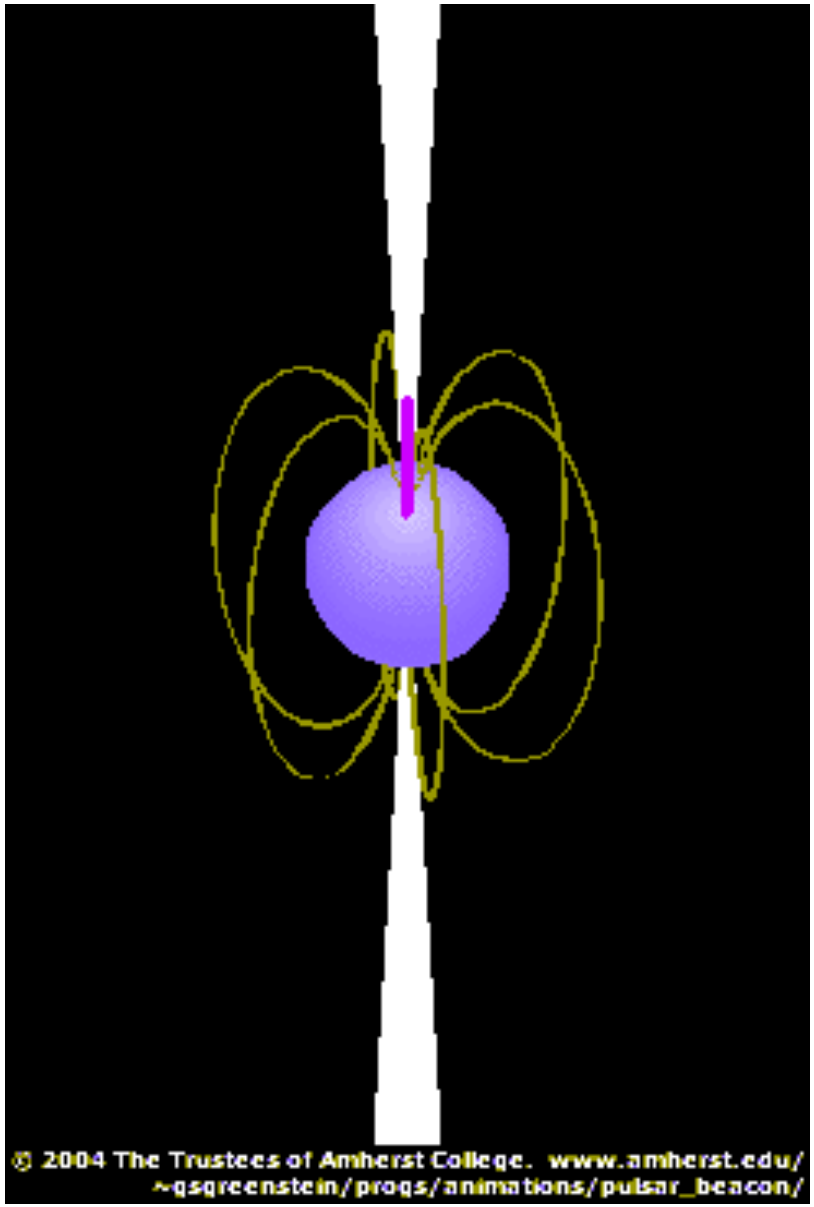
(Atoll sources; ms pulsars)

(magnetars)

Radial Structure of a Neutron Star



The Lighthouse Model of Pulsars



A Pulsar's magnetic field has a dipole structure, just like Earth.

Radiation is emitted mostly along the magnetic poles.

