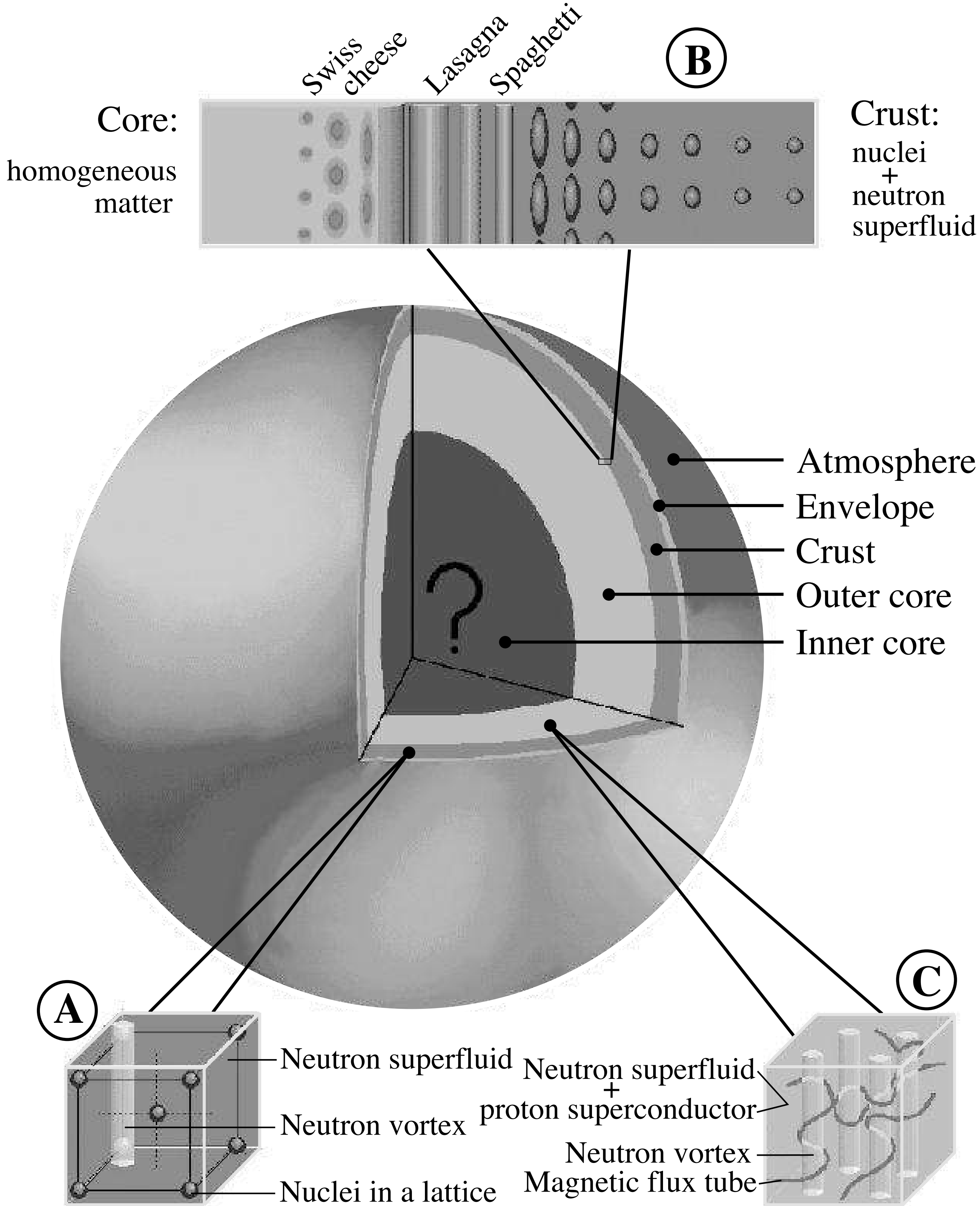
Cooling of Neutron Stars

Dany Page

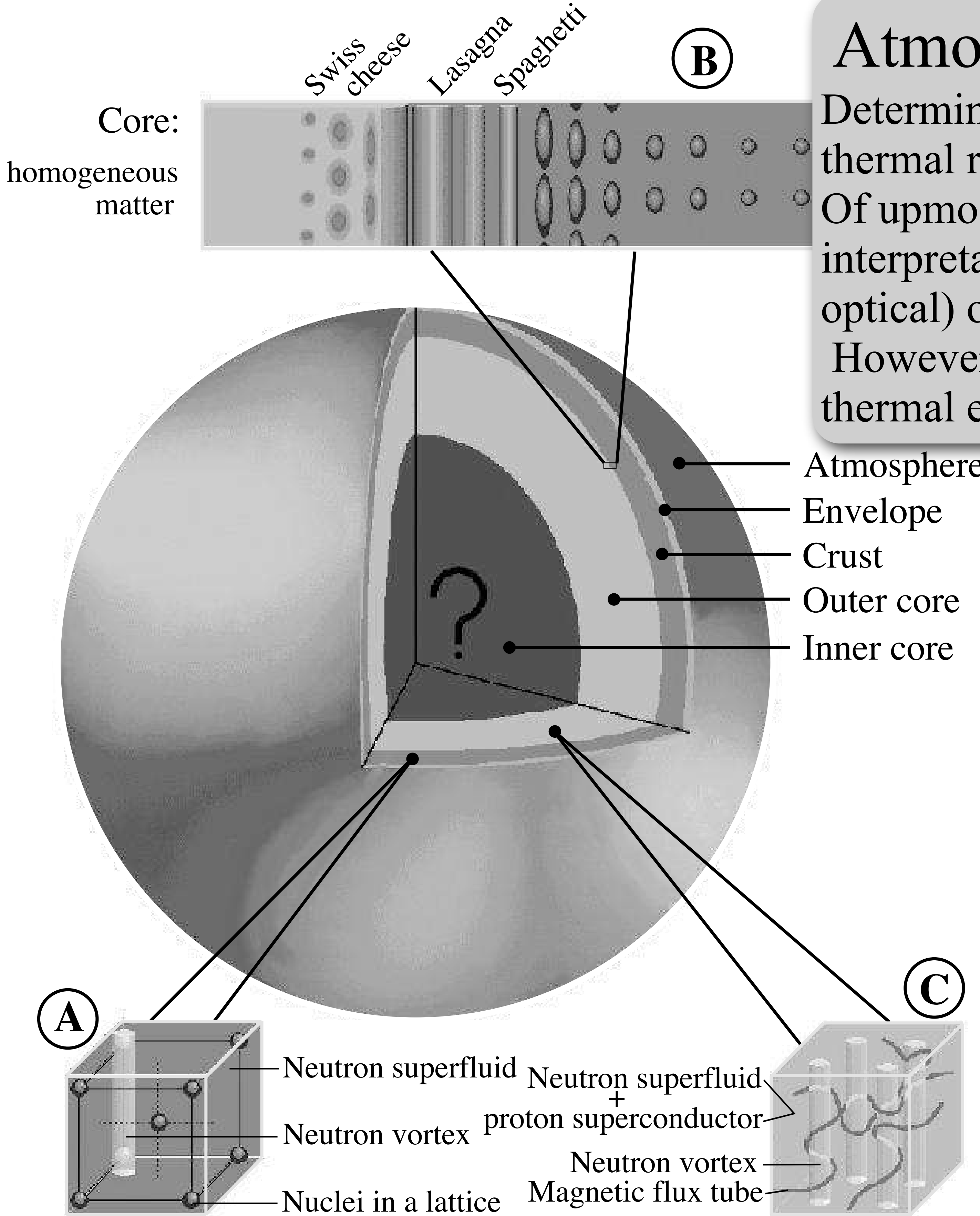
Instituto de Astronomía
Universidad Nacional Autónoma de México



Dense Matter in Compact Stars: Theoretical Developments and Observational Constraints, Page D. & Reddy S., 2006, Annu. Rev. Nucl. Part. Sci. 56, 327

Atmosphere (10 cm):

Determines the shape of the thermal radiation (the spectrum).
Of upmost importance for interpretation of X-ray (and optical) observation.
However it as NO effect on the thermal evolution of the star.



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Envelope (100 m):

Contains a huge temperature gradient: it determines the relationship between T_{int} and T_e. Extremely important for the cooling, strongly affected by magnetic fields and the presence of "polluting" light elements.

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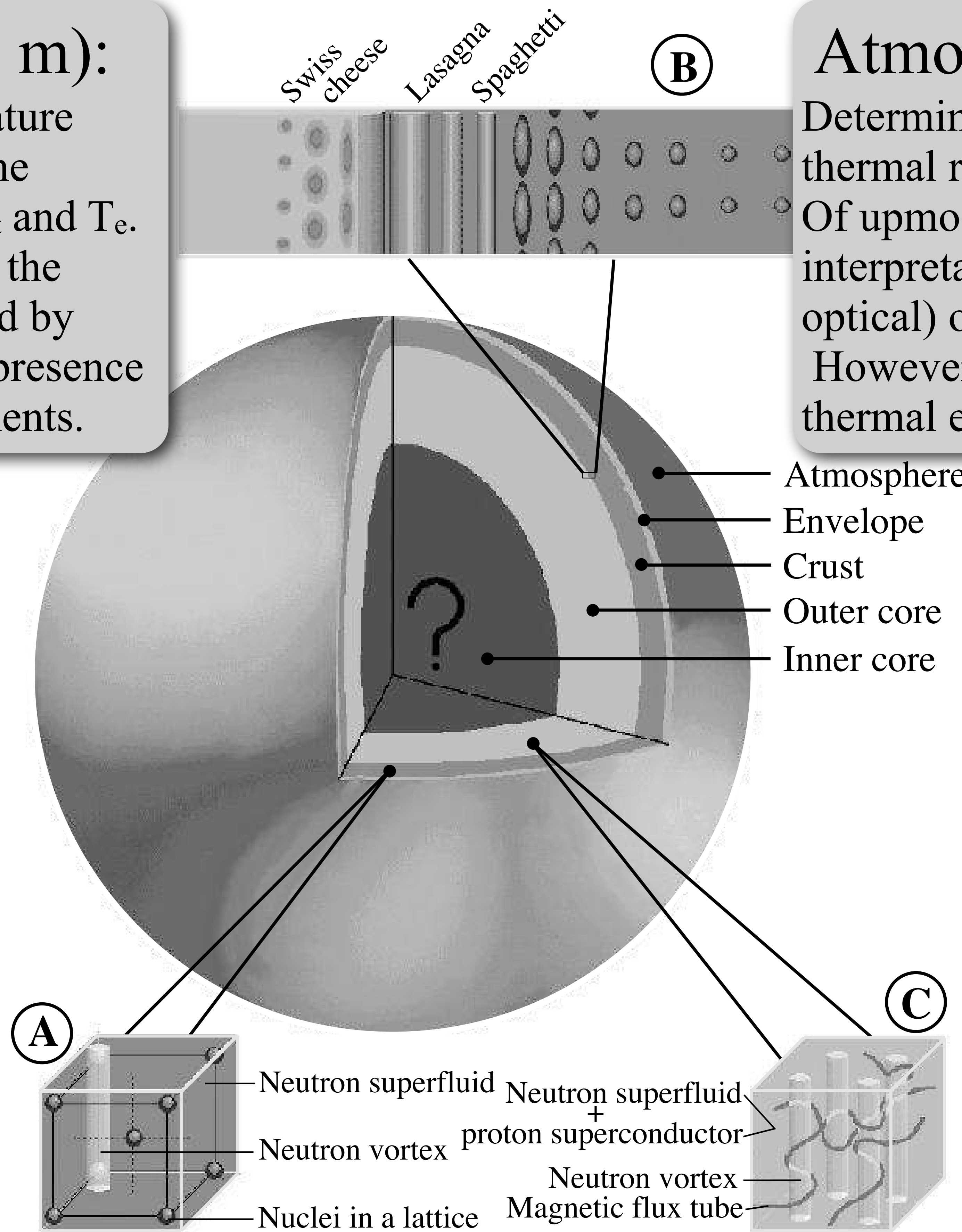

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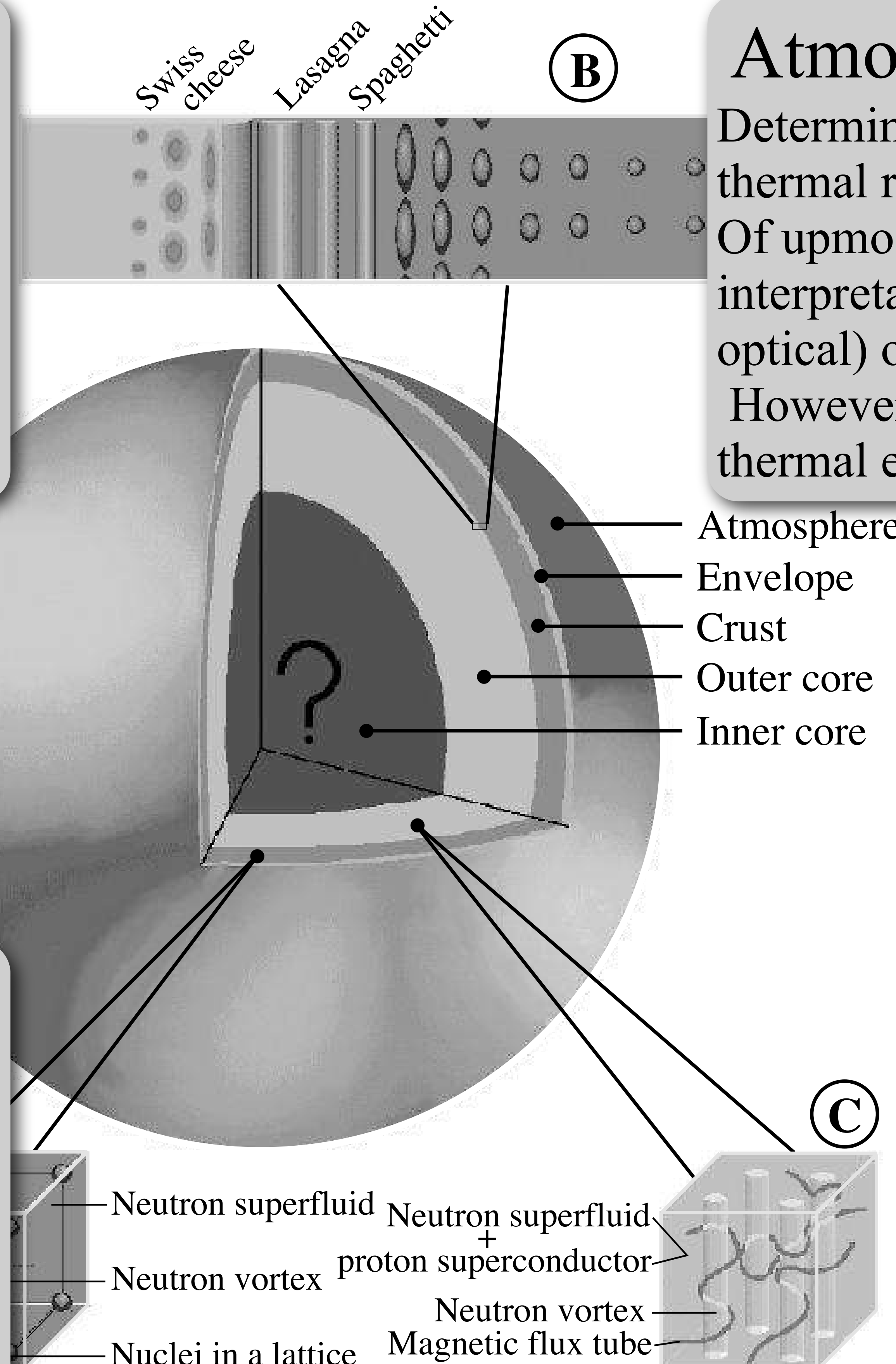

Crust (1 km):

Little effect on the long term cooling. BUT: may contain heating sources (magnetic/ rotational, pycnonuclear under accretion). Its thermal time is important for very young star and for quasi-persistent accretion

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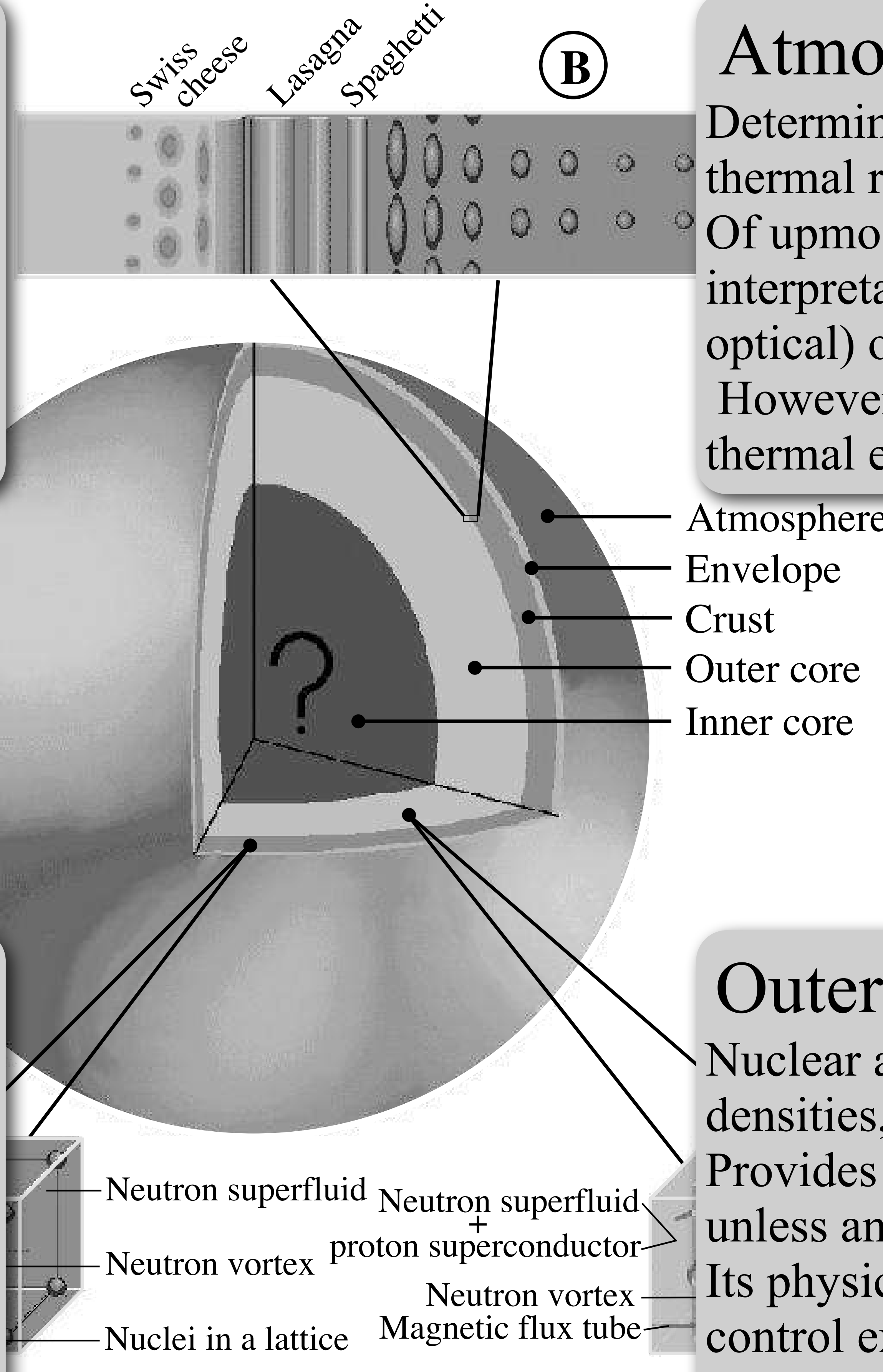

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Outer Core (10-x km):

Nuclear and supranuclear densities, containing n, p, e & μ. Provides about 90% of c_v and ε_v unless an inner core is present. Its physics is basically under control except pairing T_c which is essentially unknown.

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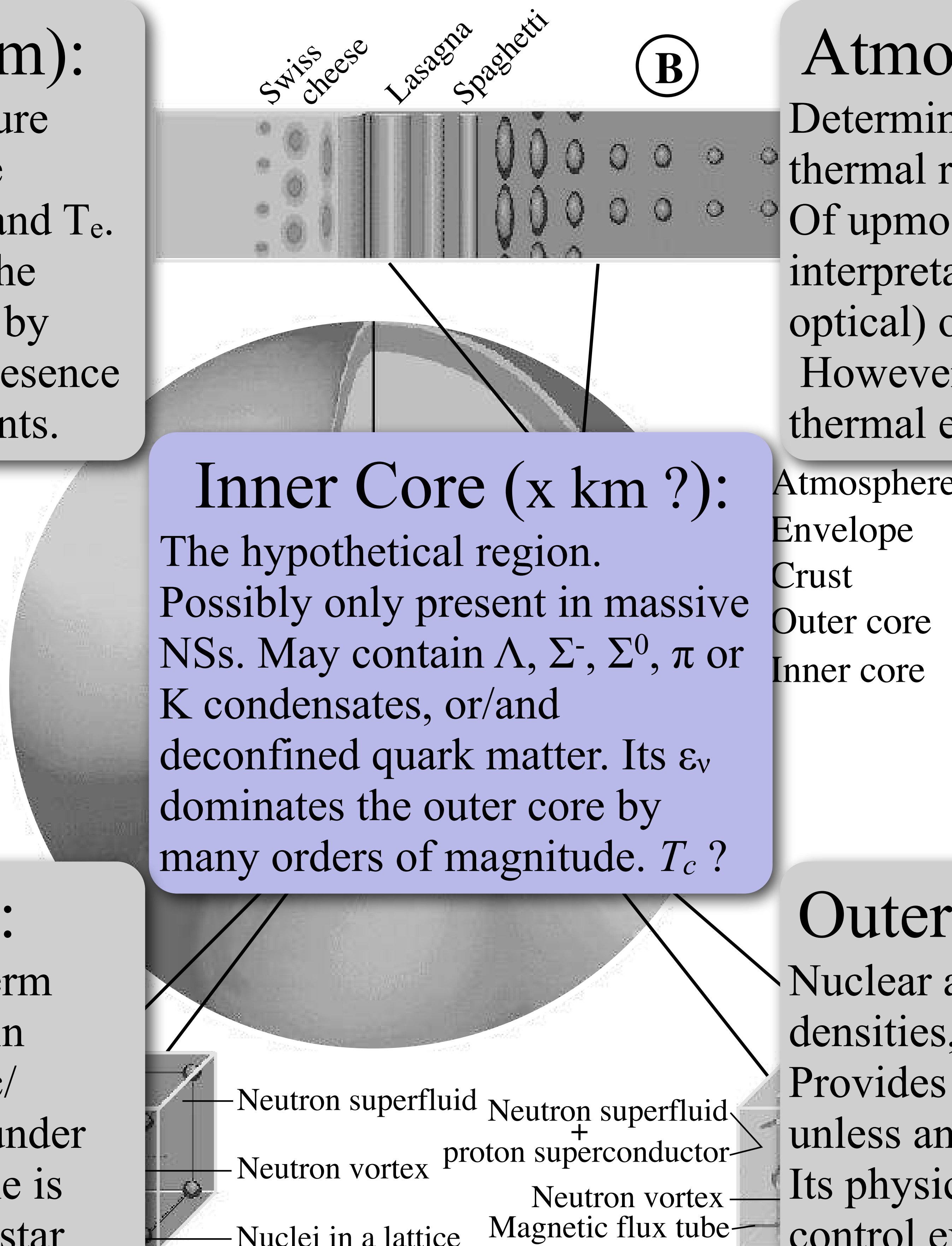

Inner Core (x km ?):

The hypothetical region. Possibly only present in massive NSs. May contain Λ, Σ^-, Σ^0, π or K condensates, or/and deconfined quark matter. Its ε_ν dominates the outer core by many orders of magnitude. T_c ?

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Neutron star cooling on a napkin

Assume the star's interior is isothermal and neglect GR effects.

Thermal Energy, E_{th} ,balance:

$$\frac{dE_{th}}{dt} = C_v \frac{dT}{dt} = -L_\gamma - L_\nu + H$$

⇒ 3 essential ingredients are needed:

- C_v = total stellar specific heat
- L_γ = total surface photon luminosity
- L_ν = total stellar neutrino luminosity

H = "heating", from B field decay, friction, etc ...

Specific Heat

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Specific heat on a napkin

Sum over all degenerate fermions: $C_V = \sum_i C_{V\ i}$ $c_{V\ i} = N(0) \frac{\pi^2}{3} k_B^2 T$ with $N(0) = \frac{m^* p_F}{\pi^2 \hbar^3}$

$$\boxed{C_v = \iiint c_v dV \simeq 10^{38} - 10^{39} \times T_9 \ \text{erg K}^{-1} \equiv C_9 T_9}$$

(lowest value corresponds to the case where extensive pairing of baryons in the core suppresses their c_v and only the leptons, e & μ, contribute)

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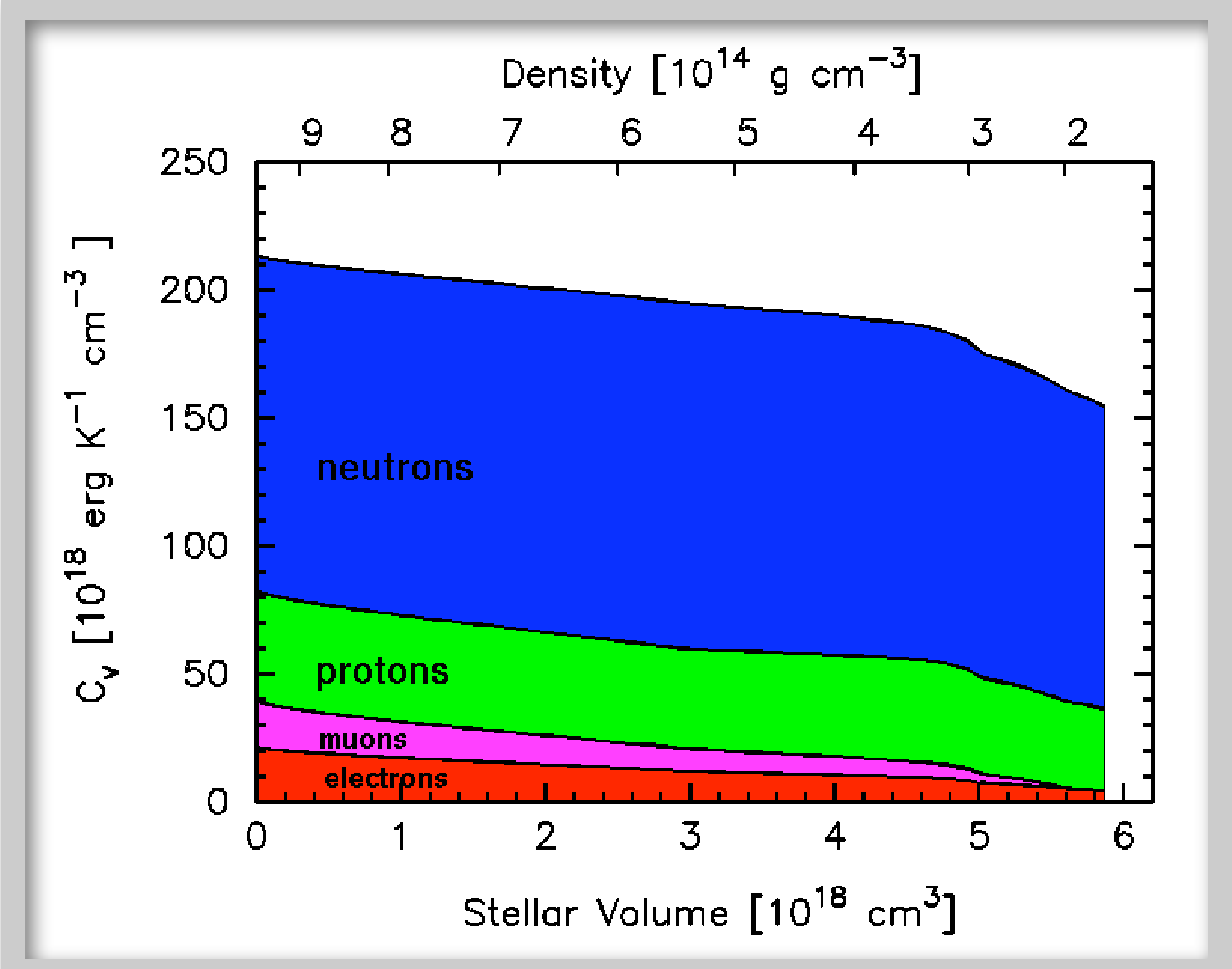

Distribution of c_V in the core of a 1.4 M_{Sun} neutron star build with the APR EOS (Akmal, Pandharipande, & Ravenhall, 1998), at

$T = 10^9$ K

Neutrinos

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The direct Urca process

Basic mechanism: β and inverse β decays:

$$n \longrightarrow p + e^- + \overline{\nu}_e \quad \text{and} \quad p + e^- \longrightarrow n + \nu_e$$

"Direct URCA process in neutron stars", JM Lattimer, CJ Pethick, M Prakash & P Haensel, 1991 PhRvL 66, 2701

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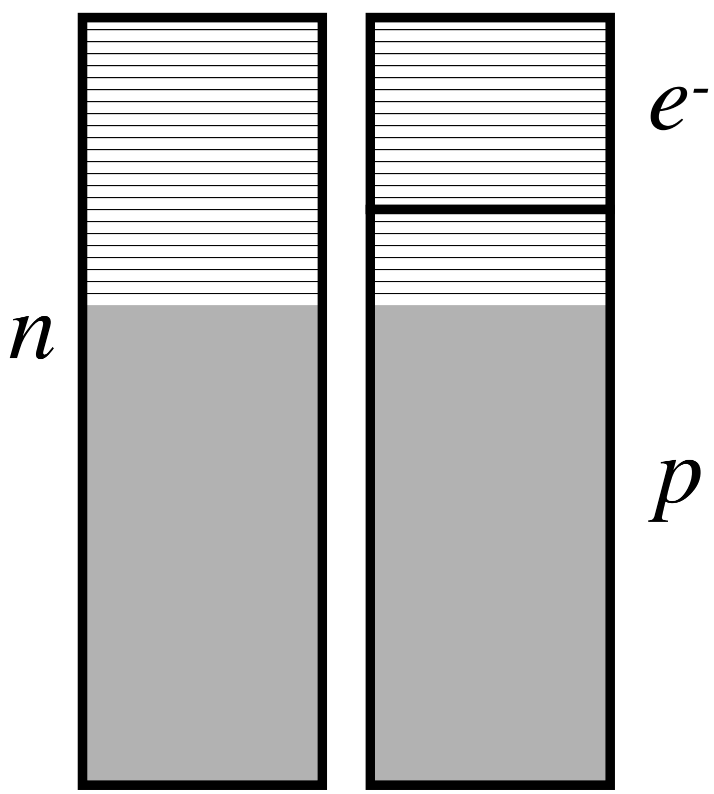
Energy conservation:

$E_{Fn} = E_{Fp} + E_{Fe}$



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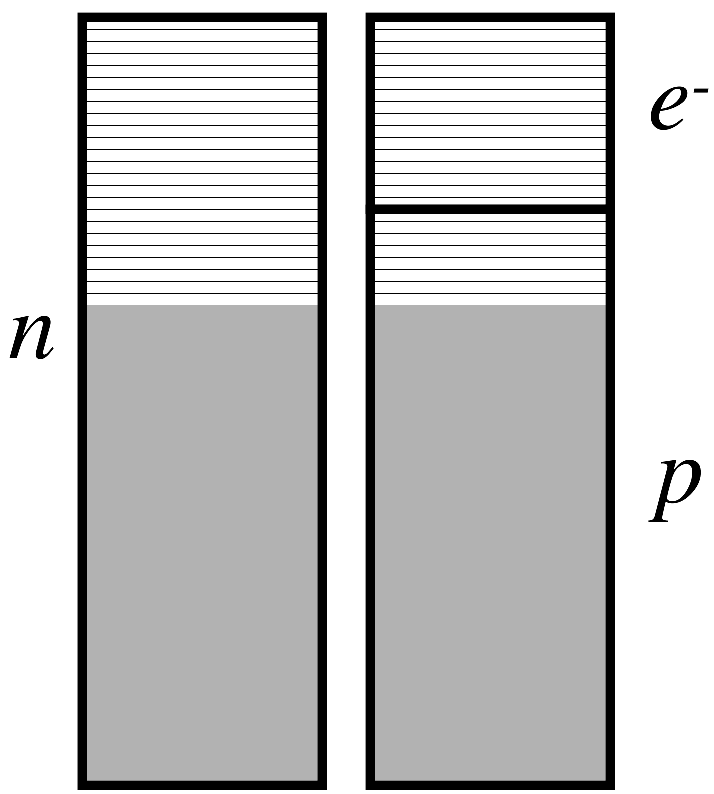

Momentum conservation:

"Triangle rule": $p_{Fn} < p_{Fp} + p_{Fe}$

$$n_i = \frac{k_{F\,i}^3}{3\pi^2} \Rightarrow n_n^{1/3} \leq n_p^{1/3} + n_e^{1/3} = 2n_p^{1/3}$$

$$\boxed{x_p \equiv \frac{n_p}{n_n + n_p} \geq \frac{1}{9} \approx 11\%}$$

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The modified URCA process

If the direct Urca process: $$\begin{cases} n & \longrightarrow \quad p + e^- + \overline{\nu}_e \\ p + e^- & \longrightarrow \quad n + \nu_e \end{cases}$$

is forbidden because of momentum conservation, add a spectator neutron:

Modified Urca process: $$\begin{cases} n + n' & \longrightarrow \quad p + n' + e^- + \overline{\nu}_e \\ p + n' + e^- & \longrightarrow \quad n + n' + \nu_e \end{cases}$$

Momentum conservation is automatic, but the price to pay is:

3 vs 5 fermions phase space: $$\left(\frac{k_B T}{E_F}\right)^2 \sim \left(\frac{0.1\,\mathrm{MeV} \cdot T_9}{100\,\mathrm{MeV} \cdot E_{F\,100}}\right)^2 \sim 10^{-6} \cdot T_9^2$$

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Neutrino emission on a napkin (I)

The Murca-Bremsstrahlung family and Durca

Name	Process	Emissivity (erg cm^{-3} s^{-1})	
Modified Urca cycle (neutron branch)	$n + n \to n + p + e^- + \bar{\nu}_e$ $n + p + e^- \to n + n + \nu_e$	$\sim 2\times10^{21}\ R\ T_9^8$	Slow
Modified Urca cycle (proton branch)	$p + n \to p + p + e^- + \bar{\nu}_e$ $p + p + e^- \to p + n + \nu_e$	$\sim 10^{21}\ R\ T_9^8$	Slow
Bremsstrahlung	$n + n \to n + n + \nu + \bar{\nu}$ $n + p \to n + p + \nu + \bar{\nu}$ $p + p \to p + p + \nu + \bar{\nu}$	$\sim 10^{19}\ R\ T_9^8$	Slow
Direct Urca cycle	$n \to p + e^- + \bar{\nu}_e$ $p + e^- \to n + \nu_e$	$\sim 10^{27}\ R\ T_9^6$	Fast

Hyperons in neutron stars (I)

Hyperons, as Λ and Σ⁻ can be produced through reactions as, e.g.

$$\begin{cases} p + e^- & \longrightarrow & \Lambda + \nu_e \\ \Lambda & \longrightarrow & p + e^- + \overline{\nu}_e \end{cases}$$

$$\begin{cases} n + e^- & \longrightarrow & \Sigma^- + \nu_e \\ \Sigma^- & \longrightarrow & n + e^- + \overline{\nu}_e \end{cases}$$

Energy conservation requires: $\mu_\Lambda = \mu_n$ and $\mu_{\Sigma^-} = \mu_n + \mu_e$

Momentum conservation: very easily satisfied for Λ
and not very difficult to satisfy for Σ^-

Hyperons will result in DUrca processes if they can be present

Charged mesons (π^- & K^-) in neutron stars

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Hyperons, as π^- and K^- can be produced through reactions as, e.g.

$$\begin{cases} n + e^- & \longrightarrow \quad n + \pi^- + \nu_e \\ n + \pi^- & \longrightarrow \quad n + e^- + \overline{\nu}_e \end{cases}$$

$$\begin{cases} n + e^- & \longrightarrow \quad n + K^- + \nu_e \\ n + K^- & \longrightarrow \quad n + e^- + \overline{\nu}_e \end{cases}$$

Energy conservation requires: $m^*_\pi = \mu_e$ or $m^*_K = \mu_e$

Momentum conservation: trivially satisfied because mesons condense (they are bosons) and the condensate can absorb *any* extra needed momentum

Charged mesons will result in DUrca processes if they can be present

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Neutrino emission on a napkin (III)

Name	Process	Emissivity (erg cm^{-3} s^{-1})	
Modified Urca cycle (neutron branch)	$n+n \to n+p+e^- + \bar{\nu}_e$ $n+p+e^- \to n+n+\nu_e$	$\sim 2\times10^{21}\ R\ T_9^8$	Slow
Modified Urca cycle (proton branch)	$p+n \to p+p+e^- + \bar{\nu}_e$ $p+p+e^- \to p+n+\nu_e$	$\sim 10^{21}\ R\ T_9^8$	Slow
Bremsstrahlung	$n+n \to n+n+\nu+\bar{\nu}$ $n+p \to n+p+\nu+\bar{\nu}$ $p+p \to p+p+\nu+\bar{\nu}$	$\sim 10^{19}\ R\ T_9^8$	Slow
Cooper pair formations	$n+n \to [nn]+\nu+\bar{\nu}$ $p+p \to [pp]+\nu+\bar{\nu}$	$\sim 5\times10^{21}\ R\ T_9^7$ $\sim 5\times10^{19}\ R\ T_9^7$	Medium
Direct Urca cycle (nucleons)	$n \to p+e^- + \bar{\nu}_e$ $p+e^- \to n+\nu_e$	$\sim 10^{27}\ R\ T_9^6$	Fast
Direct Urca cycle (Λ hyperons)	$\Lambda \to p+e^- + \bar{\nu}_e$ $p+e^- \to \Lambda+\nu_e$	$\sim 10^{27}\ R\ T_9^6$	Fast
Direct Urca cycle (Σ^- hyperons)	$\Sigma^- \to n+e^- + \bar{\nu}_e$ $n+e^- \to \Sigma^- +\nu_e$	$\sim 10^{27}\ R\ T_9^6$	Fast
π^- condensate	$n+ < \pi^- > \to n+e^- + \bar{\nu}_e$	$\sim 10^{26}\ R\ T_9^6$	Fast
K^- condensate	$n+ < K^- > \to n+e^- + \bar{\nu}_e$	$\sim 10^{25}\ R\ T_9^6$	Fast

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Bremsstrahlung	$n+n \to n+n$ $n+p \to$ p		Slow
Cooper pair formations		$10^{21}\ R\ T_9^7$ $\sim 5\times 10^{19}\ R\ T_9^7$	Medium
Direct Urc (nucle	e	$\sim 10^{27}\ R\ T_9^6$	Fast
Direct U (Λ hyper	$e^- + \bar{\nu}_e$ $+ e^- \to \Lambda + \nu_e$	$\sim 10^{27}\ R\ T_9^6$	Fast
Direct Urc (Σ^- hyperon	$\Sigma^- \to n+e^- + \bar{\nu}_e$ $n+e^- \to \Sigma^- + \nu_e$	$\sim 10^{27}\ R\ T_9^6$	Fast
π^- condensate	$n+ <\pi^-> \to n+e^- + \bar{\nu}_e$	$\sim 10^{26}\ R\ T_9^6$	Fast
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Anything beyond just neutrons and protons
(and only a small amount of them)
results in enhanced neutrino emission

Dominant neutrino processes in the crust

Plasmon decay process

$$\Gamma \longrightarrow \nu + \overline{\nu}$$

Bremsstrahlung processes:

$$e^- + {}^A Z \longrightarrow e^- + {}^A Z + \nu + \overline{\nu}$$

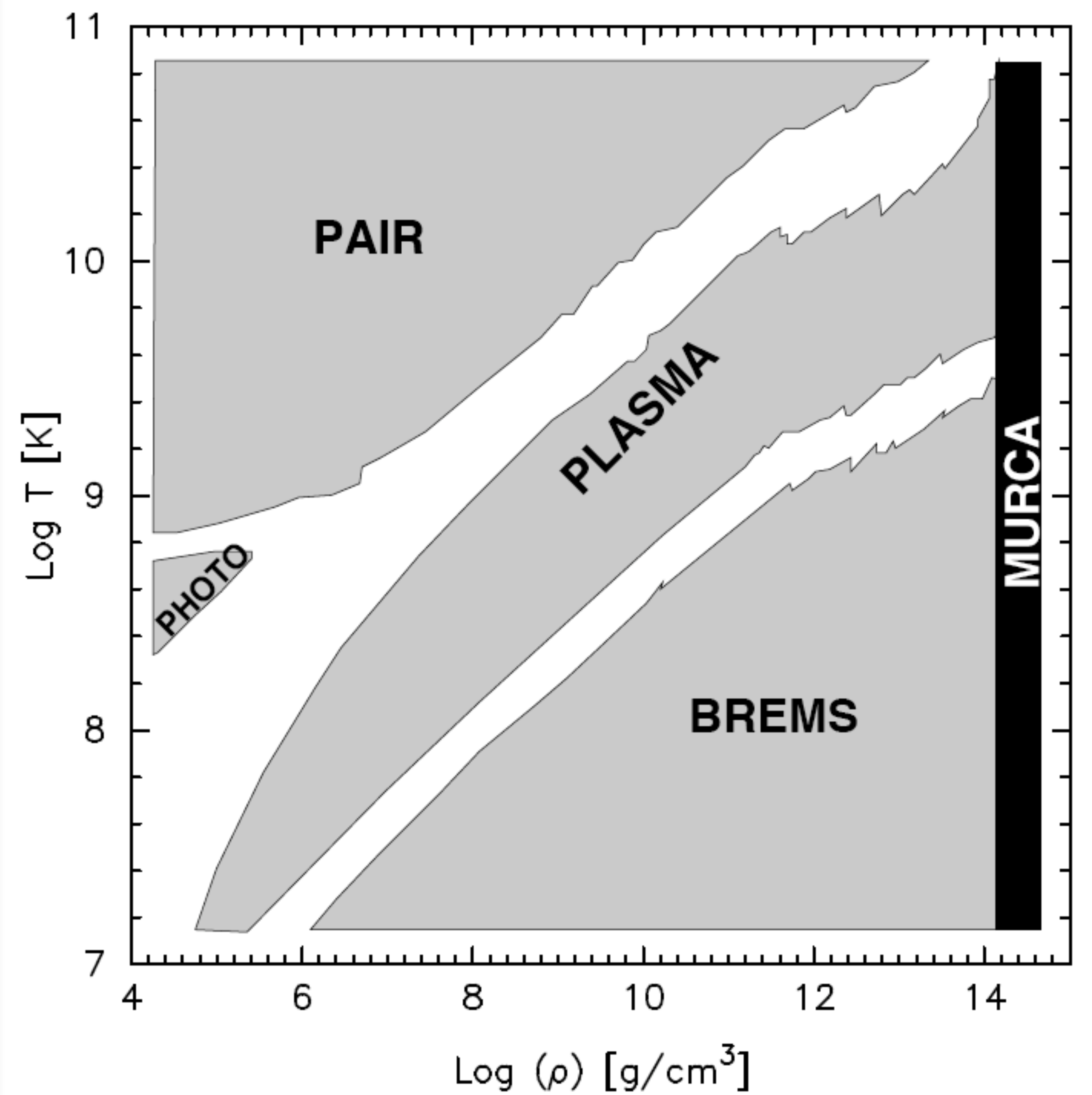
$$n + n \longrightarrow n + n + \nu + \overline{\nu}$$

Pair annihilation process:

$$\gamma + \gamma \leftrightarrow e^- + e^+ \longrightarrow \nu + \overline{\nu}$$

Photo-neutrino process:

$$\gamma + e^- \longrightarrow e^- + \nu + \overline{\nu}$$



Regions where the indicated neutrino emission process contributes more than 90% of the total.

