A one and a half hours walk through the physics of neutron stars

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CFisUC

Rewriting Nuclear Physics Textbooks: Basic nuclear interactions and their applications to nuclear processes in the Cosmos and on Earth
24th-28th July, Pisa (Italy)

The final message of this lecture

Neutron stars are excellent observatories to test fundamental properties of matter under extreme conditions and offer an interesting interplay between nuclear processes and astrophysical observables

Road Map of this Walk

- Very general introduction to neutron stars
- A brush-stroke on the role of hyperons in neutron stars

Neutron stars are different things for different people

- For astronomers are very little stars "visible" as radio pulsars or sources of X- and γ-rays.

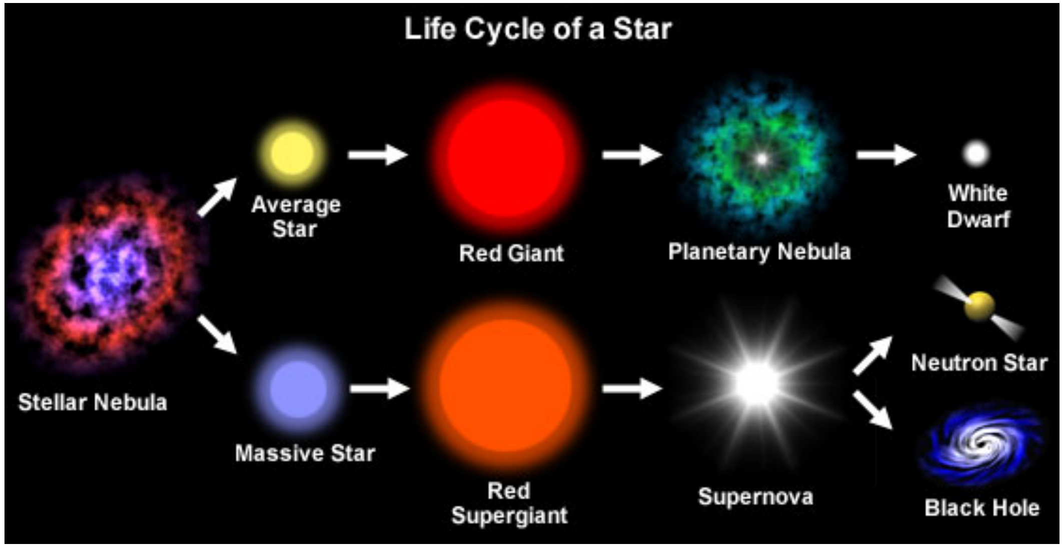
- For particle physicists are neutrino sources (when they born) and probably the only places in the Universe where deconfined quark matter may be abundant.

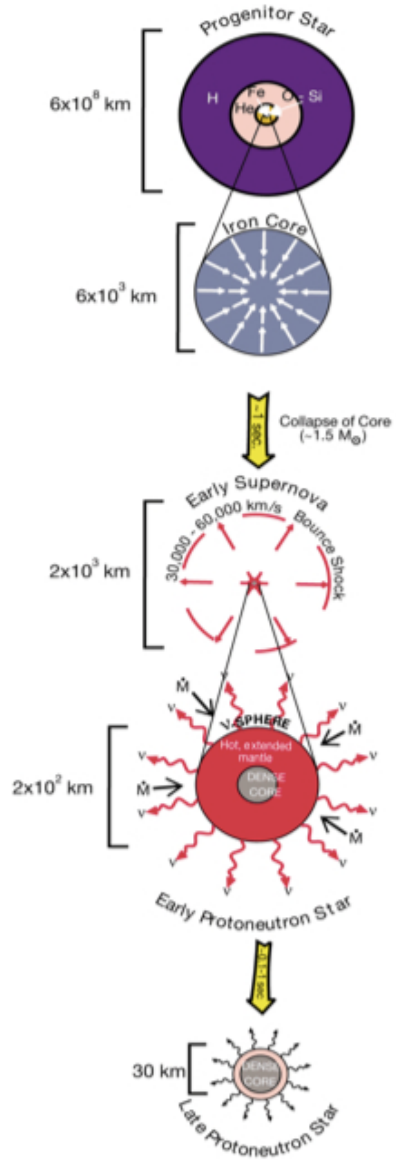
- For cosmologists are "almost" black holes.