Rapid rotation along axis not aligned with magnetic field axis

→ Light house model of pulsars

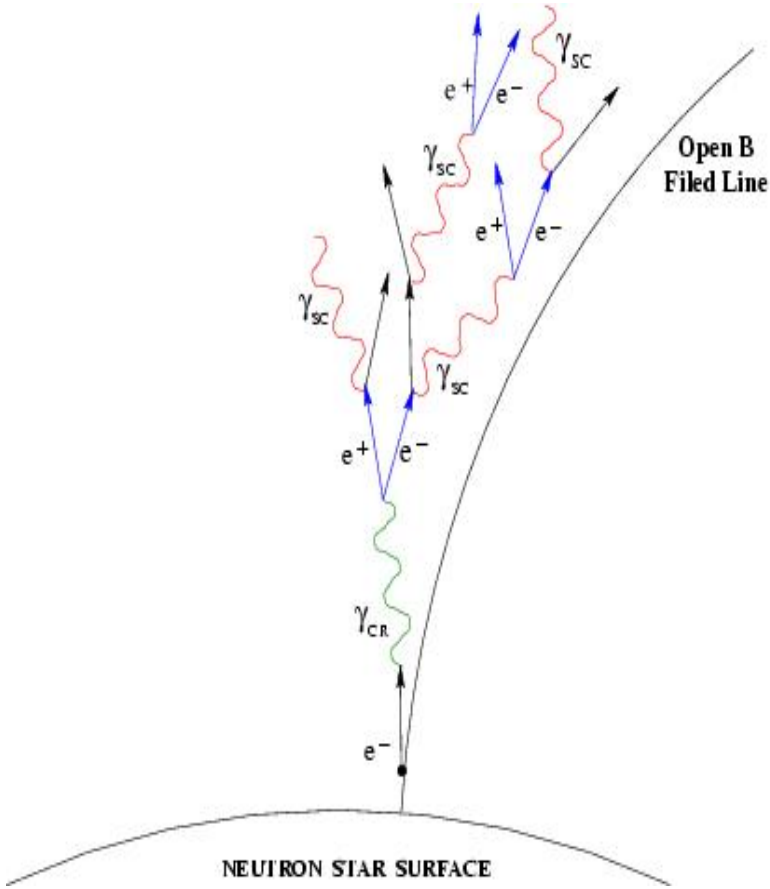
Pulses are not perfectly regular

→ gradual build-up of average pulse profiles

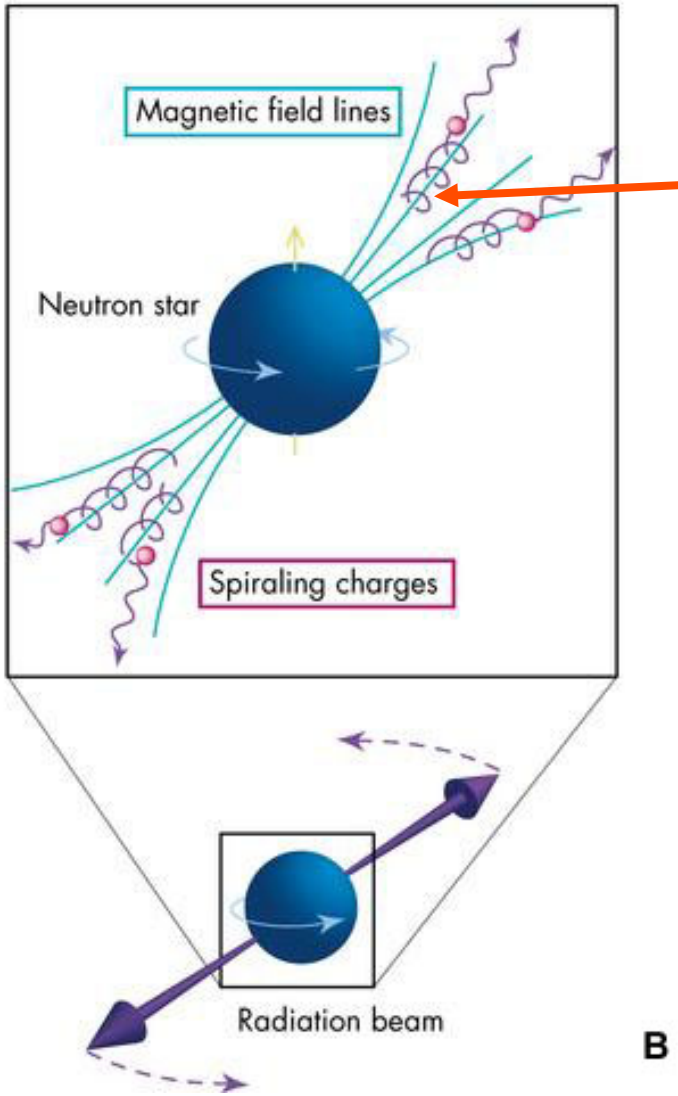
e+e- production in neutron stars

Mechanism: the spinning B of the pulsar strips e^-, that are accelerated at the polar cap or at the outer gap, emit γ that make production of $e^\pm$ that are trapped in the cloud, further accelerated and later released at $\tau \sim 10^5$ years.

In both models of polar gap and outer gap, electrons can be accelerated in different regions of the pulsar magnetosphere due to induced electric fields and produce an electromagnetic cascade through the emission of curvature radiation, which in turn results in production of photons which are above threshold for pair production in the strong pulsar magnetic field. This process results in lower energy electrons and positrons that can escape the magnetosphere either through the open field lines [25] or after joining the pulsar wind. In this second case, the electrons and positrons lose part of their energy adiabatically because of the expansion of the wind.