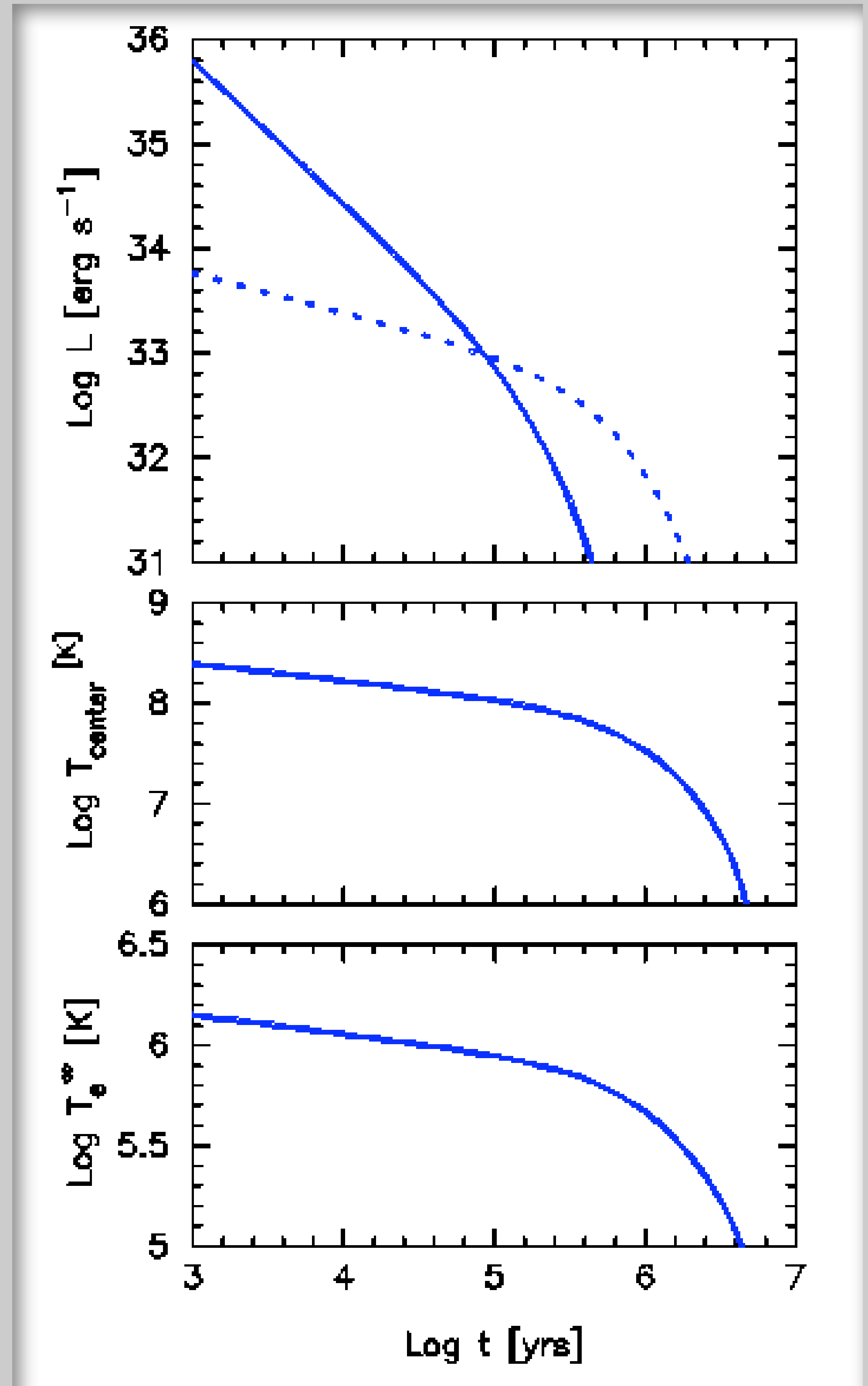
Simple
Analytical Solutions

Analytical Solution

$$\frac{dE_{th}}{dt} = C_v \frac{dT}{dt} = -L_\gamma - L_\nu$$

$$C_v = CT \quad L_\nu = NT^8 \quad L_\gamma = ST^{2+4\alpha}$$

$$L_\gamma = 4\pi R^2 \sigma T_e^4 \;\; \text{using} \;\; T_e \propto T^{0.5+\alpha} \;\; \text{with} \;\; \alpha \ll 1$$



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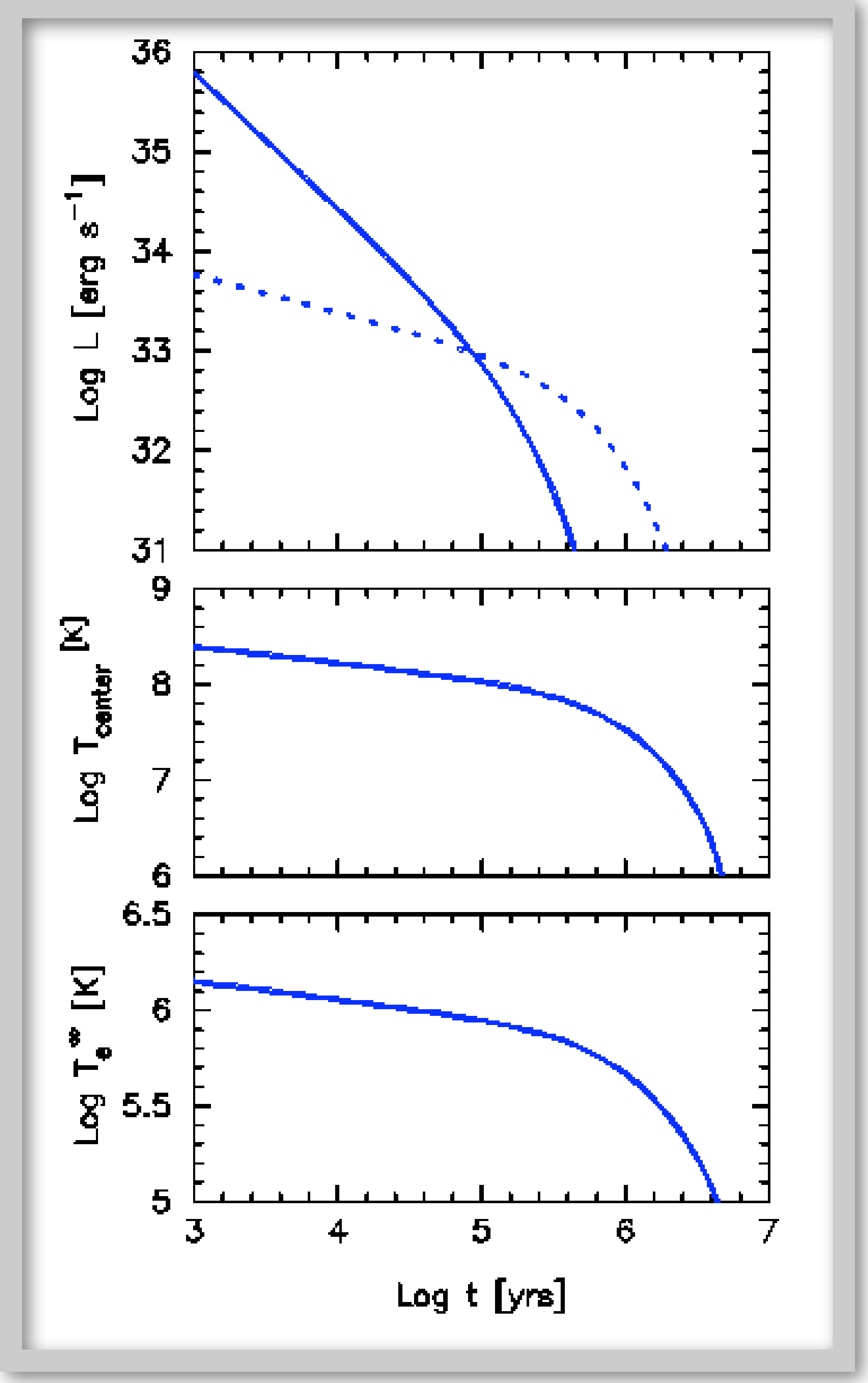
- Neutrino Cooling Era: $L_\nu >> L_\gamma$

$$\frac{dT}{dt} = -\frac{N}{C} T^7 \Rightarrow t - t_0 = A \left[\frac{1}{T^6} - \frac{1}{T_0^6} \right]$$

$$\boxed{T \propto t^{-1/6} \ \text{ and } \ T_e \propto t^{-1/12}}$$



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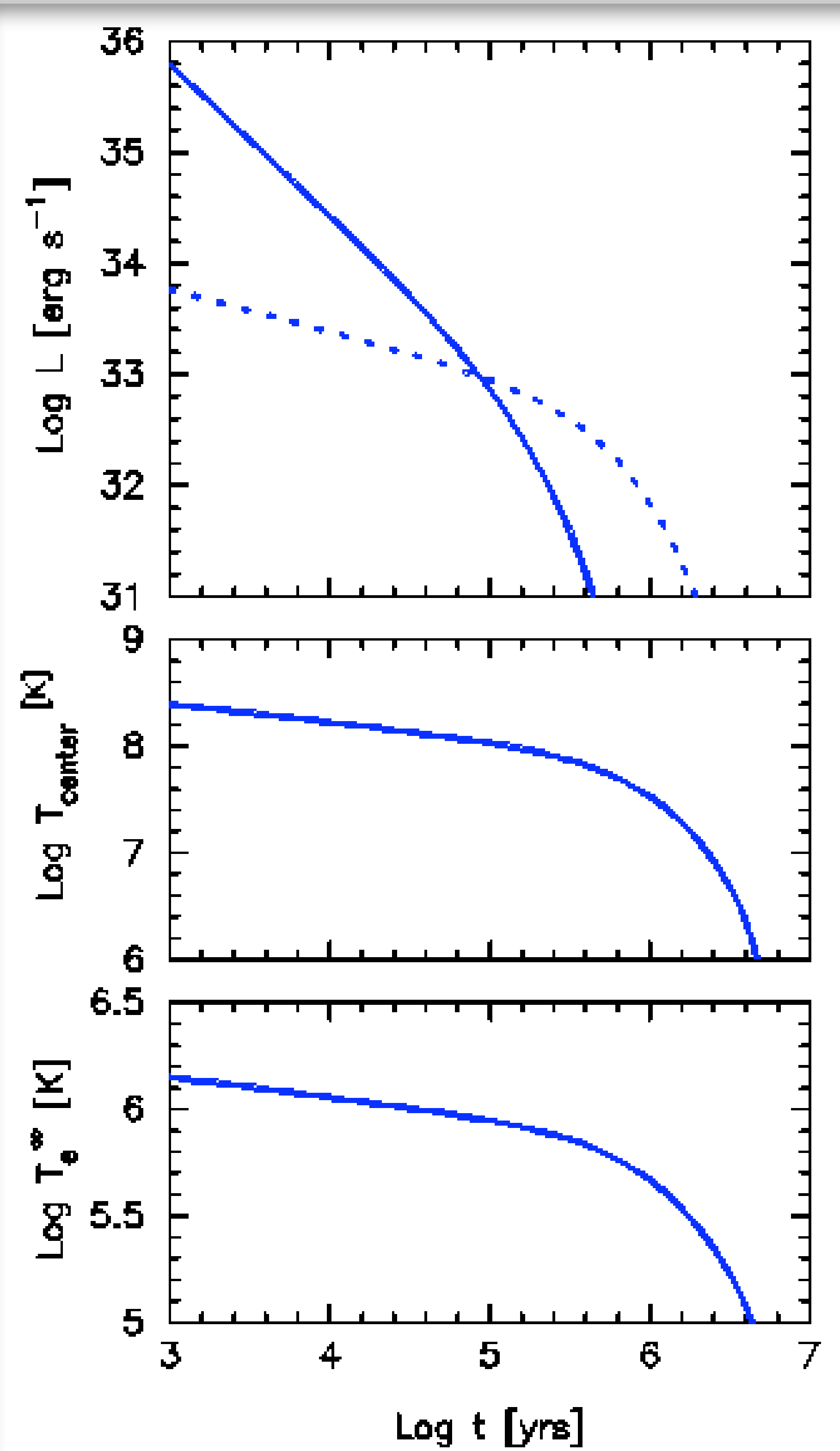
- Photon Cooling Era: $L_\gamma >> L_\nu$

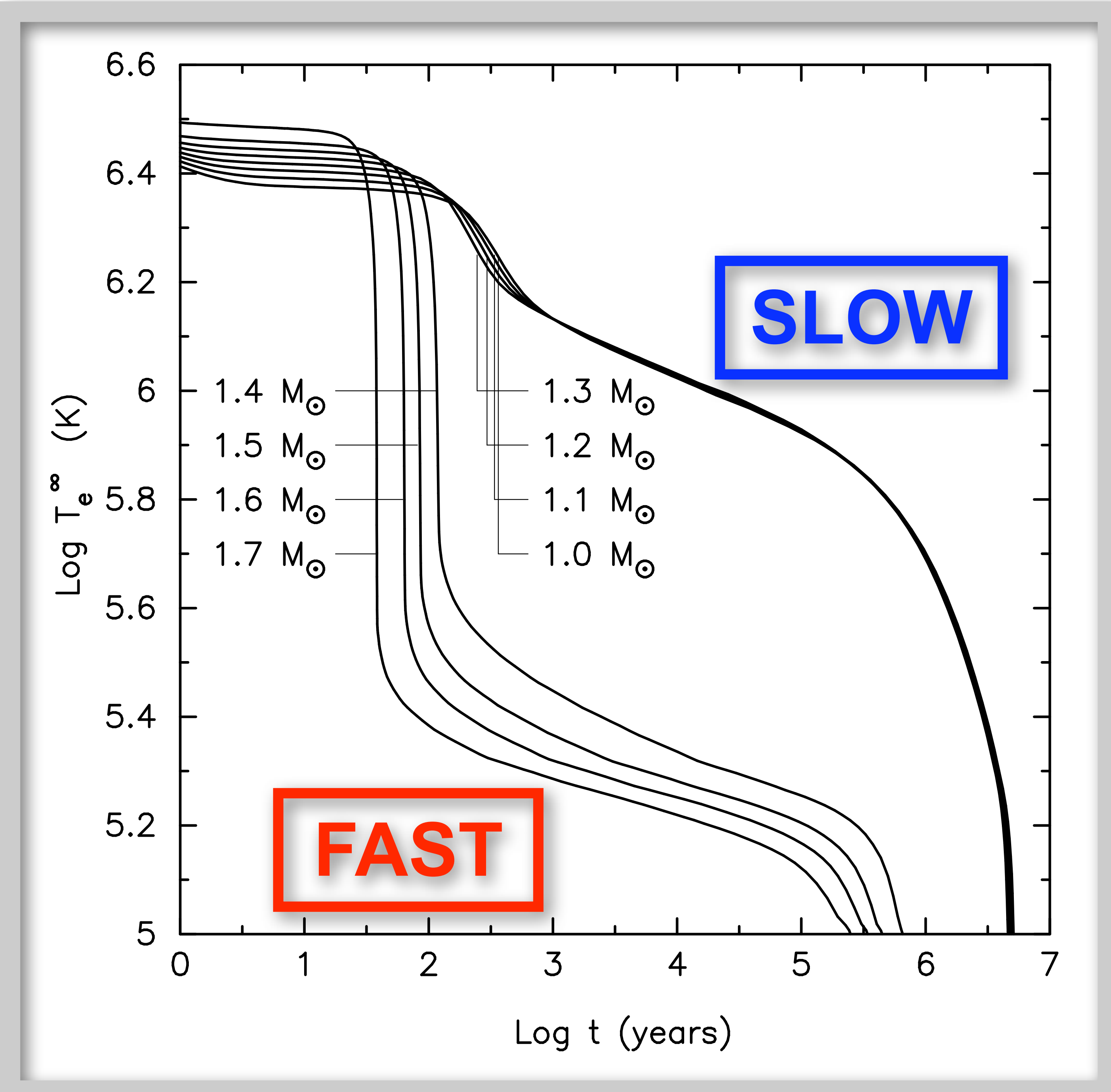
$$\frac{dT}{dt} = -\frac{N}{S} T^{1+\alpha} \Rightarrow t - t_0 = A \left[\frac{1}{T^\alpha} - \frac{1}{T_0^\alpha} \right]$$

$$T \propto t^{-1/\alpha} \ \text{ and } \ T_e \propto t^{-1/2\alpha}$$



MUrca vs DUrca

Direct vs modified Urca cooling

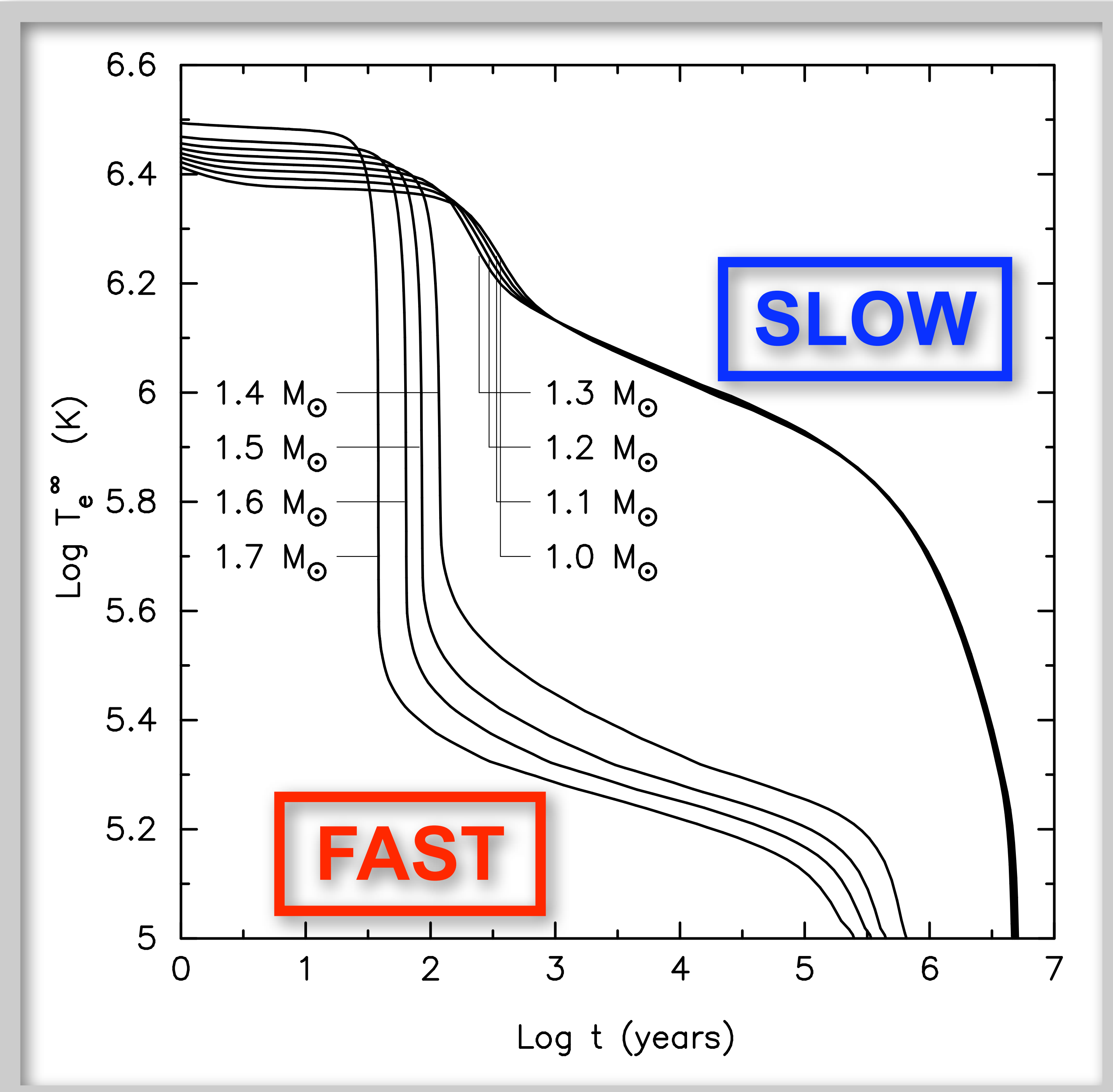


Models based on the PAL EOS:

adjusted (by hand) so that DURCA becomes allowed (triangle rule !) at M > 1.35 M_{Sun}.

“The Cooling of Neutron Stars by the Direct Urca Process”, Page & Applegate, ApJ 394, L17 (1992)

Direct vs modified Urca cooling



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This value is arbitrary:
we DO NOT know the value of this critical mass, and hopefully observations will, some day, tell us what it is !

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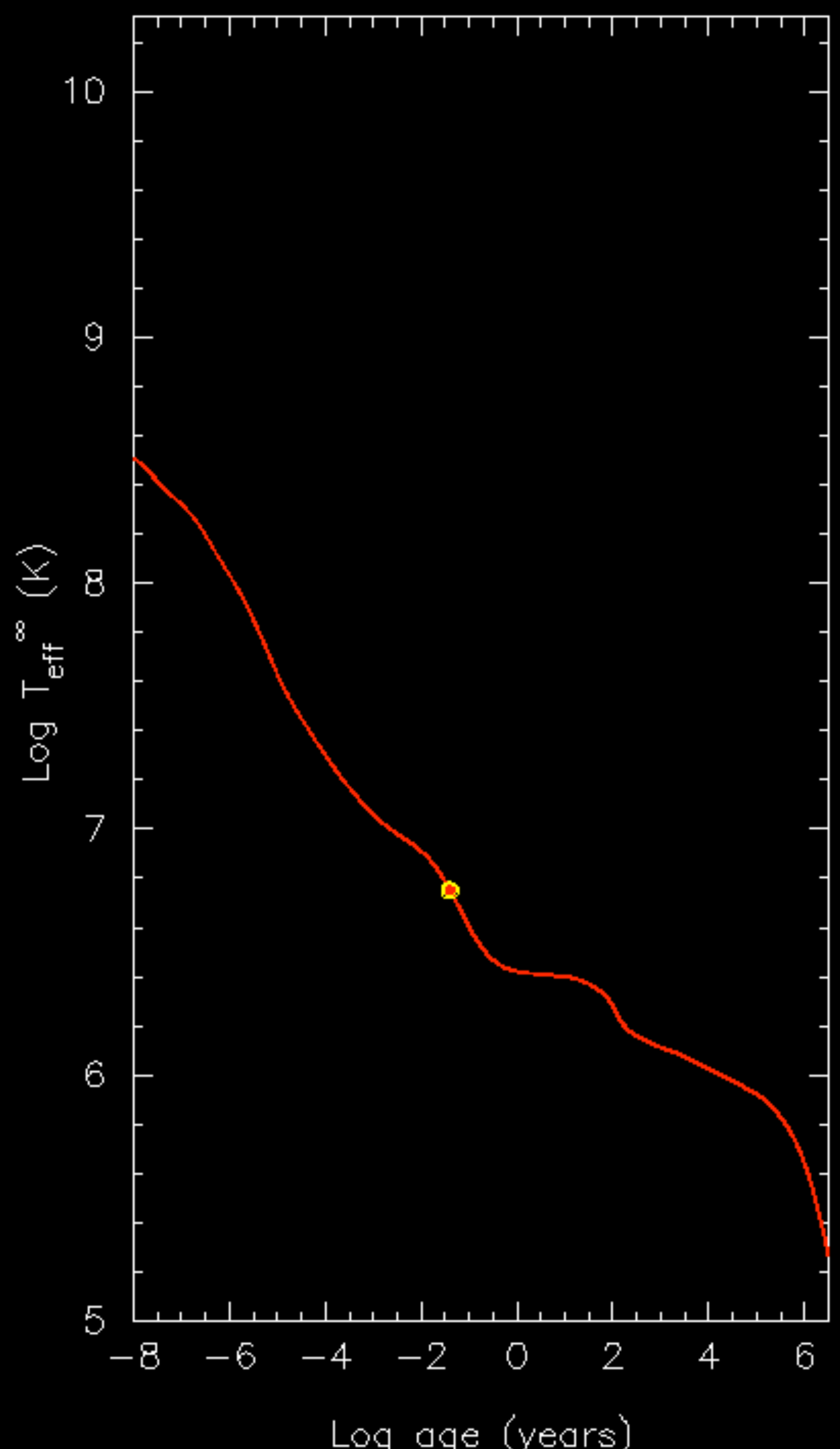
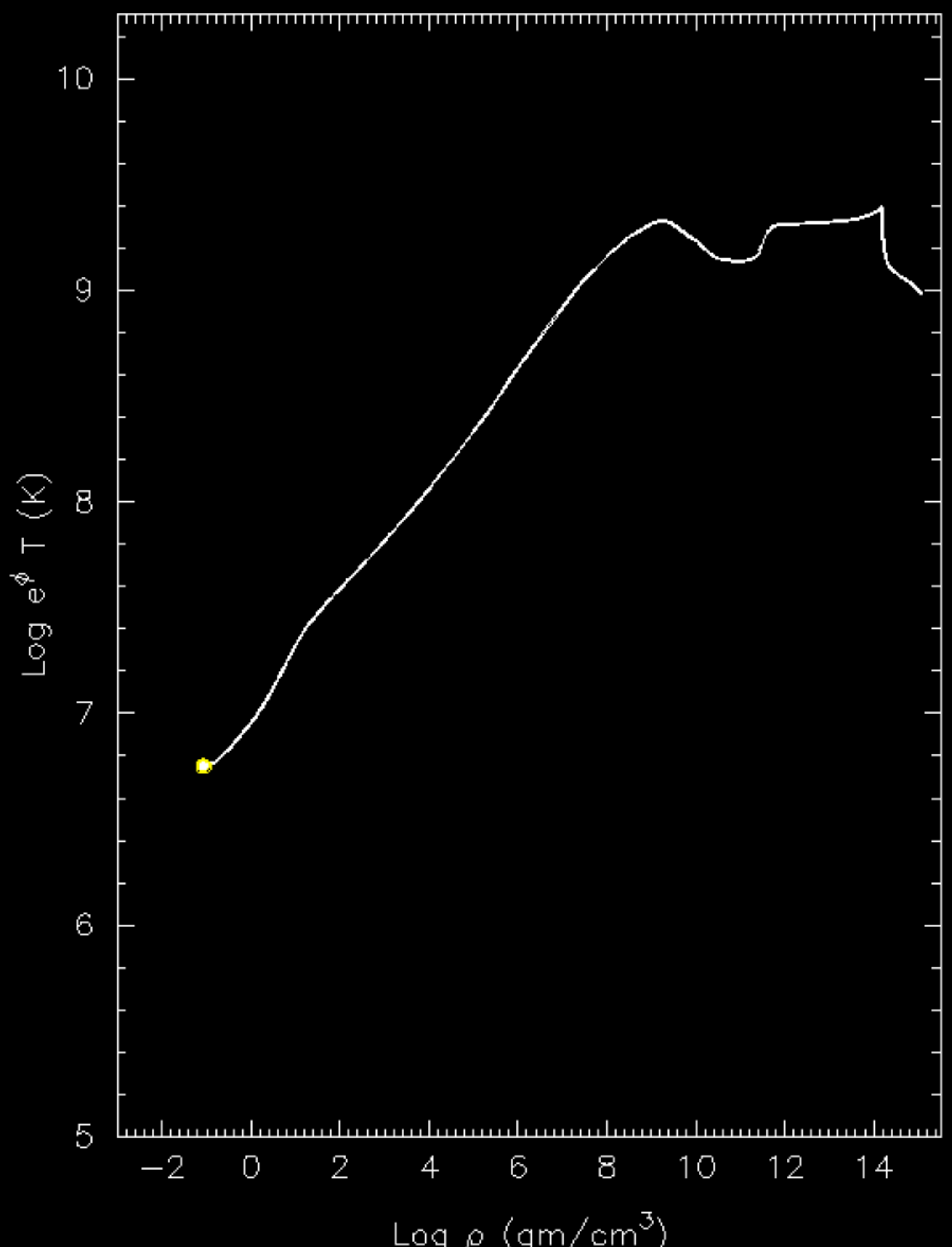
Standard cooling of a 1.3 M⊙ neutron star





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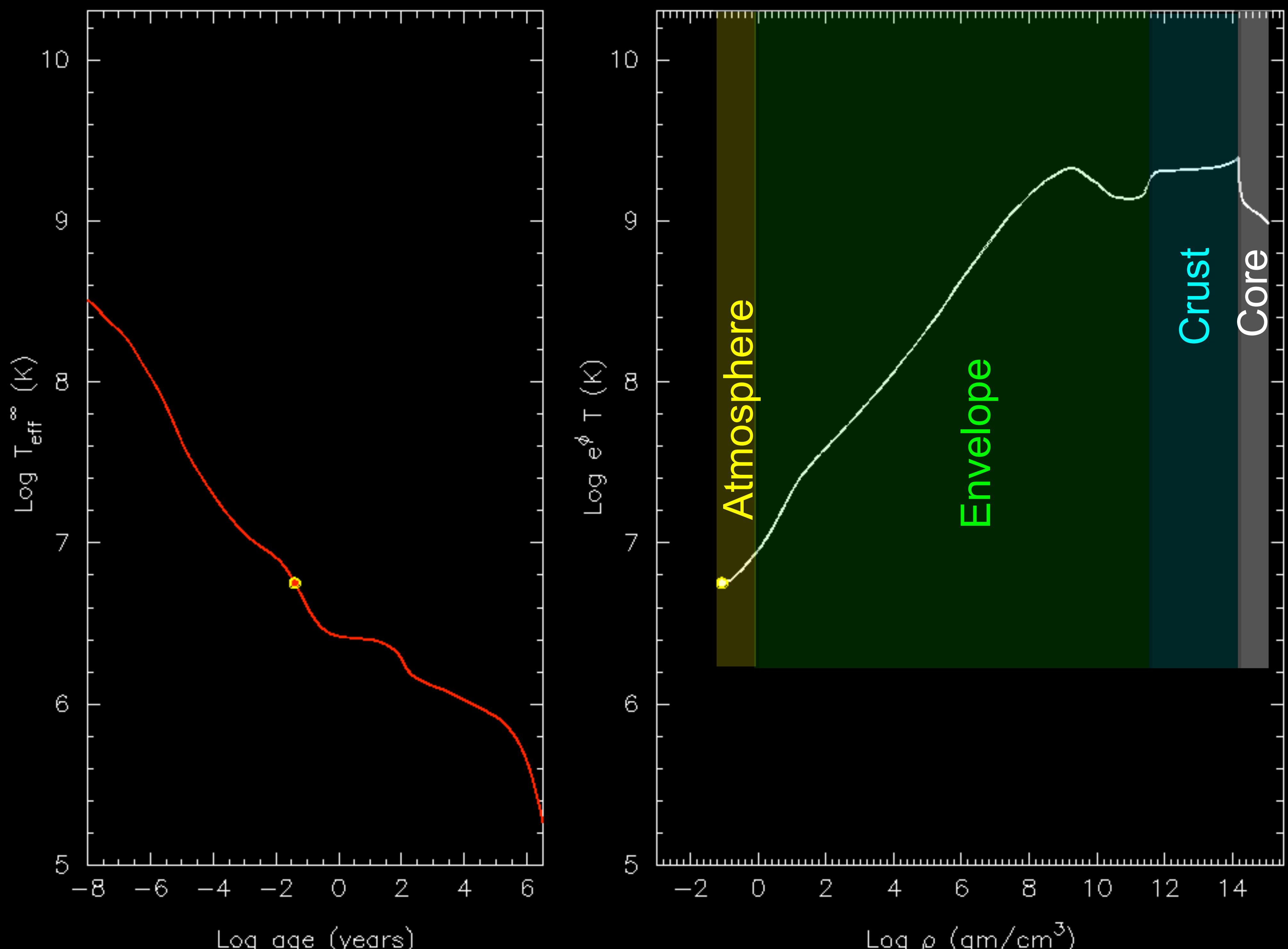
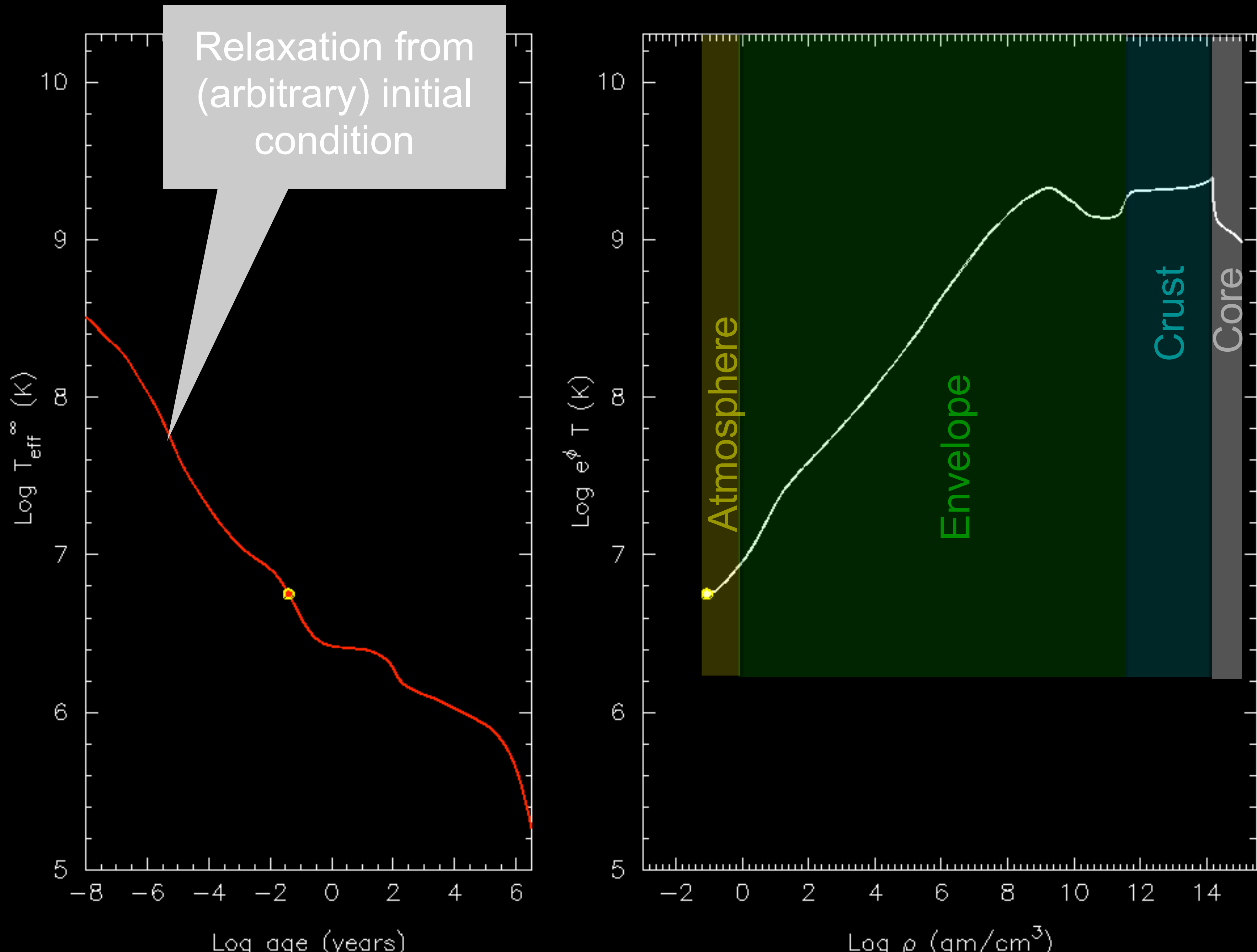


Standard cooling of a 1.3 M☉ neutron star



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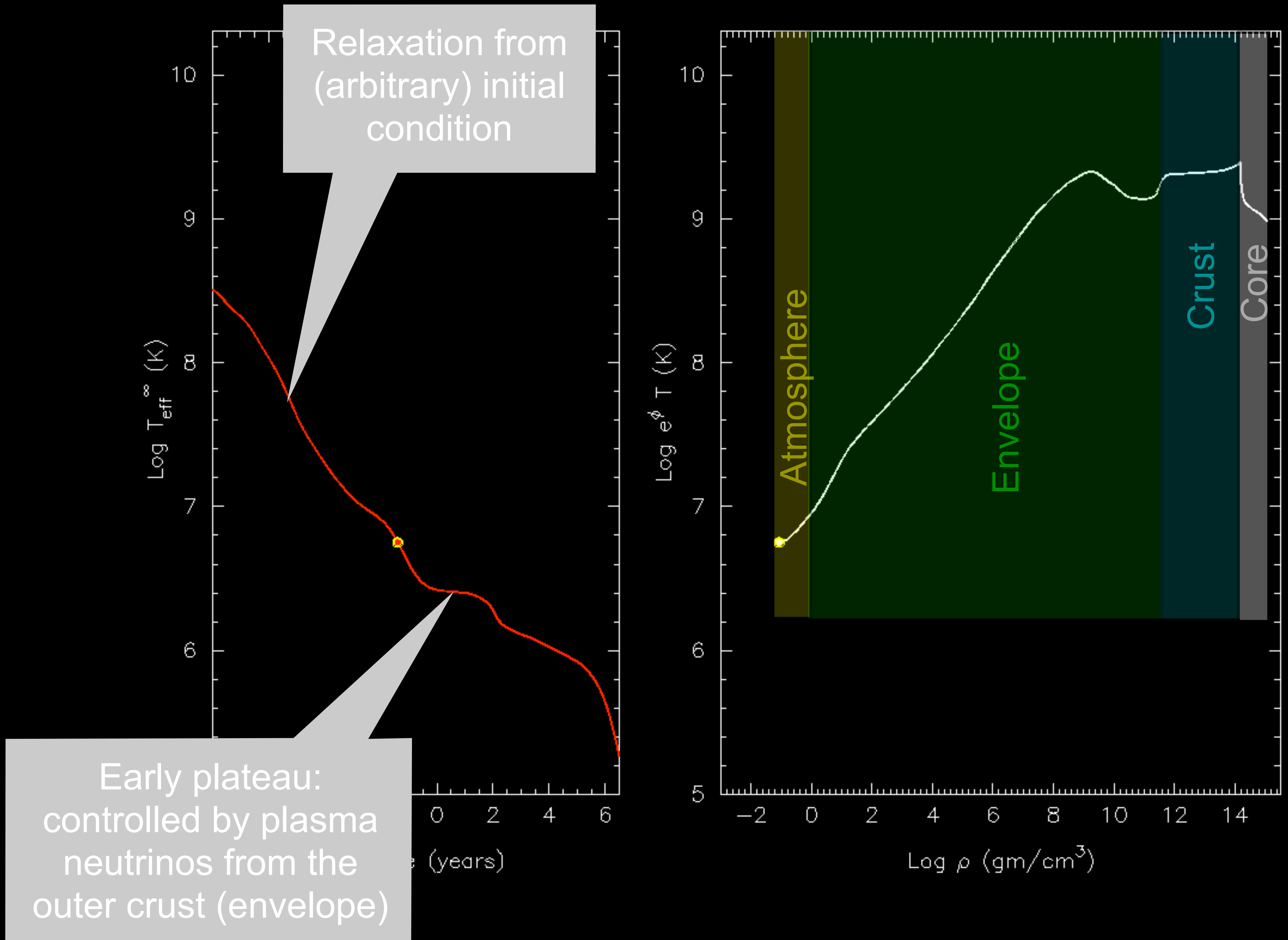

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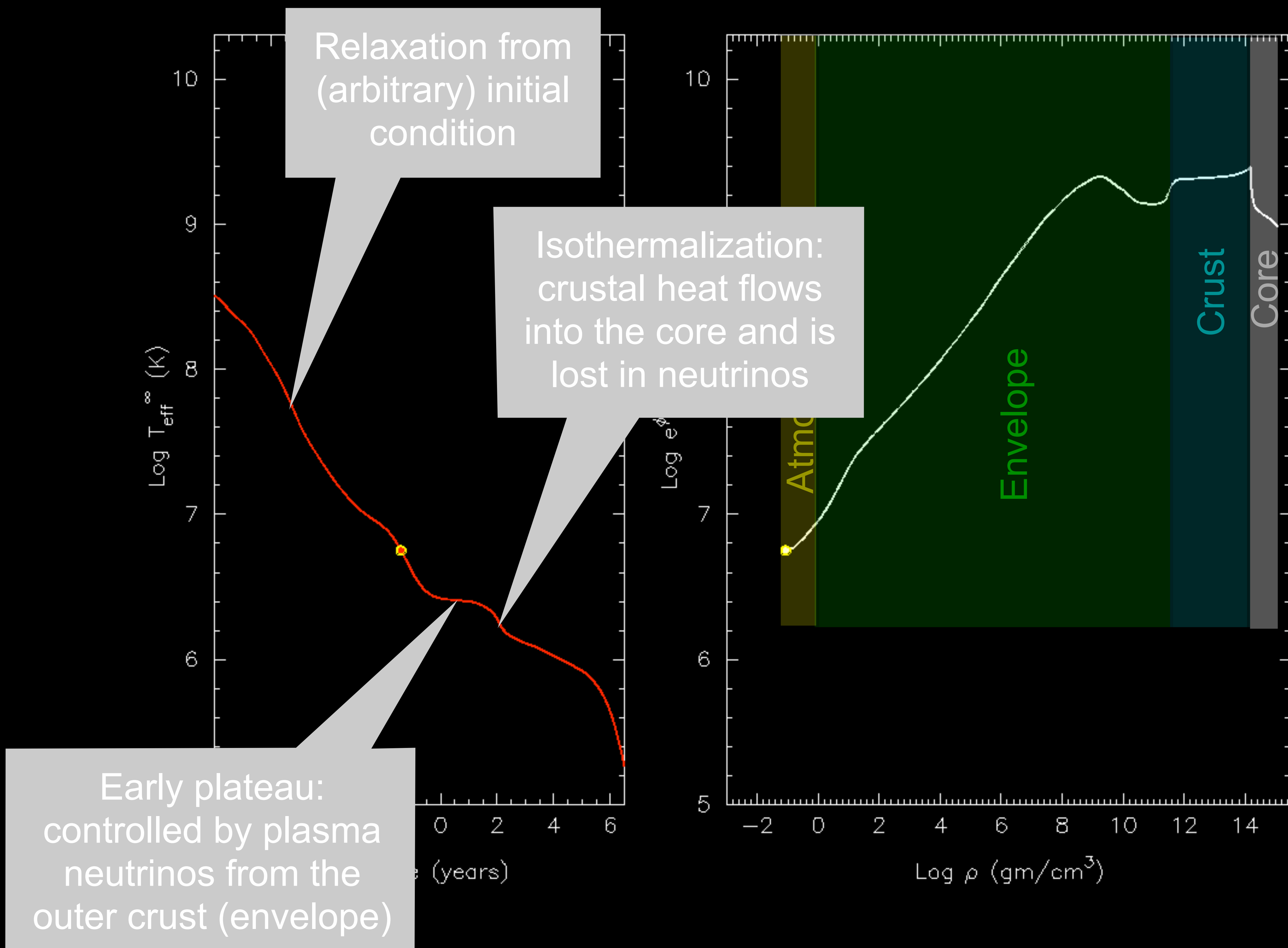

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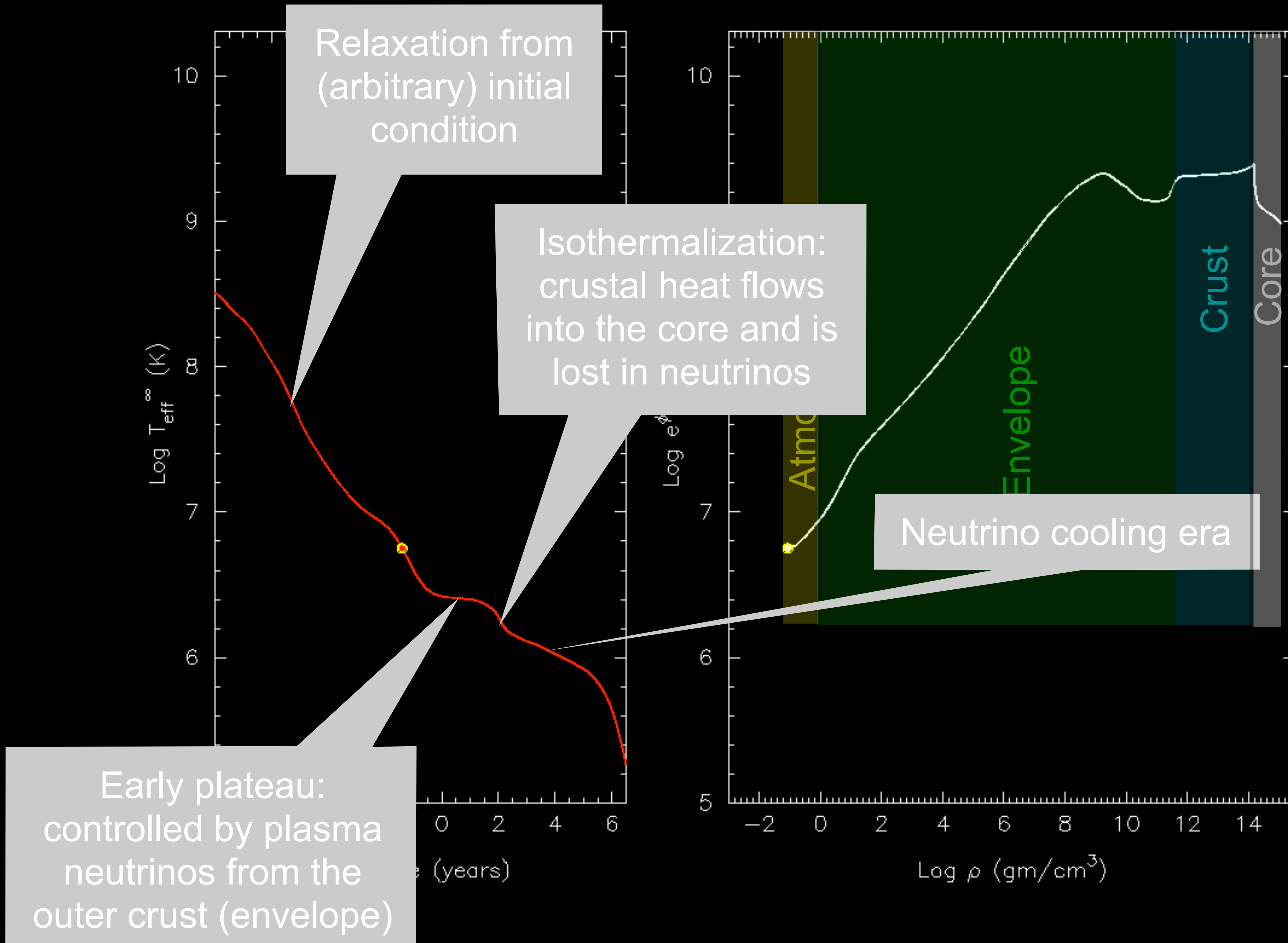