- For nuclear physicists are the biggest neutron-rich nuclei of the Universe ($A \sim 10^{56}$-10^{57}, $R \sim 10$ km, $M \sim 1$-$2\ M_\odot$).

But everybody agrees that …

Neutron stars are a type of stellar compact remnant that can result from the gravitational collapse of a massive star (8 $M_{\odot}$< M < 25 $M_{\odot}$) during a Type II, Ib or Ic supernova event.



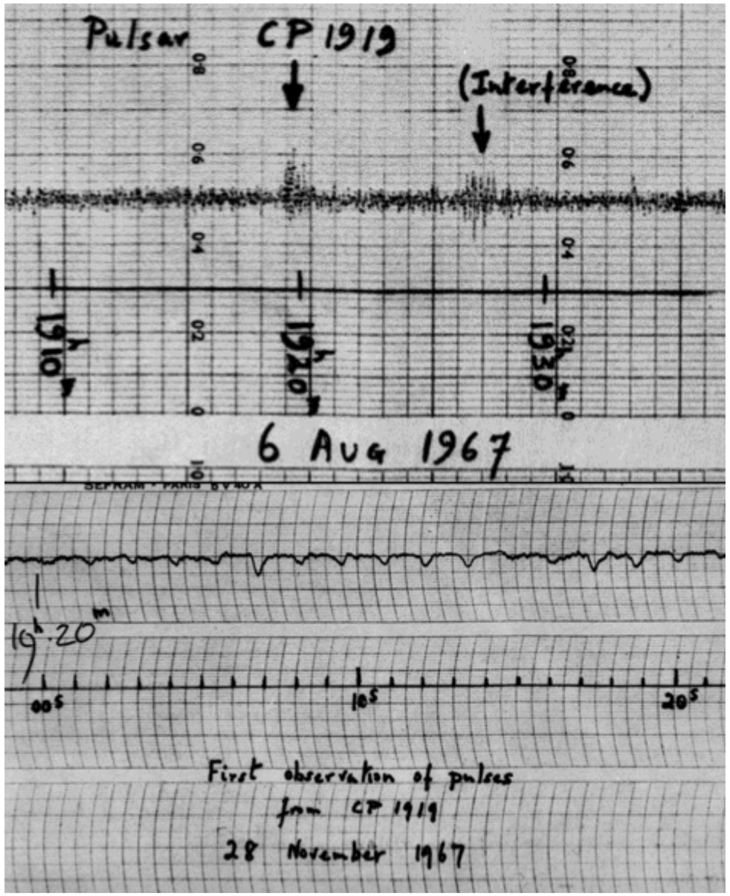


50 years of the discovery of the first radio pulsar

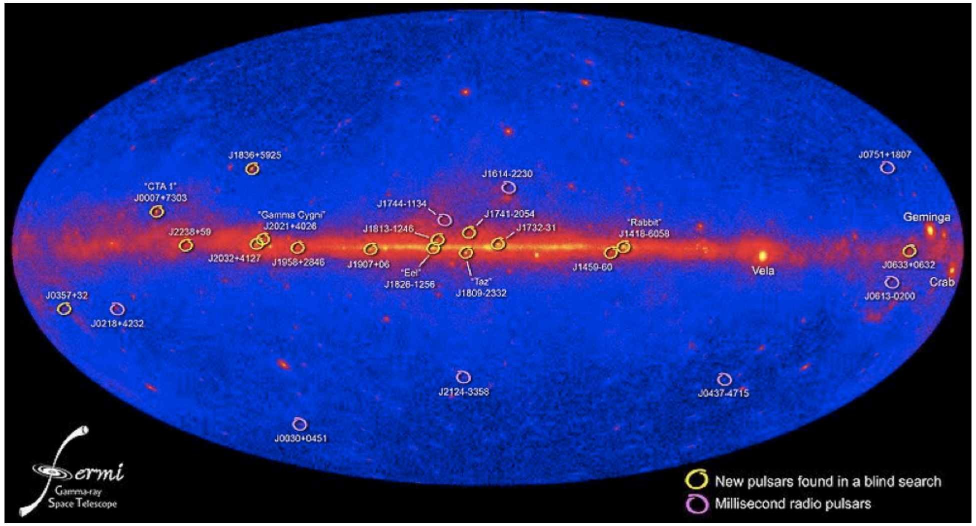
✧ radio pulsar at 81.5 MHz

✧ pulse period P=1.337 s