Pulsar Emission Models:Polar Cap model



Particle acceleration along magnetic field lines

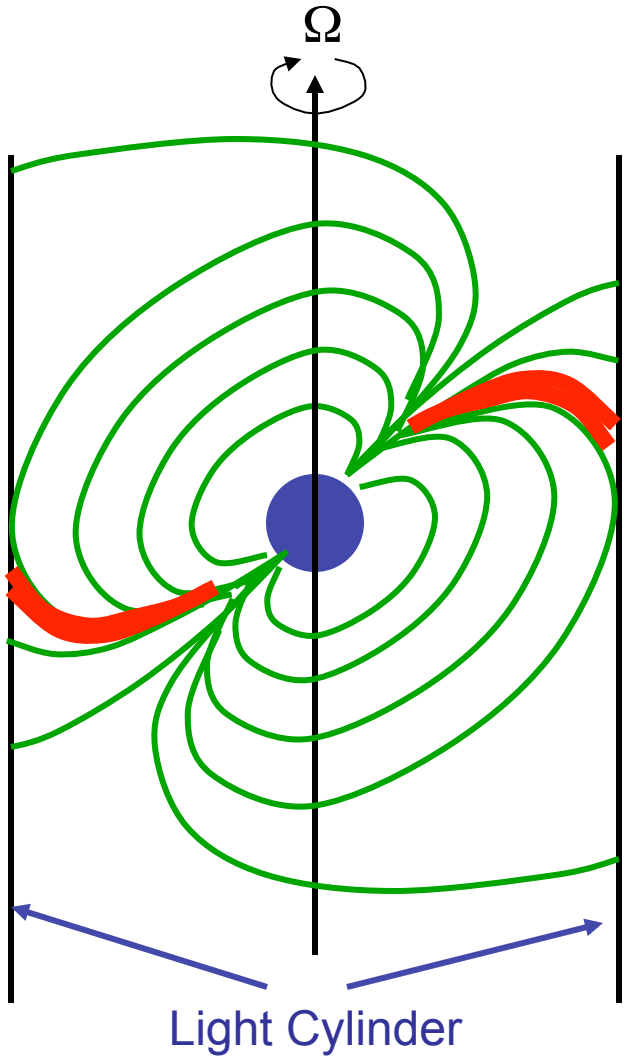
Synchrotron emission

Curvature radiation

Pair production

Electromagnetic cascades

Pulsar Emission Models: Outer Gap model



Electrons are bound to magnetic fields co-rotating with the pulsar

At a radial distance $r = c/\Omega$ co-rotation at the speed of light → "light cylinder"

→ Particles ripped off magnetic fields

Synchrotron emission

Curvature radiation

- La velocità di rotazione delle pulsar è nota sperimentalmente, e può essere facilmente stimata:

$$\frac{GM^2}{R^2} = M\omega^2{}_{NS}R \longrightarrow \omega_{NS} \approx 10^3 - 10^4 s^{-1}$$

La Pulsar fornisce energia tramite induzione EM.

L'energia massima fornita ad una particella di carica Ze:

$$\nabla \times \varepsilon = -\frac{1}{c}\frac{\partial B}{\partial t} \longrightarrow \frac{\varepsilon}{L} = \frac{B\omega_o}{c}$$

(ε = campo elettrico indotto su una regione lineare L)

$$E_{Max} = \int Ze \cdot \varepsilon \cdot dx = Ze \cdot B\omega_o L \cdot L / c$$

PULSAR: Massima energia acquisita dalla particella accelerata nel caso di p (Z=1), B=10^8 T, ω=10^3 s^{-1}, R=10 km

$$E_{Max} = Ze \cdot B\omega_o L^2 / c =$$

$$= 1.6 \times 10^{-19}(C) \cdot 10^8(T) \cdot 10^3(s^{-1}) \cdot \left[10^4(m)\right]^2$$

$$= 1.6(J) = 10^{19} eV$$

.

The energy spectrum injected by a single pulsar depends on the environmental parameters of the pulsar, but some attempts to calculate the average spectrum injected by a population of mature pulsars suggest that the spectrum may be relatively hard, having a slope of ~1.5- 1.6.

This spectrum, however, results from a complex interplay of individual pulsar spectra, of the spatial and age distributions of pulsars in the Galaxy, and on the assumption that the chief channel for pulsar spin down is magnetic dipole radiation. Due to the related uncertainties, variations from this injection spectra cannot be ruled out.

$$\frac{dN_e}{dE_e} \approx 8.6 \times 10^{38} \dot{N}_{100} \left(E_e/\mathrm{GeV}\right)^{-1.6} \exp\left(-E_e/\ E_c\ GeV\right) \mathrm{GeV}^{-1}\ \mathrm{s}^{-1}\ ,$$

With E_c ~ 10-1000 GeV and N_{100} is rate of pulsar generation per century

This expression corresponds to an average energy output in electron-positron pairs of approximately 6×10^{46} erg per pulsar, to be compared eith the energy injected by the pulsar of the order of 10^{49} erg

- For details of derivation: Cosmic-ray positrons from mature gamma-ray pulsars, L. Zhang And K.S. Cheng, Astronomy&Astrophysics, 368, 1063-1070 (2001),

Pulsars as the sources of high energy cosmic ray positrons, Dan Hooper,a,b Pasquale Blasia,c,d and Pasquale Dario Serpico, Journal of cosmology and astrophysics, 2009

- This cutoff is determined by the details of the development of the electromagnetic cascade in the pulsar magnetosphere and, even more importantly, by the distribution of periods, magnetic fields, and radii of mature pulsars. The exact value of the cutoff energy should, therefore, not be considered to be a robust prediction of the theory, although it represents a good estimate of the order of magnitude of the cutoff energy. For instance, in some models it is argued that the typical energy of electrons and positrons in the cascade associated with the polar gap is $E_e \approx 8B_{12}^{5/7} P_o^{-17/7}$ GeV which, for a period of 200 ms, would yield a maximum energy of the resulting pairs of ~ 400 GeV. On the other hand, other models lead to different maximum energies (~80 GeV) and even different scalings of the cutoff with the parameters of the pulsar.