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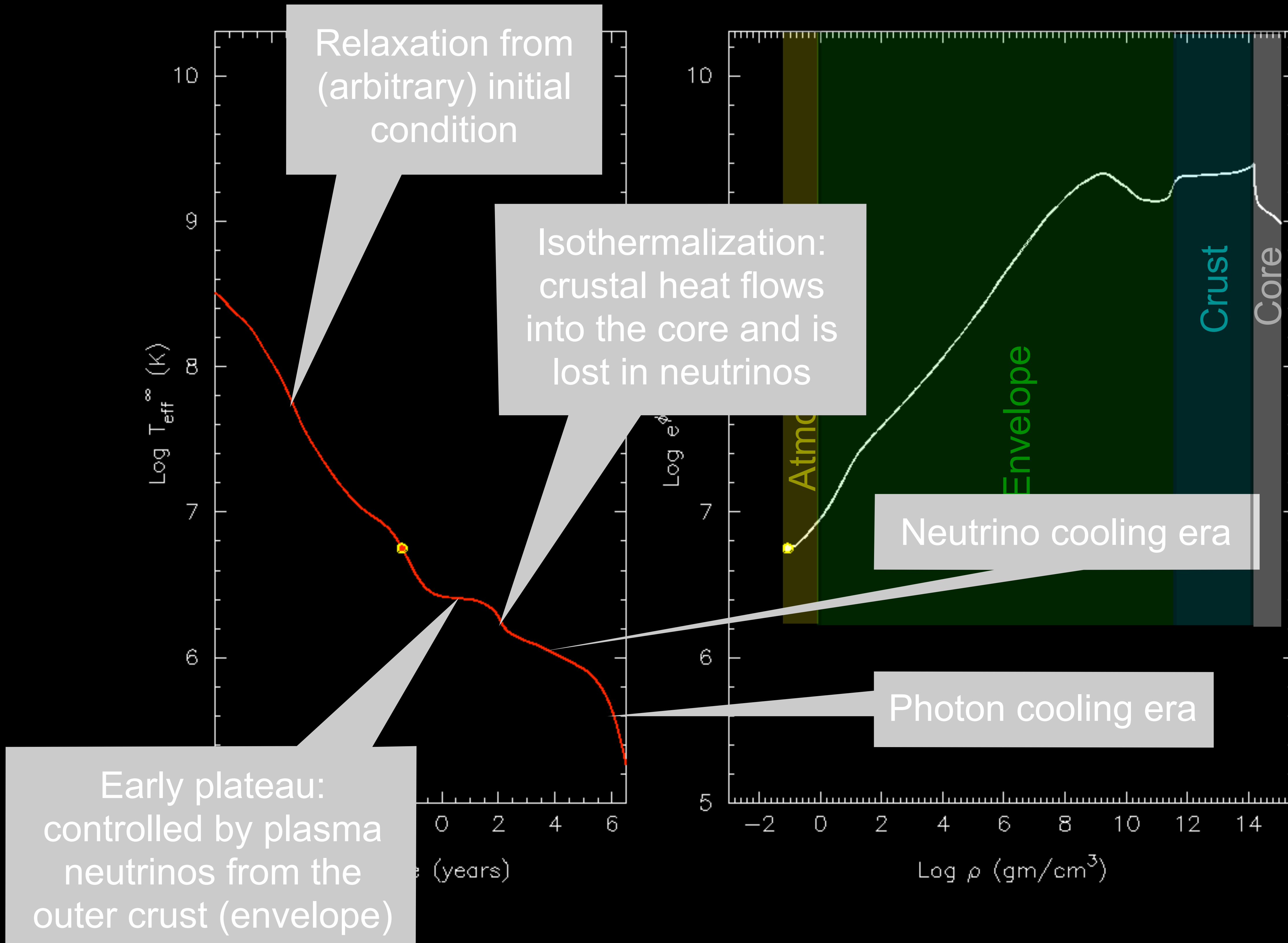
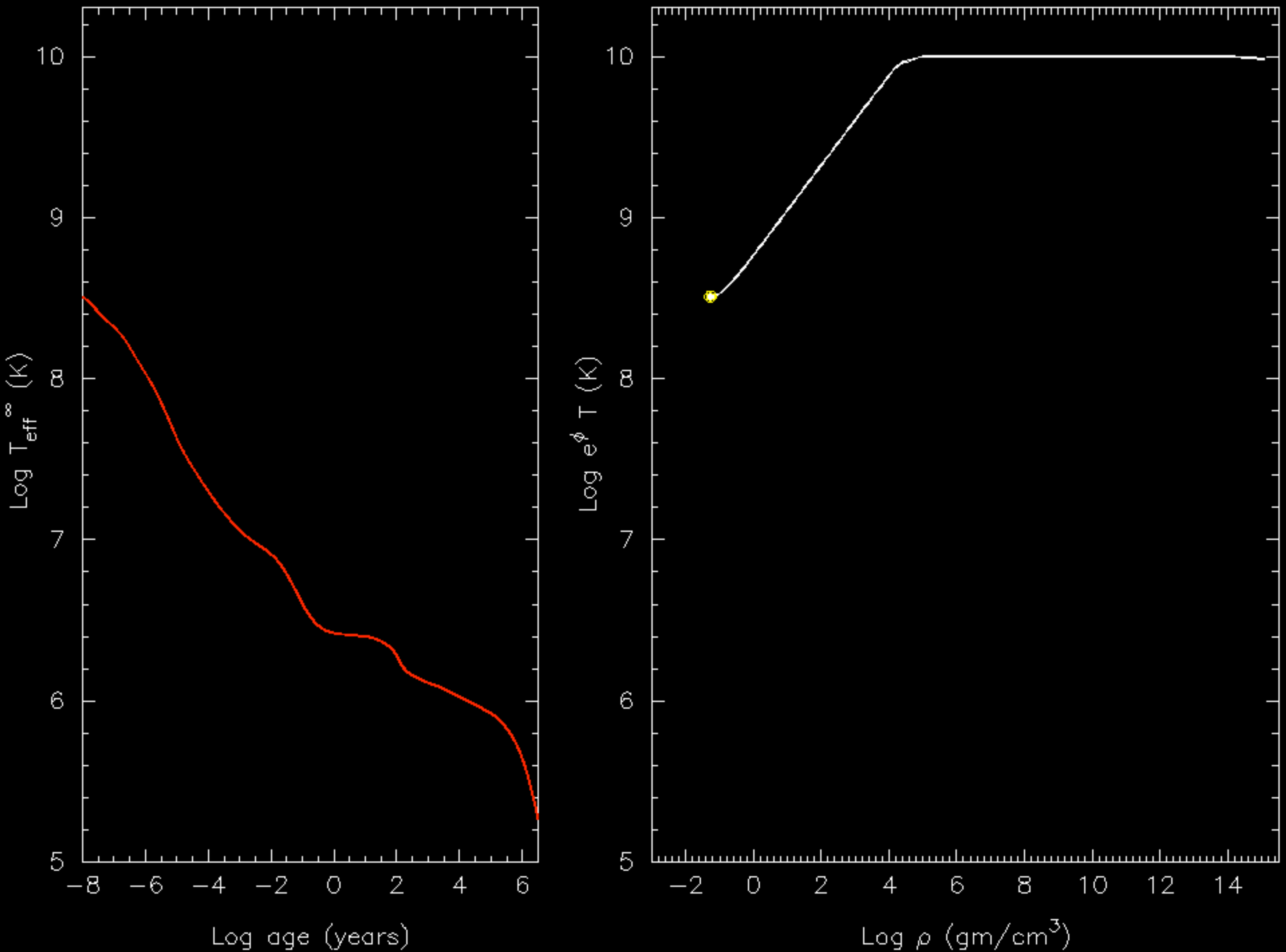
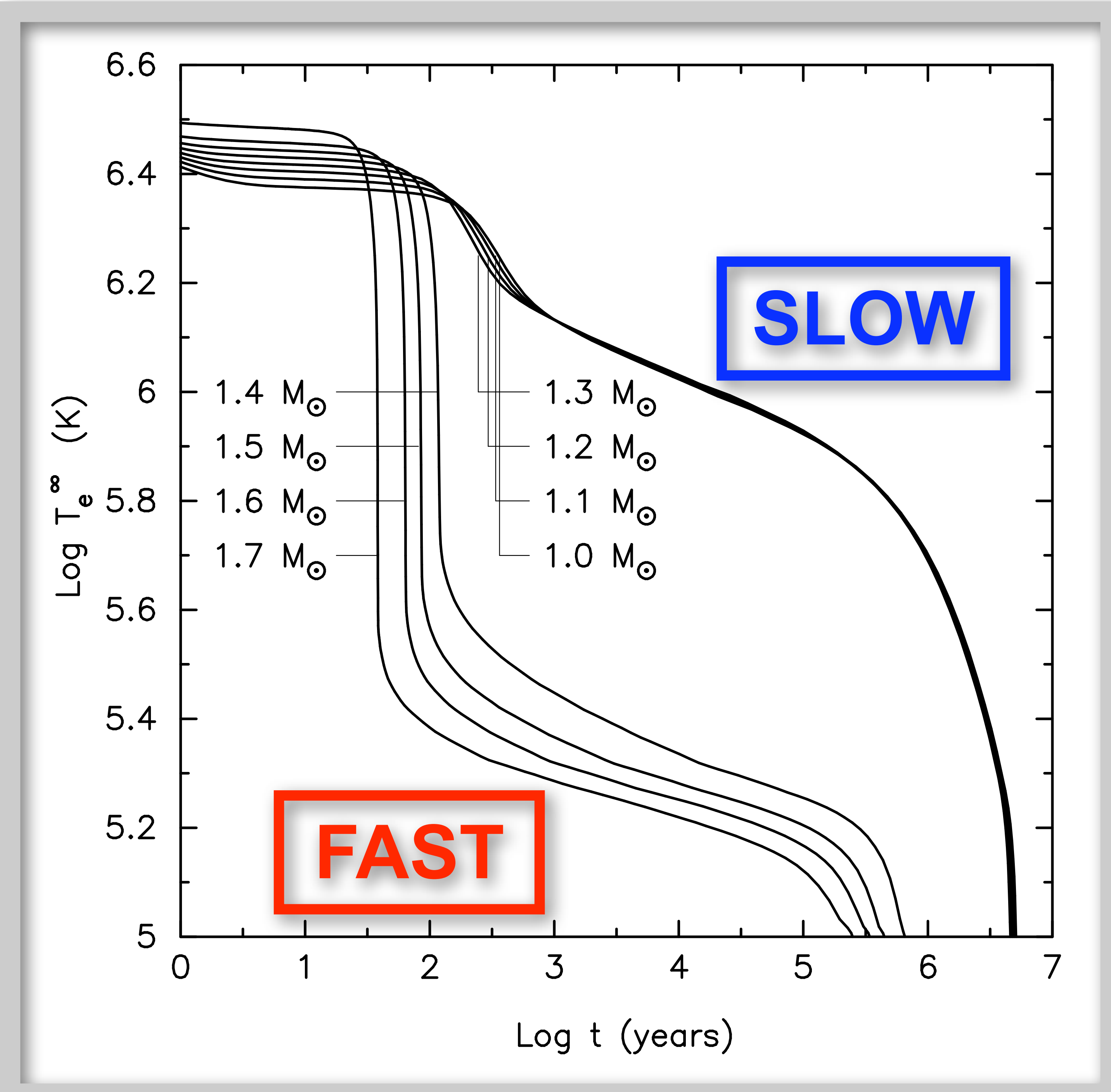


Standard cooling of a 1.3 M☉ neutron star [Animation file: 1.3_N.mov]



Direct vs modified Urca cooling



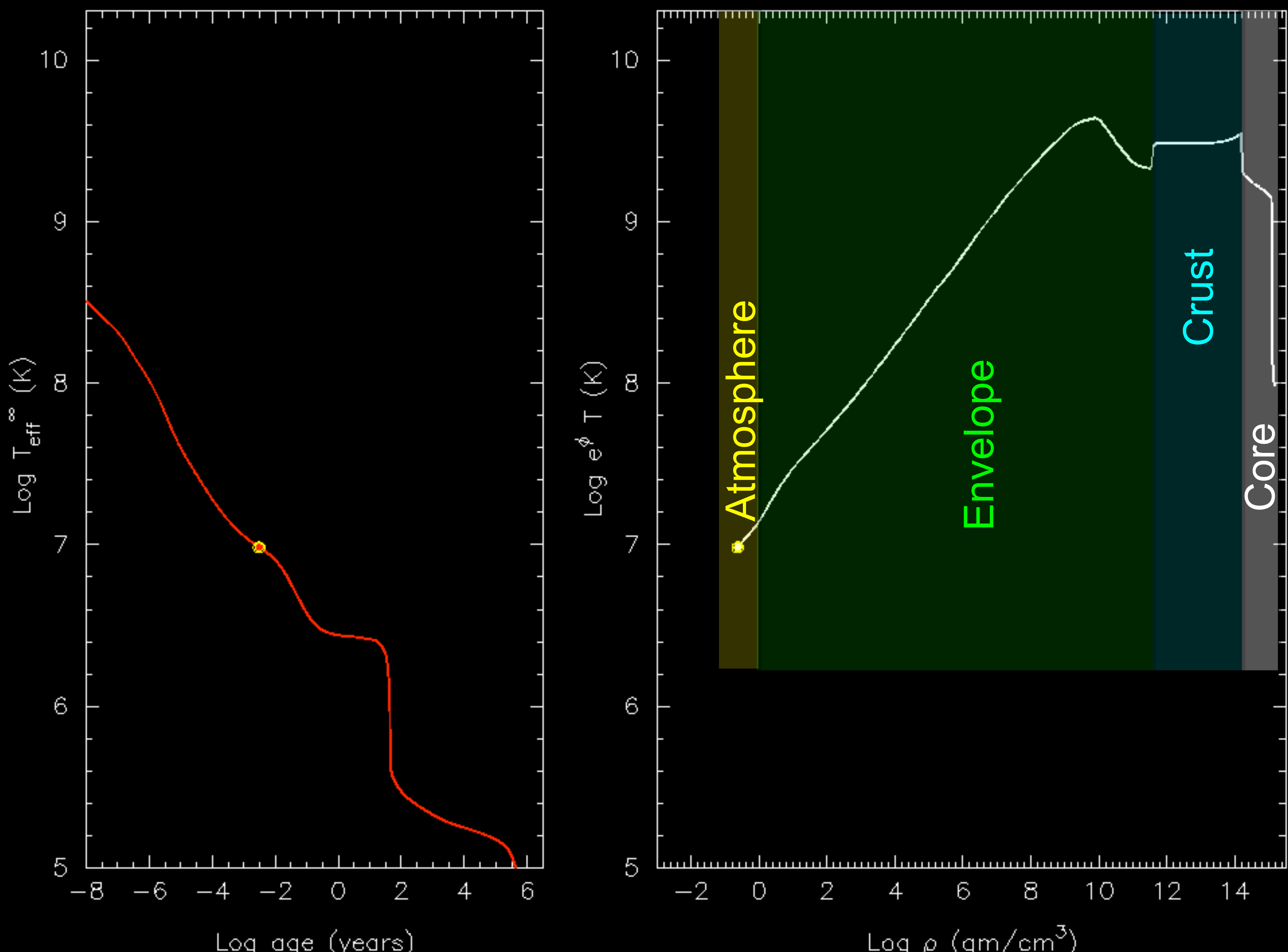
Models based on the PAL EOS:

adjusted (by hand) so that DURCA becomes allowed (triangle rule !) at M > 1.35 M_{Sun}.

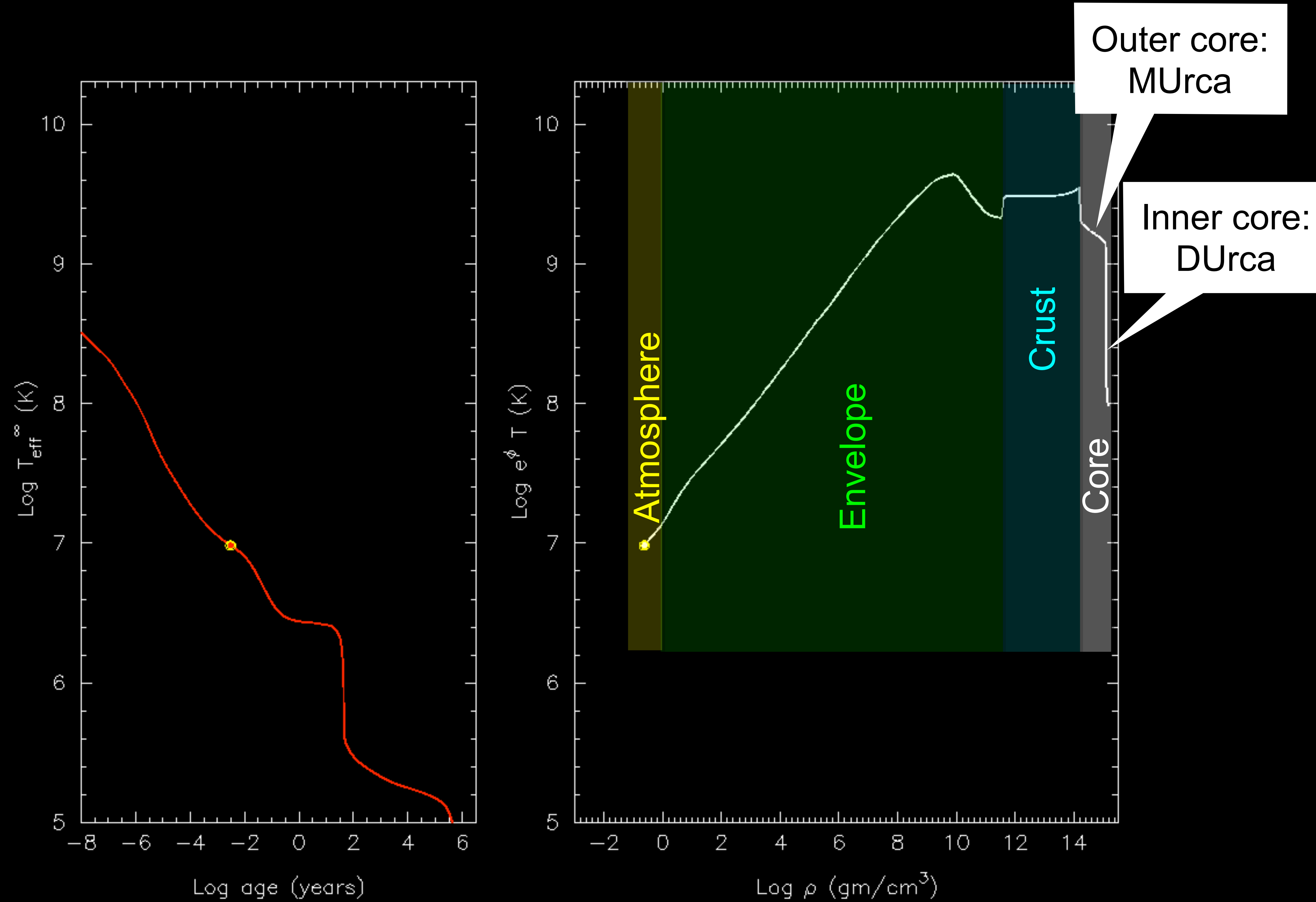
This value is arbitrary:
we DO NOT know the value of this critical mass, and hopefully observations will, some day, tell us what it is !

"The Cooling of Neutron Stars by the Direct Urca Process", Page & Applegate, ApJ 394, L17 (1992)

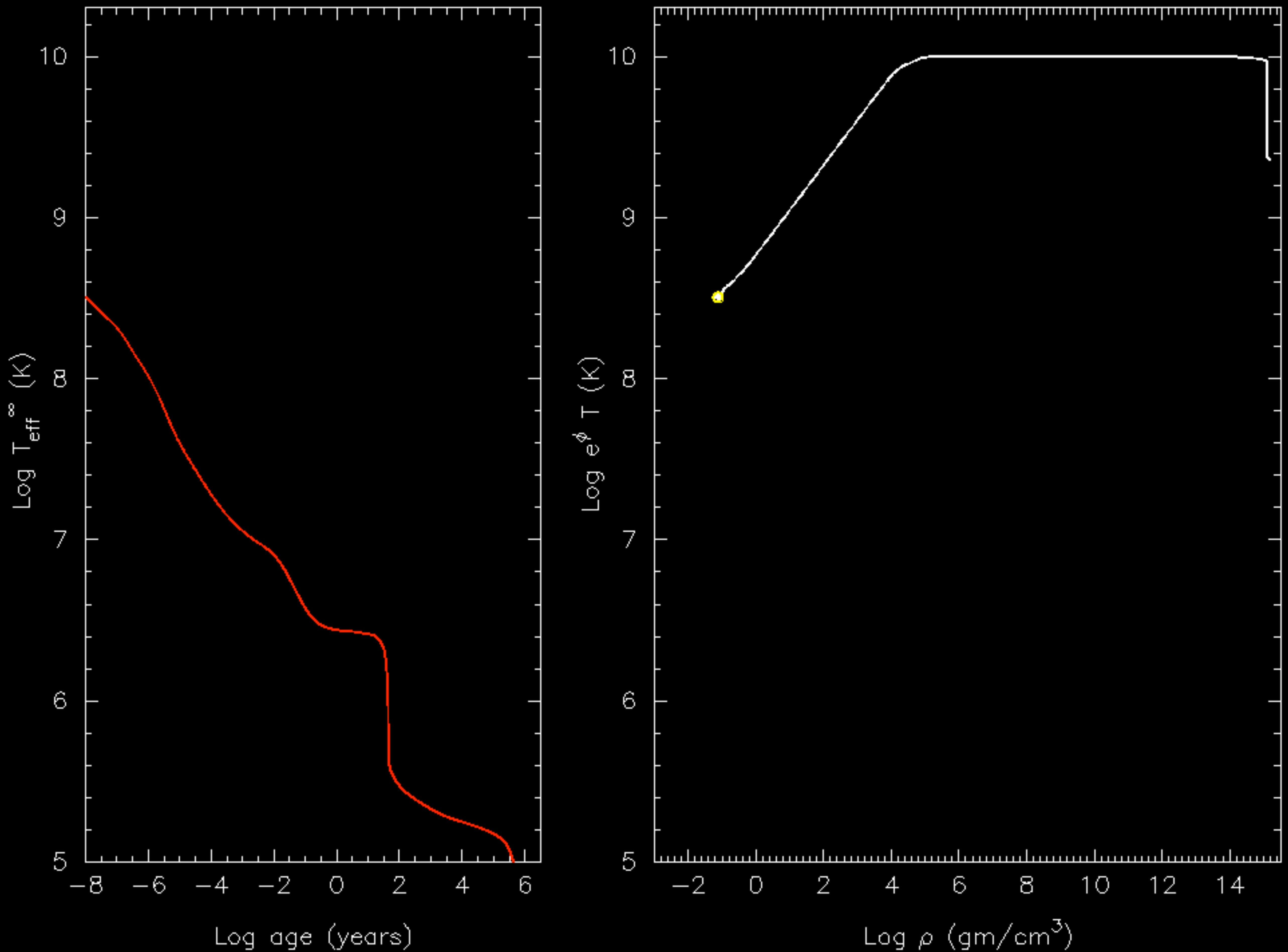
Enhanced cooling of a 1.5 M☉ neutron star



Enhanced cooling of a 1.5 M$_\odot$ neutron star



Enhanced cooling of a 1.5 M⊙ neutron star [Animation file: 1.5_N.mov]



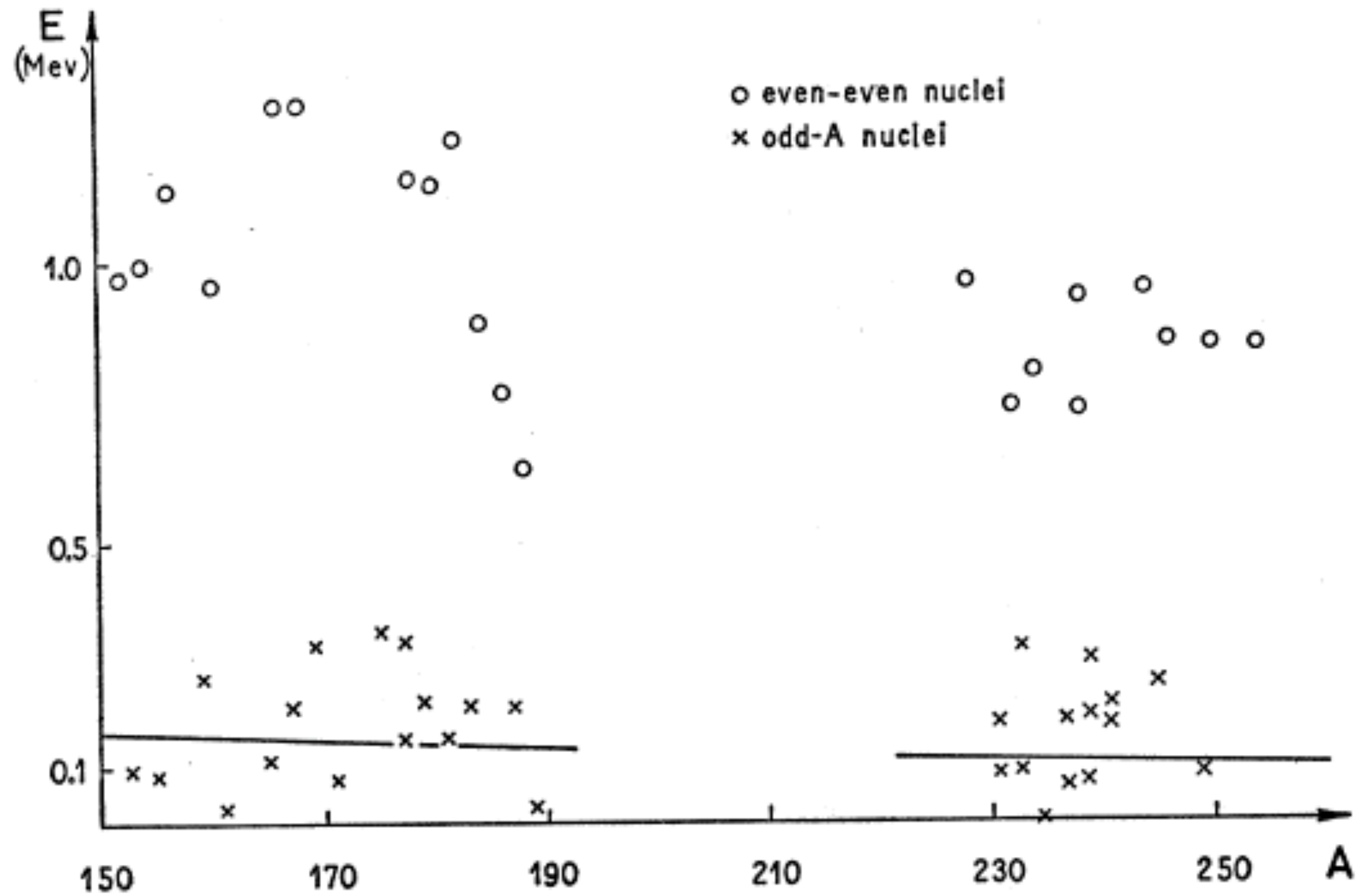
Pairing

Pairing in nuclei

FIG. 1. Energies of first excited intrinsic states in deformed nuclei, as a function of the mass number. The experimental data may be found in *Nuclear Data Cards* [National Research Council, Washington, D. C.] and detailed references will be contained in reference 1 above. The solid line gives the energy $\delta/2$ given by Eq. (1), and represents the average distance between intrinsic levels in the odd-A nuclei (see reference 1).

The figure contains all the available data for nuclei with $150<A<190$ and $228<A$. In these regions the nuclei are known to possess nonspherical equilibrium shapes, as evidenced especially by the occurrence of rotational spectra (see, e.g., reference 2). One other such region has also been identified around $A=25$; in this latter region the available data on odd-A nuclei is still represented by Eq. (1), while the intrinsic excitations in the even-even nuclei in this region do not occur below 4 Mev.

We have not included in the figure the low lying $K=0$ states found in even-even nuclei around Ra and Th. These states appear to represent a collective odd-parity oscillation.



Possible Analogy between the Excitation Spectra of Nuclei and Those of the Superconducting Metallic State
Bohr, A.; Mottelson, B. R.; Pines, D. (1958), Phys. Rev. 110, p.936

Pairing in nuclei

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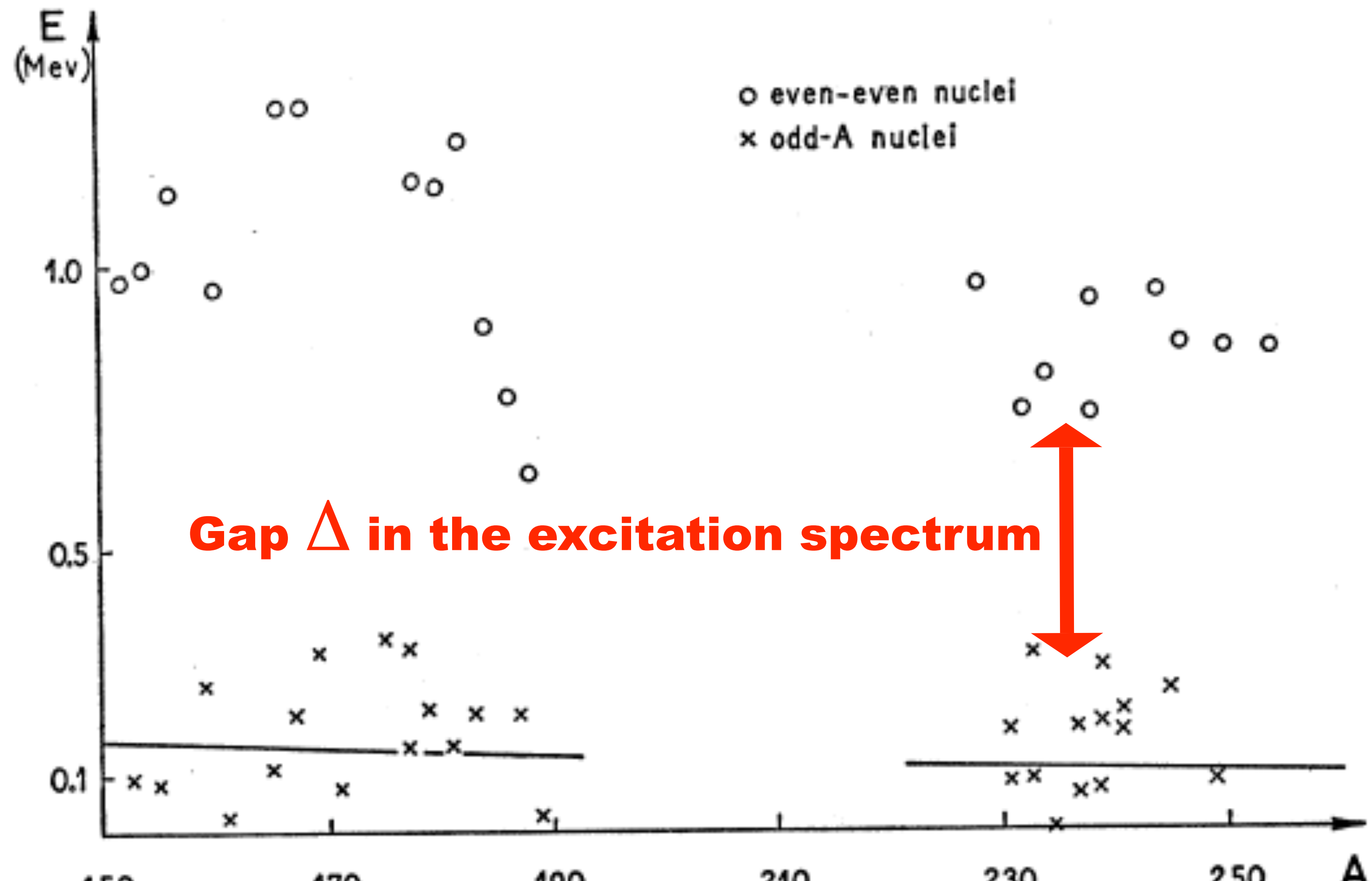


⇒ Excitations are suppressed by a factor ~ exp(- Δ/kT)

Possible Analogy between the Excitation Spectra of Nuclei and Those of the Superconducting Metallic State
Bohr, A.; Mottelson, B. R.; Pines, D. (1958), Phys. Rev. 110, p.936

Pairing of nucleons ?

Cooper theorem: the Fermi surface is unstable (at low enough temperature) to the formation of Cooper pairs if there is *any* attractive interaction in some channel.

Question 1: which attractive interaction and which channel ?
Question 2: at which temperature ?

Pairing of nucleons ?

Cooper theorem: the Fermi surface is unstable (at low enough temperature) to the formation of Cooper pairs if there is *any* attractive interaction in some channel.

Question 1: which attractive interaction and which channel ?
Question 2: at which temperature ?

"Answer" 1: look at phase shifts for in-vacuum nucleon-nucleon interactions (positive phase-shift means attraction):

- **at low energy the $^1S_0$ channel is attractive**
- **at higher energies the $^3P_2$ channel is attractive**

Other attractive channels exist but lead to very small gaps: apparently the $^3P_2$ gap dominates.

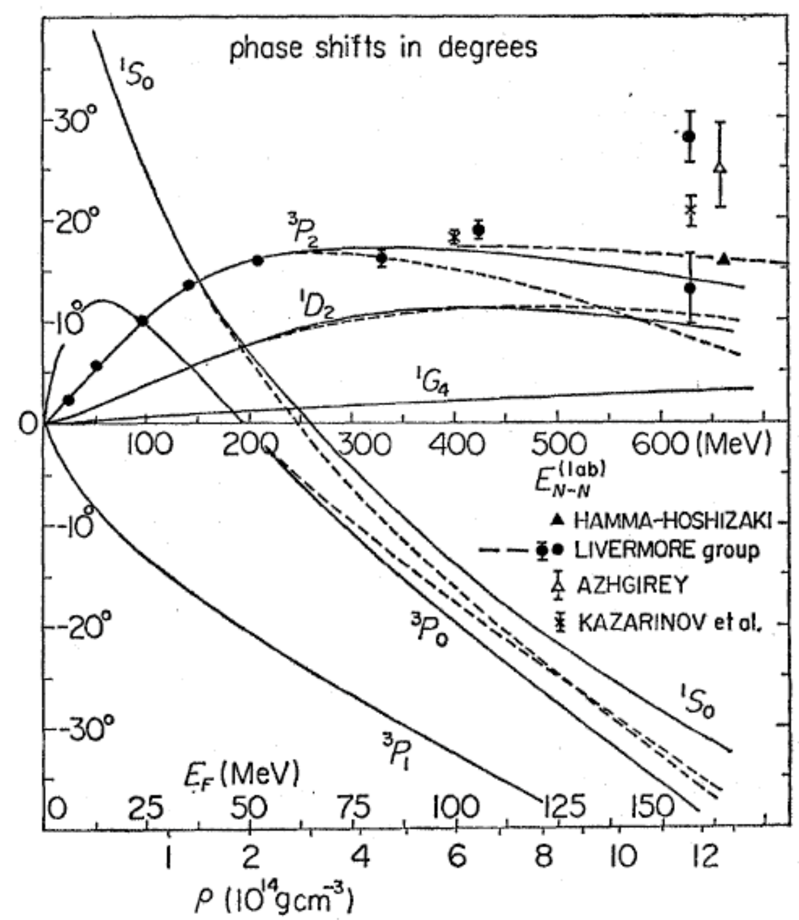
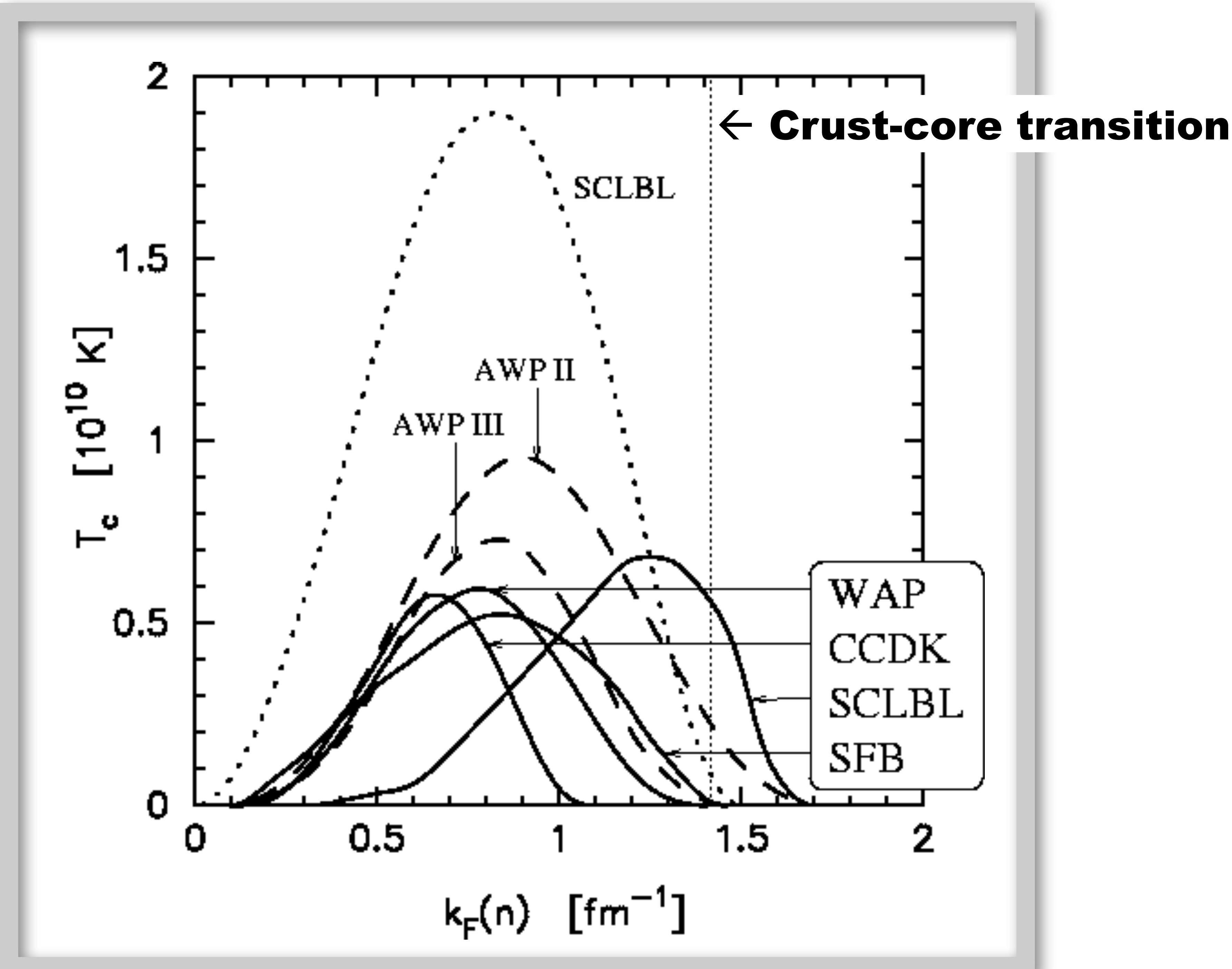


Fig. 1. Nucleon-nucleon scattering phase shifts versus $E_{N-N}^{(LAB)} = 4E_F$. Solid (dotted) lines represent the phase shifts calculated from the OPEG potential with 2 GeV soft core (the OPEH potential with the hard core radius = 0.42 fm).[9] For the $^3P_2$ phase shifts, solutions of the phase shift analysis are shown.[10]

Superfluid State in Neutron Star Matter. I Generalized Bogoliubov Transformation and Existence of $^3P_2$ Gap at High Density
Tamagaki, R., 1970PThPh..44..905T

Prediction for the neutron $^1S_0$ T_c



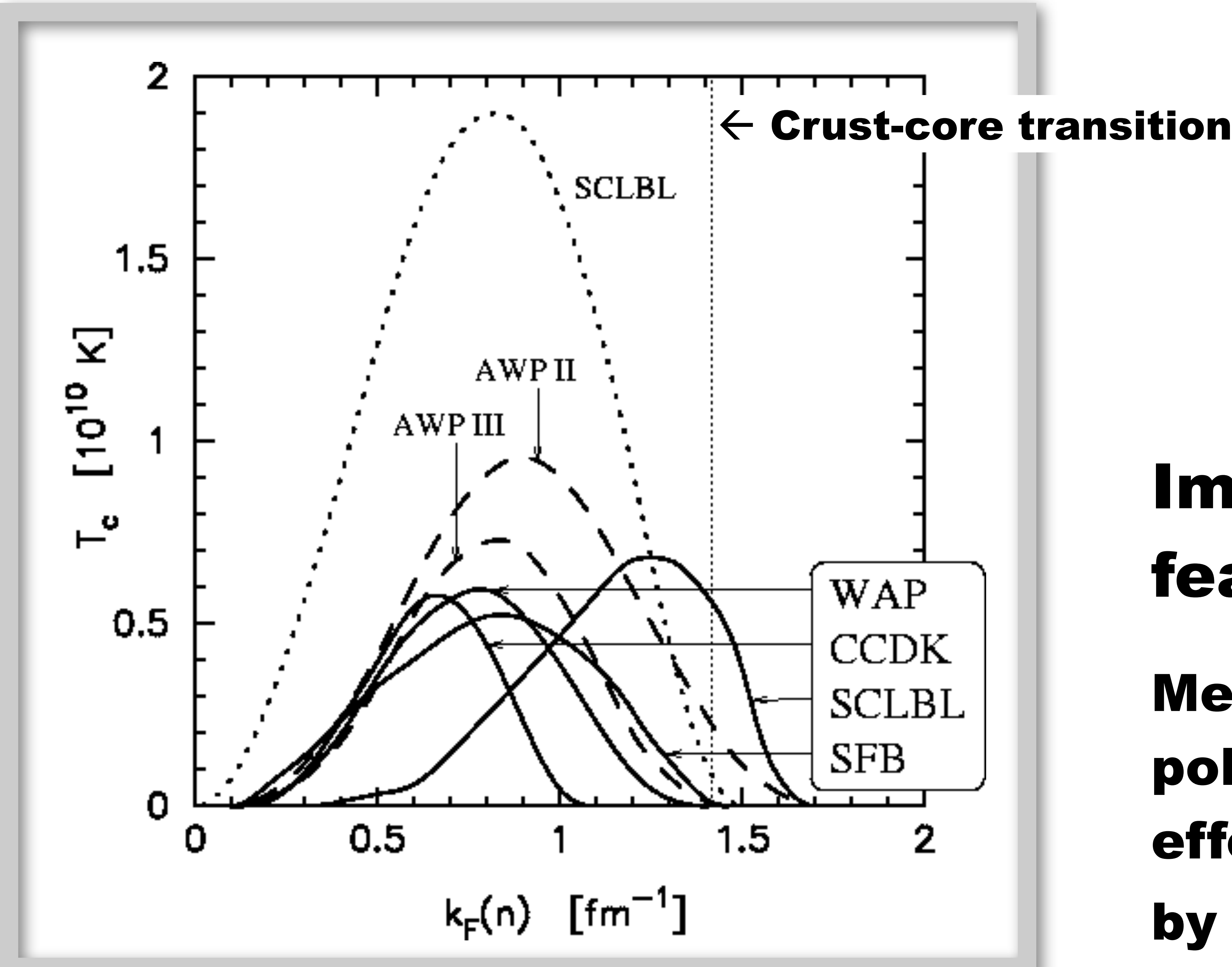
WAP: Wambach, Ainsworth & Pines, Nulc. Phys. A555 (1993), 128

CCDK: Chen, Clark, Davé & Khodel, Nucl. Phys. A555 (1993), 59

SCLBL: Schulze, Cugnon, Lejeune, Baldo & Lombardo, Phys. Lett. B375 (1996), 1

SFB: Schwenk, Friman & Brown, Nucl. Phys. A717 (2003), 191

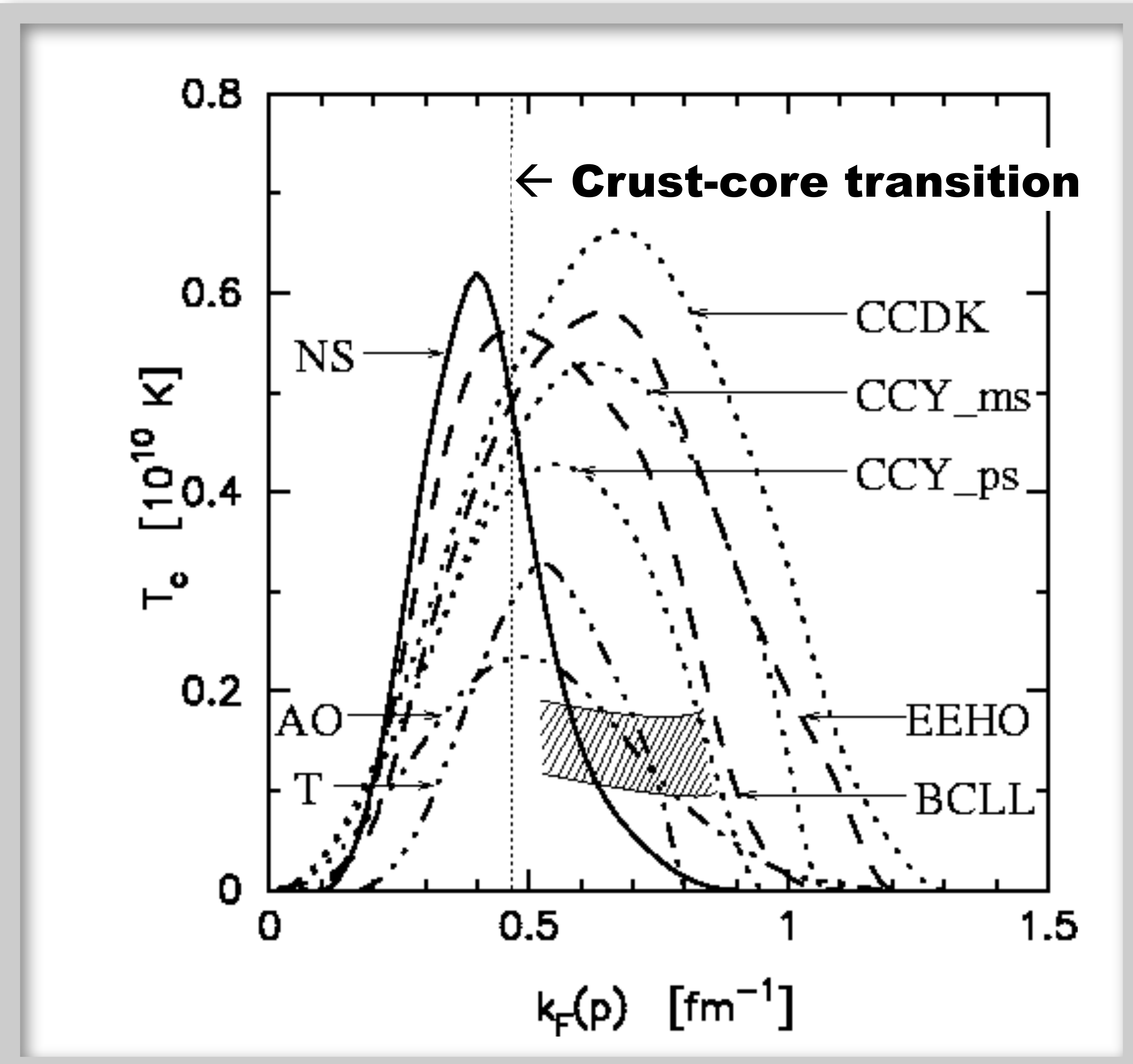
Prediction for the neutron $^1S_0$ T_c



Important feature:

Medium polarization effects reduce T_c by a factor three

Prediction for the proton $^1S_0$ T_c



T: Takatsuka, Prog. Thero. Phys. 50 (1970), 905

CCY: Chao, Clark & Yang, Nucl. Phys. A179 (1972), 320

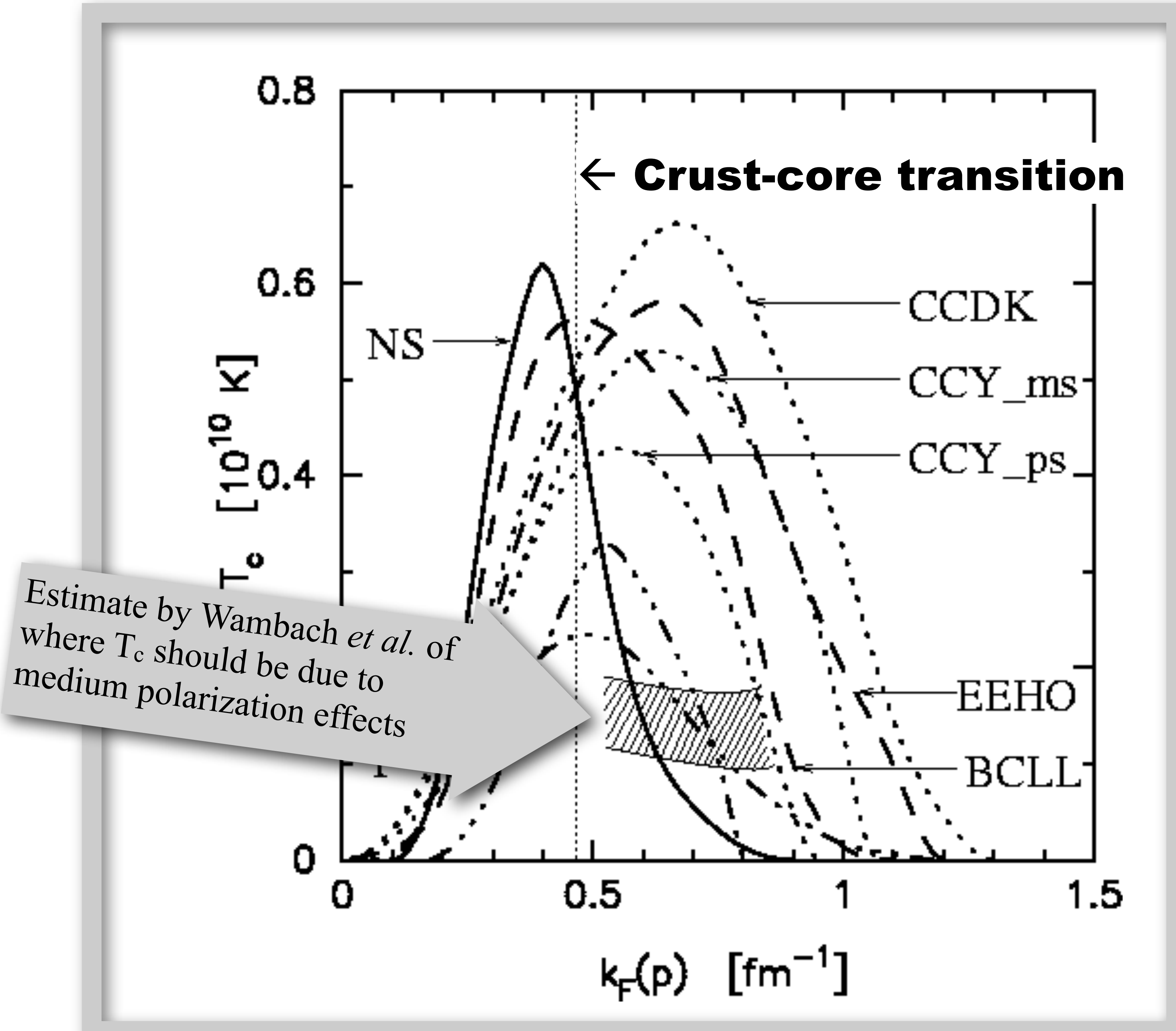
AO: Amundsen & Osgaard, Nucl. Phys. A437 (1985), 487

BCLL: Baldo, Cugnon, Lejeune & Lombardo, Nucl. Phys. A536 (1992), 349

CCDK: Chen, Clark, Davé & Khodel, Nucl. Phys. A555 (1993), 59

EEHO: Elgaroy, Engvik, Horth-Jensen & Osnes, Nucl. Phys. A604 (1996), 466

Prediction for the proton $^1S_0$ T_c



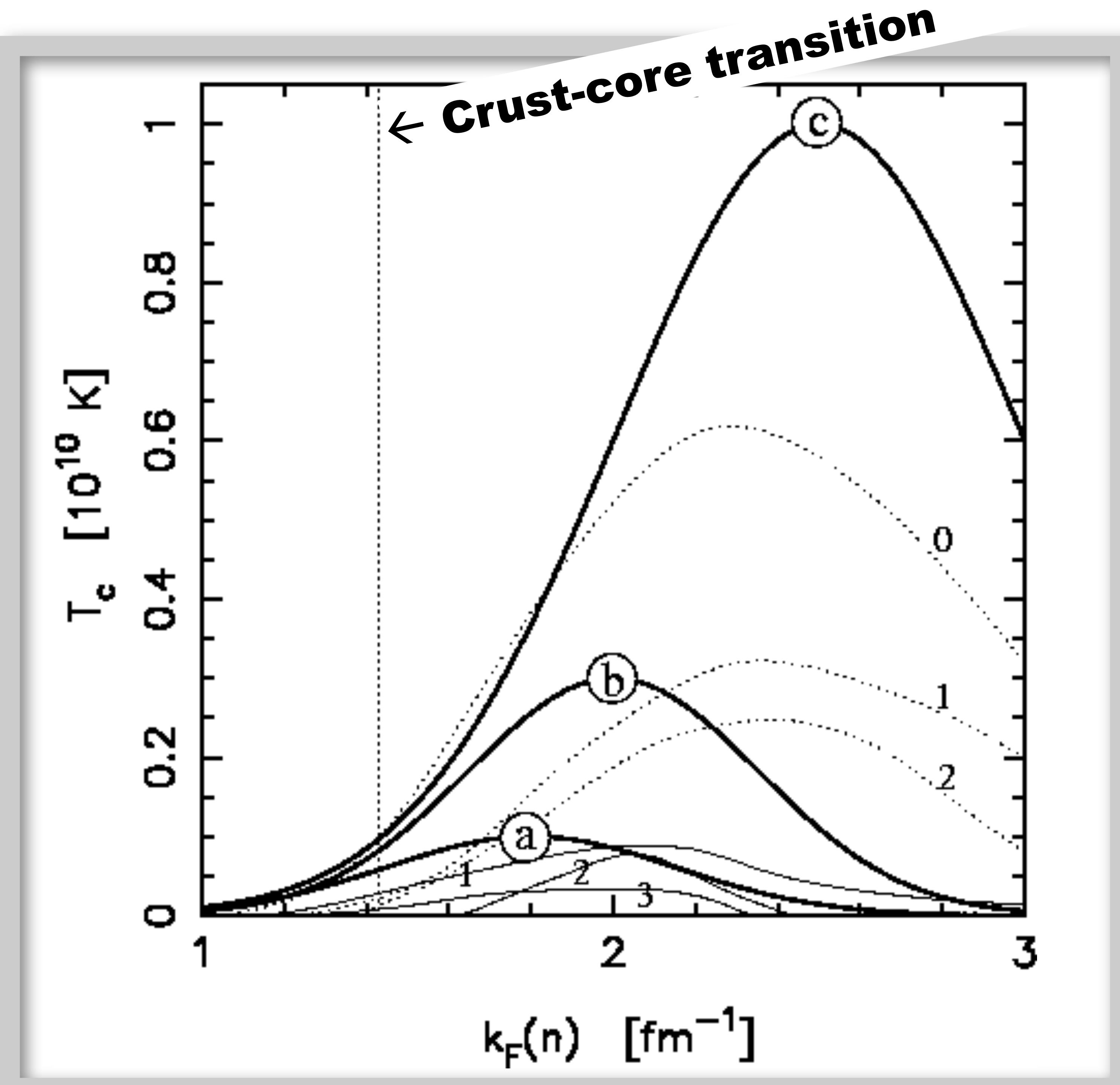
Important features:

All vanish at p_F >1.3 fm^{-1} and most at p_F > 1 fm^{-1}

Expected maximum T_c ~ 1 - 2 x 10^9 K

Medium polarization effects seem to reduce T_c by a factor three

Prediction for the neutron $^3P_2$ T_c



0: Hoffberg, Glassgold, Richardson & Ruderman, Phys. Rev. Lett. 24 (1970), 775

1: Amundsen & Osgaard, Nucl. Phys. A442 (1985), 4163

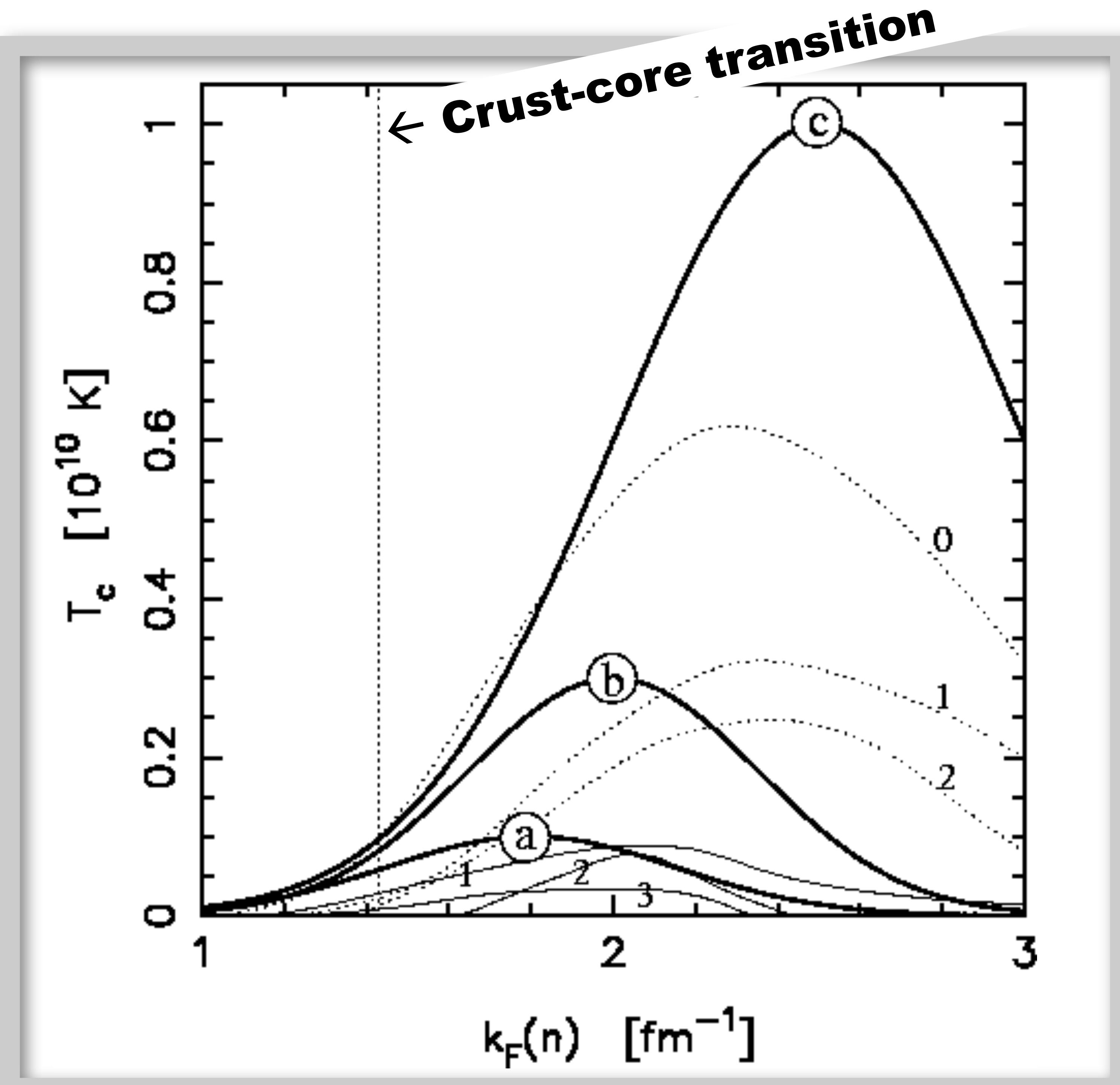
2: Takatsuka, Prog. Theor. Phys. 48 (1972), 1517

a, b, c:

Baldo, Elgaroy, Engvik, Horth-Jensen & Schulze,

Phys. Rev. C58 (1998), 1921

Prediction for the neutron $^3P_2$ T_c



Important feature:

WE DO NOT REALLY KNOW WHAT IT IS

Medium polarization effects were expected to increase the $^3P_2$ gap while they probably strongly suppress it.

Problem with the $^3P_2$ phase-shift

No model of the nucleon-nucleon interaction reproduces the experimental $^3P_2$ phase-shift above 300 MeV.

The T_c curves a, b, & c on the previous slide reflect this uncertainty:
any one of them is possible

$E_{lab} \leftrightarrow k_F$ conversion: $E_{lab} = (2k_F)^2/2m$

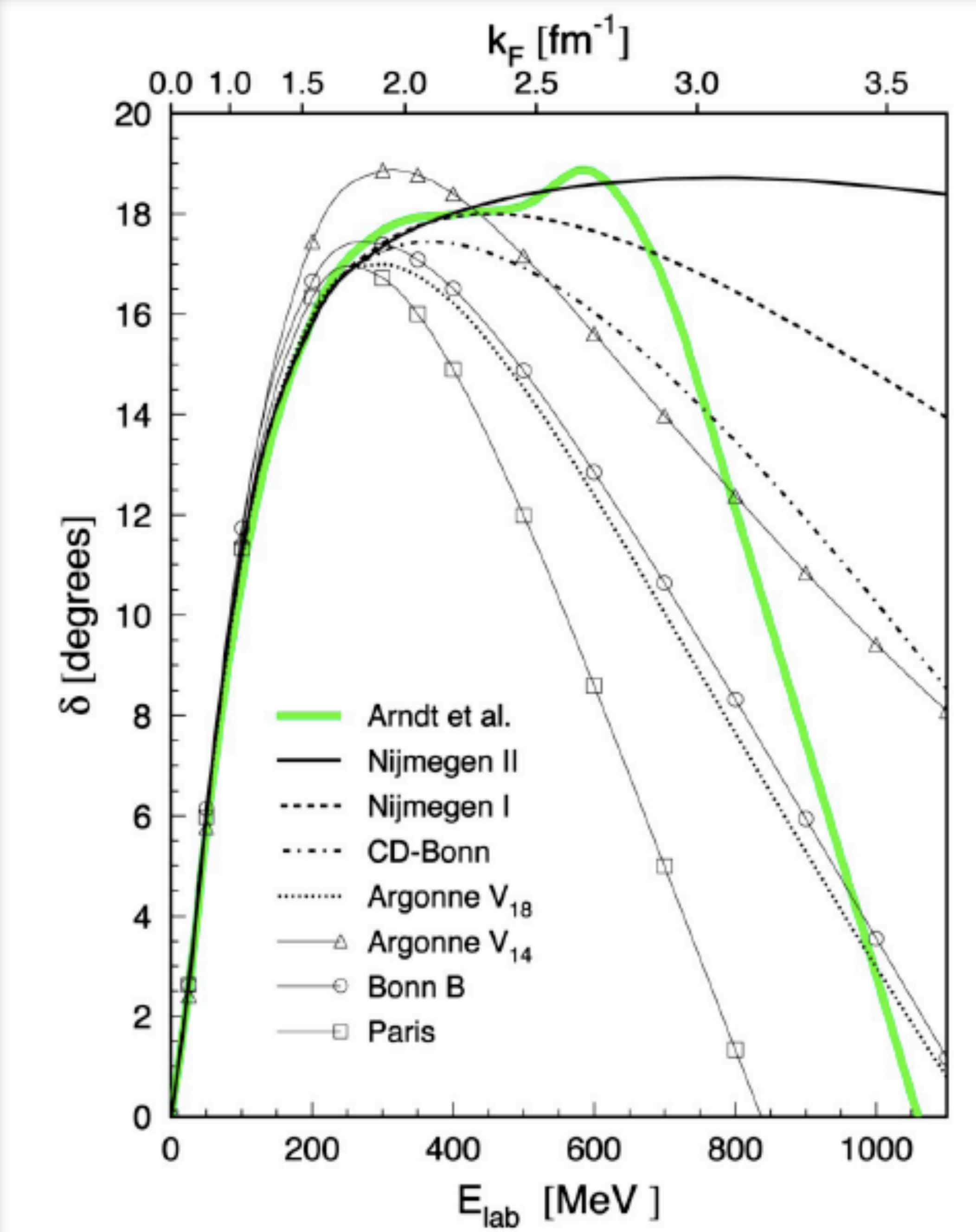


FIG. 4. $^3P_2$ phase-shift predictions of different potentials up to $E_{\rm lab}$=1.1 GeV, compared with the phase-shift analysis of Arndt *et al.* (1997) (Color in online edition).

$^3P_2$-$^3F_2$ pairing in neutron mater with modern nucleon-nucleon potentials
M. Baldo, Ø. Elgarøy, L. Engvik, M. Hjorth-Jensen, and H.-J. Schulze
Phys. Rev. C 58, 1921-1928 (1998)

Suppression of c_v and ε_ν by pairing

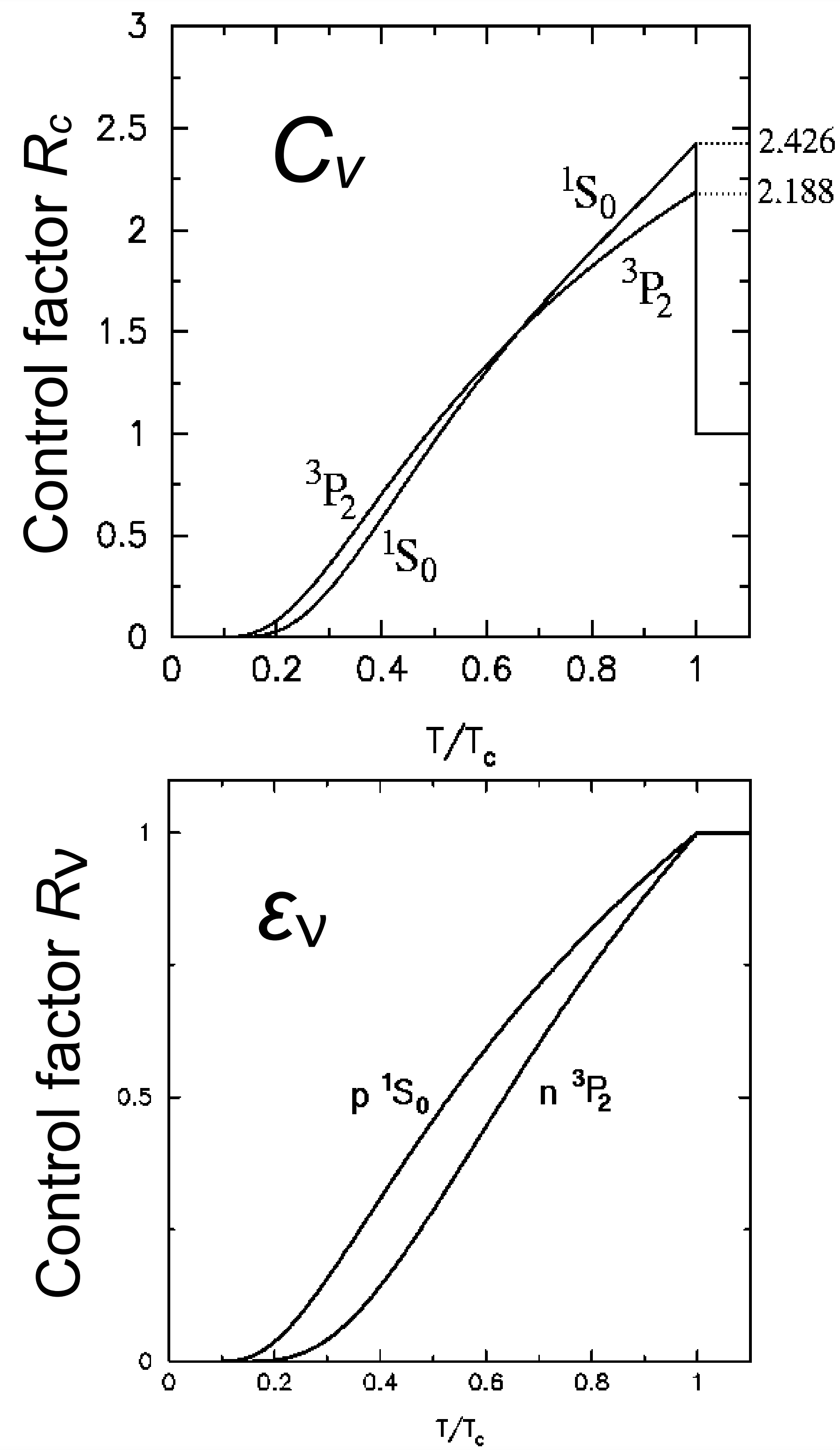
The presence of a pairing gap in the single particple excitation spectrum results in a Boltzmann-like

$\sim \exp(-\Delta/k_B T)$

suppression of c_v and ε_ν:

$$c_v \to c_v^{\text{Paired}} = R_c \, c_v^{\text{Normal}}$$

$$\epsilon_\nu \to \epsilon_\nu^{\text{Paired}} = R_\nu \epsilon_\nu^{\text{Normal}}$$



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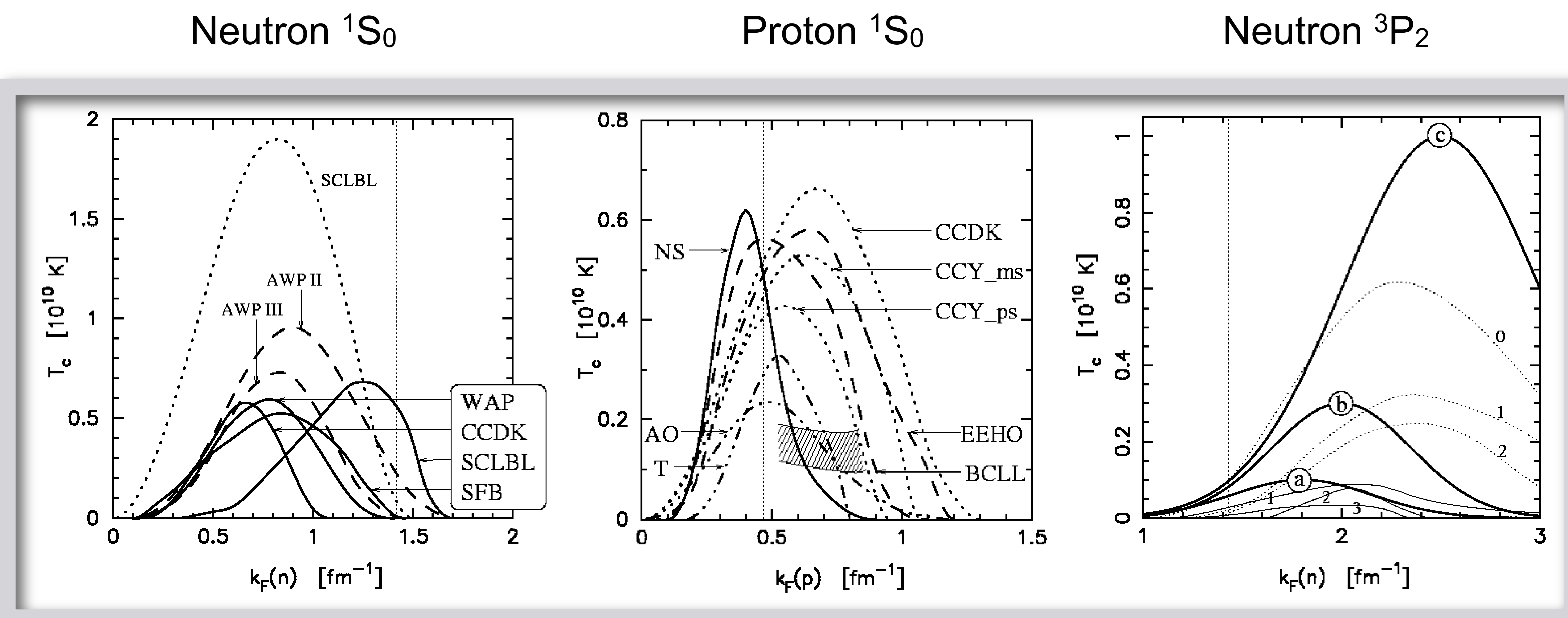
Pairing T_c models



Size and extent of pairing gaps is highly uncertain

Slow vs fast cooling with pairing

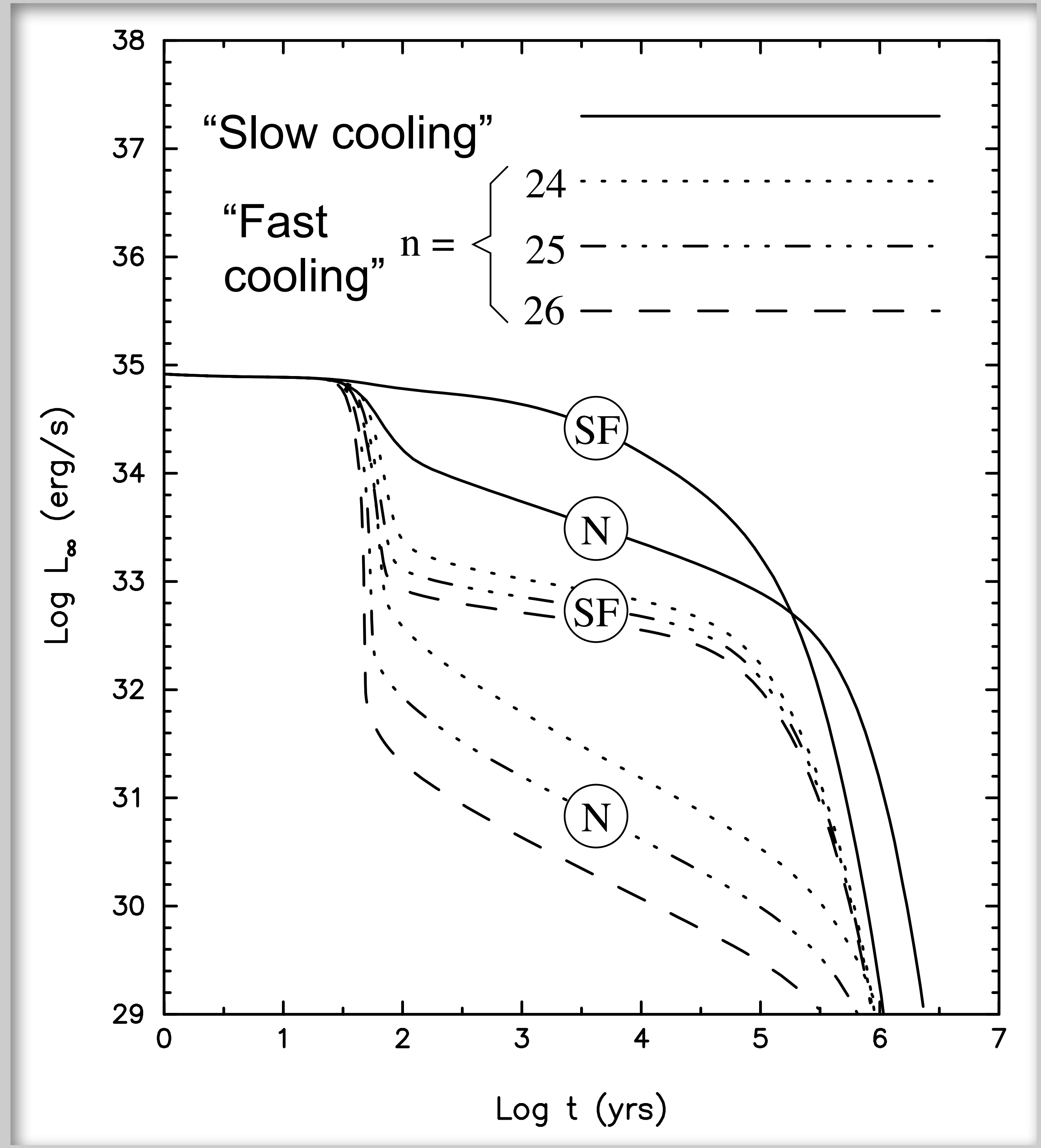
Slow neutrino emission
(modified URCA process)

$$\epsilon_\nu^{\rm slow} \sim 10^{21}\ T_9^8\ \ {\rm erg\ cm^{-3}\ s^{-1}}$$

Fast neutrino emission
(almost anything else)

$$\epsilon_\nu^{\rm fast} \sim 10^{n}\ T_9^6\ \ {\rm erg\ cm^{-3}\ s^{-1}}$$

- n = 24 ~ Kaon condensate
- n = 25 ~ Pion condensate
- n = 26 ~ Direct Urca



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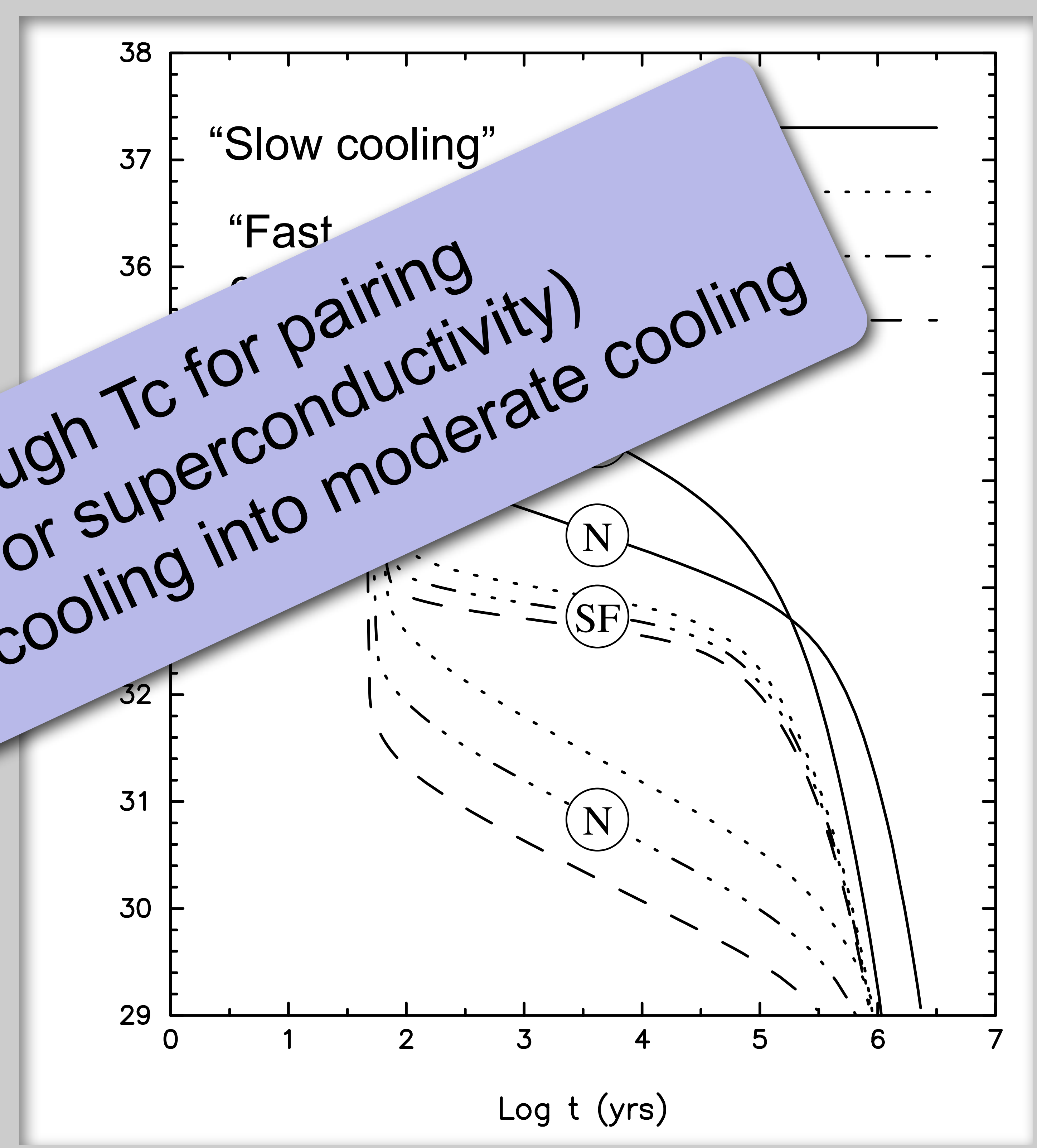
$$\epsilon_\nu^{\text{slow}} \sim 10^{21}\ T_9^8\ \ \text{erg cm}^{-3}\ \text{s}^{-1}$$

Fast neutrino emission
(almost anything else)

$$\epsilon_\nu^{\text{fast}} \sim 10^{n}\ T_9^6\ \ \text{erg c}$$

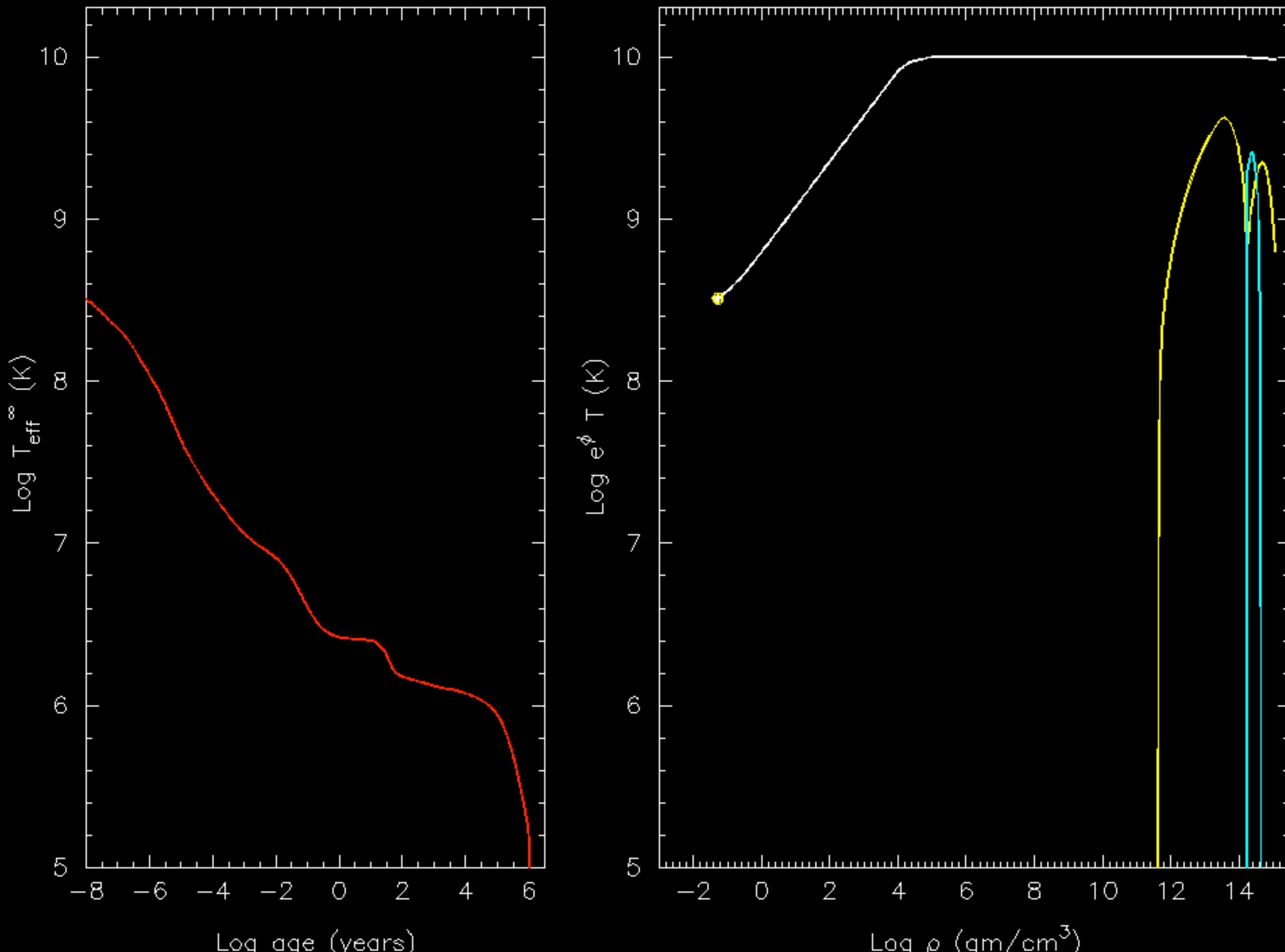
- n = 2
- n = 25
- n = 26

A large enough Tc for pairing (superfluidity or superconductivity) can transform fast cooling into moderate cooling

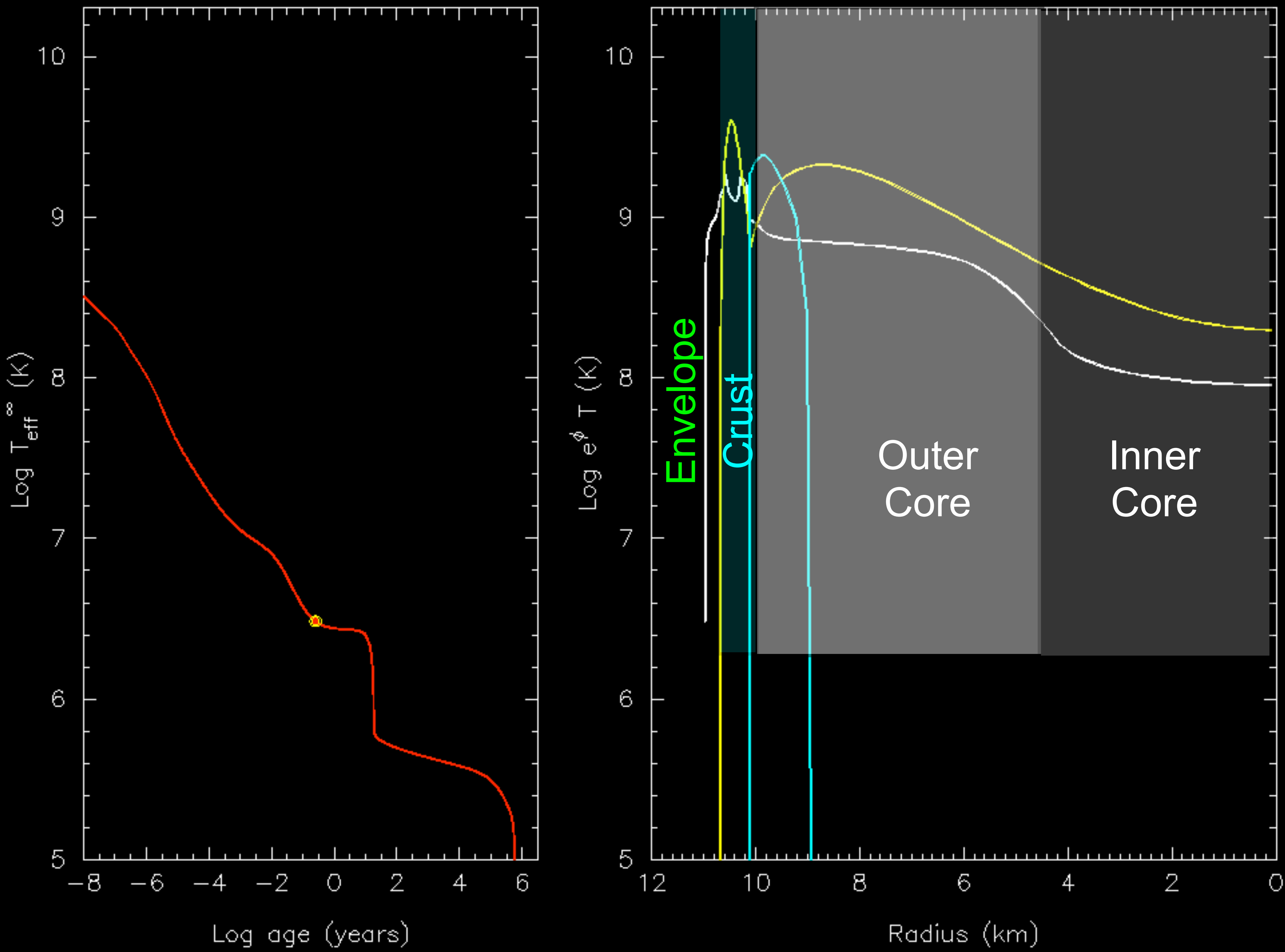


Wednesday, July 7, 2010

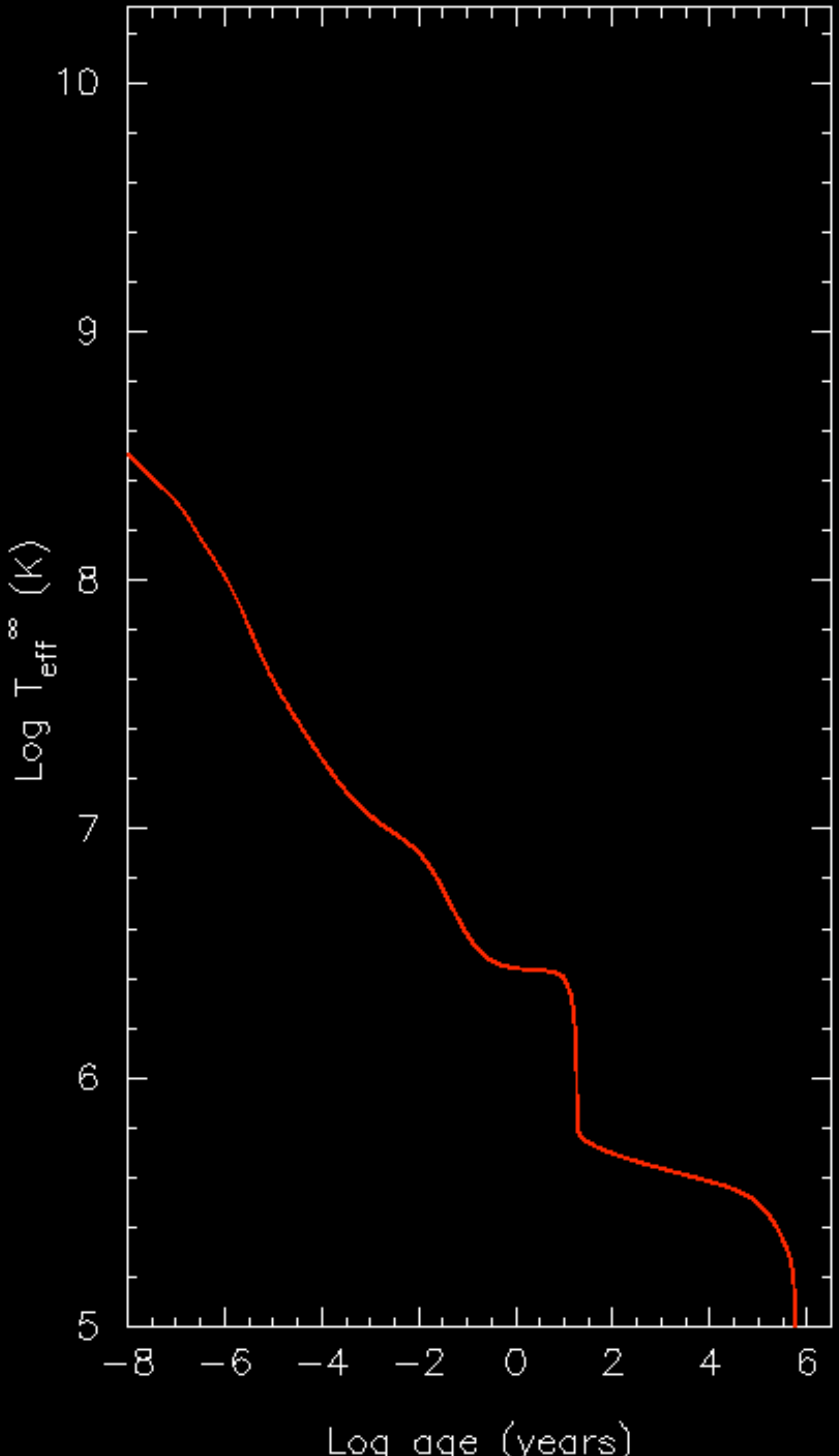
Standard cooling of a 1.3 M☉ neutron star with pairing [Animation file: 1.3_P.mov]

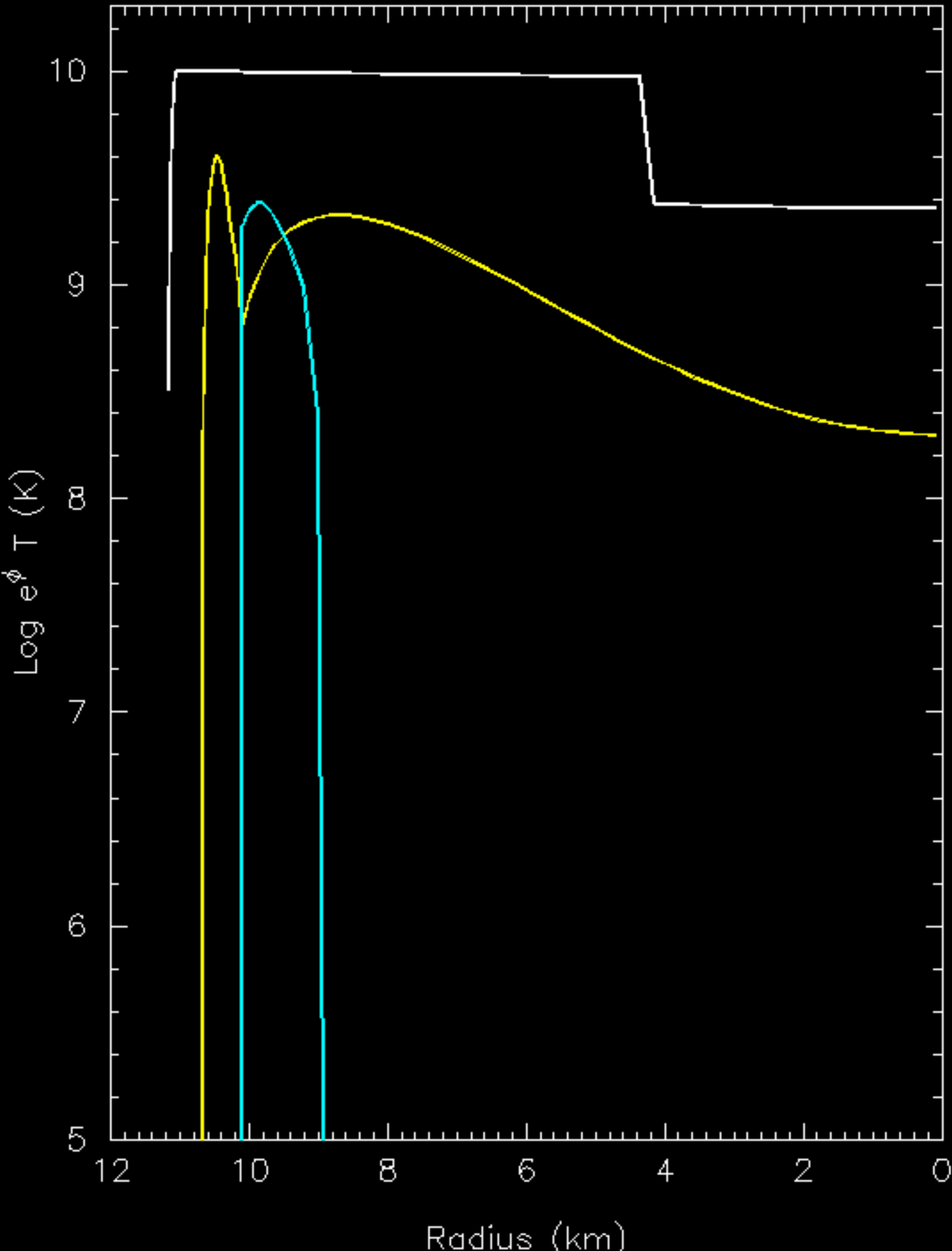


Enhanced cooling of a 1.5 M⊙ neutron star with pairing



Enhanced cooling of a 1.5 M⊙ neutron star moderated pairing [Animation file: 1.5_P.mov]





Envelopes:

Heavy vs light elements
Magnetic fields

Neutron star cooling on a napkin

Assume the star's interior is isothermal and neglect GR effects.