Most NS are observed as pulsars. In 1967 Jocelyn Bell & Anthony Hewish discover the first radio pulsar, soon identified as a rotating neutron star (1974 Nobel Prize for Hewish but not for Jocelyn)



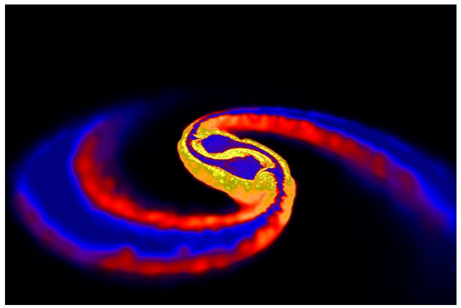
Nowadays more than 2000 pulsars are known (~ 1900 Radio PSRs (141 in binary systems), ~ 40 X-ray PSRs & ~ 60 γ-ray PSRs)



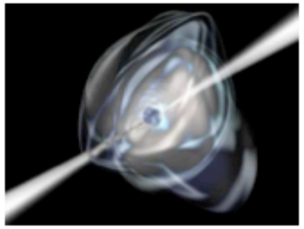
http://www.phys.ncku.edu.tw/~astrolab/mirrors/apod_e/ap090709.html

Observables

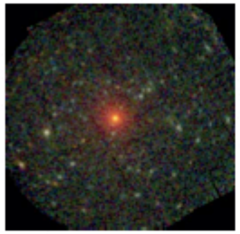
- Period (P, dP/dt)
- Masses
- Luminosity
- Temperature
- Magnetic Field
- Gravitational Waves (NS-NS, BH-NS mergers, NS oscillation modes)



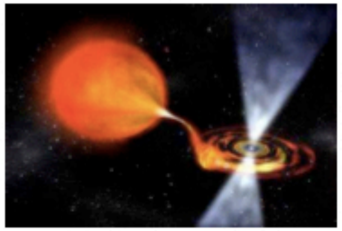
The 1001 Astrophysical Faces of Neutron Stars