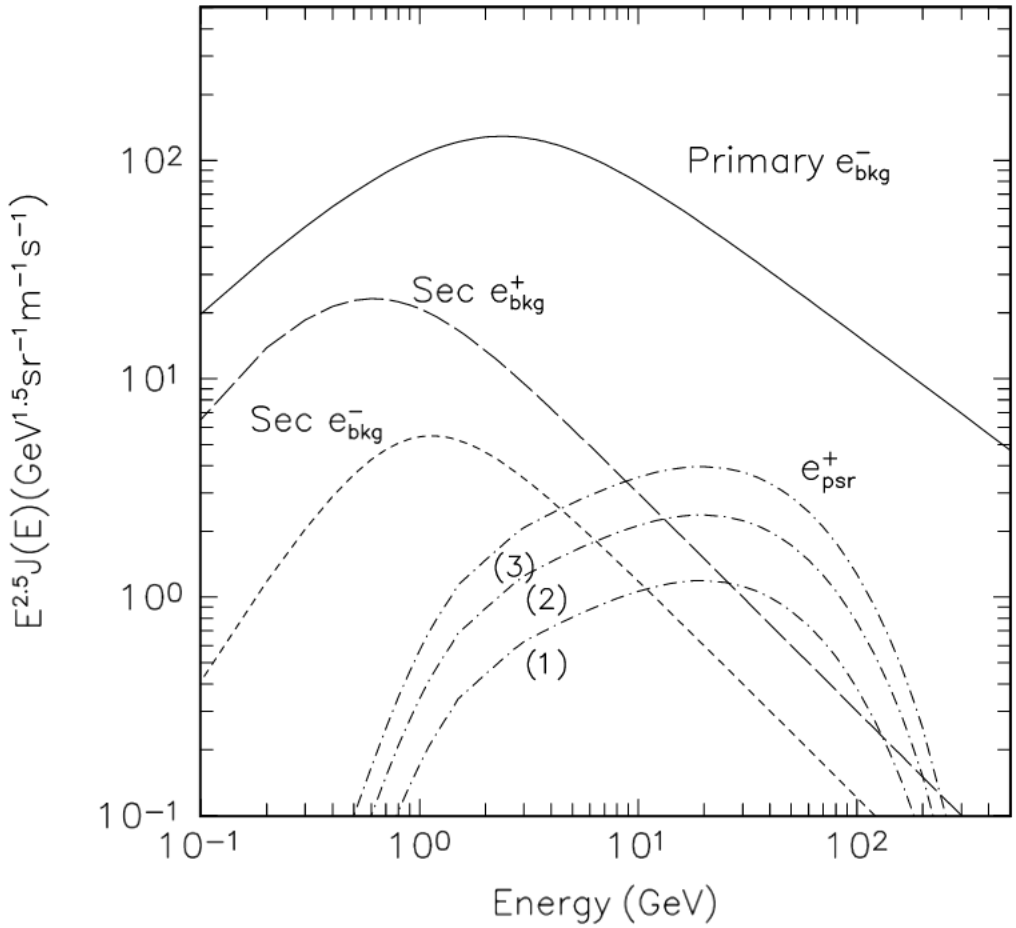


Fig. 2. The spectra of cosmic-ray electrons and positrons. The background of cosmic-ray electrons and positrons are approximated by those given by Moskalenko & Strong (1998): primary e^-_{bkg} (solid curve), secondary e^-_{bkg} (short-dashed curve) and secondary e^+_{bkg} (long-dashed curve). The positron spectra from the mature γ-ray pulsars for various birth rate of the neutron stars in the Galaxy are labeled as (1) ($\dot{N}_{100} = 1$), (2) ($\dot{N}_{100} = 2$) and (3) ($\dot{N}_{100} = 10/3$)