Thermal Energy, E_{th} ,balance:

$$\frac{dE_{th}}{dt} = C_v \frac{dT}{dt} = -L_\gamma - L_\nu + H$$

⇒ 3 essential ingredients are needed:

- C_v = total stellar specific heat
- L_γ = total surface photon luminosity
- L_ν = total stellar neutrino luminosity

H = "heating", from B field decay, friction, etc ...

Envelope models

GUDMUNDSSON, PETHICK, AND EPSTEIN

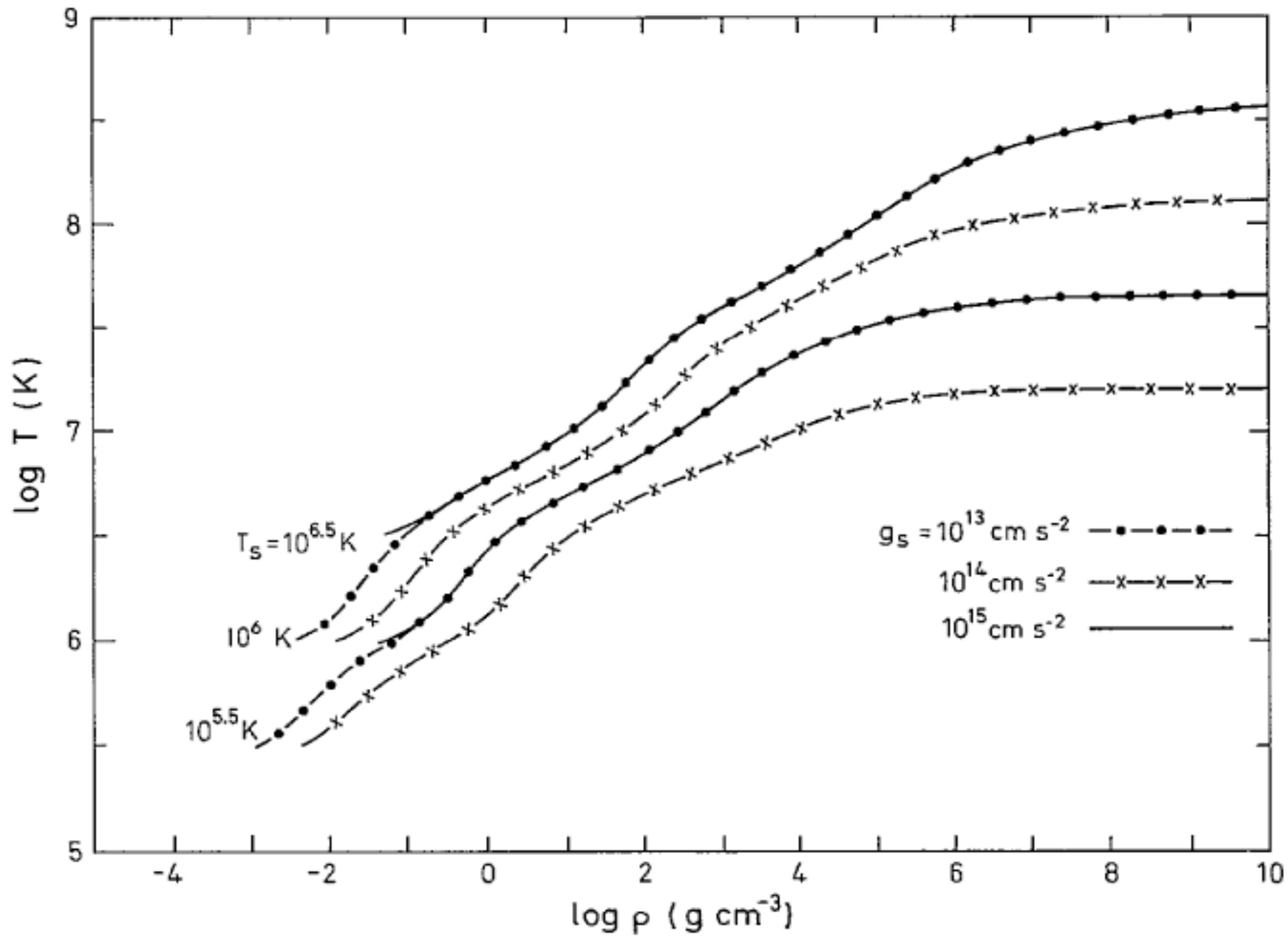


FIG. 2.—Temperature profiles for three values of the surface temperature T_s and various values of the surface gravity g_s

Neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I., 1982ApJ...259L..19G

Structure of neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I. 1983ApJ...272..286G

The Eddington "Atmosphere"

Eddington approximation:

- F is uniform
- Diffusion approximation:

$$F = -\frac{c}{3\overline{\kappa}\rho}\frac{d\epsilon^P}{dr} = \frac{c}{3}\frac{d(aT^4)}{d\tau}$$

$$(d\tau = -\overline{\kappa}\rho dr)$$

- No incident radiation at $\tau = 0$

Integrate to get $aT^4 = 3\tau F/c + Constant$ and $Constant = aT_0^4$ (at $\tau = 0$).

No incident radiation at τ=0 gives:

$$F = \frac{1}{2}c\epsilon_0^P = \frac{1}{2}acT_0^4$$

$$aT^4 = \frac{3\tau}{c}\frac{1}{2}acT_0^4 + aT_0^4 = aT_0^4\left(1 + \frac{3}{2}\tau\right)$$

Effective temperature:

$$F = \sigma T_e^4 = \frac{1}{2}acT_0^4 \implies T_e^4 = 2T_0^4$$

$$\boxed{T^4 = \frac{1}{2}T_e^4\left(1 + \frac{3}{2}\tau\right)}$$

Photosphere at: $\boxed{\tau = \frac{2}{3}}$

$$\kappa\rho\Delta z|_P = \kappa\frac{P_P}{g} = \frac{2}{3}$$ (g=surface gravity) gives:

$$\boxed{P_P = \frac{2g_s}{3\overline{\kappa}_P}}$$ **Eddington condition**

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$$(d\tau = -\overline{\kappa}\rho dr)$$

- No incident radiation at $\tau = 0$

In ... *stant*

a

N

Self-inconsistent:
diffusion approximation breaks down near the photosphere because I_ν becomes highly anisotropic !

$$aT \;\ldots\; \frac{\;}{c}\;2 \;\ldots\; 0 \;\ldots\; 0 \;\ldots\; 0\left(\ldots + \frac{3}{2}\tau\right)$$

Effective temperature:

$$F = \sigma T_e^4 = \frac{1}{2}acT_0^4 \implies T_e^4 = 2T_0^4$$

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$$\boxed{P_P = \frac{2g_s}{3\overline{\kappa}_P}}$$ **Eddington condition**

Atmosphere Temperature Profiles

The simple results of the Eddington approximation give a honest qualitative description of the photosphere when compared with detailed models:

$$T^4 = \frac{1}{2} T_e^4 \left(1 + \frac{3}{2} \tau \right)$$

$$\tau = \frac{2}{3}$$

$$P_P = \frac{2 g_s}{3 \overline{\kappa}_P}$$

GR effects: $$g_s = \frac{GM}{R^2} \frac{1}{\sqrt{1 - 2GM/c^2 R}}$$

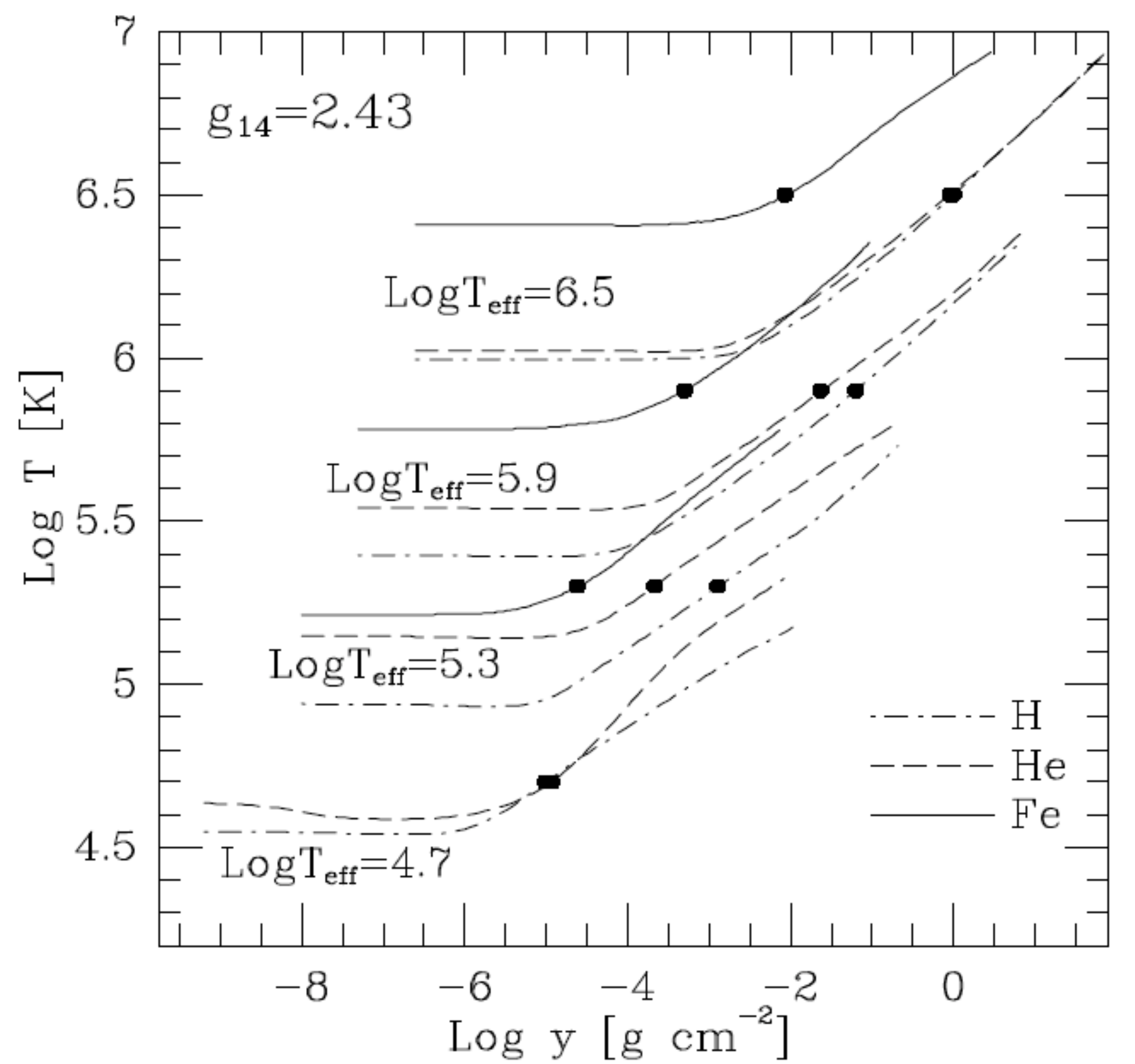


Fig. 2. Temperature profiles $T(y)$ of the models with different effective temperatures and chemical compositions. The filled circles mark the points where $T(y) = T_{\rm eff}$; the corresponding $y = y_{\rm eff}$ can be considered as a characteristic depth of the atmosphere.

Model neutron star atmospheres with low magnetic fields. I. Atmospheres in radiative equilibrium. Zavlin, V. E.; Pavlov, G. G.; Shibanov, Yu. A. 1996A&A...315..141Z

Opacities

Free electron scattering (Thomson):

$$\kappa_{es} = \frac{\sigma_T n_e}{\rho} = 0.40 f_e \ \mathrm{cm}^2 \ \mathrm{g}^{-1}$$

$$\sigma_T = \frac{8\pi}{3} \left(\frac{e^2}{m_e c^2} \right)^2 = 0.665 \times 10^{-24} \ \mathrm{cm}^2$$

$$n_e = f_e \rho / m_p \quad \text{or} \quad f_e = \frac{Z}{A}$$

Free-free absorption (inverse bremsstrahlung) and **bound-free absorption** (photoionization):

κ_ν has an overall ν^{-3} frequency dependence and

$$\overline{\kappa} \propto \rho T^{-3.5}$$

this type of ρ T dependence is usually called a **Kramer opacity**

For more details see: Shapiro & Teukolsky´s book, Appendix I

A simple envelope model

In a thin envelope: $m=M$, $r=R$, and L (and F) are uniform:

$$\frac{dT}{dr} = -\frac{1}{\lambda}F = -\frac{3\kappa\rho}{4acT^3}\frac{L}{4\pi R^2}$$

$$\frac{dP}{dr} = -\frac{GM}{R^2}\rho$$

$$\Longrightarrow \quad \frac{\partial T}{\partial P} = \frac{3L}{16\pi acGM}\frac{\kappa}{T^3}$$

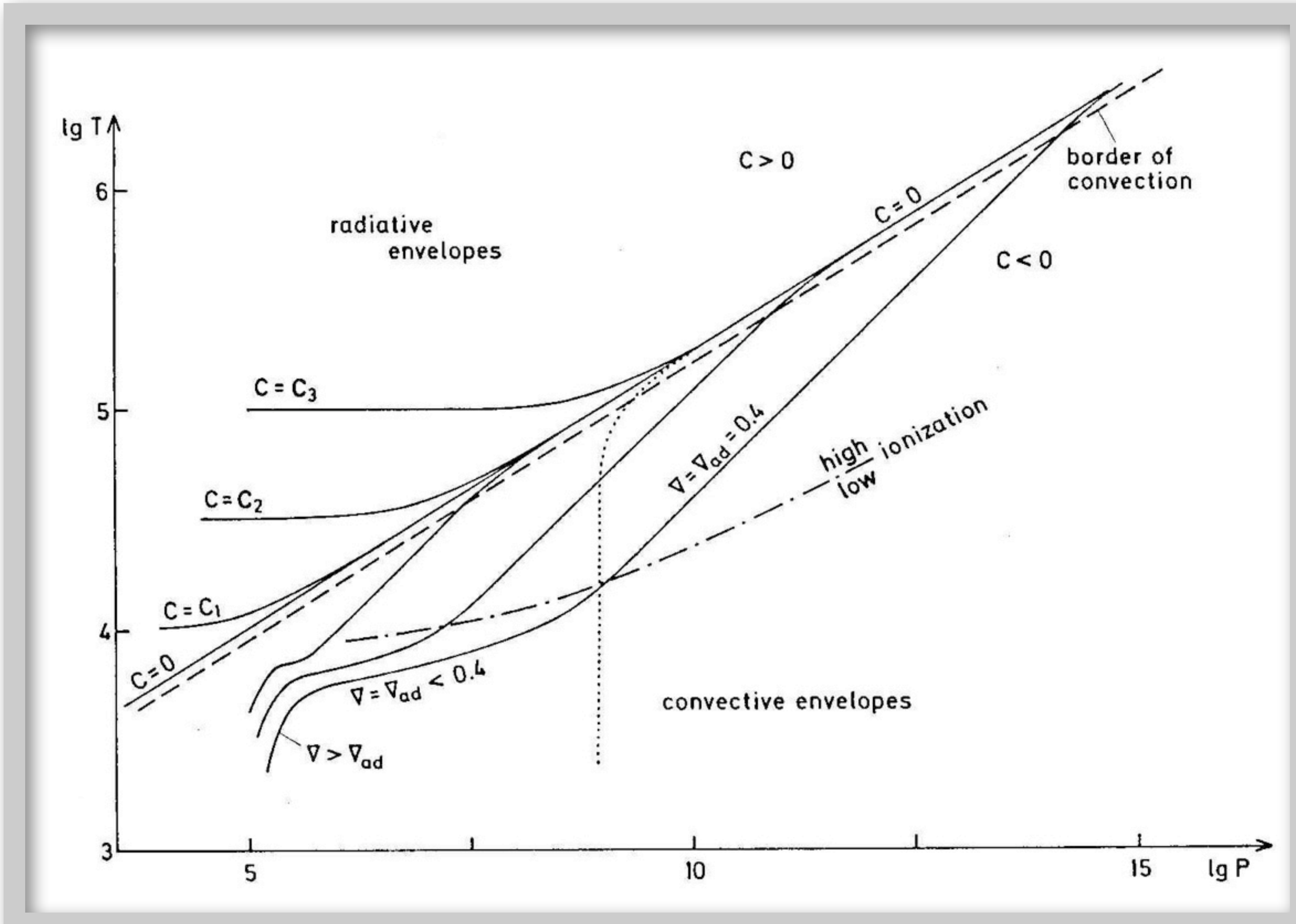
Kramer opacity: $\kappa = \kappa_0\, \rho\, T^{-3.5}$ and ideal gas $P = (R/\mu)\, \rho T \Rightarrow \kappa = \kappa_P\, P\, T^{-4.5}$ and thus:

$$\frac{T^{7.5}}{P}\frac{\partial T}{\partial P} = \frac{3\kappa_P}{16\pi acG}\frac{L}{M} \Longrightarrow \boxed{T^{8.5} = B(P^2 + C)}$$

C=0 : "zero solution" gives P=T=ρ=0 at the surface

C = integration constant: has to be determined by an appropriate photospheric boundary condition

Convergence to Zero Solution



CONCLUSION: all solutions (for various choices of C) converge toward the "zero solution": P=T=ρ at the "surface"

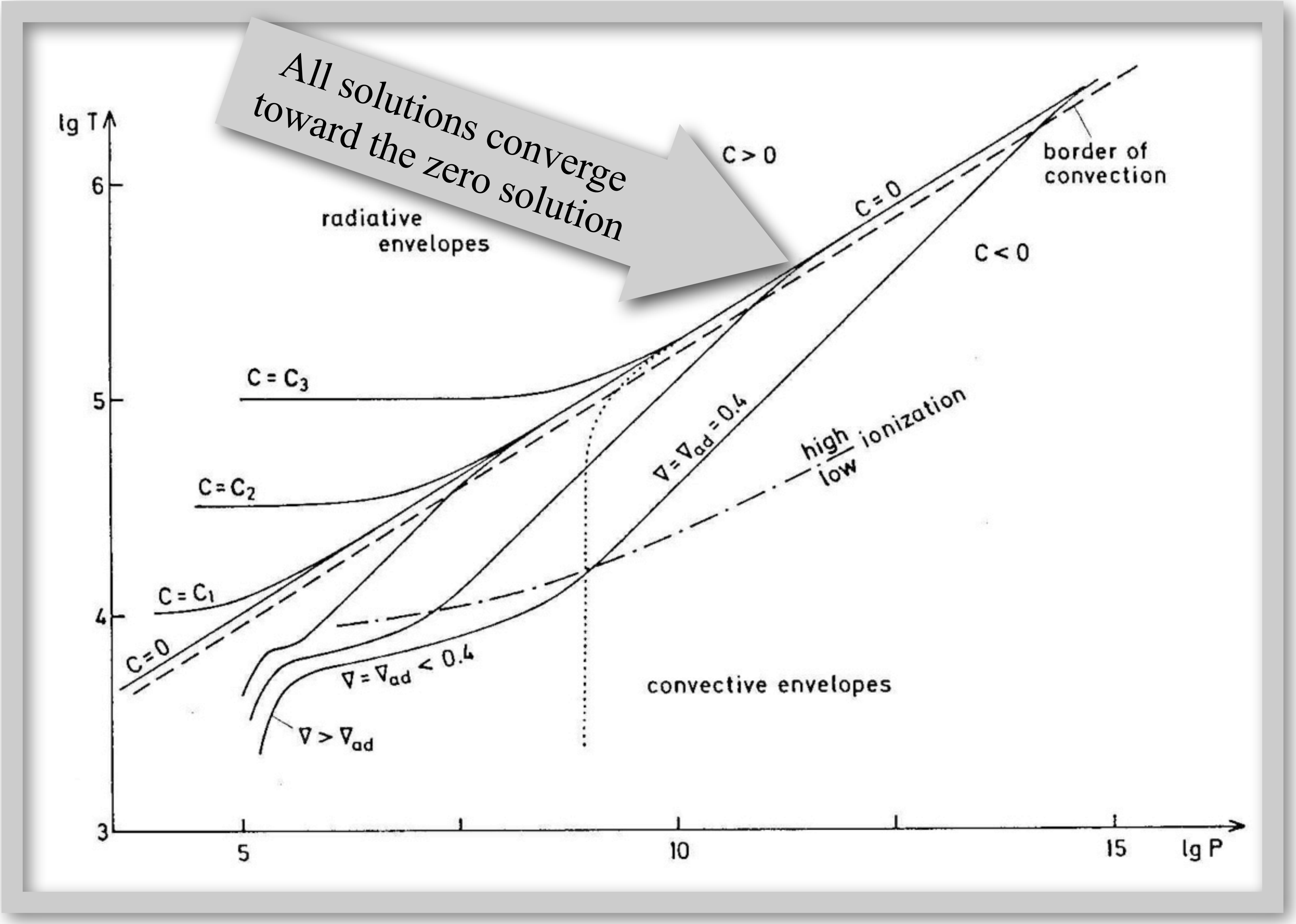
$$T^{8.5} = B(P^2 + C)$$

One can be sloppy at the surface: no effect deeper inside the star !

In practice: the Eddington boundary condition ($P_P = 2g_s/3\kappa_P$) is OK

Stellar Structure and Evolution
R. Kippenhahn & A. Weigert, A&A Library, 1990

Convergence to Zero Solution



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Stellar Structure and Evolution
R. Kippenhahn & A. Weigert, A&A Library, 1990

Conclusion from all this:

At low densities (near the photosphere) things are very complicated and one must do a complete detailed (frequency dependent) radiative transfer calculation to model the atmosphere correctly and obtain the spectrum.

Deeper into the envelope the diffusion approximation is reliable, one can use frequency averaged Rosseland means, and the resulting temperature profile is essentially independent of the assumed T profile in the atmosphere (one can safely use the Eddington approximation for the atmosphere)

The GPE envelope models

Ingredients:

Thin plane parallel layer with

m=M, r=R

$L=4\pi R^2\,\sigma T_e^4$ uniform in the envelope

Neglect neutrinos

$$\frac{dT}{dP} = \frac{3}{16}\frac{\kappa}{T^3}\frac{T_e^4}{g_s}$$

$$g_s = \frac{GM}{R^2}\frac{1}{\sqrt{1-2GM/c^2R}}$$

Los Alamos opacity tables
and equation of state for pure iron

GUDMUNDSSON, PETHICK, AND EPSTEIN

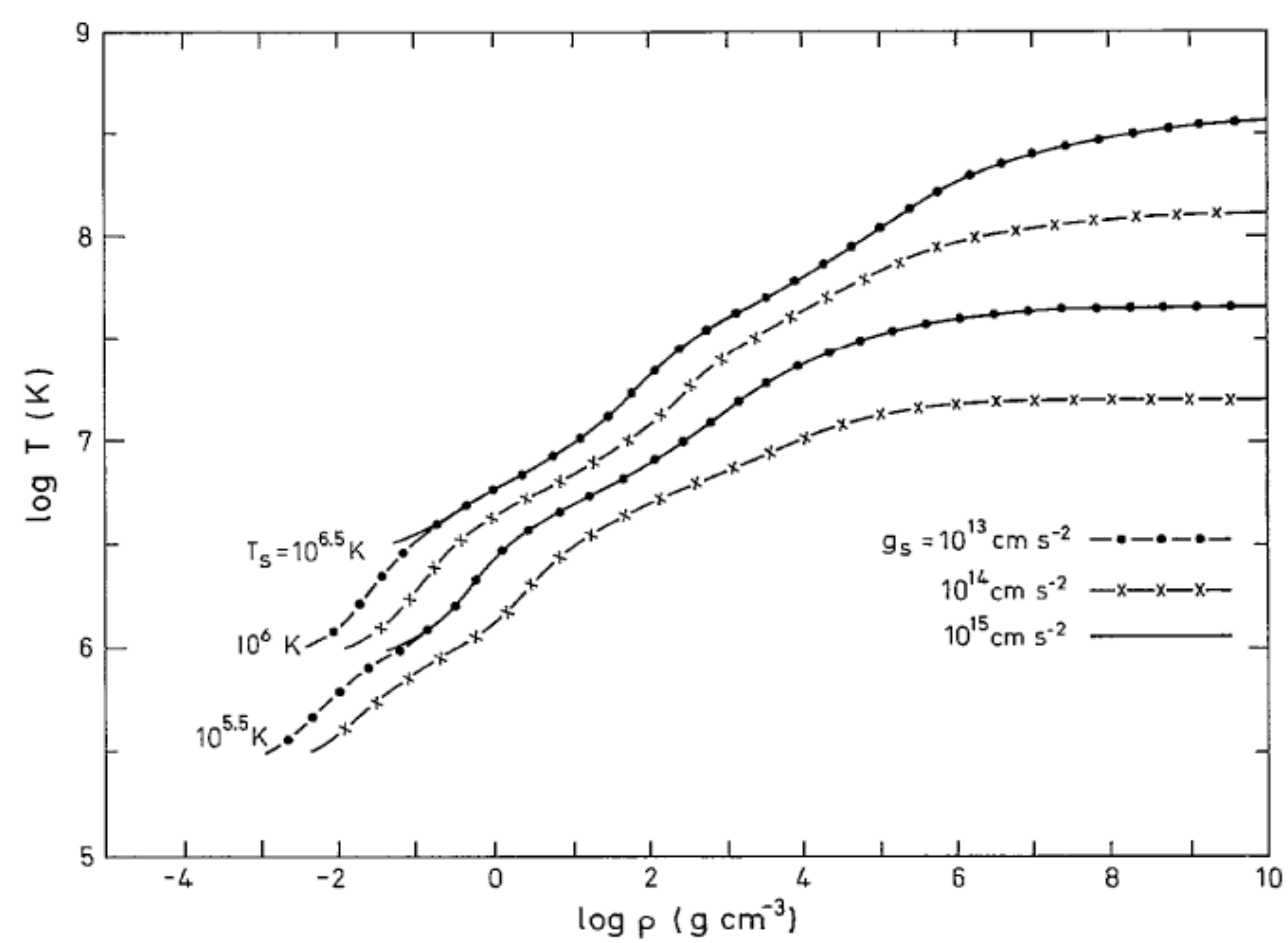

FIG. 2.—Temperature profiles for three values of the surface temperature T_s and various values of the surface gravity g_s

RESULT: "T_b - T_e" relationship. T_b = T at $\rho_B = 10^{10}$ g cm^{-3}

Neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I., 1982ApJ...259L..19G

Structure of neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I. 1983ApJ...272..286G

The GPE envelope models

Ingredients:

$$T_{b\,8} = 1.288 \left(\frac{T_{s\,6}^4}{g_{s\,14}} \right)^{0.455}$$

$$T_{s\,6} = 0.87\, g_{s\,14}^{1/4}\, T_{b\,8}^{0.55}$$

Neglect neutrinos

$$\frac{dT}{dP} = \frac{3}{16} \frac{\kappa}{T^3} \frac{T_e^4}{g_s}$$

$$g_s = \frac{GM}{R^2} \frac{1}{\sqrt{1 - 2GM/c^2R}}$$

Los Alamos opacity tables
and equation of state for pure iron

GUDMUNDSSON, PETHICK, AND EPSTEIN

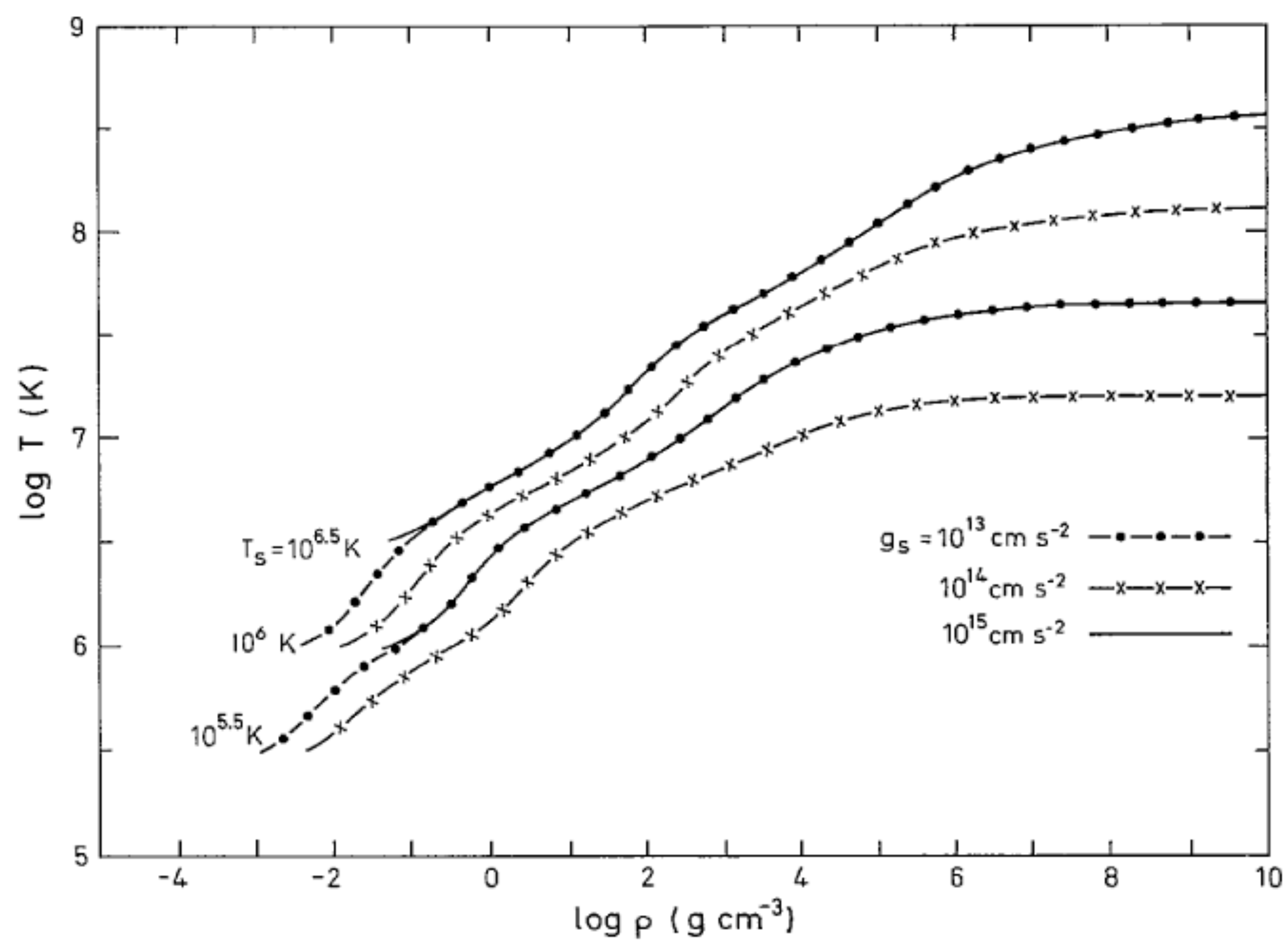

FIG. 2.—Temperature profiles for three values of the surface temperature T_s and various values of the surface gravity g_s

RESULT: "T_b - T_e" relationship. T_b = T at $\rho_B = 10^{10}$ g cm^{-3}

Neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I., 1982ApJ...259L..19G

Structure of neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I. 1983ApJ...272..286G

Physical conditions in the GPE model

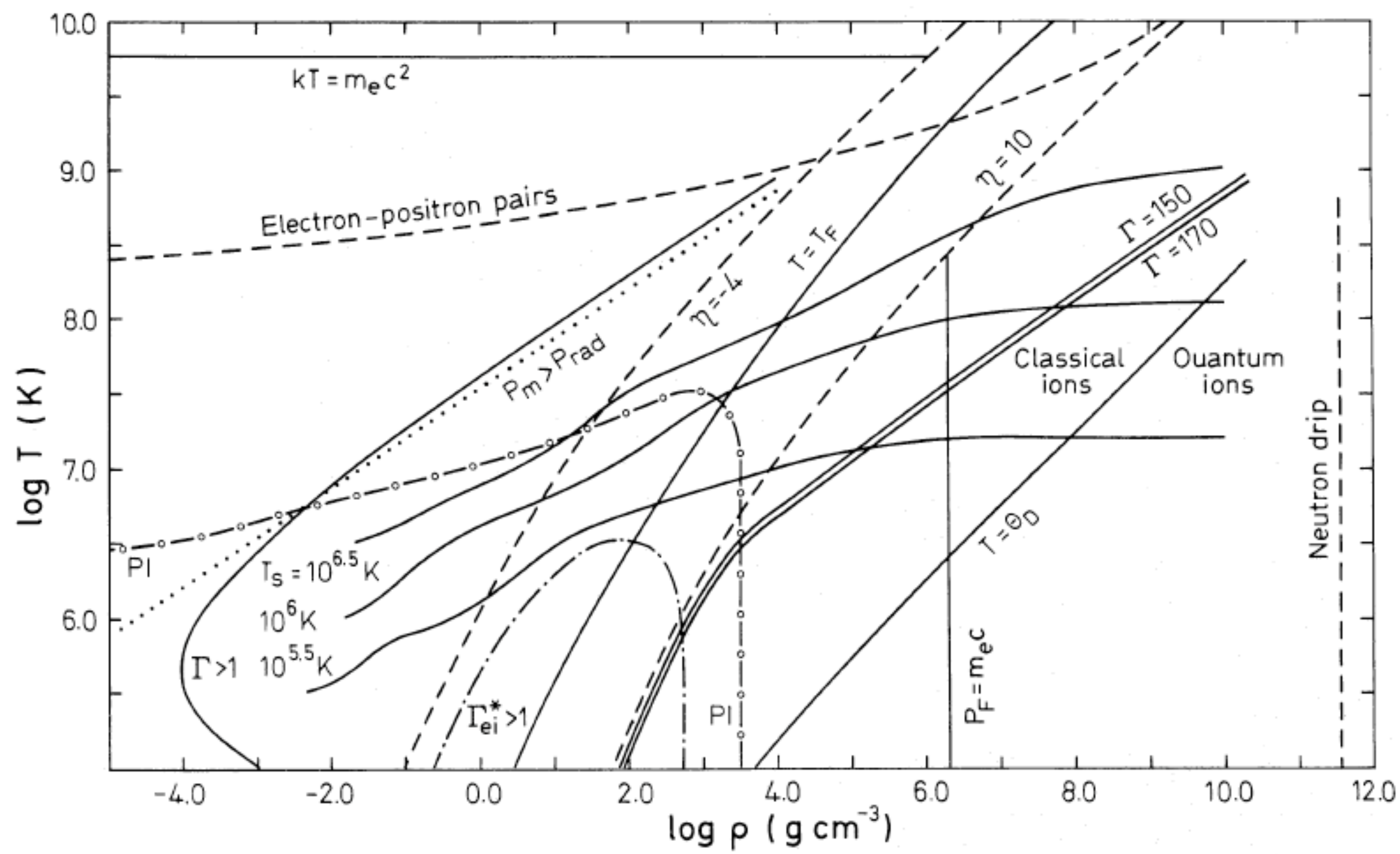

FIG. 1.—Physical conditions at densities and temperatures of interest in the study of neutron star envelopes. The various regions are identified in the text. Also shown are temperature-density profiles for envelopes for three values of the surface temperature and a surface gravity of 10^{14} cm s^{-2}.

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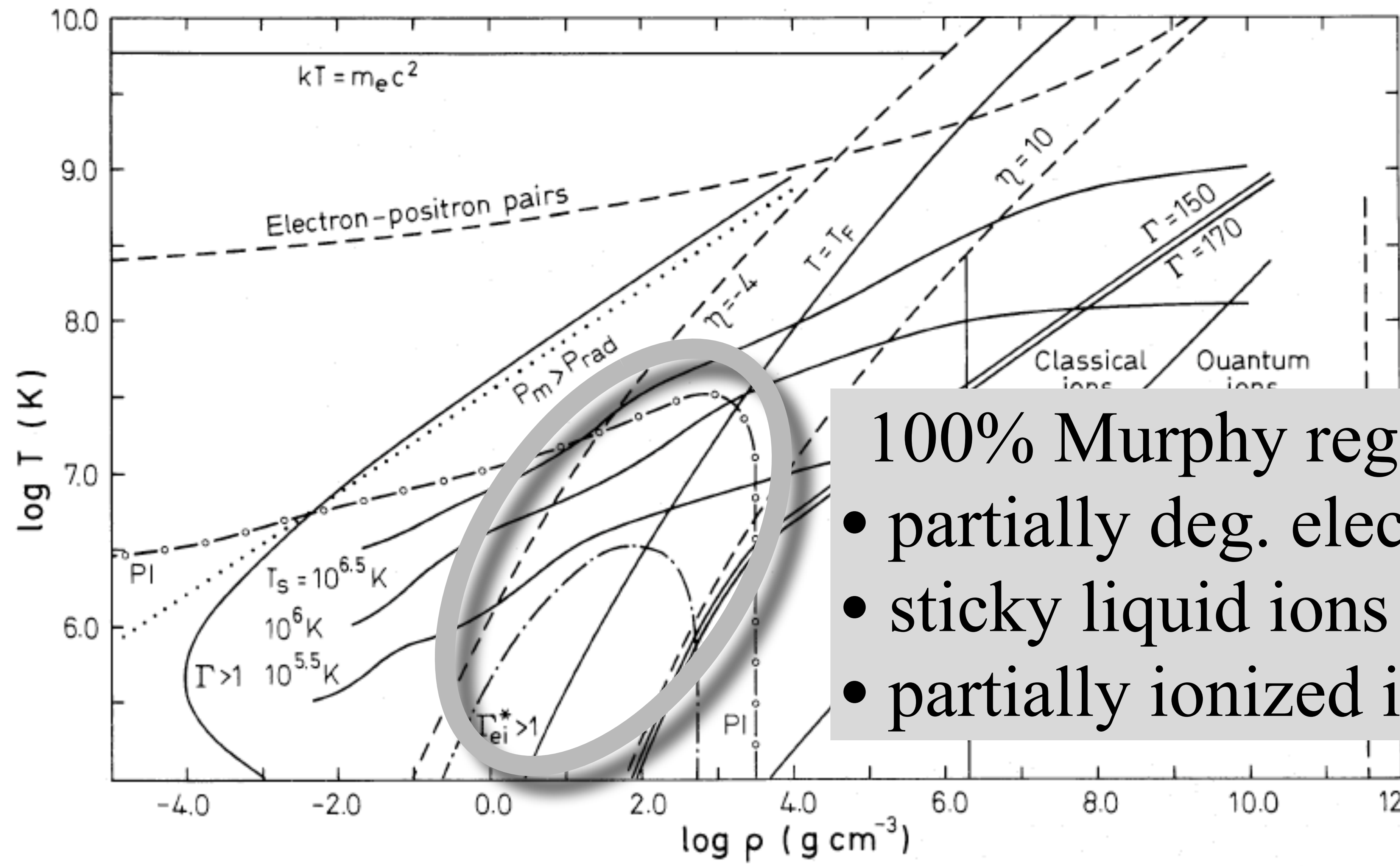

100% Murphy regime:

- partially deg. electrons
- sticky liquid ions
- partially ionized ions

FIG. 1.—Physical conditions at densities and temperatures of interest in the study of neutron star envelopes. The various regions are identified in the text. Also shown are temperature-density profiles for envelopes for three values of the surface temperature and a surface gravity of 10^{14} cm s^{-2}.

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Sources of opacity in the GPE model

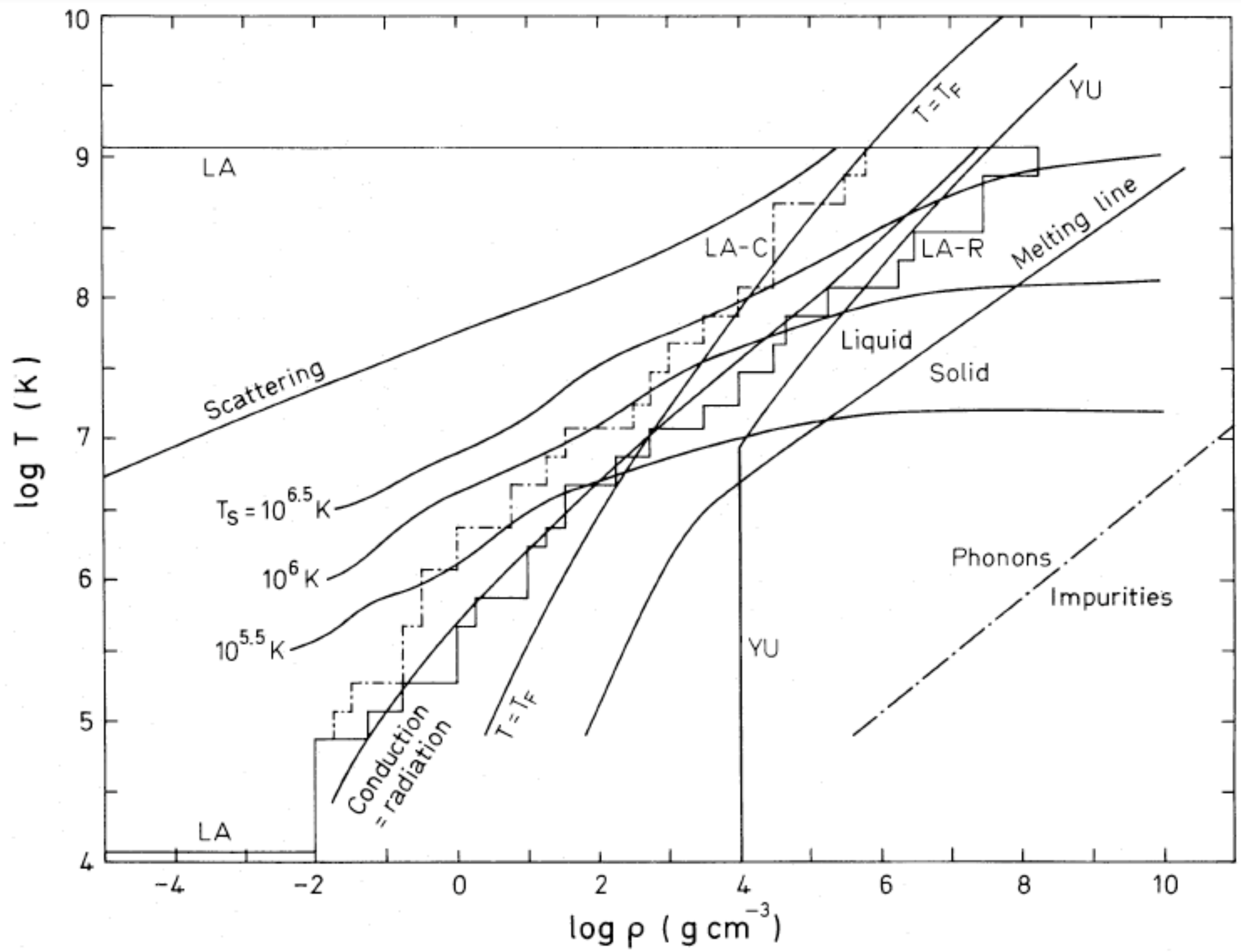

FIG. 2.—The dominant sources of opacity at various densities and temperatures. Also shown are temperature-density profiles for neutron star envelopes for three values of the surface temperature and a surface gravity of 10^{14} cm s^{-2}. See text for further explanations.