Anomalous X-ray Pulsars

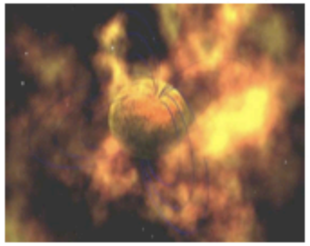


dim isolated neutron stars



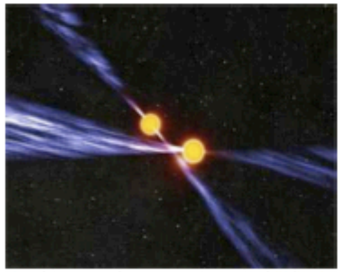
X-ray binaries

bursting pulsars

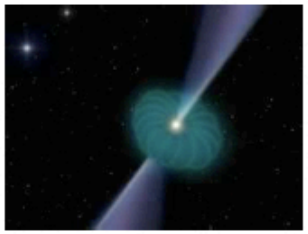


Soft Gamma Repeaters

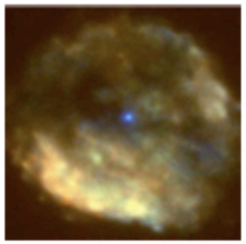
pulsars



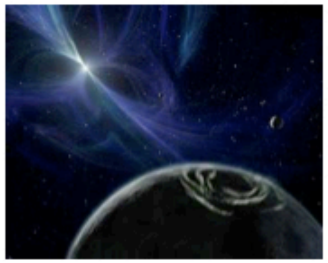
binary pulsars



Rotating Radio Transients



Compact Central Objects



planets around pulsar

Observation of Neutron Stars

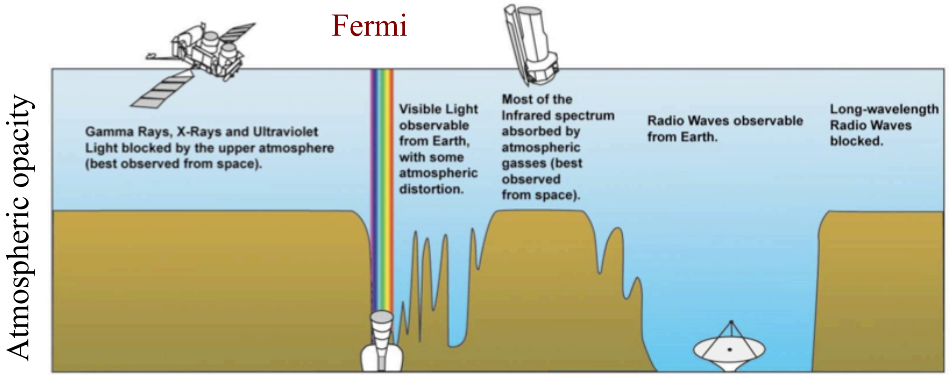
X- and γ-ray telescopes

Chandra

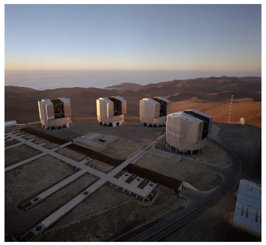
Fermi

Space telescopes

HST (Hubble)



Optical telescopes



VLT (Atacama, Chile)