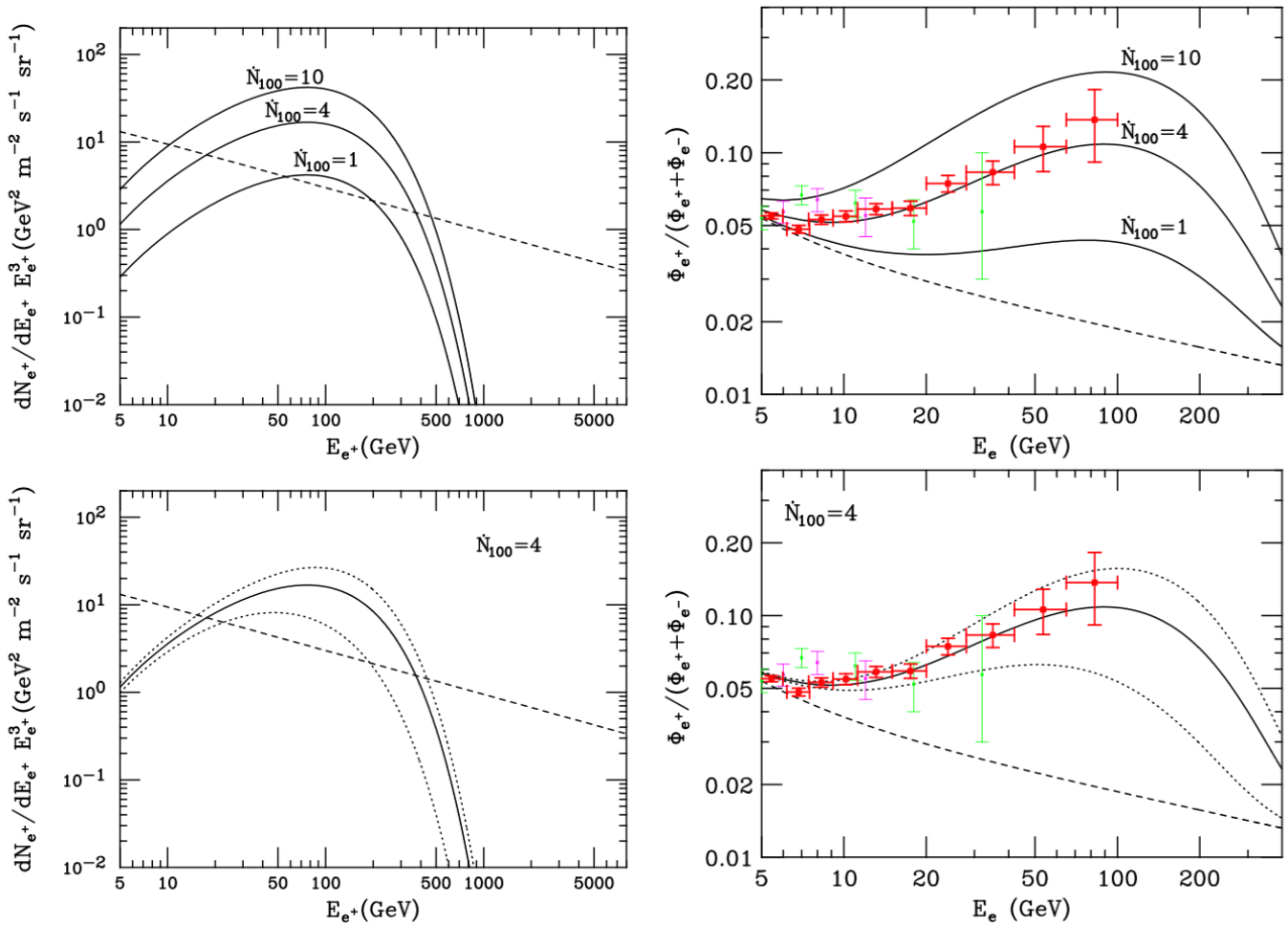


Figure 1: The spectrum of cosmic ray positrons (left) and the positron fraction (right) resulting from the sum of all pulsars throughout the Milky Way. Also shown as a dashed line is the prediction for secondary positrons (and primary and secondary electrons in the right frames) as calculated in ref. [27]. In the right frames, the measurements of HEAT [3] (light green and magenta) and measurements of PAMELA [2] (dark red) are also shown. We have used the injected spectrum reported in eq. (2.7). In the lower frames, the upper (lower) dotted line represents the case in which the injection rate within 500 parsecs of the Solar System is doubled (neglected), providing an estimate the variance resulting from the small number of nearby pulsars contributing to the spectrum.

Case study: Crab nebula

- La luminosita' totale della Crab e' L=5x 10^{38} $ergs^{-1}$

- Da dove viene l'energia irraggiata?

- Se si considera la pulsar come una sfera rigida di raggio R»10 km e massa M=1 massa solare=1.988x10^{33} gr, l'energia rotazionale e'

$$I = \frac{3}{5}MR^2 = 1.2 \times 10^{45}\ grcm^2 \qquad E = \frac{1}{2}I\omega^2 = 4.8 \times 10^{49}\ erg$$

- La frequenza di rotazione misurata e' ω = 190 s^{-1}

- Il tasso di diminuzione di velocita' angolare $\frac{d\omega}{dt} = -2.4 \times 10^{-9}\ s^{-2}$