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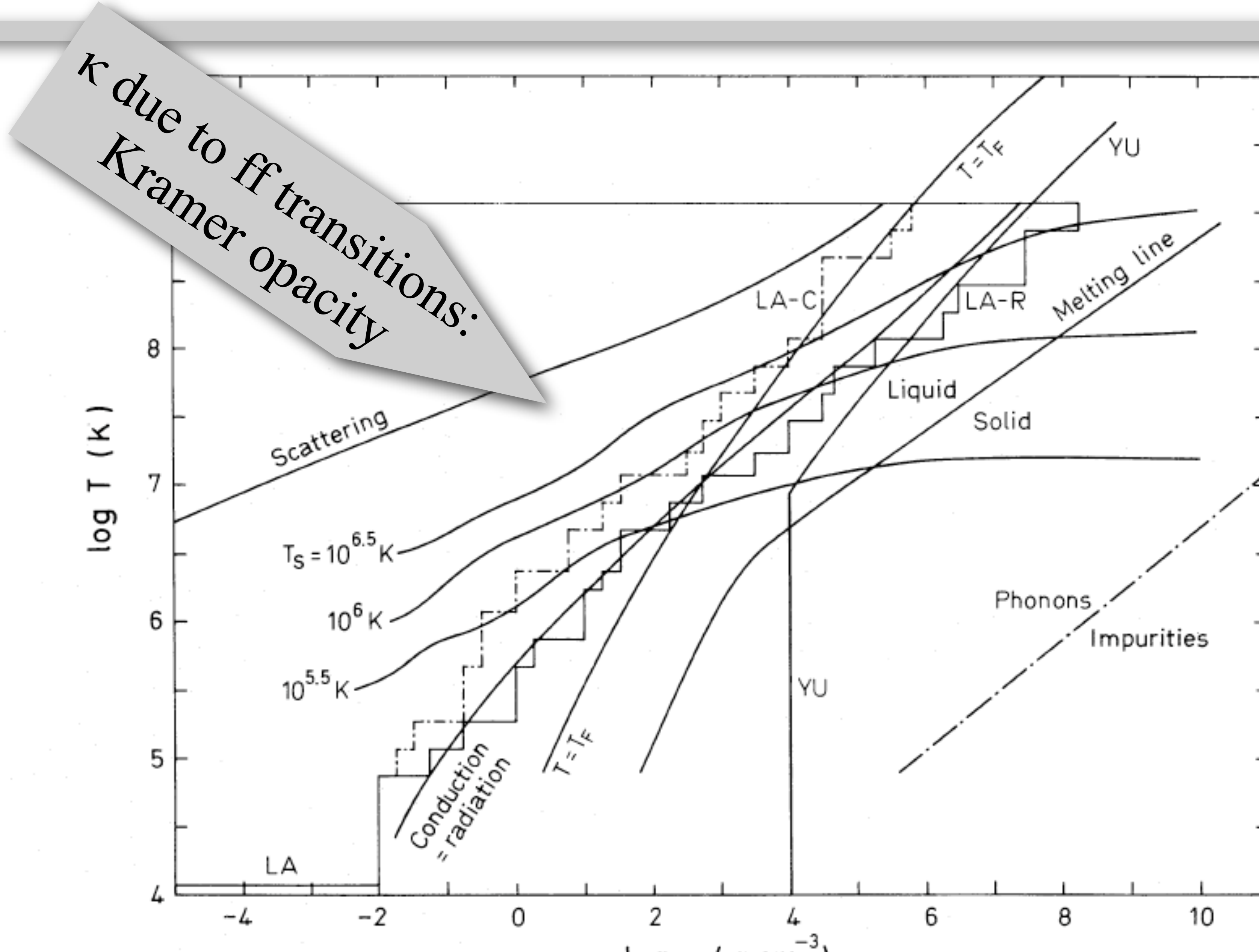

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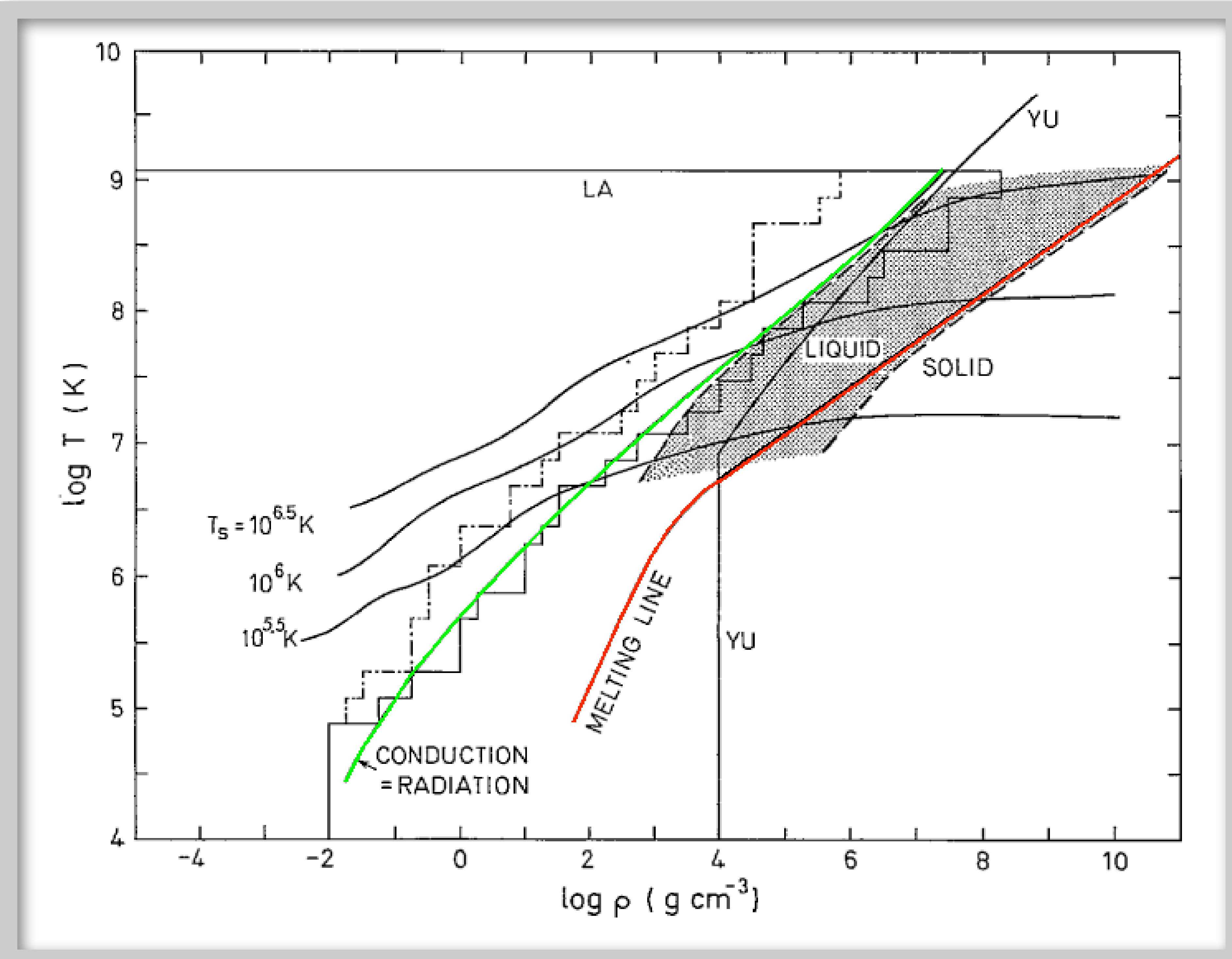
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I. 1983ApJ...272..286G

The Sensitivity Strip



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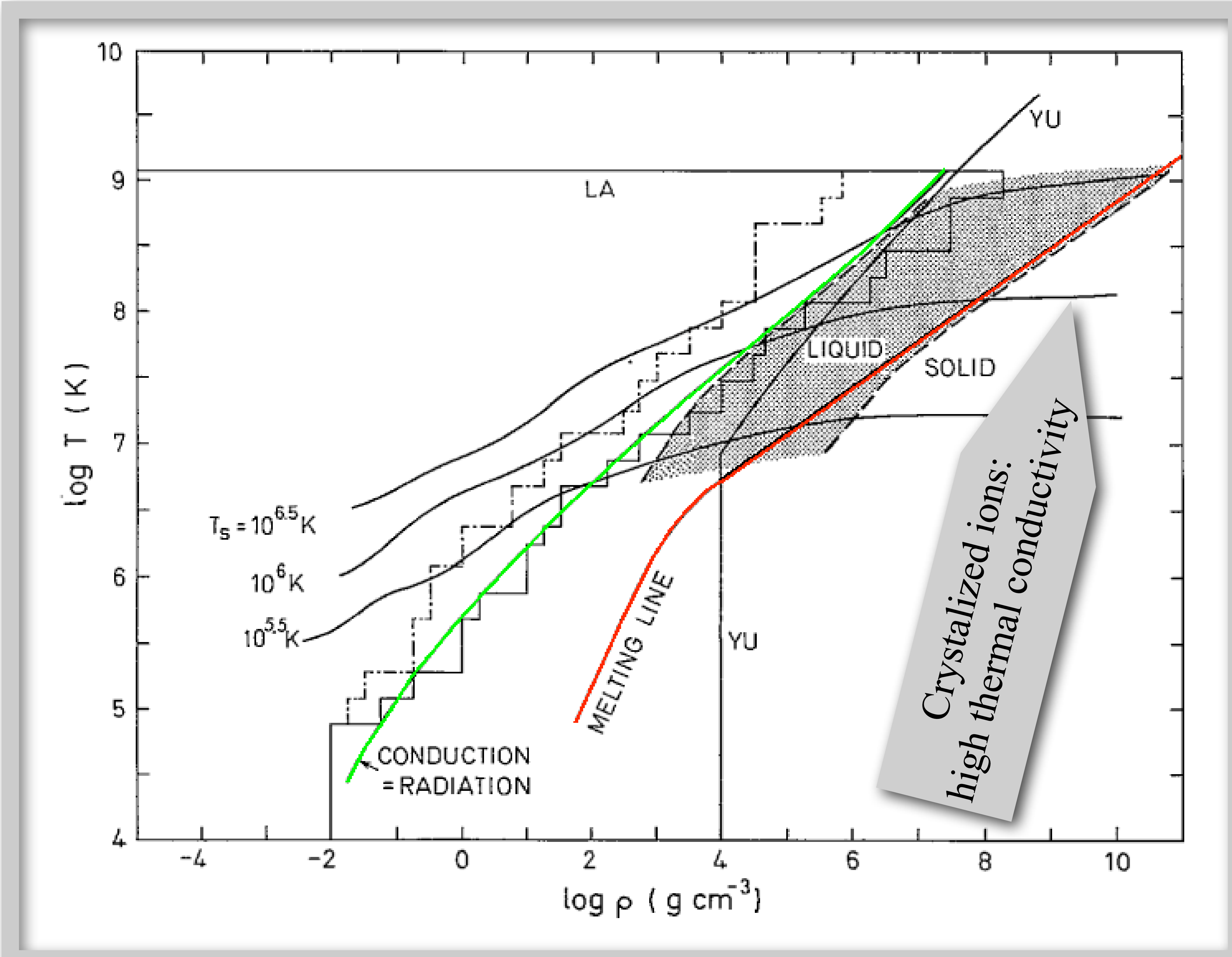

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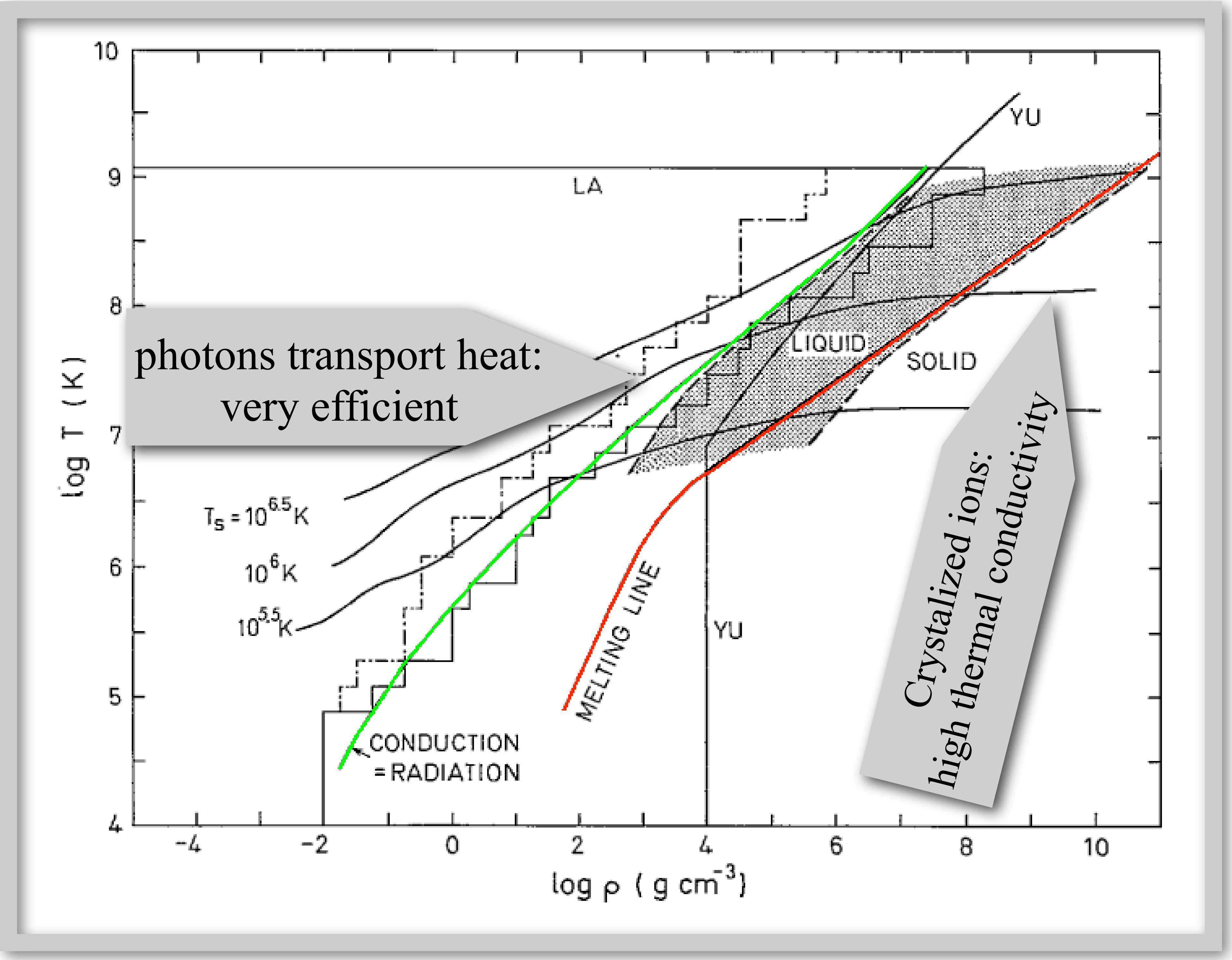

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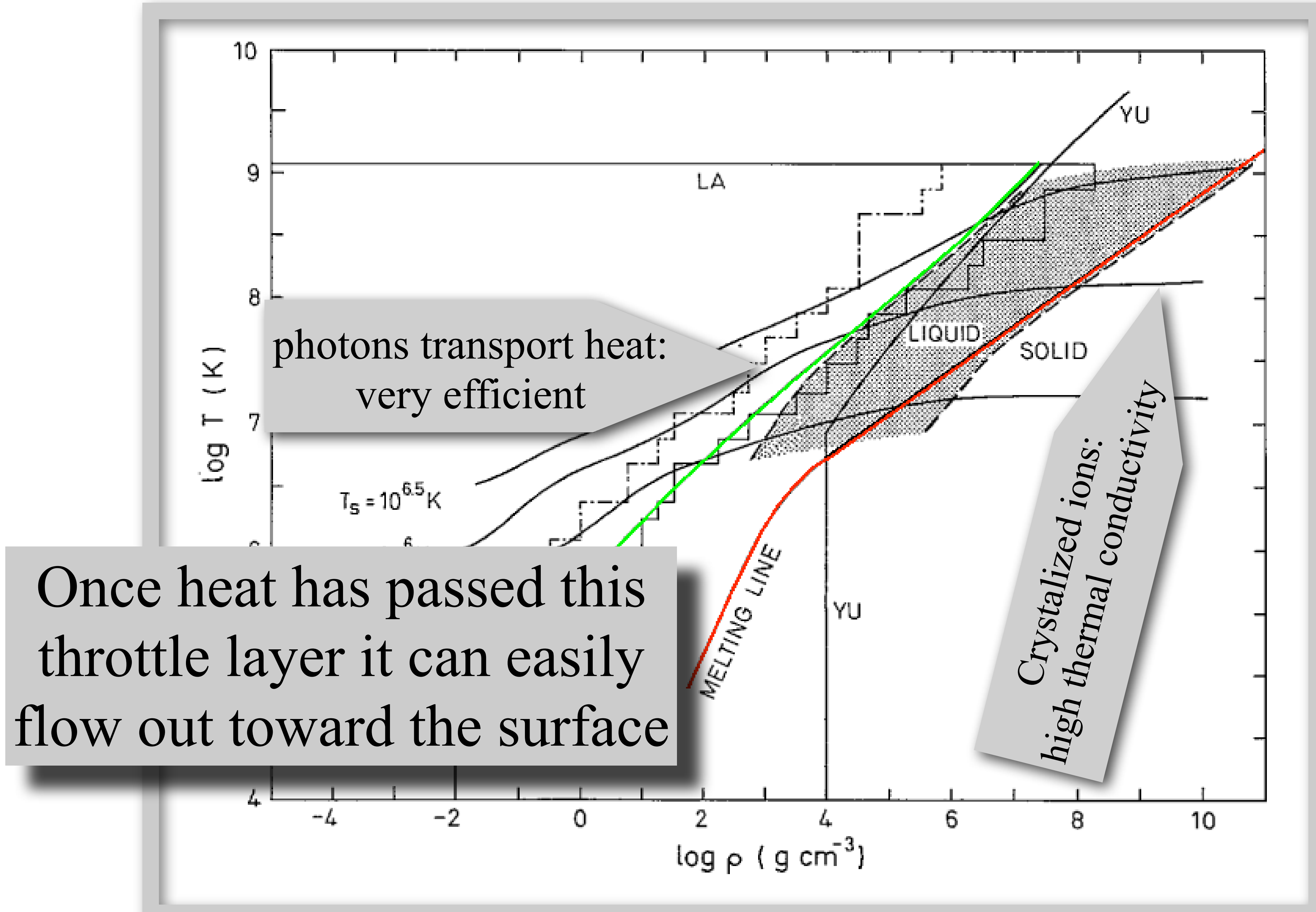
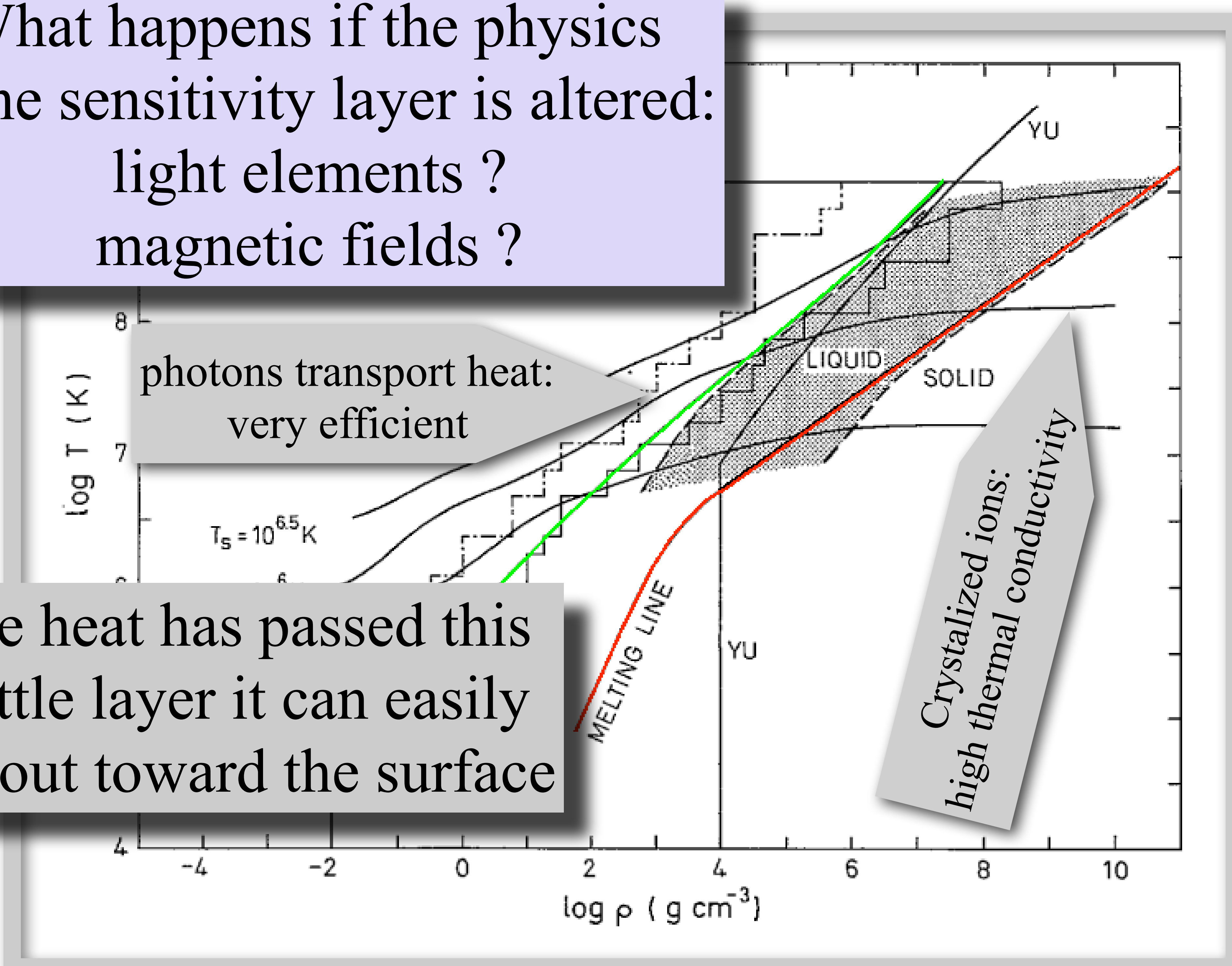


Structure of neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I. 1983ApJ...272..286G

instituto de astronomía
unam

The Sensitivity Strip

What happens if the physics
in the sensitivity layer is altered:
light elements ?
magnetic fields ?



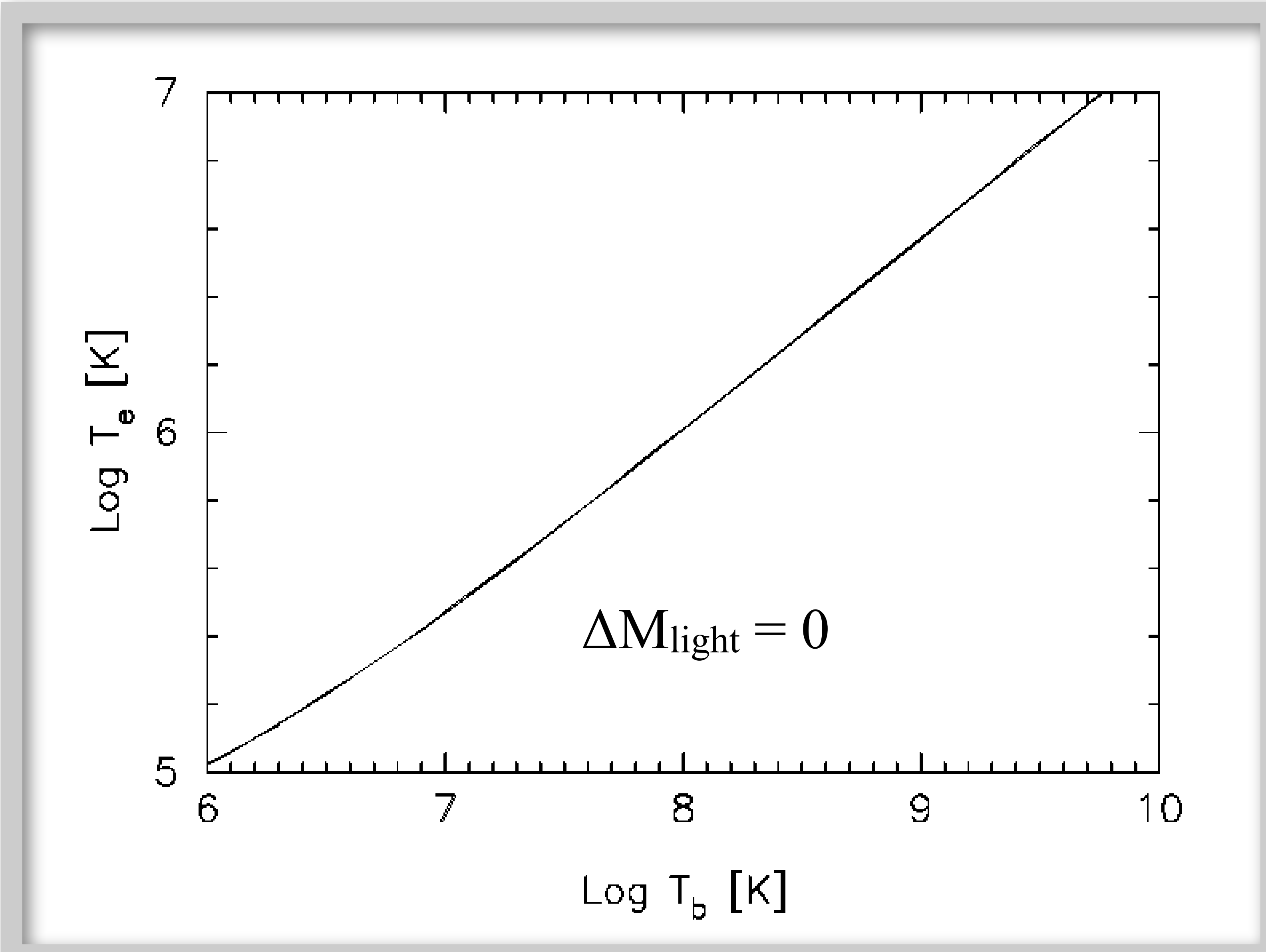
photons transport heat:
very efficient

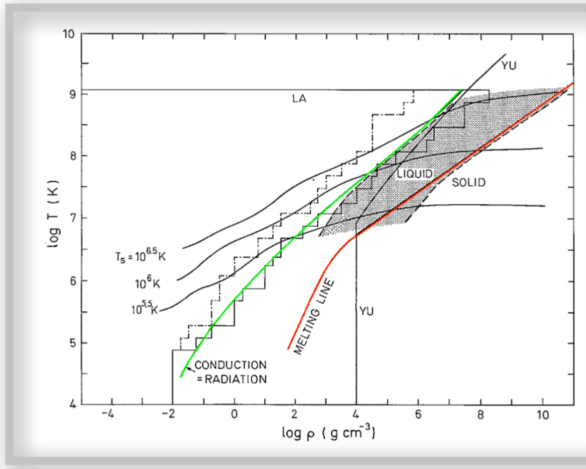
Crystalized ions:
high thermal conductivity

Once heat has passed this
throttle layer it can easily
flow out toward the surface

Structure of neutron star envelopes
Gudmundsson, E. H.; Pethick, C. J.; Epstein, R. I. 1983ApJ...272..286G

Light Element Envelopes



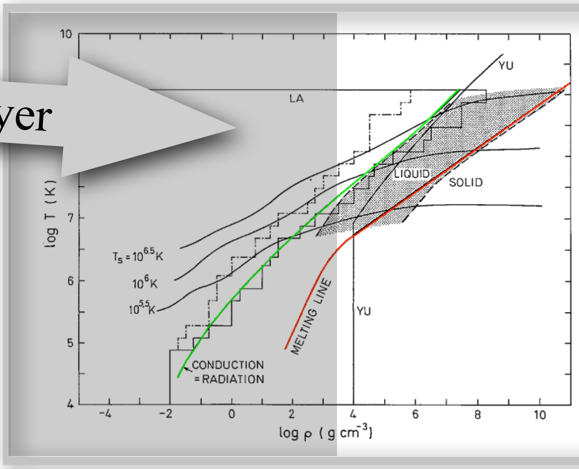


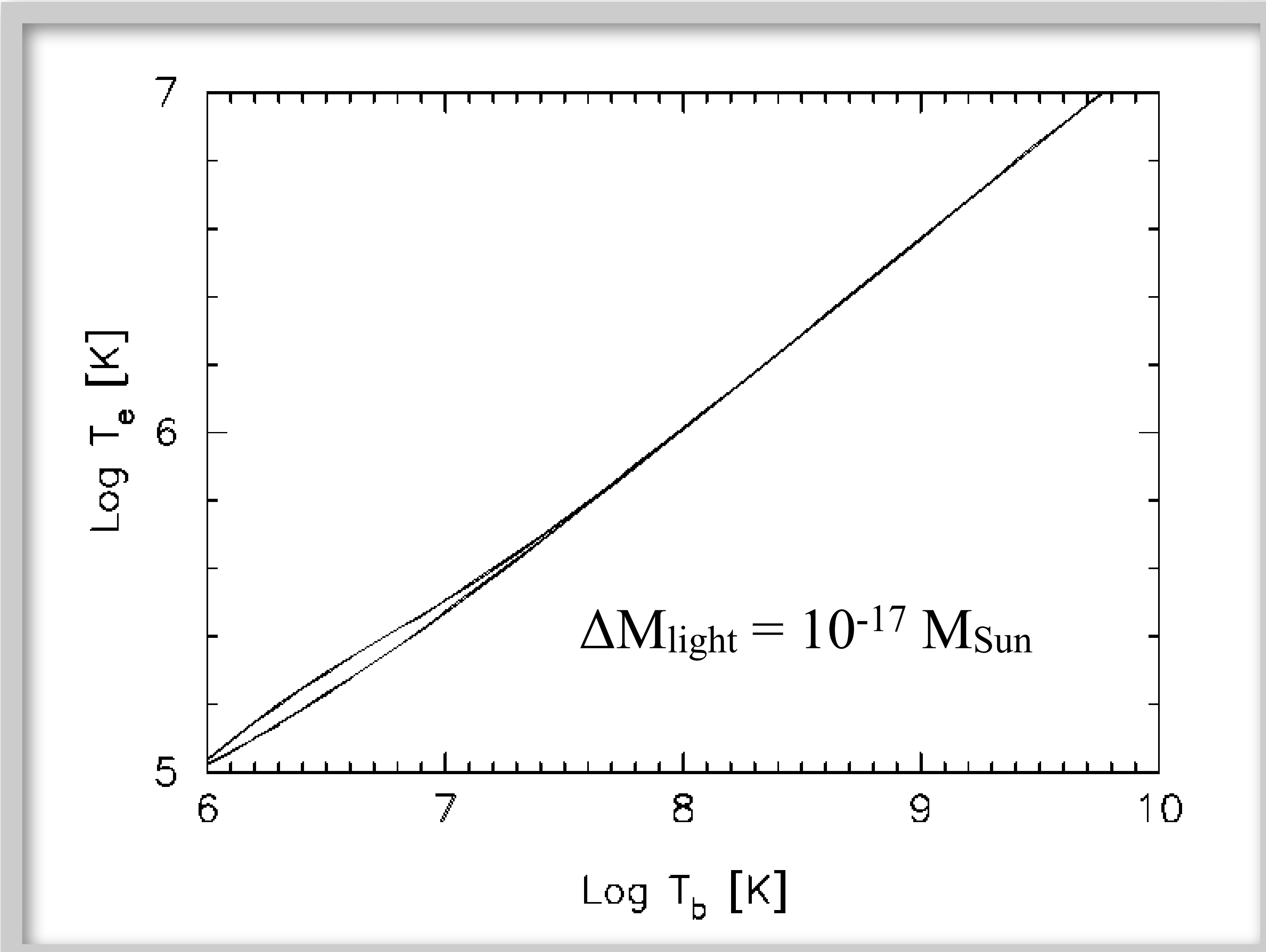
ΔM_{light} = mass of light in the upper envelope

Cooling Neutron Stars with Accreted Envelopes
Chabrier, Gilles; Potekhin, Alexander Y.; Yakovlev, Dmitry G., 1997ApJ...477L..99C

Light Element Envelopes

Thickness of light elements layer





Electron thermal conductivity, due to e-ion scattering in the liquid sensitivity layer:

$$\lambda_{\text{liquid}} \propto \frac{1}{Z}$$

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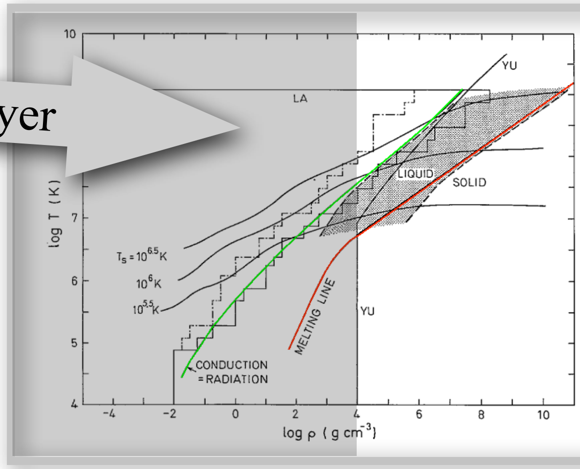

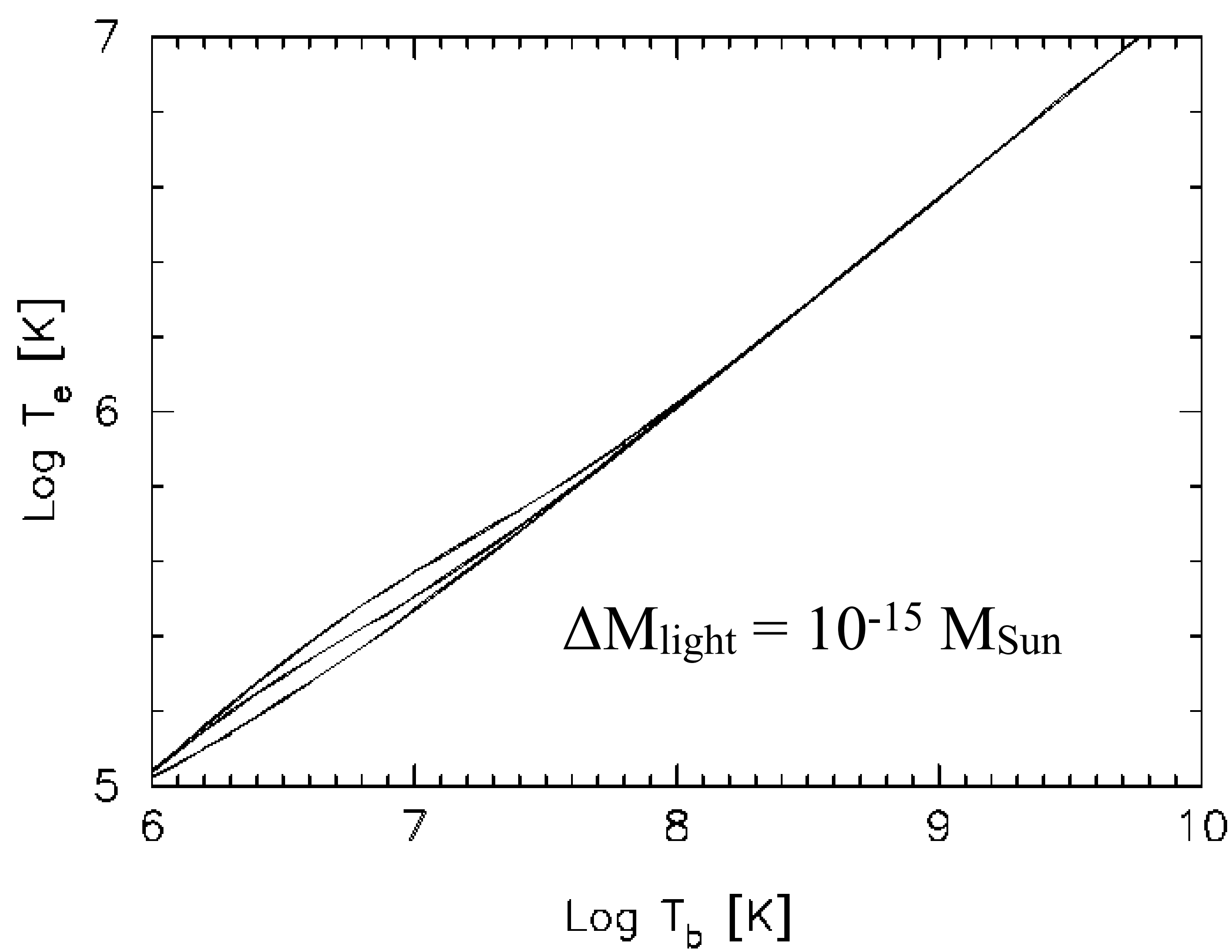

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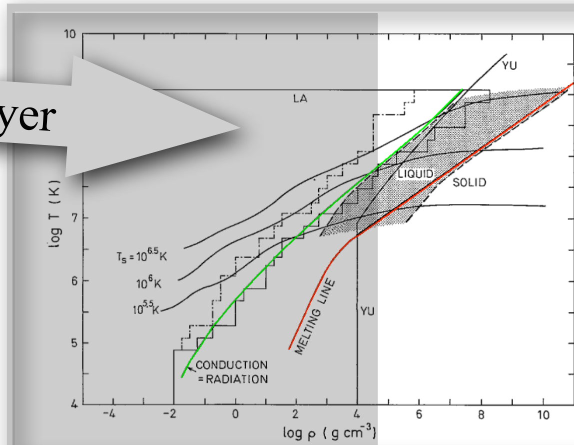

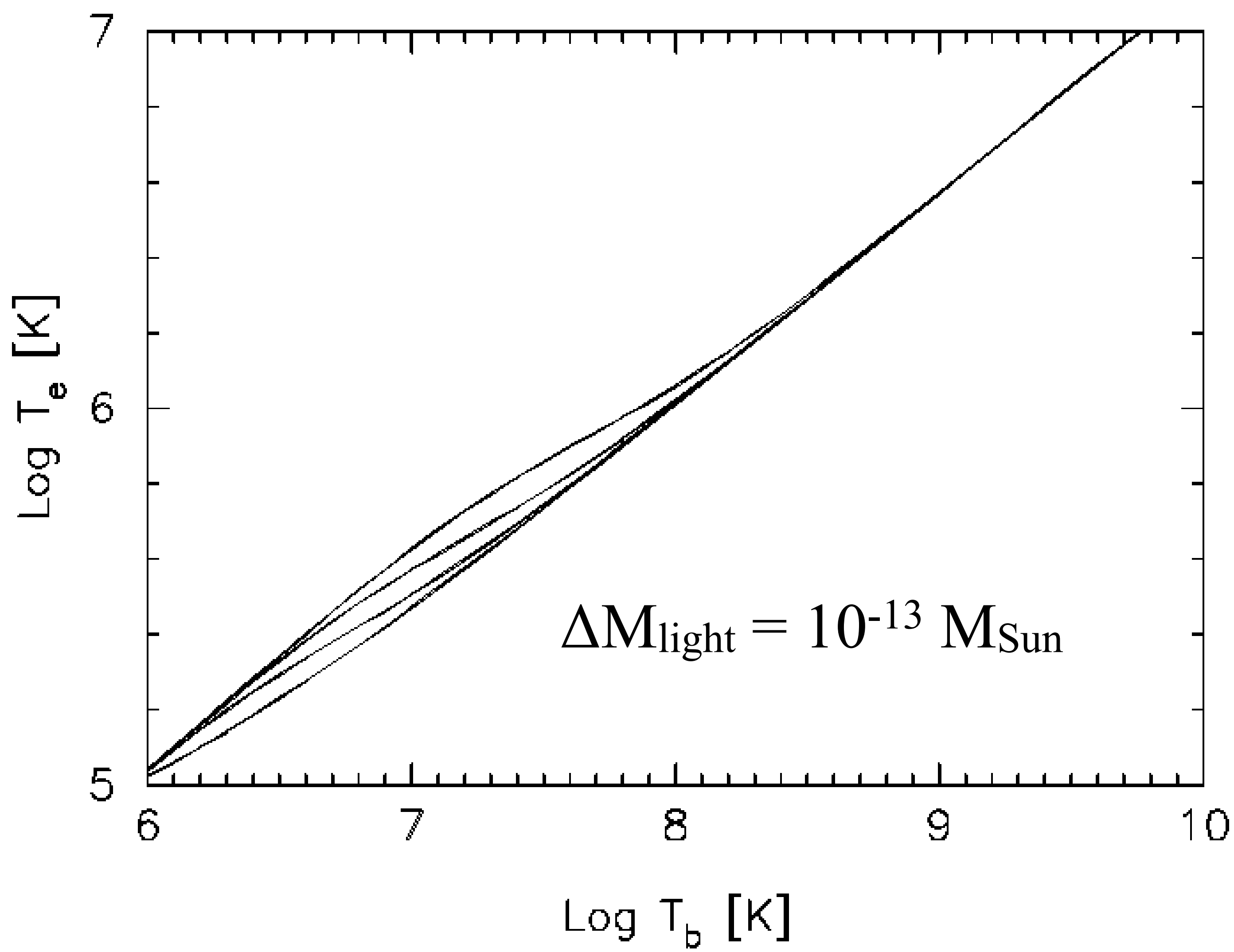

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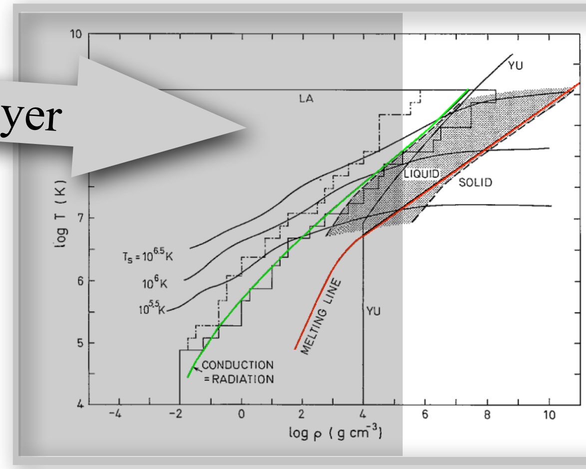

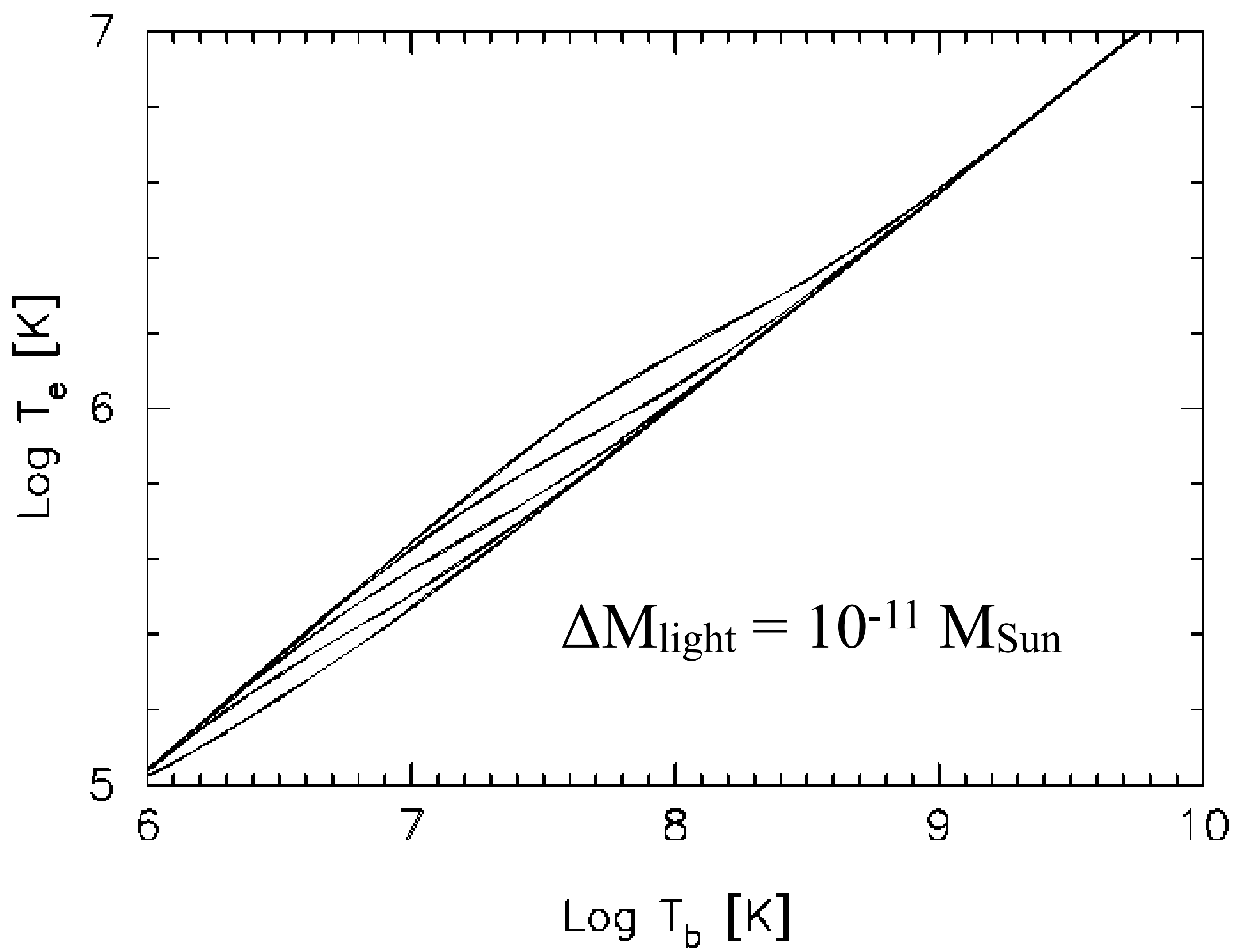

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Cooling Neutron Stars with Accreted Envelopes
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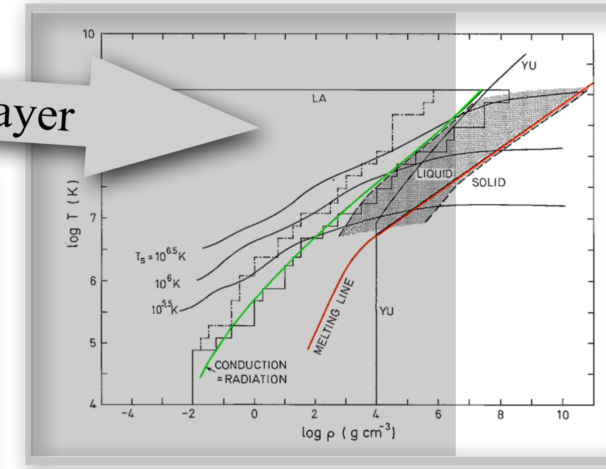

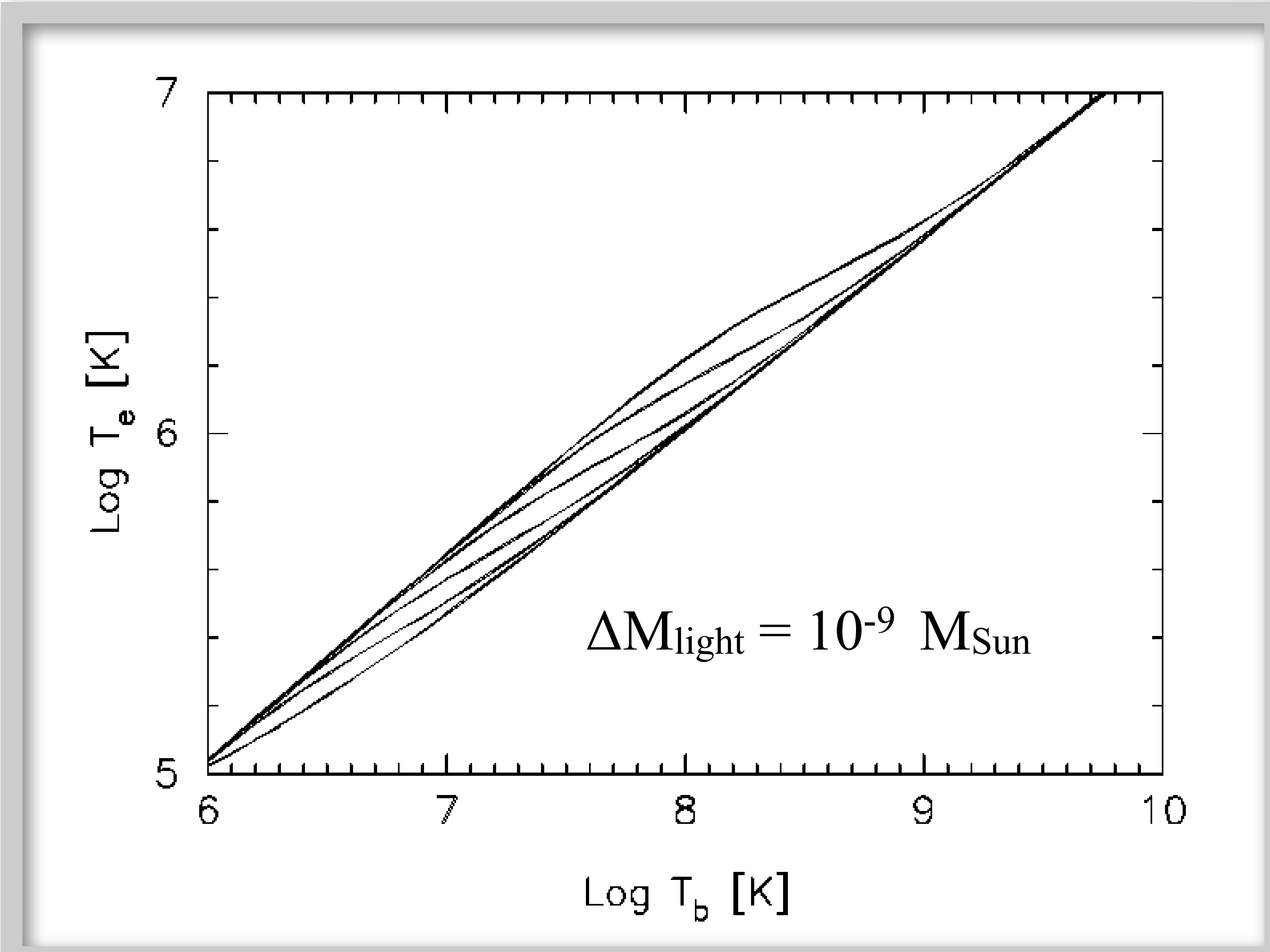

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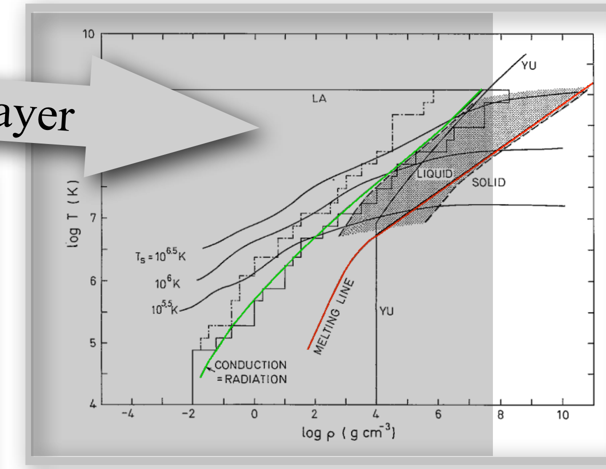