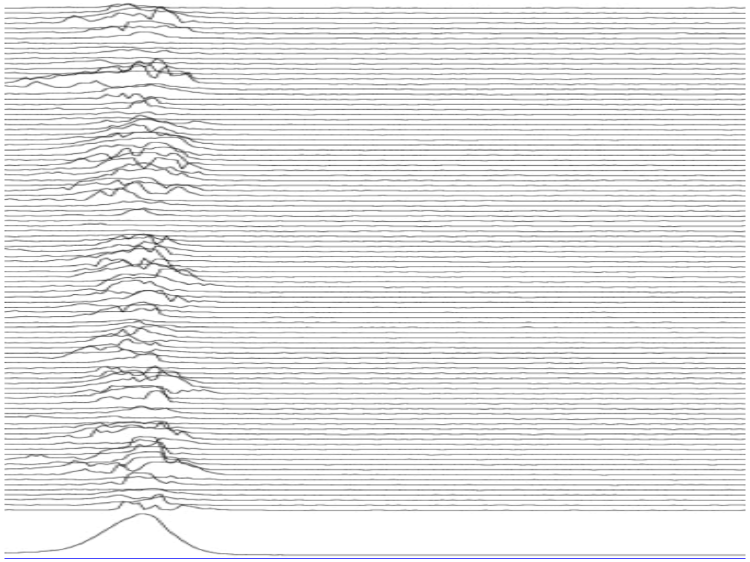
Arecibo (Puerto Rico): d= 305 m

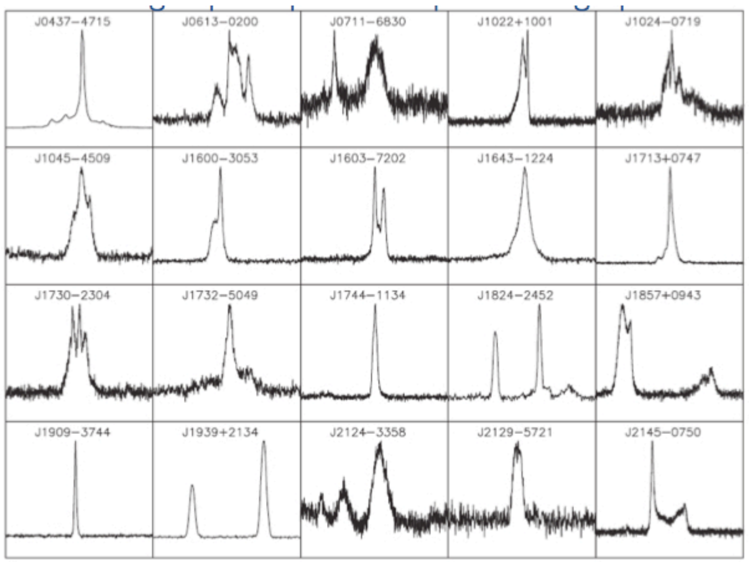
Radio telescopes

Green Banks (USA): d= 100 m

Nançay (France): d ~ 94 m

The Fingerprint of a Pulsar



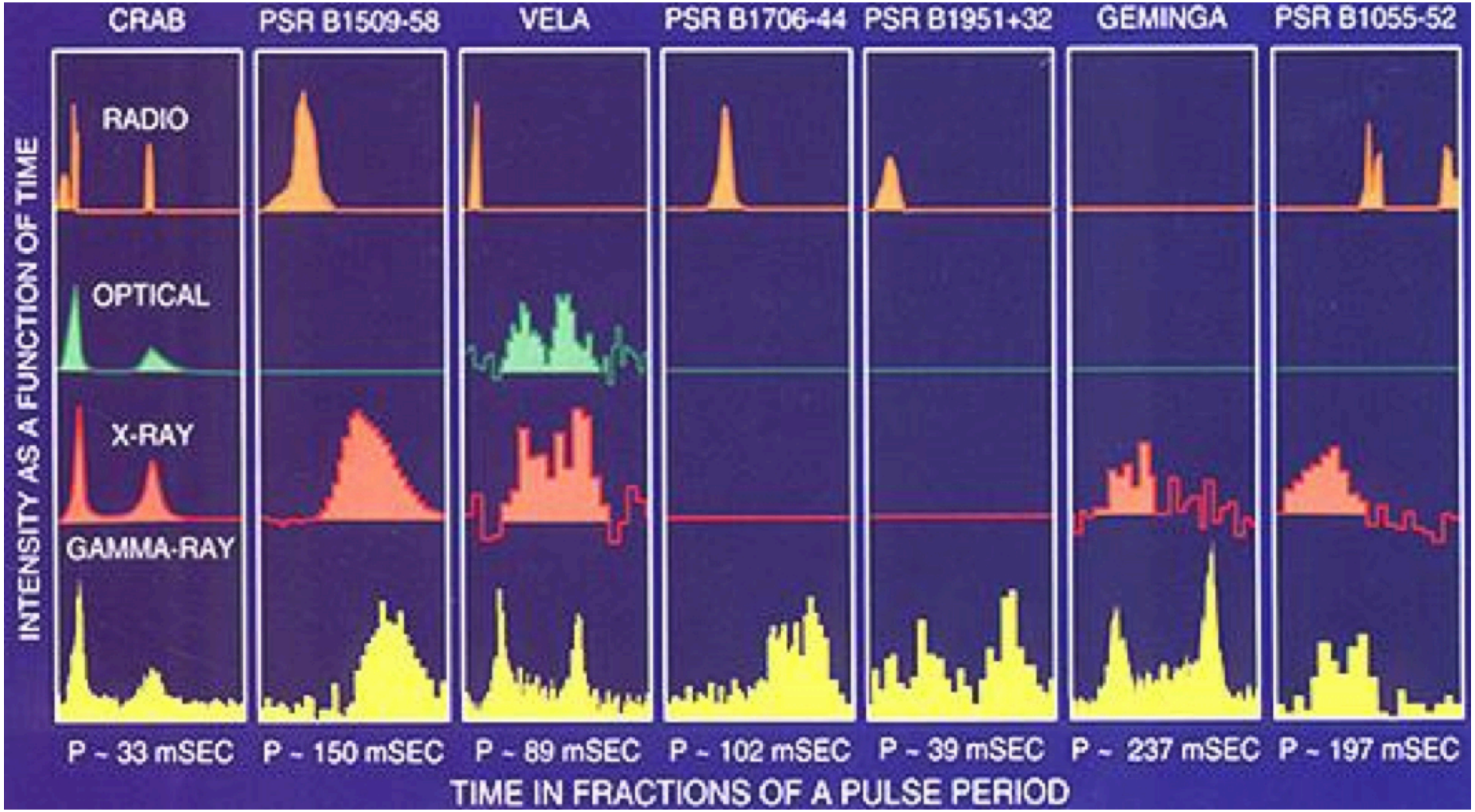


Individual pulses are very different. But the average over 100 or more pulses is extremely stable and specific of each pulsar

- Top: 100 single pulses from the pulsar PSR B0950+08 (P=0.253 s) showing the pulse-to-pulse variability in shape and intensity