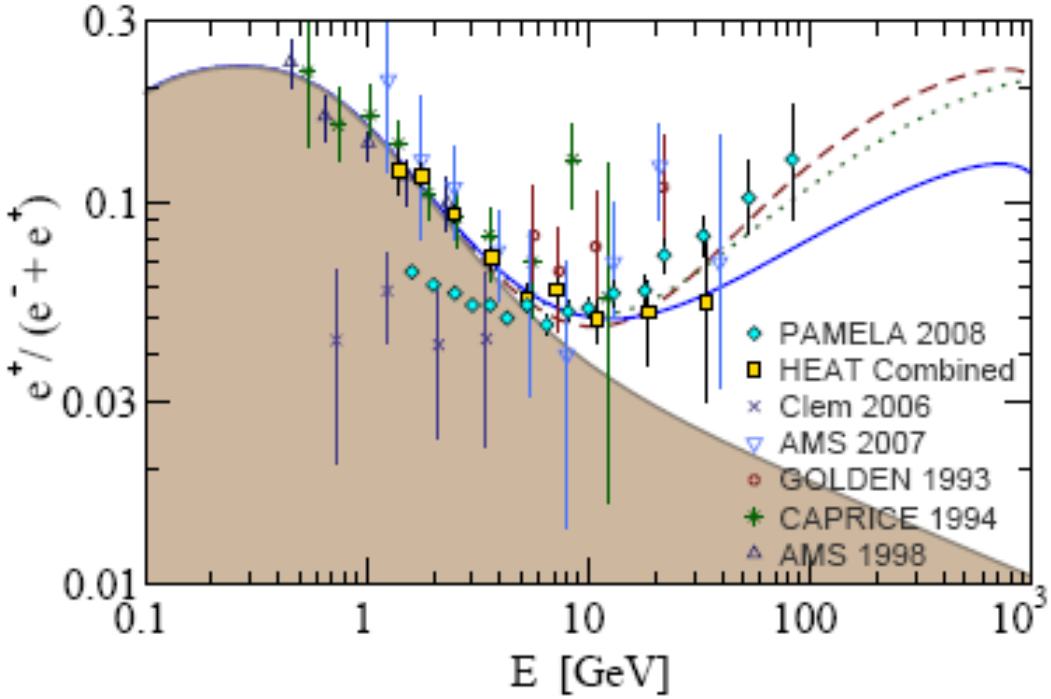
- L'energia dissipata (che viene trasferita al mezzo circostante) e' quindi $\frac{dE}{dt} = I\omega\frac{d\omega}{dt}$

- tramite processi che coinvolgono campi magnetici molto intensi e rapidamente variabili in prossimita' della superficie della pulsar

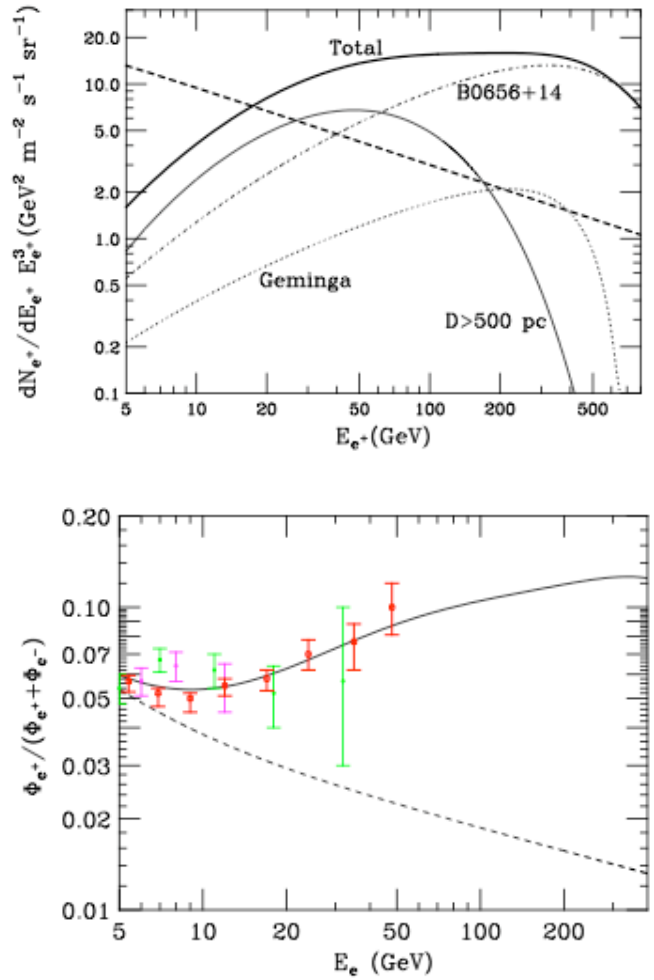
$$= (1.2\times10^{45})\times(190)\times(2.4\times10^{-9})\ ergs^{-1} \approx 5.5\times10^{38}\ ergs^{-1}$$

- Valore molto vicino alla potenza totale irraggiata dalla pulsar → la sorgente di energia e' proprio la rotazione della pulsar che rallenta il suo moto