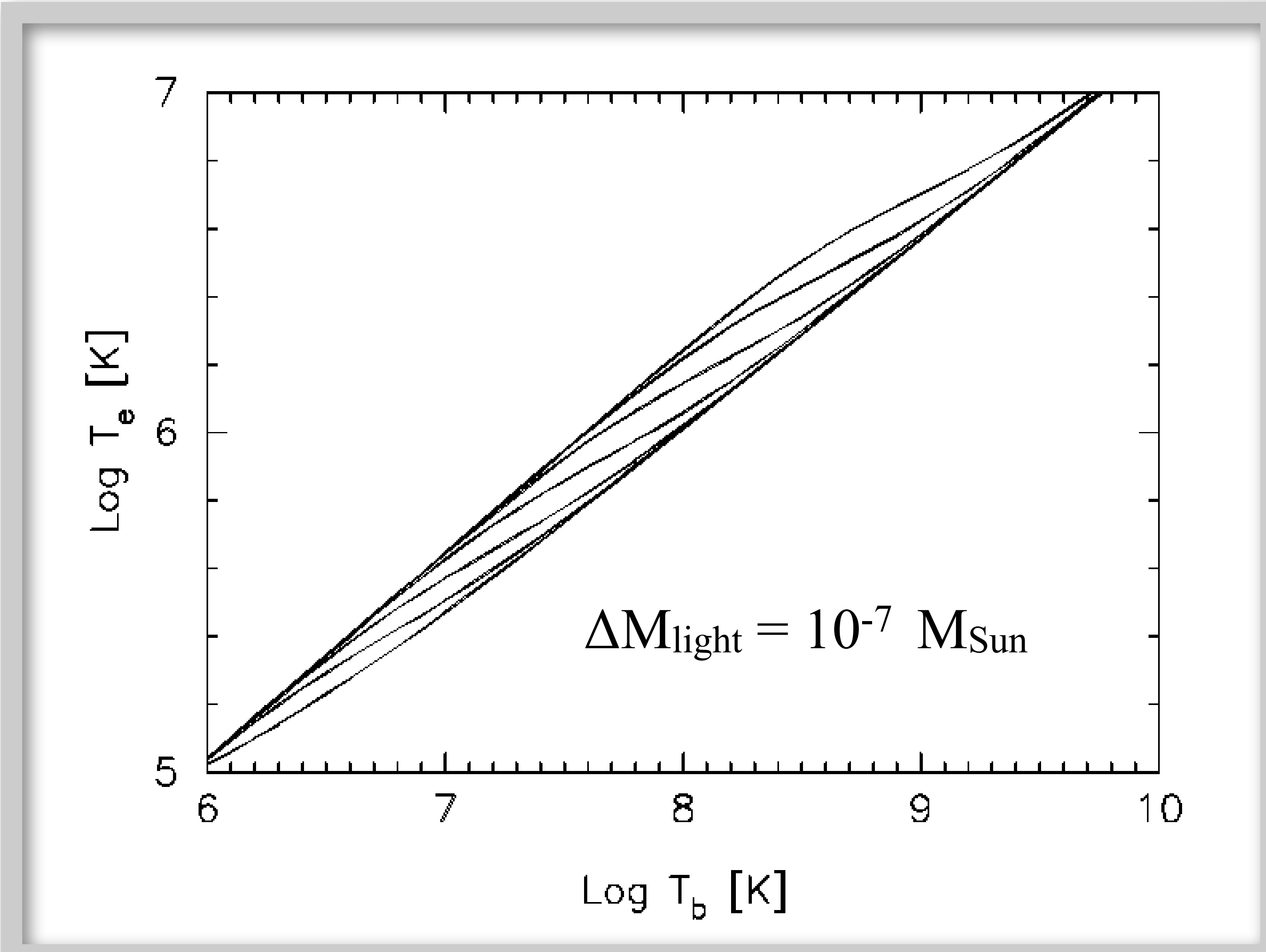

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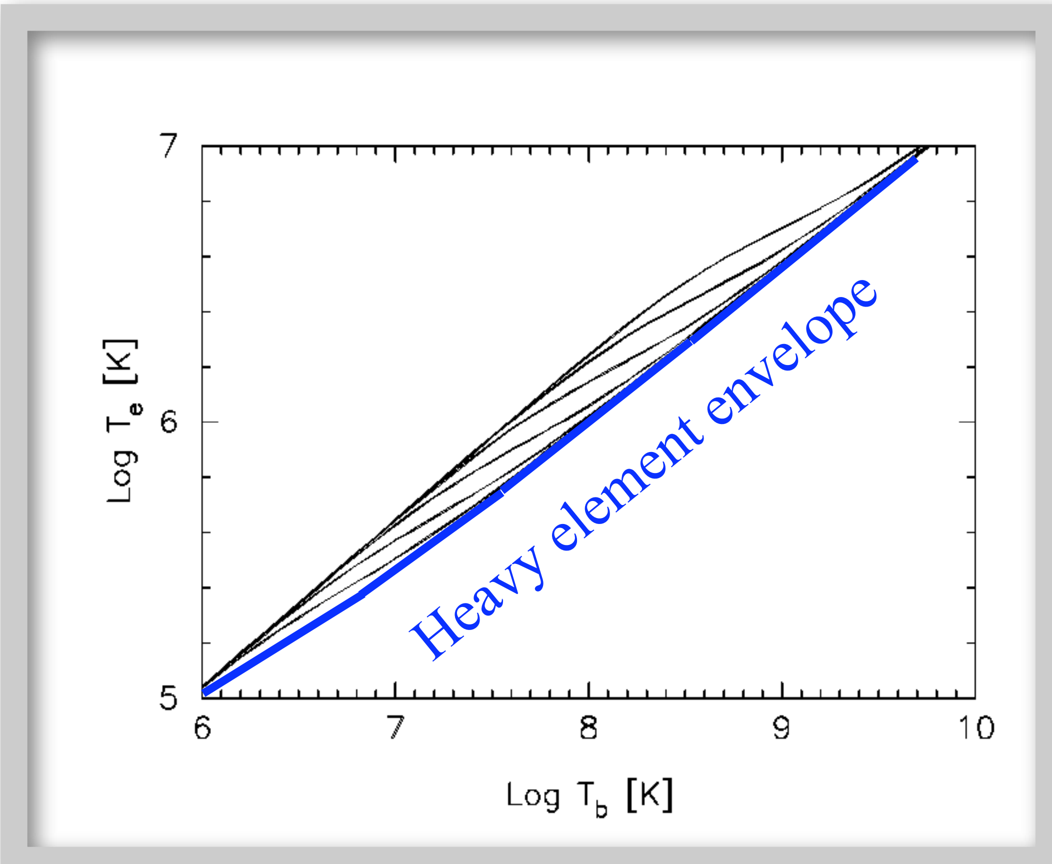
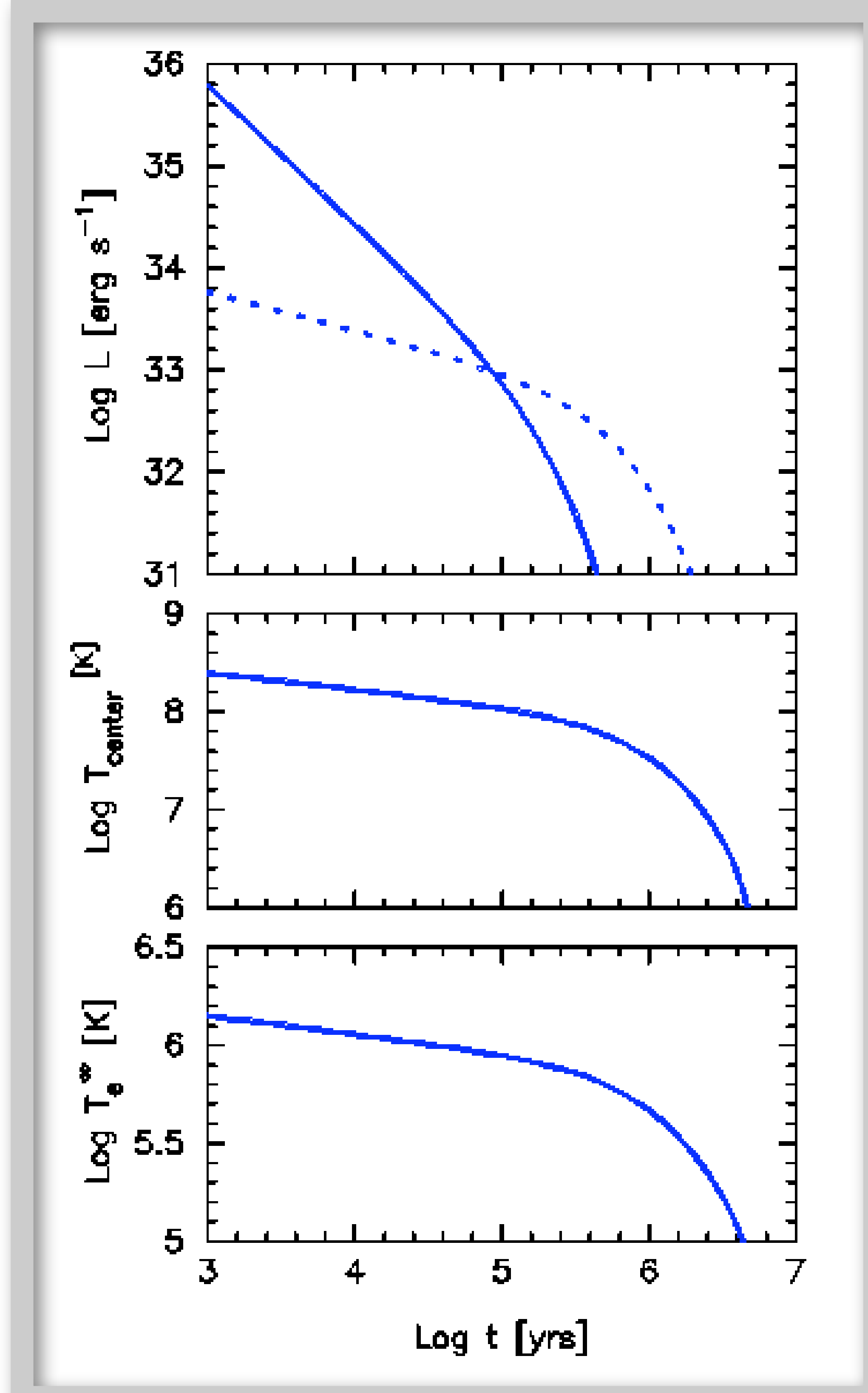
Neutron Star Cooling on a napkin





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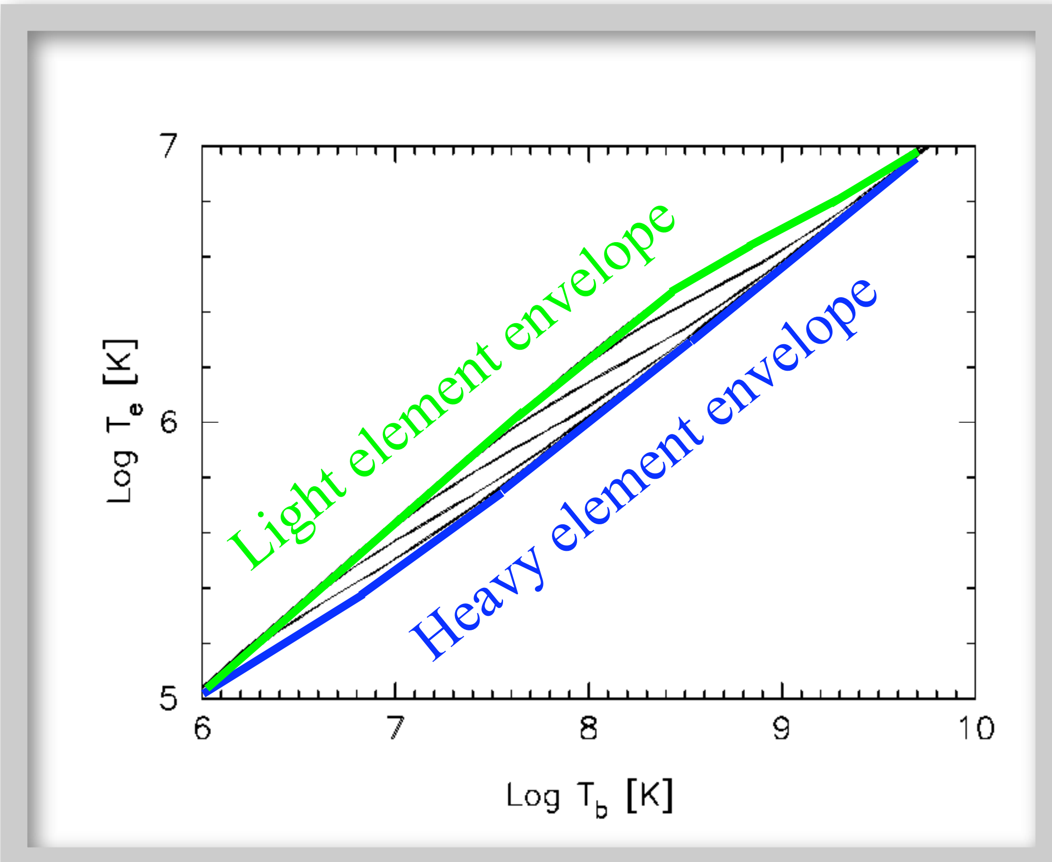

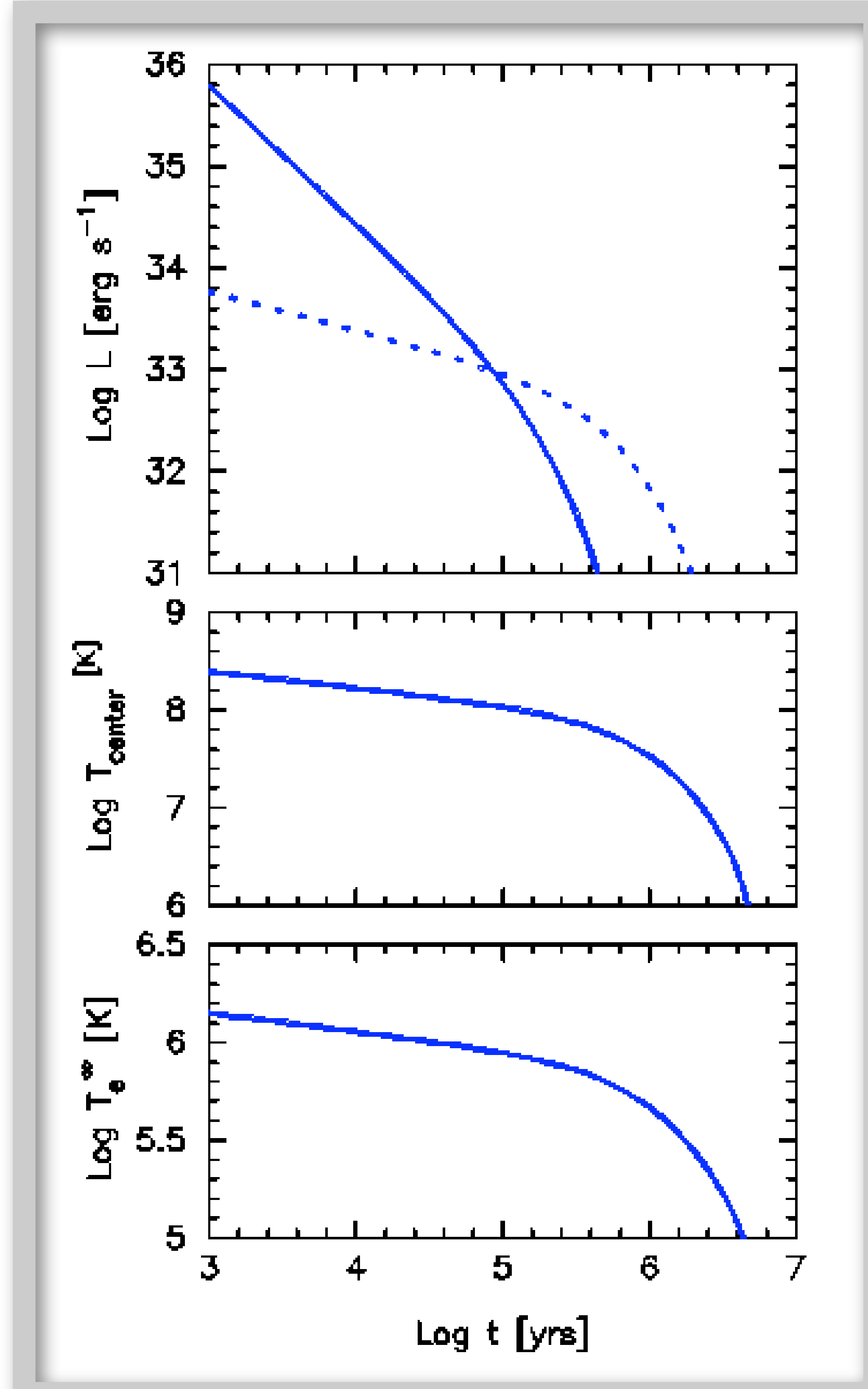

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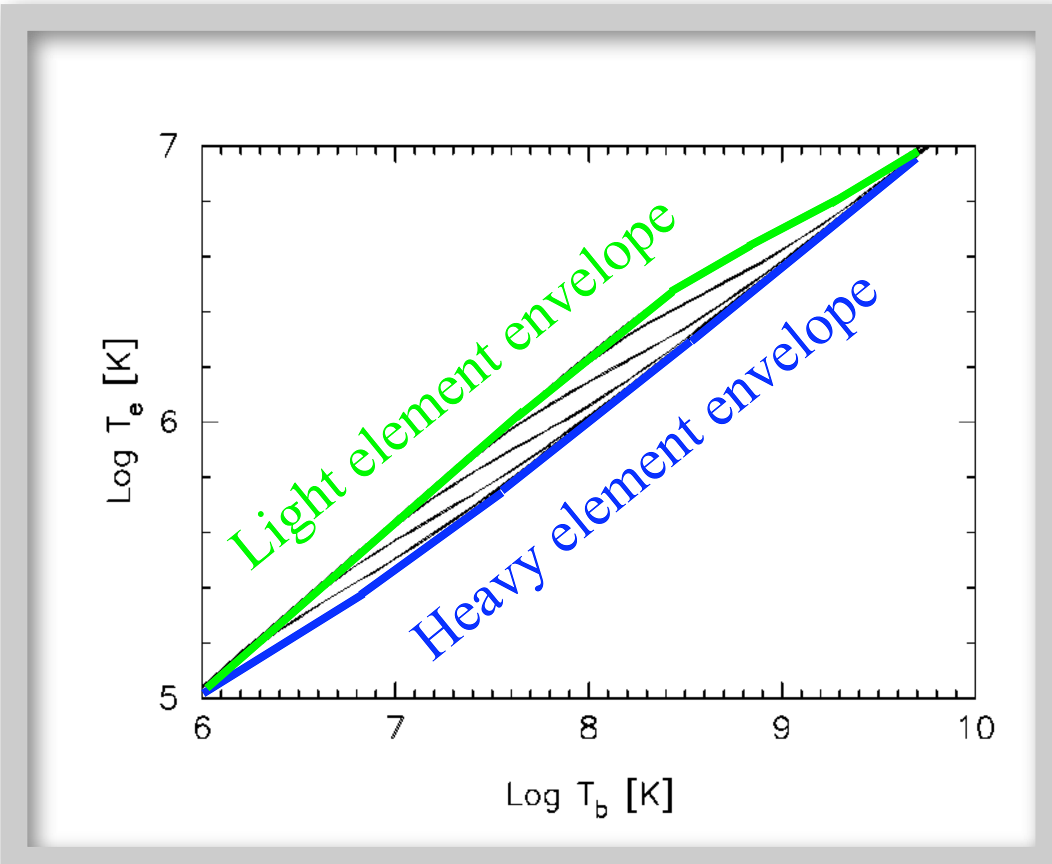

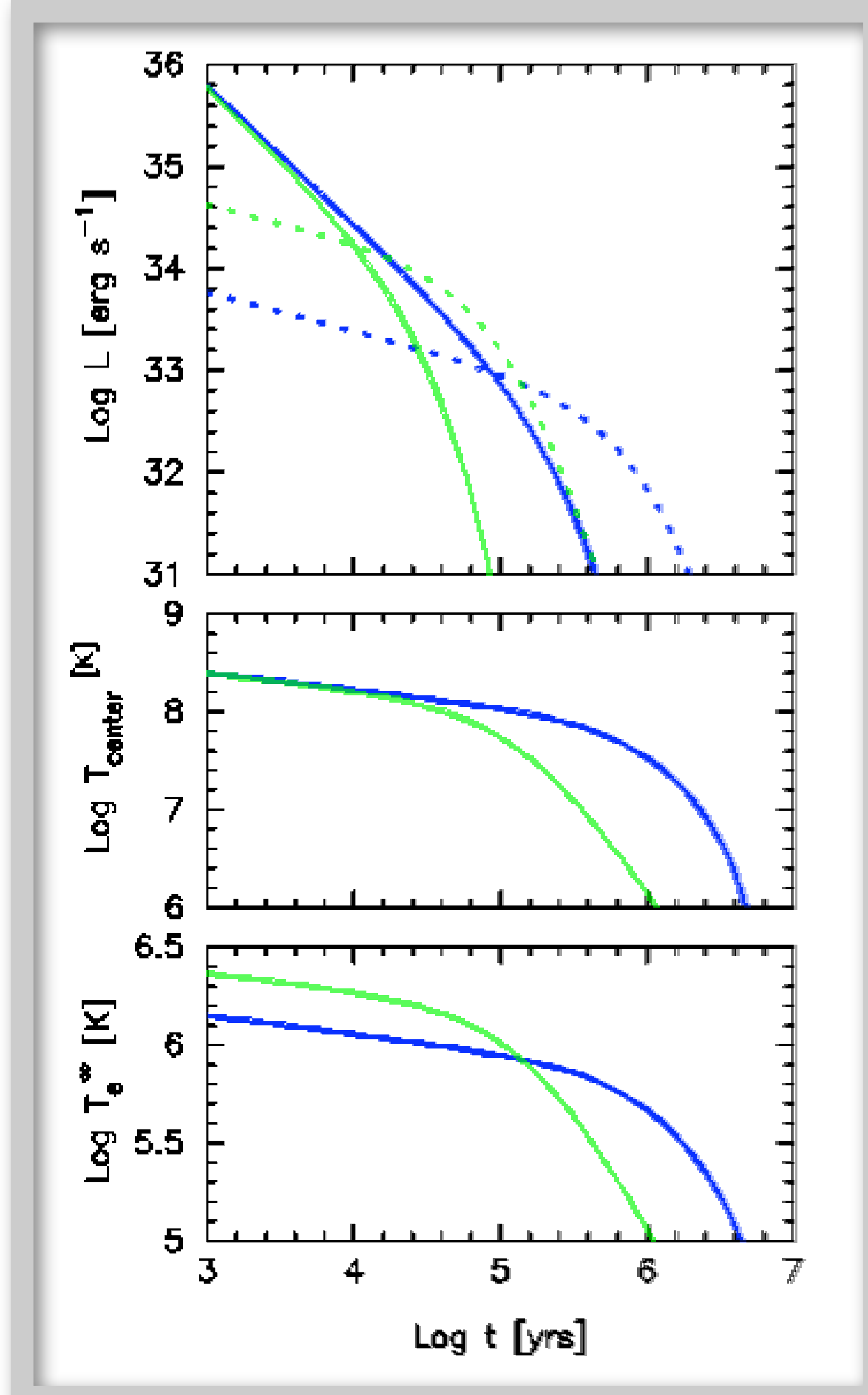

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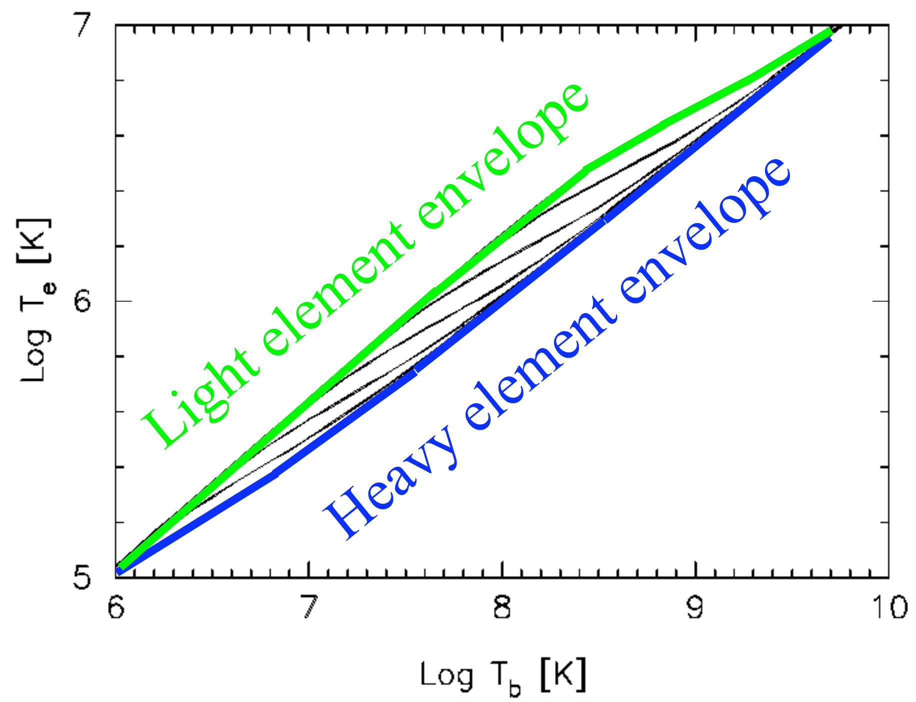

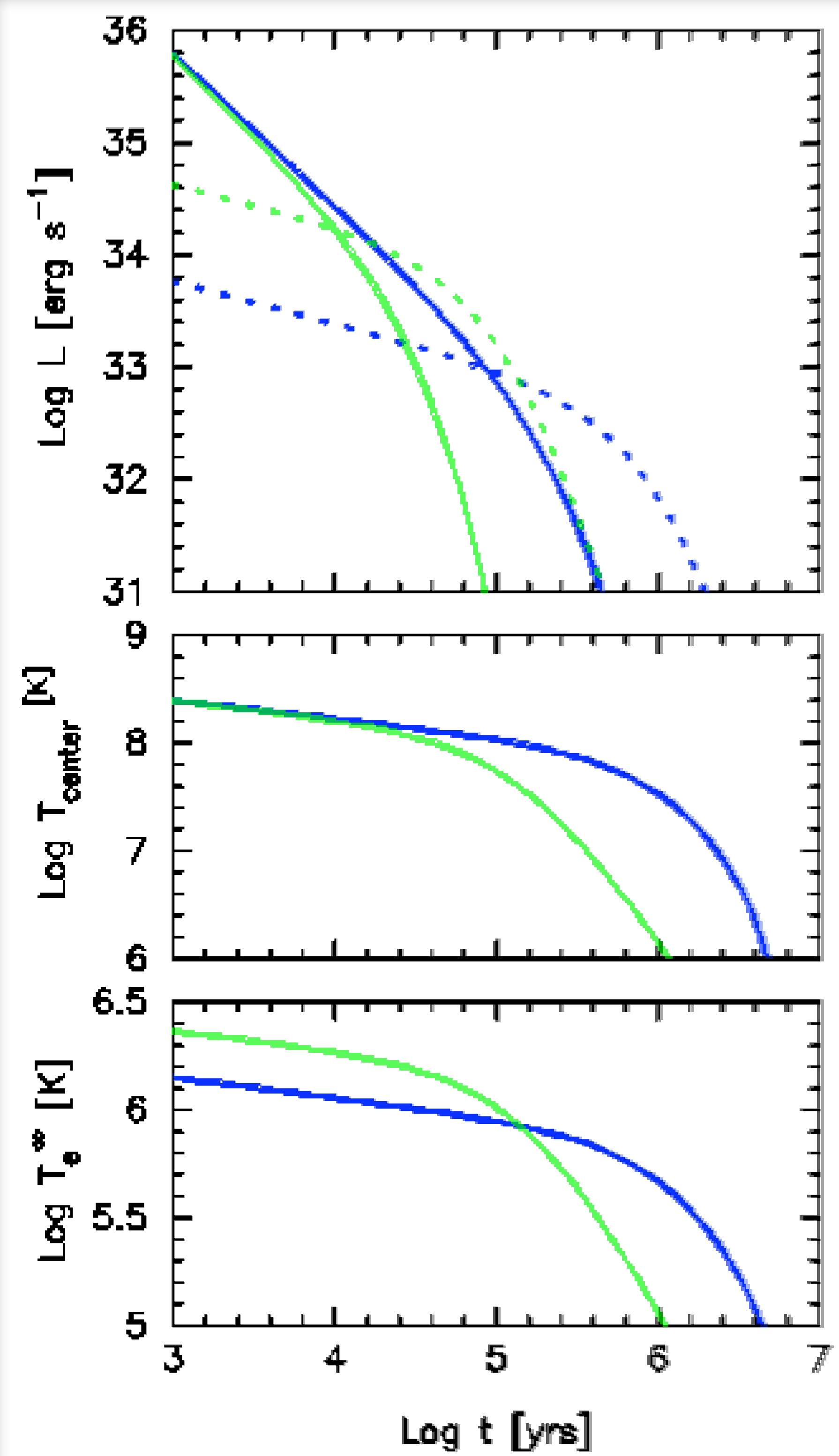

Light element envelopes: star looks warmer during neutrino cooling era and then cools faster during photon cooling era

Heat transport with magnetic field

$$\vec{F} = -\kappa \cdot \vec{\nabla} T$$

$$\kappa_0 = \frac{1}{3} c_{\rm v} \overline{\rm v}^2 \tau \quad = \quad \frac{\pi^2 k_{\rm B}^2 T n_e}{3 m_e^*} \tau$$

τ = electron relaxation time

Temperature distribution in magnetized neutron star crusts
U Geppert, M Kueker & D Page, 2004A&A...426..267G

Heat transport with magnetic field

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In the presence of a strong magnetic field κ becomes a tensor:

$$\kappa = \begin{pmatrix} \kappa_\perp & \kappa_\wedge & 0 \\ -\kappa_\wedge & \kappa_\perp & 0 \\ 0 & 0 & \kappa_{||} \end{pmatrix}$$

$$\kappa_{||} = \kappa_0$$

$$\kappa_\perp = \frac{\kappa_0}{1 + (\omega_B \tau)^2}$$

$$\kappa_\wedge = \frac{\kappa_0 \; \omega_B \tau}{1 + (\omega_B \tau)^2}$$

$$\omega_B = \frac{eB}{m_e^* c}$$ = electron cyclotron frequency

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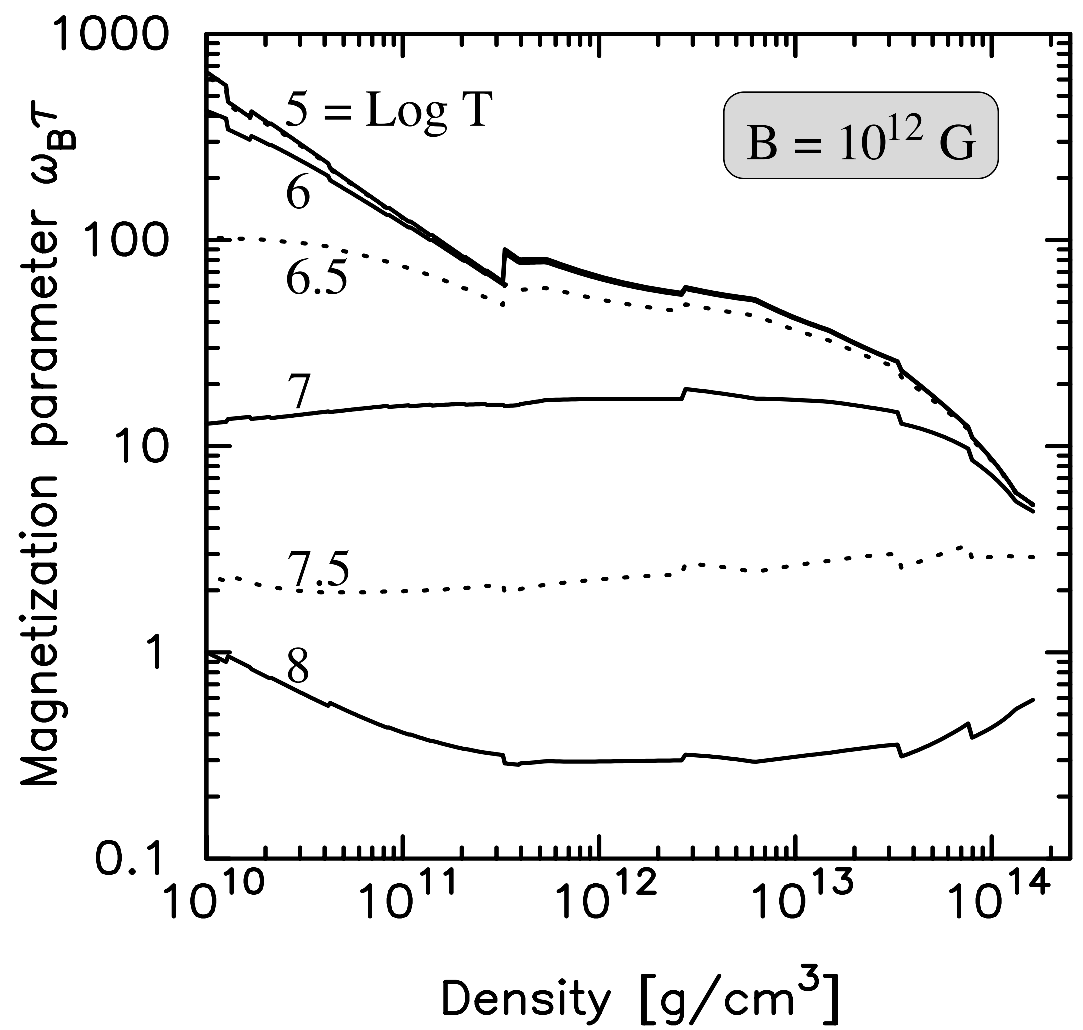


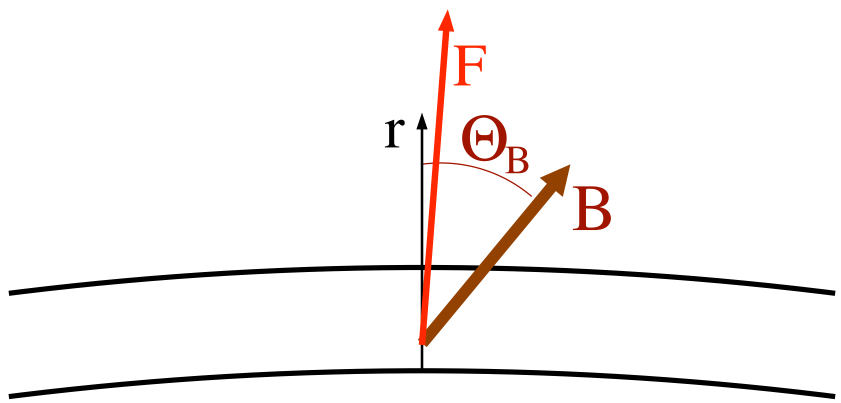
Temperature distribution in magnetized neutron star crusts
U Geppert, M Kueker & D Page, 2004A&A...426..267G

Magnetized Envelopes

Thinness of envelope $\Rightarrow F$ essentially radial: only need to calculate F_r.

Thermal conductivity in the radial direction:

$$\kappa(\Theta_B) = \cos^2\Theta_B \times \kappa_{||} + \sin^2\Theta_B \times \kappa_{\perp}$$



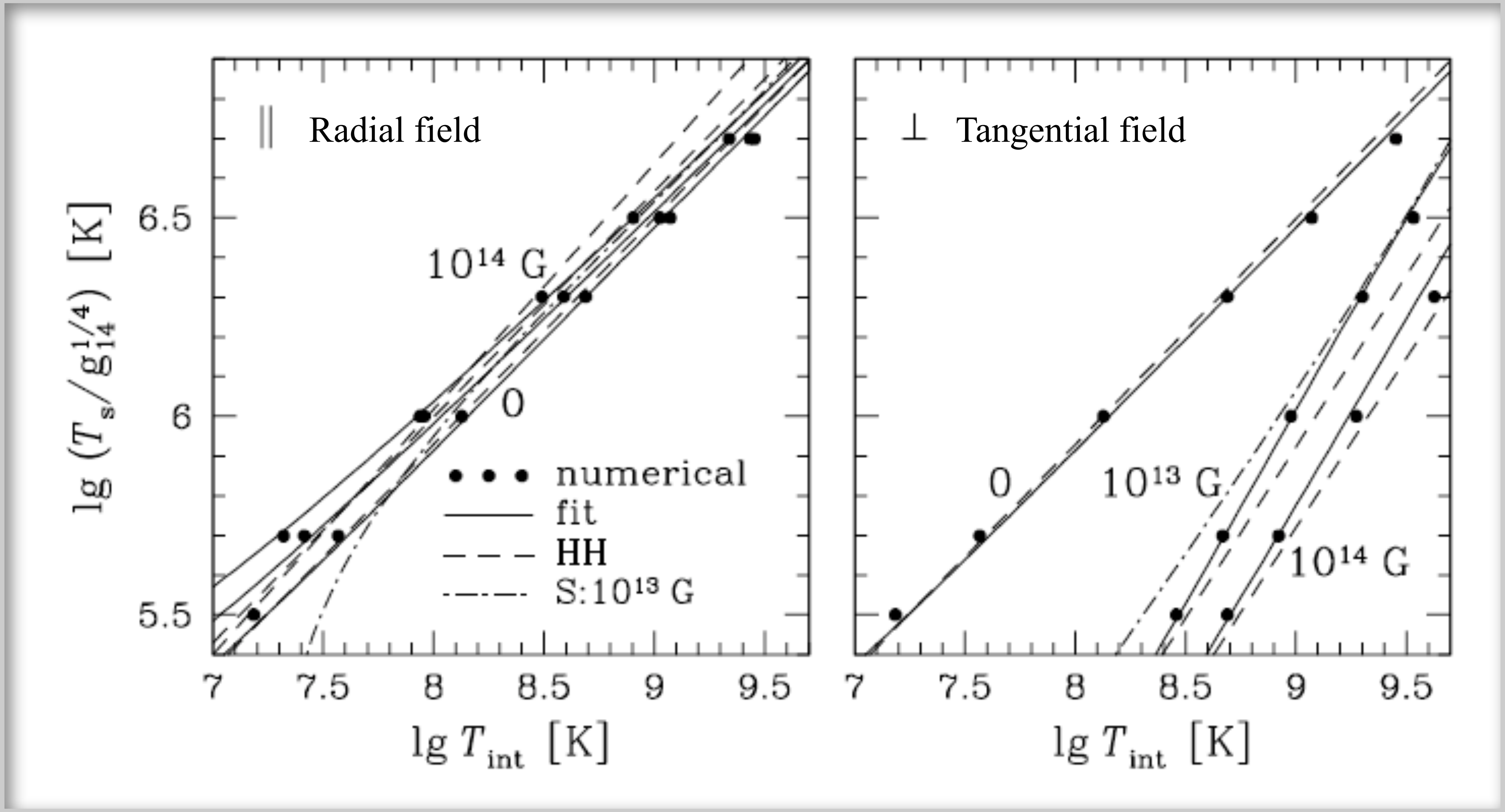
Greenstein & Hartke (1983) proposed to approximate $T_s(\Theta_B)$ by:

$$T_s^4(\Theta_B) = T_{s\,||}^4 \cos^2\Theta_B + T_{s\,\perp}^4 \sin^2\Theta_B$$

where $T_{s\,||} = T_s(\Theta_B = 0^o)$ and $T_{s\,\perp} = T_s(\Theta_B = 90^o)$:
only need to calculate two cases, radial and tangential field, and use the formula to interpolate.

T_s here means the local effective temperature such that $F = \sigma\, T_s^4$ at each point on the surface

Modern calculations of $T_{||}$ & $T_{\perp}$



Thermal structure and cooling of neutron stars with magnetized envelopes
AY Potekhin & DG Yakovlev, A&A 374, 213 (2001)

Surface temperature distributions

With the Greenstein-Hartke interpolation formula one can take any field geometry at the surface (envelope) and calculate the surface temperature distribution:

Surface temperature of a magnetized neutron star and interpretation of the ROSAT data. I. Dipolar fields
D Page, ApJ 442, 273 (1995)

Surface temperature of a magnetized neutron star and interpretation of the ROSAT data. II.
D Page & A Sarmiento, ApJ 473, 1067 (1996)

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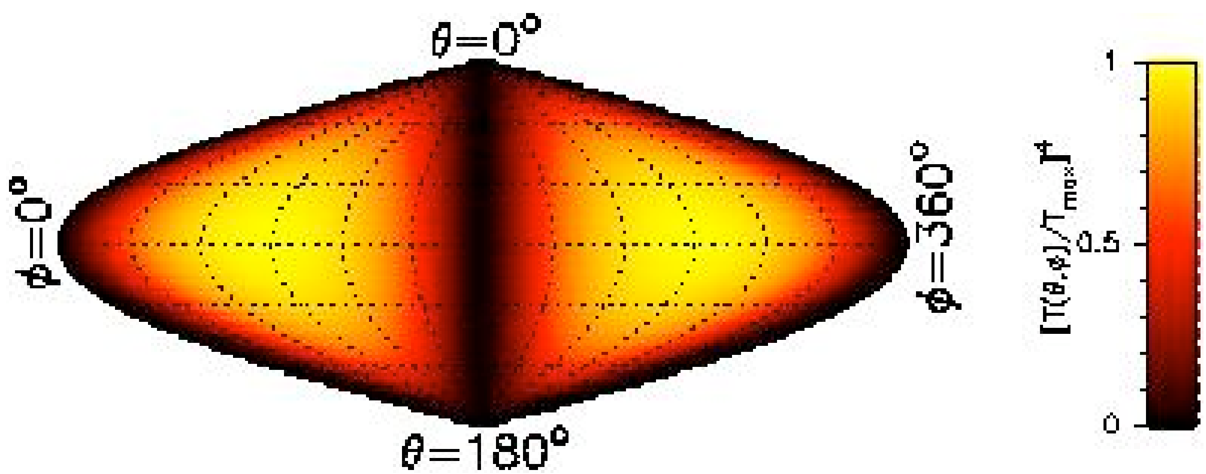

Purely dipolar field

(oriented on the equatorial plane to make a prettier picture !)