Example: pulsars



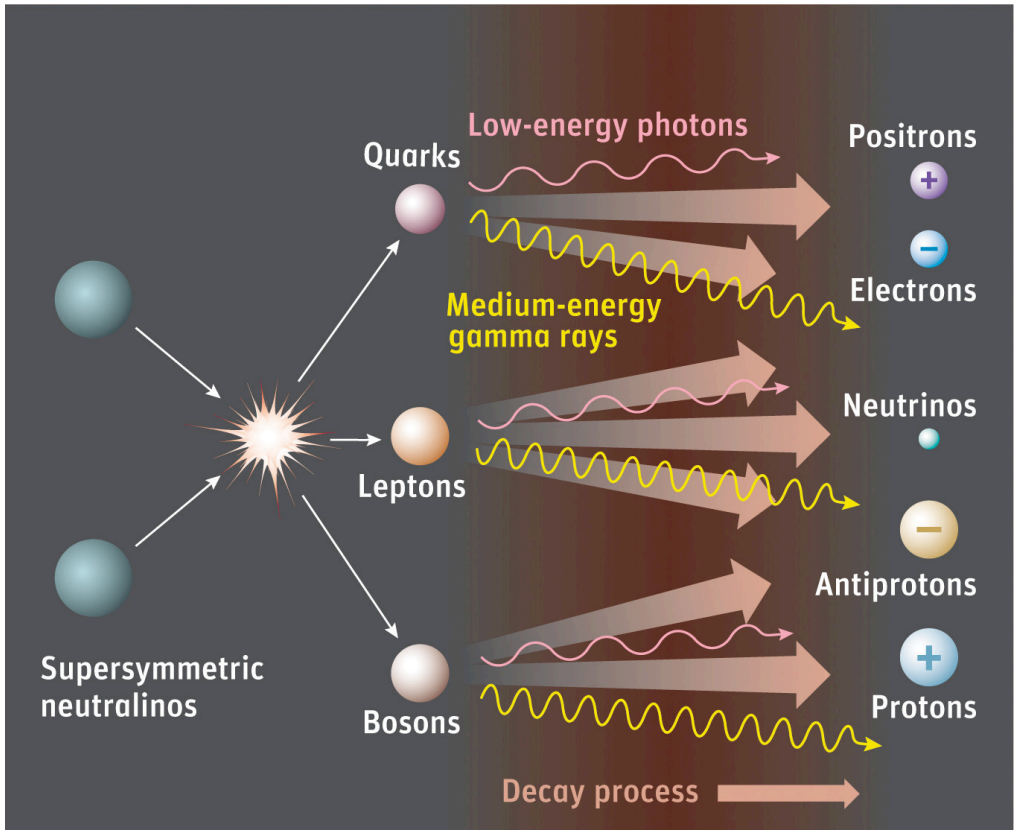
H. Yüksak et al., arXiv:0810.2784v2
Contributions of e- & e+ from Geminga assuming different distance, age and energetic of the pulsar



Hooper, Blasi, and Serpico
arXiv:0810.1527

Dark Matter

- Dark matter in the halo annihilates producing a cascade of standard particles, as p, e+,e-, gammas, neutrinos, anti-p and so on
- Depends on: halo DM distribution (unknown) rates of decays of different channels of DM particles into standard particles (SUSY, Kaluza-Klein,…)



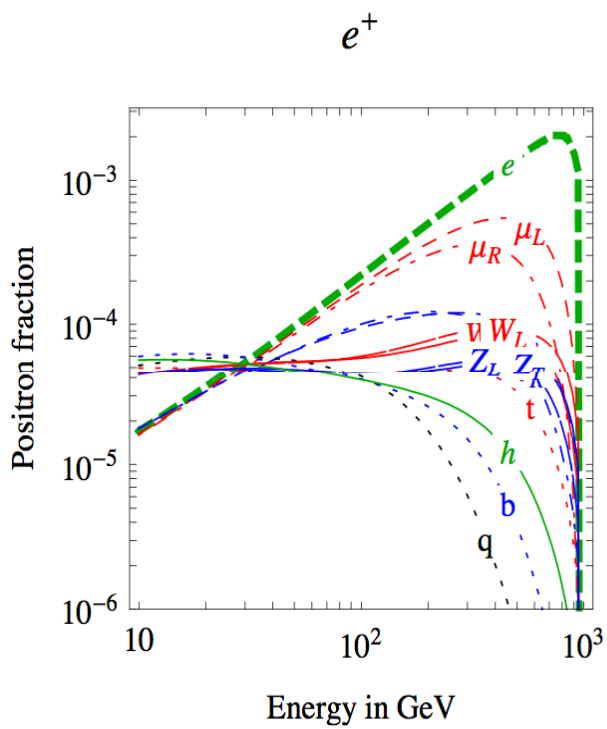
DM annihilations

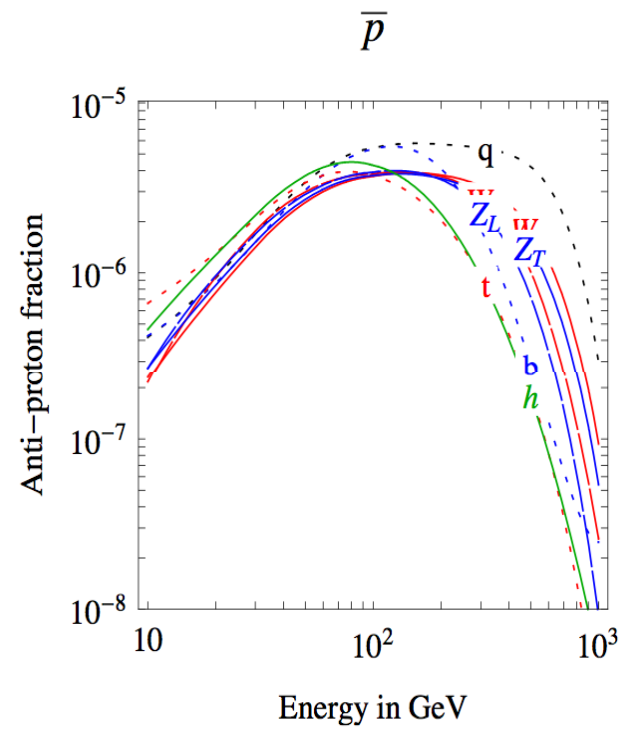
Resulting spectrum for positrons and antiprotons M= 1 TeV

The flux shape is completely determined by:

1) WIMP mass

2) Annihilations channels

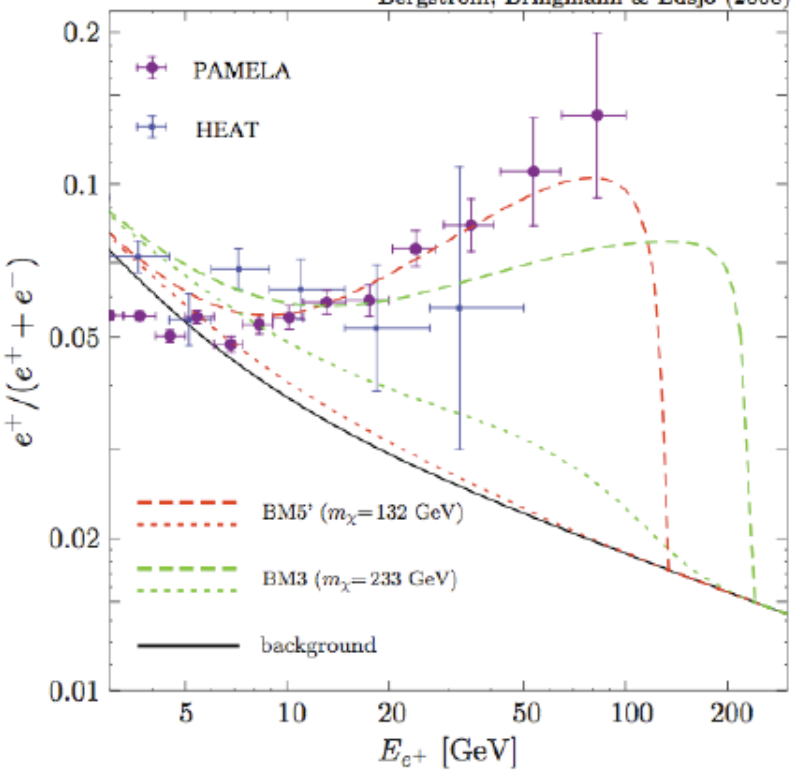




Example: Dark Matter

arXiv:0808.3725

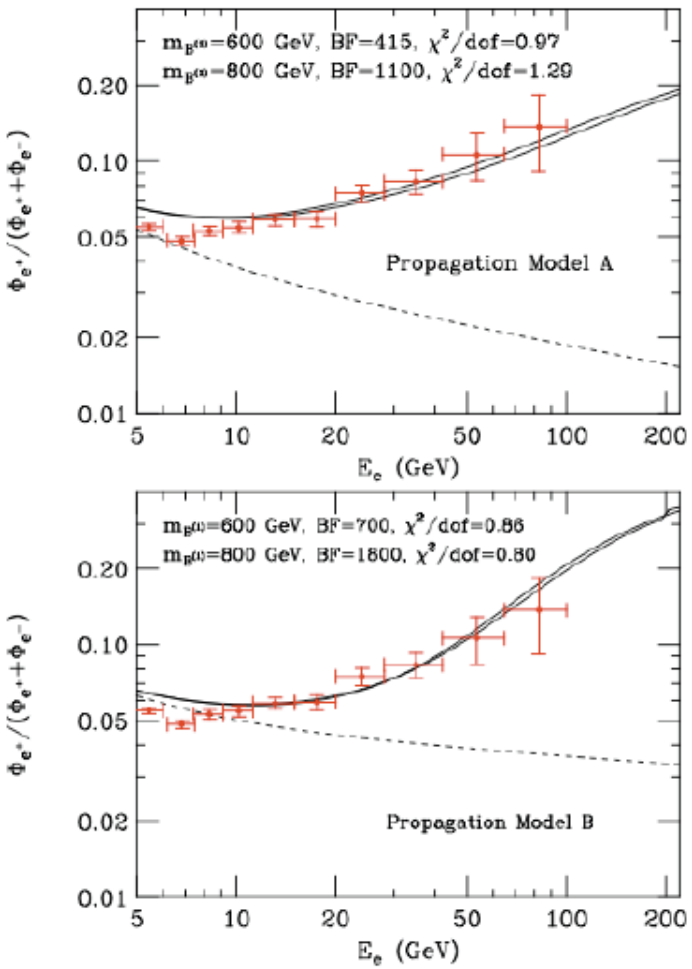
Bergström, Bringmann & Edsjö (2008)



Majorana DM with **new** internal bremsstrahlung correction. NB: requires annihilation cross-section to be 'boosted' by >1000.

Hooper and Zurek

arXiv:0902.0593v1



Kaluza-Klein dark matter