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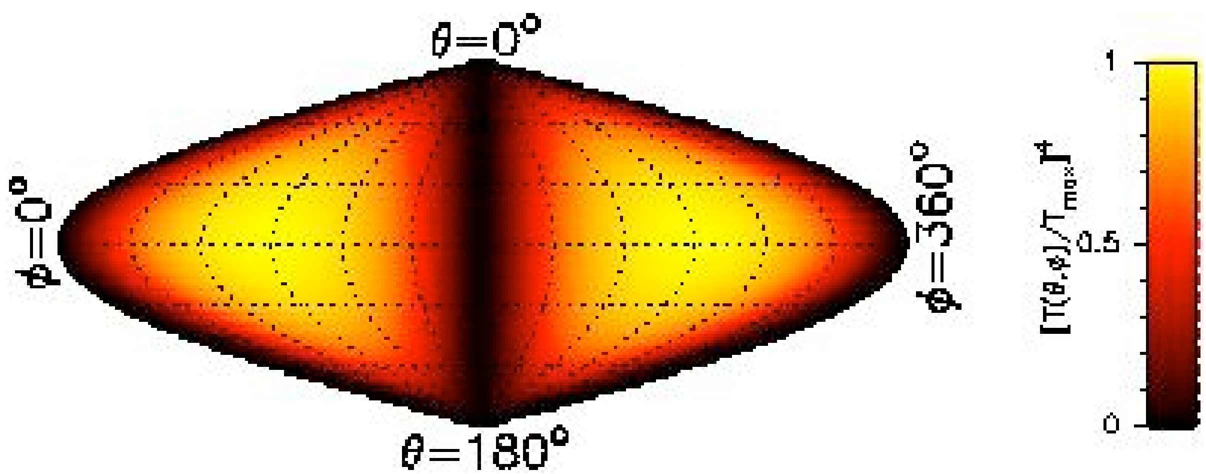

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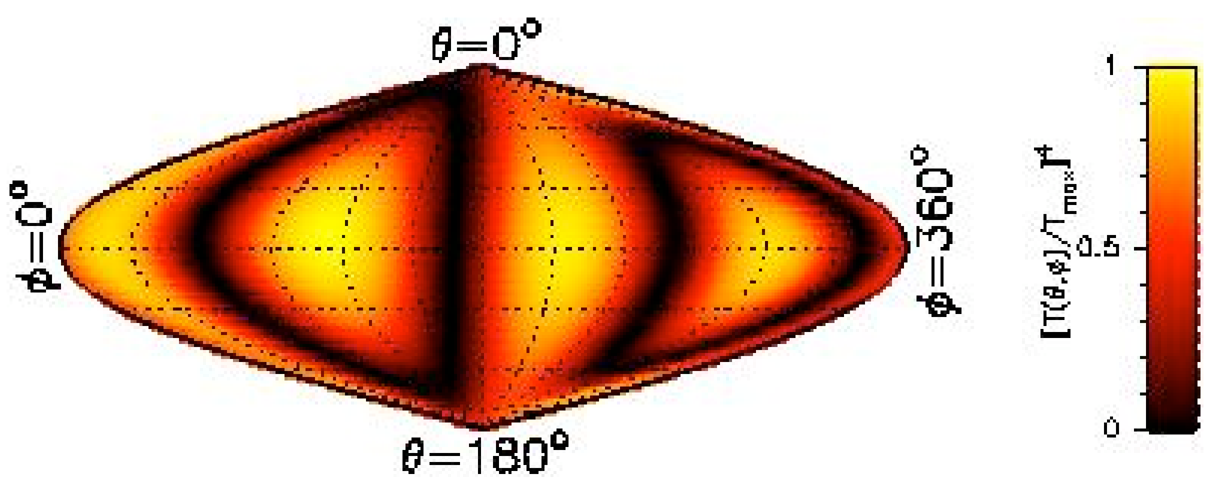

Dipolar + quadrupolar field

Surface temperature of a magnetized neutron star and interpretation of the ROSAT data. I. Dipolar fields
D Page, ApJ 442, 273 (1995)

Surface temperature of a magnetized neutron star and interpretation of the ROSAT data. II.
D Page & A Sarmiento, ApJ 473, 1067 (1996)

Magnetized T_b - T_e relationships

The star's effective temperature is then easily calculated:

$$L = \iint \sigma_B T_s(\theta,\phi)^4 dS \;=\; 4\pi R^2 \sigma T_e^4$$

$$(dS \;=\; R^2 \cdot d\Omega)$$

$$\boxed{T_e^4 = \frac{1}{4\pi}\iint T_s(\theta,\phi)^4 \, d\Omega}$$

This directly generates a T_b - T_e relationship for any surface magnetic field geometry

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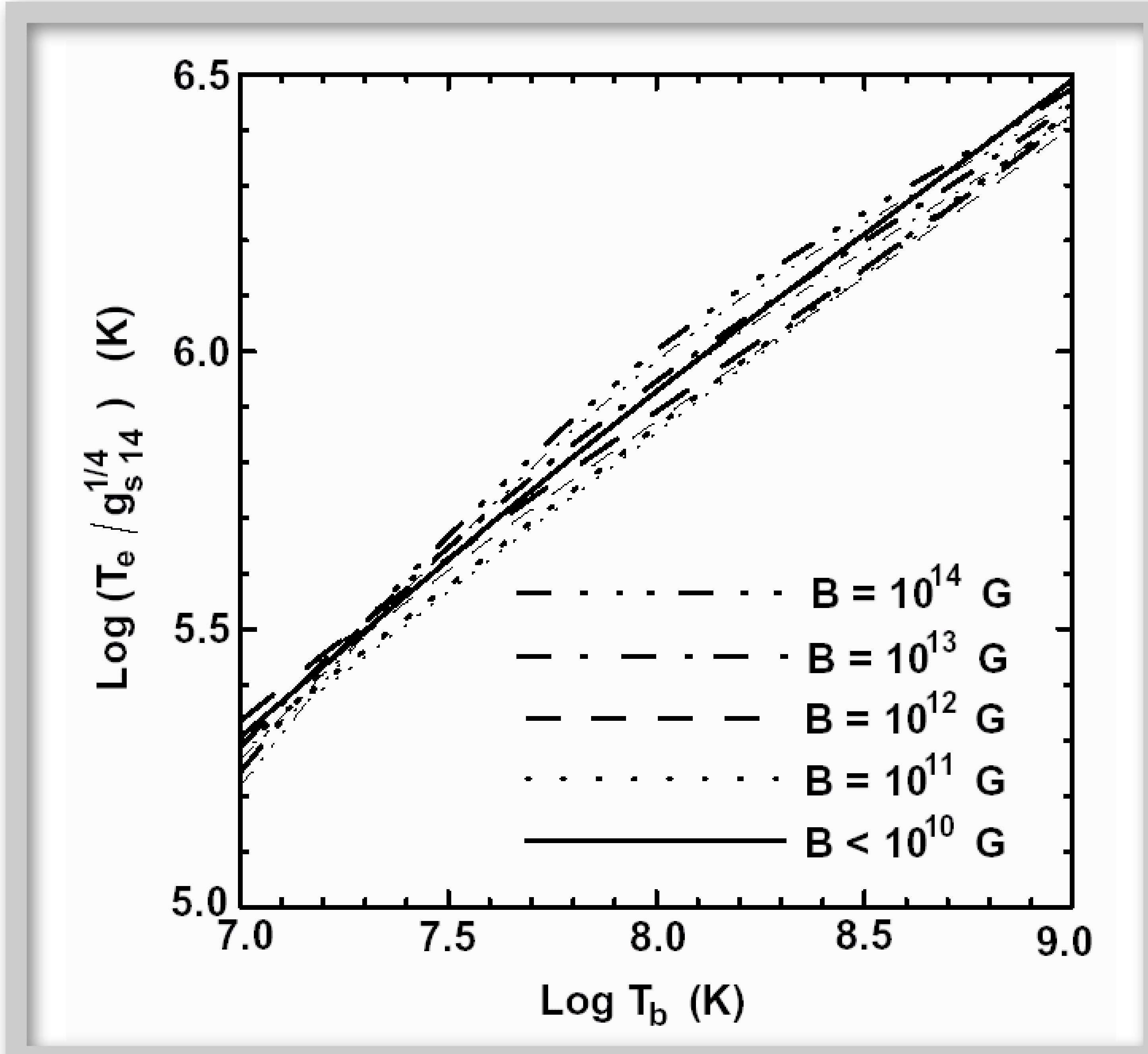
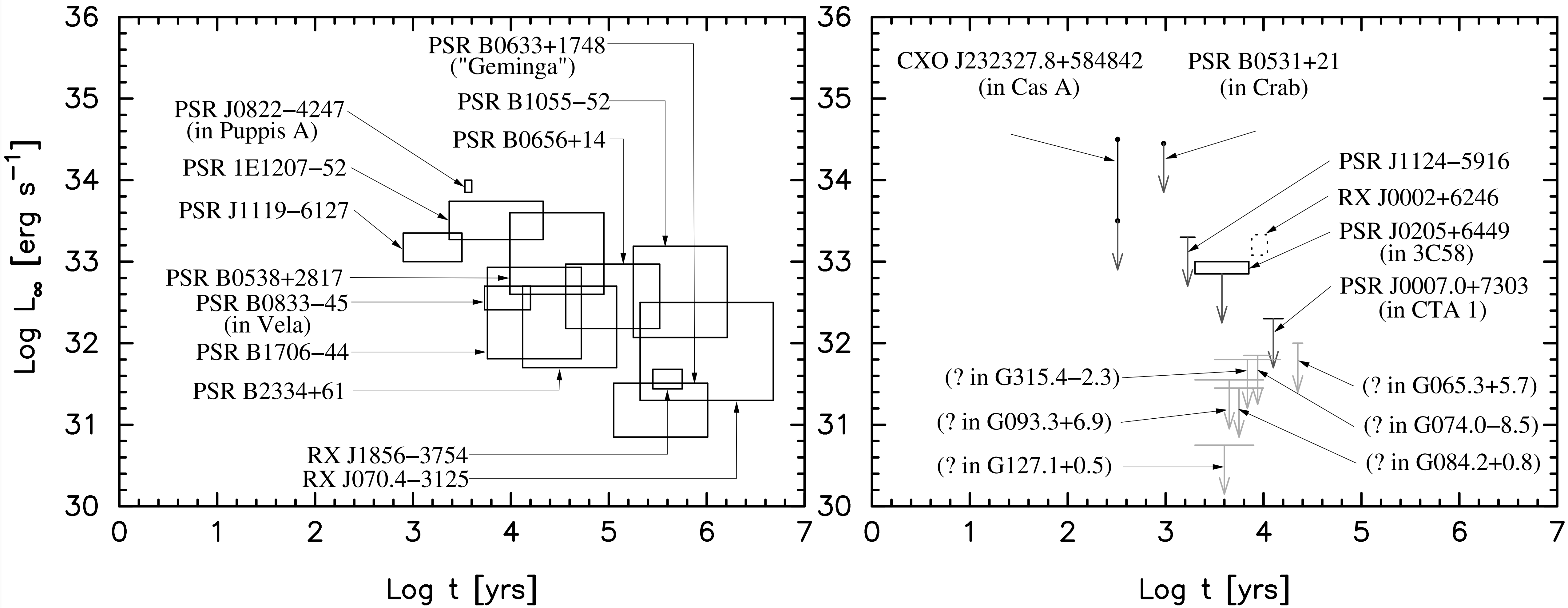


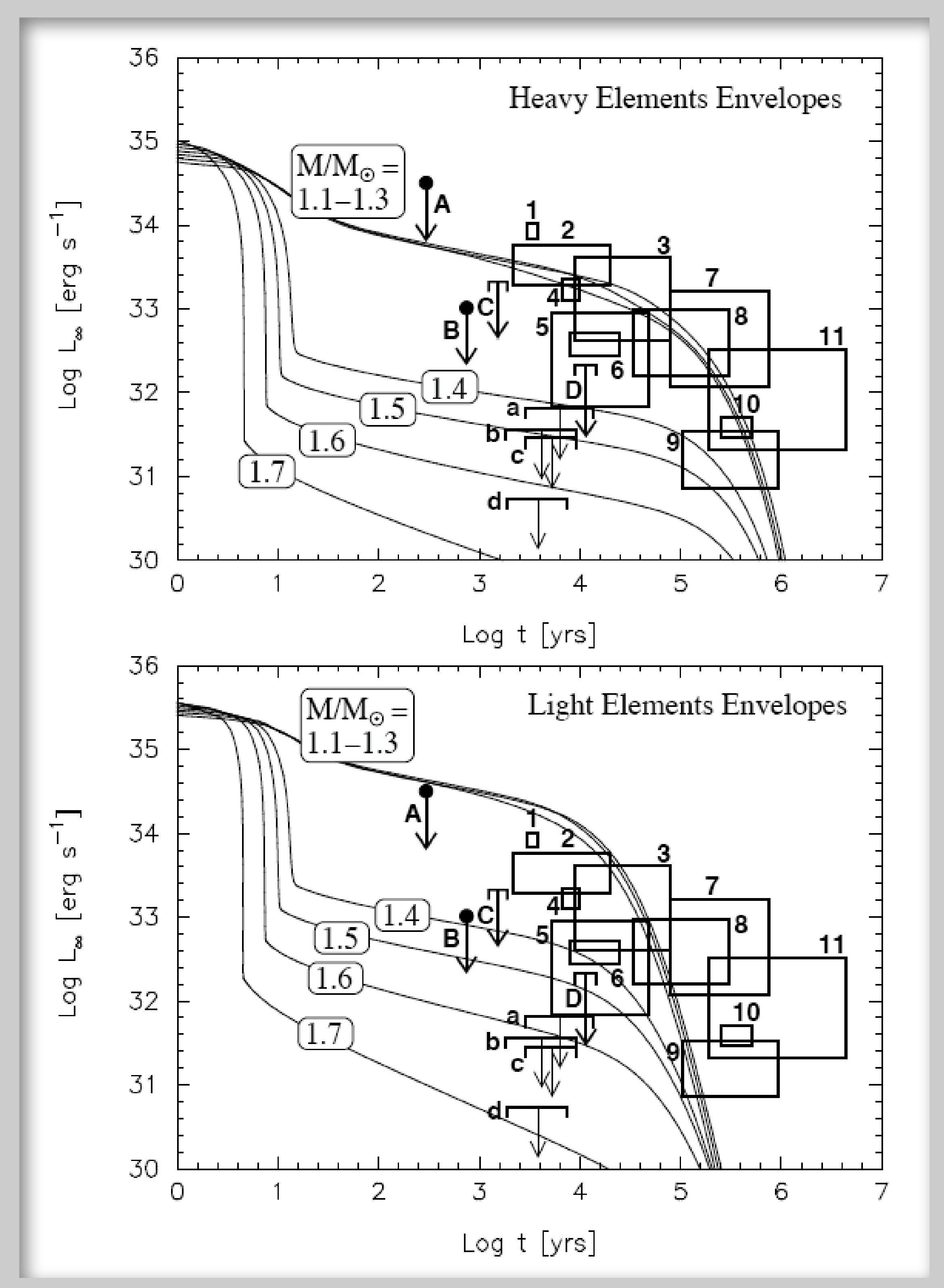
Surface temperature of a magnetized neutron star and interpretation of the ROSAT data. II. D Page & A Sarmiento, ApJ 473, 1067 (1996)

Comparison with Data

Observational data



Direct Urca with pairing vs data



EOS: PAL
$M_{cr} = 1.35\ M_{Sun}$

Pairing gaps:

Neutron $^1S_0$: “SFB”
Neutron $^3P_2$: “b”
Proton $^1S_0$: “T73”

Minimal Cooling

Minimal Cooling or, do we need fast cooling ?

Motivation:

Many new observations of cooling neutron stars
with CHANDRA and XMM-NEWTON

Do we have any strong evidence for the presence of some "exotic" component in the core of some of these neutron stars ?

Minimal Cooling of Neutron Stars: A New Paradigm
D. Page, J.M. Lattimer, M. Prakash & A.W. Steiner
2004ApJS..155..623P

Neutrino Emission from Cooper Pairs and Minimal Cooling of Neutron Stars
Page, Dany; Lattimer, James M.; Prakash, Madappa; Steiner, Andrew W.
2009ApJ...707.1131P

Minimal Cooling or, do we need fast cooling ?

Minimal Cooling assumes:
nothing special happens in the core, i.e.,
no direct URCA, no π^- or K^- condensate,
no hyperons, no deconfined quark matter, no ...

(and no medium effects enhance the
modified URCA rate beyond its standard value)

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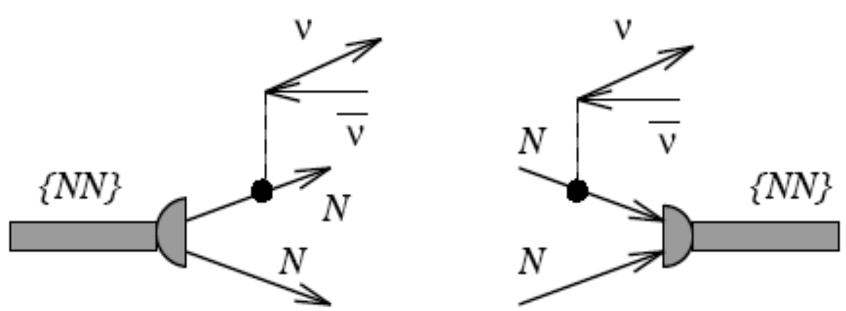
Minimal Cooling is not naive cooling:

it takes into account uncertainties due to

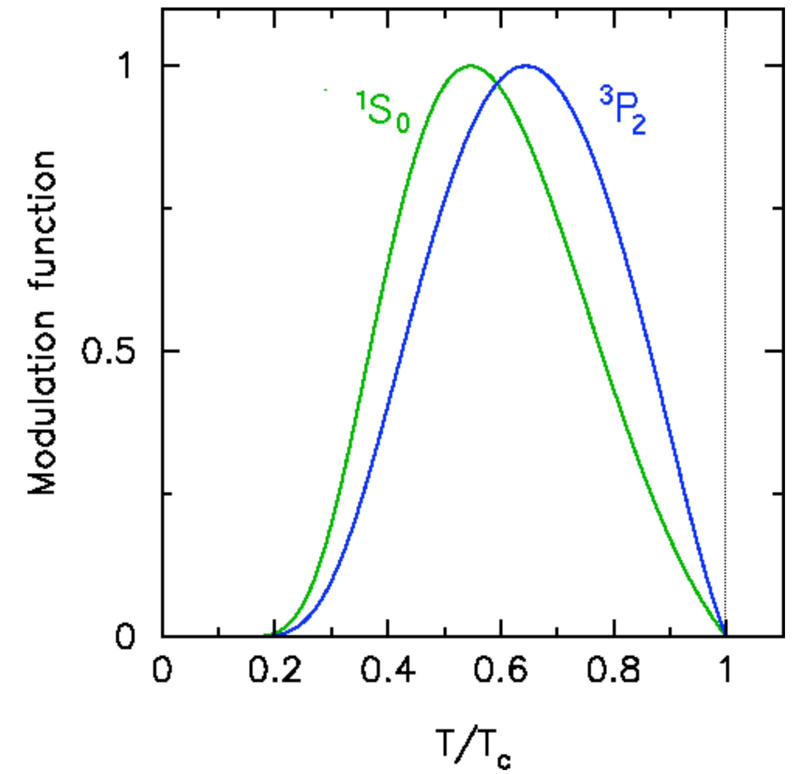
- Large range of predicted values of T_c for n & p.
- Enhanced neutrino emission at $T \leq T_c$ from the Cooper pair formation mechanism.
- Chemical composition of upper layers (envelope), i.e., iron-peak elements or light (H, He, C, O, ...) elements, the latter significantly increasing T_e for a given T_b.
- Equation of state.
- Magnetic field.

Neutrino emission from the breaking (and formation) of Cooper pairs: “PBF”



$$Q = \frac{4G_F^2 m_i^* p_{F,i}}{15\pi^5 \hbar^{10} c^6} (k_B T)^7 \mathcal{N}_\nu a_{i,j} F_j \left[\Delta_i(T)/T\right]$$

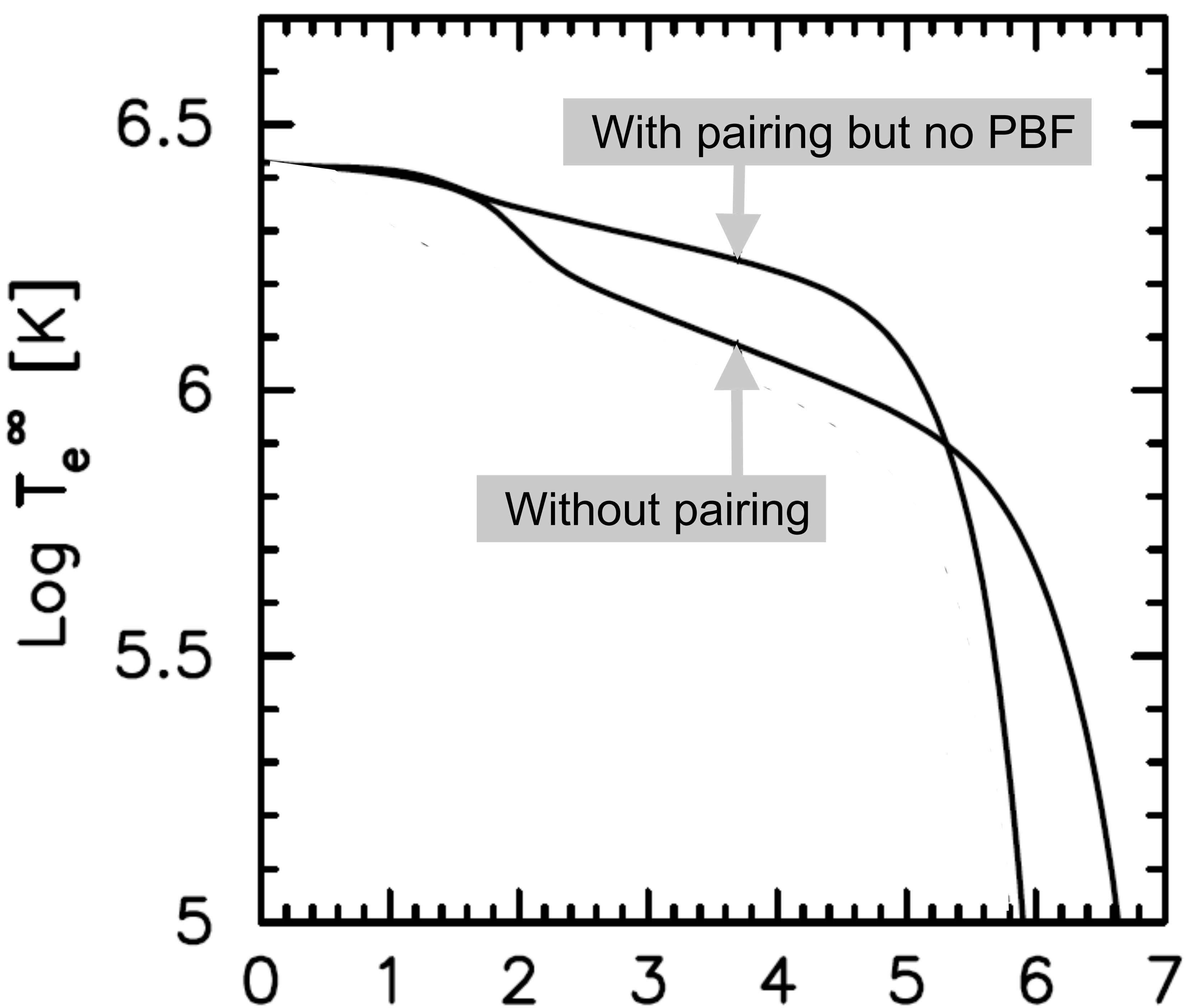
$$= 3.51 \times 10^{21} \frac{\mathrm{erg}}{\mathrm{cm^3\ s}} \left(\frac{m_i^*}{m_i}\right) \left(\frac{p_{F,i}}{m_i c}\right) \times T_9^7 a_{i,j} F_j \left[\Delta(T)/T\right]$$



Neutrino pair emission from finite-temperature neutron superfluid and the cooling of neutron stars
E Flowers, M Ruderman & P Sutherland, 1976ApJ...205..541F

Voskresensky D., Senatorov A., 1986, Sov. Phys.–JETP 63, 885

Basic effects of pairing on the cooling



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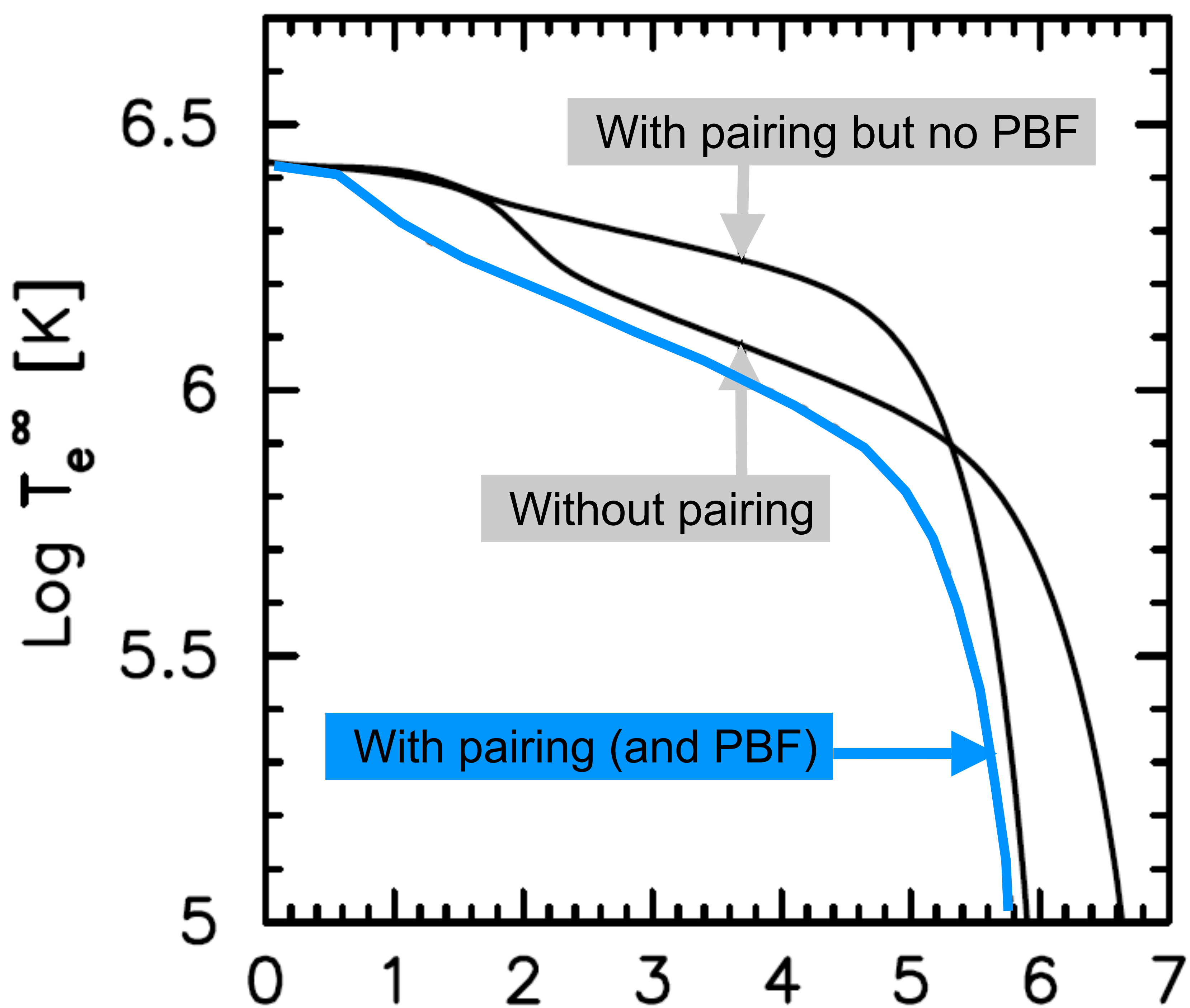
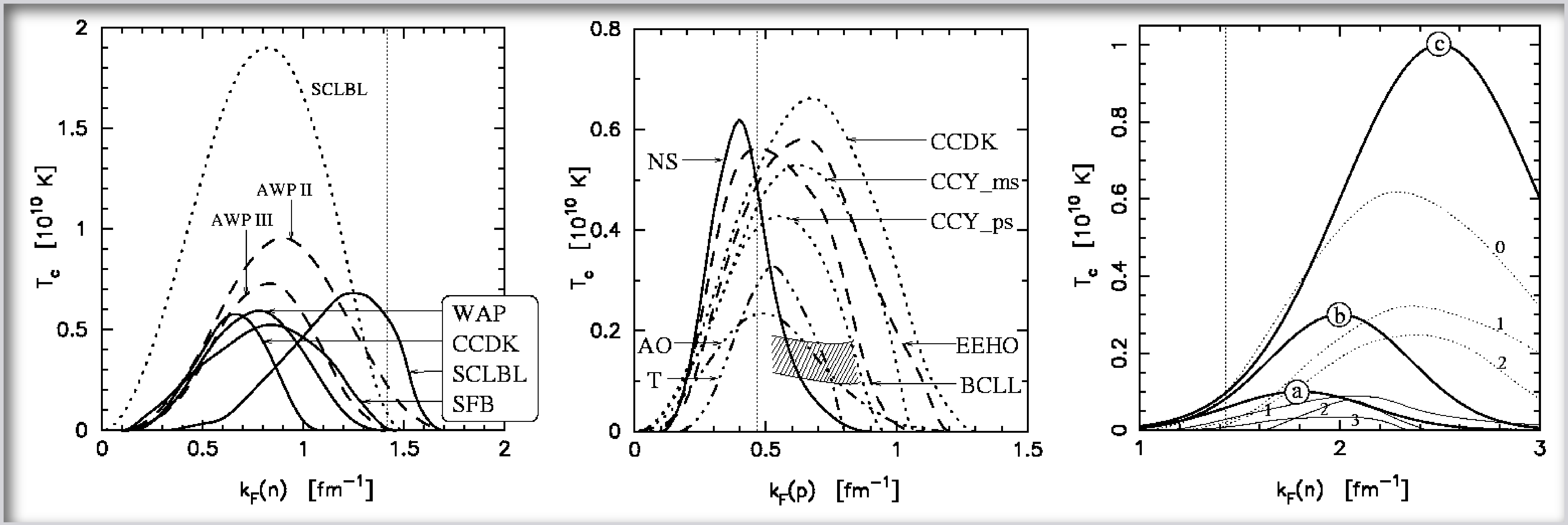


Pairing T_c models

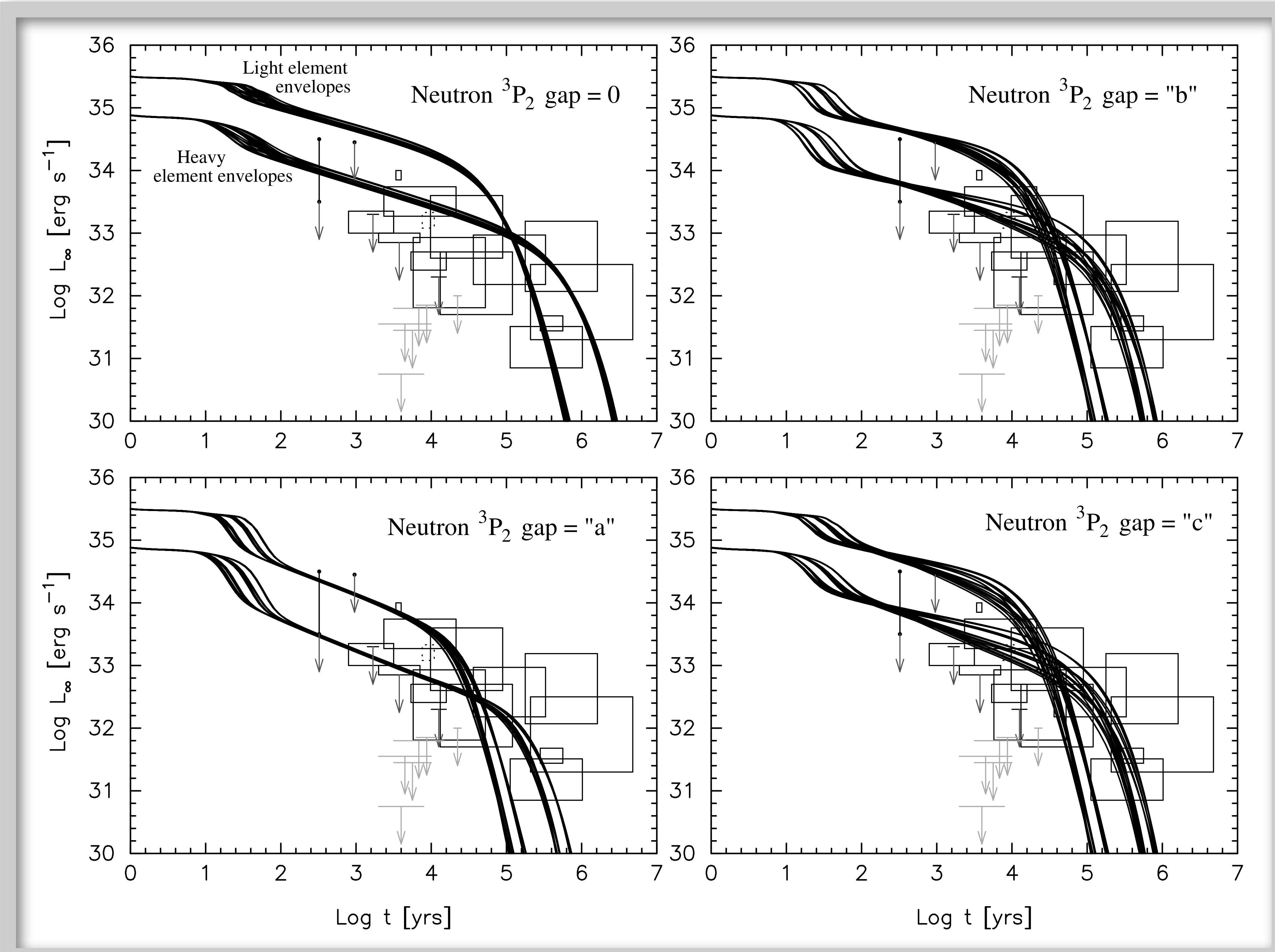
Neutron $^1S_0$ Proton $^1S_0$ Neutron $^3P_2$



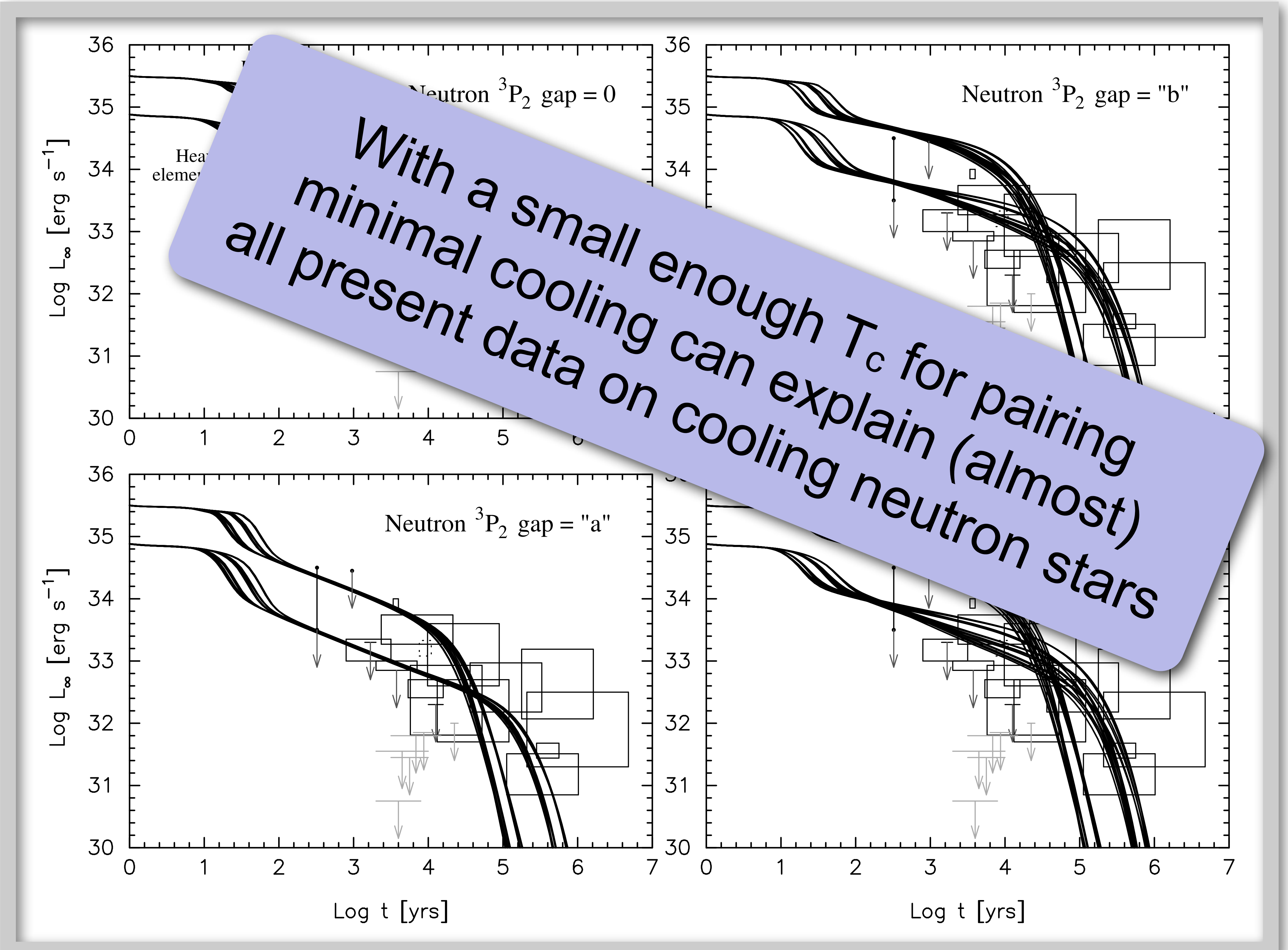
Size and extent of pairing gaps is highly uncertain

Wednesday, July 7, 2010

Minimal cooling versus data



Minimal cooling versus data



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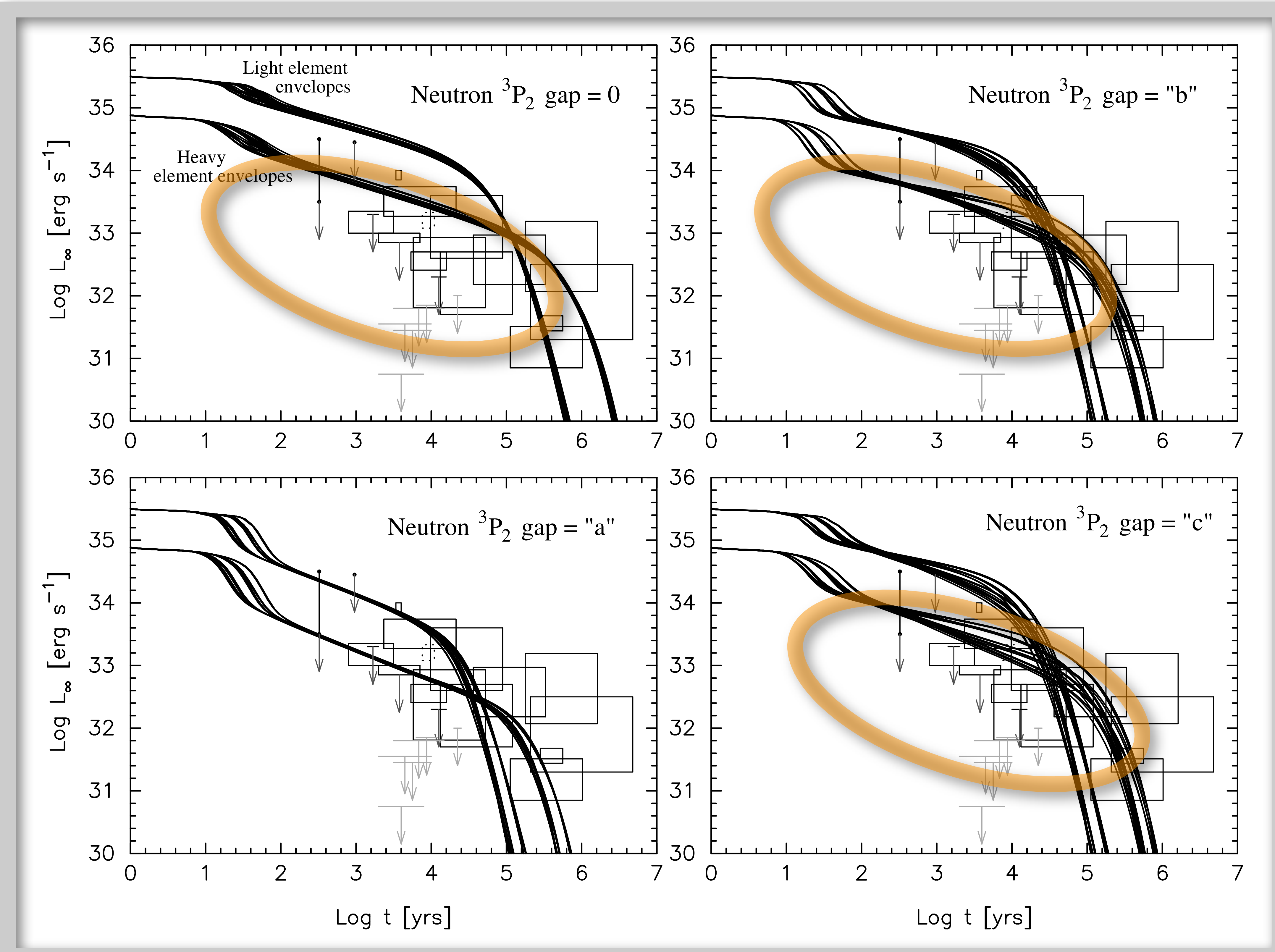

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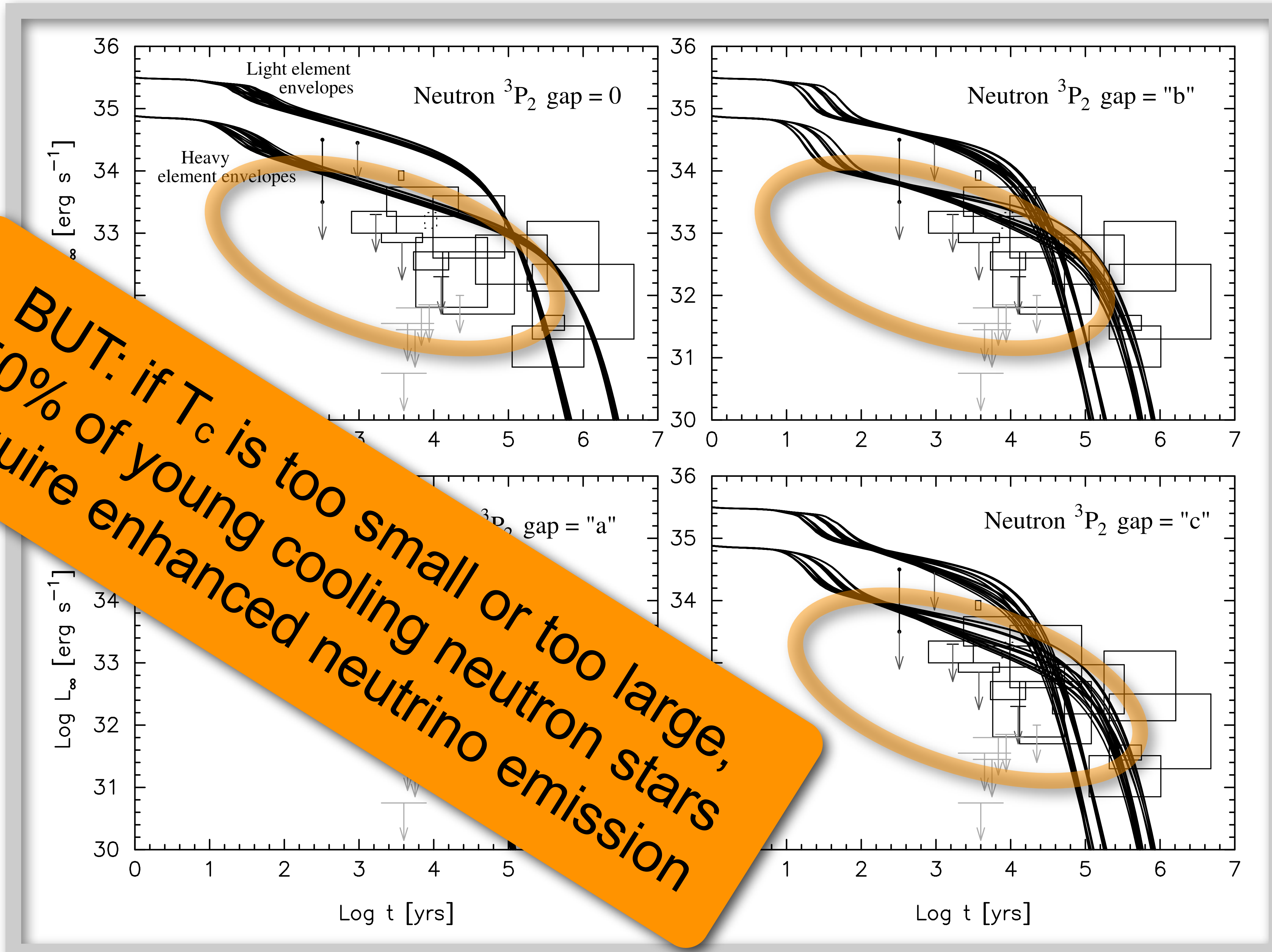
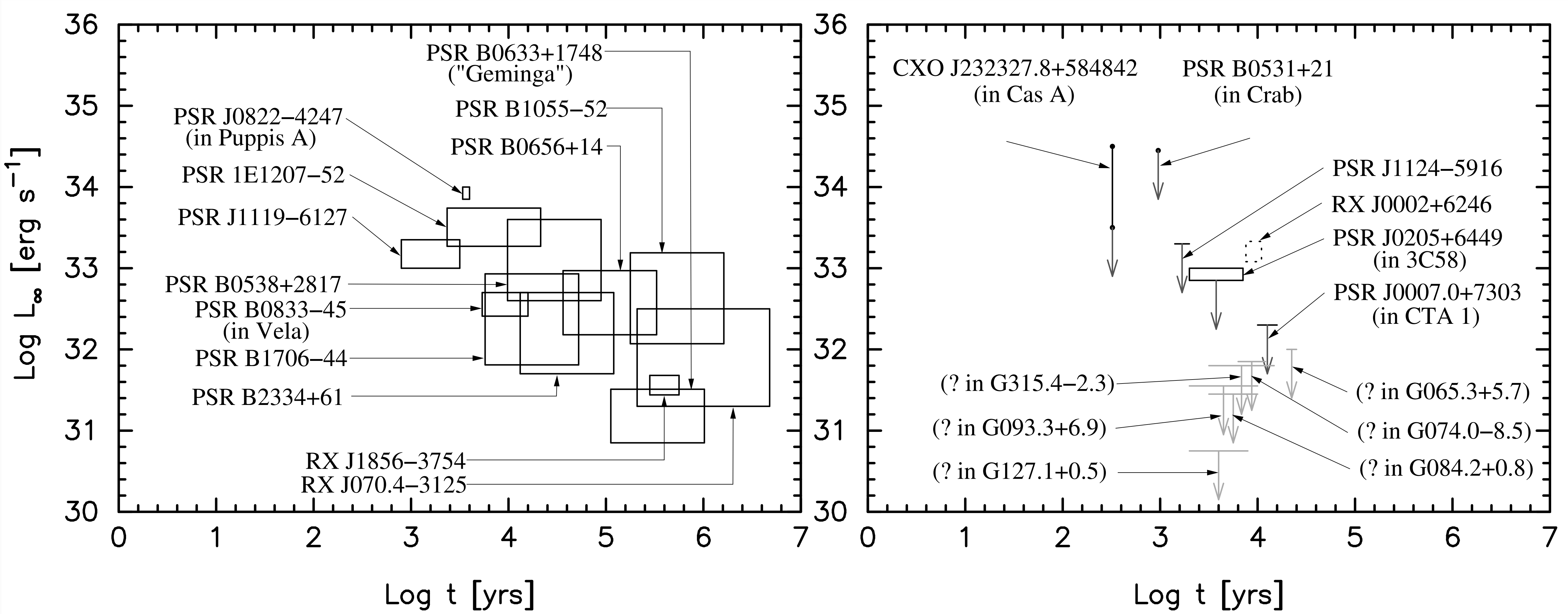
Minimal cooling versus data



BUT: if T_c is too small or too large, 50% of young cooling neutron stars require enhanced neutrino emission

Observational data



Conclusions

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- Many possibilities for fast neutrino emission.
- Neutrino emission can be strongly suppressed by pairing.
- Fast cooling scenarios are compatible with if T_c for pairing is large enough.
- Minimal Cooling: most observed isolated cooling neutron stars are OK
  if the neutron $^3P_2$ gap has the correct size.
  If not about 50% of them require some faster neutrino emission.
- A few serious candidates for neutrino cooling beyond minimal.


"That's all Folks!"