- Bottom: Average profiles of several pulsars

Hobbs et al., Pub Astr. Soc. Aust., 202, 28 (2011)

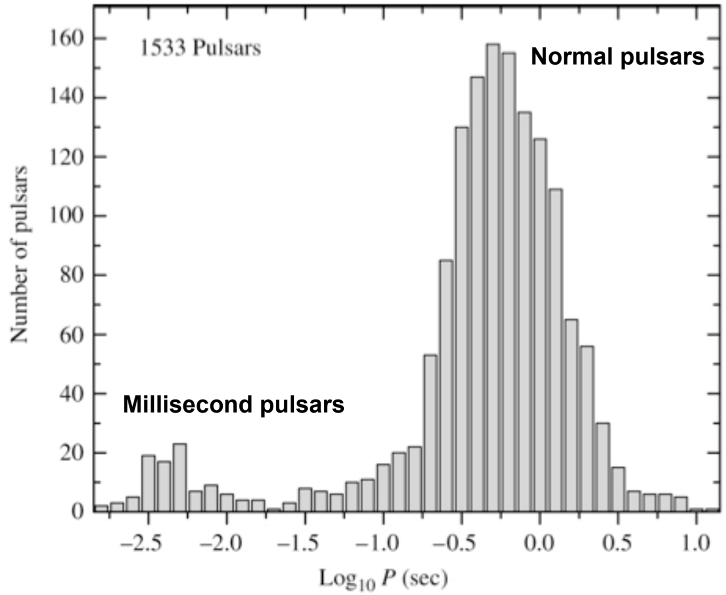
Pulsar shape at different wavelength



Pulsar Rotational Period

The distribution of the rotational period of pulsars shows two clear peaks that indicate the existence of two types of pulsars

- normal pulsars with P ~ s
- millisecond pulsars with P ~ ms

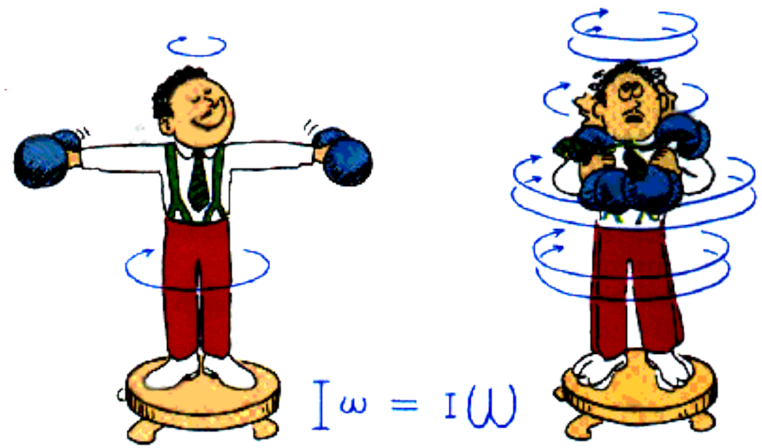


Globular cluster Terzan 5

- First millisecond pulsar discovered in 1982 (Arecibo)
- Nowadays more than 200 millisecond pulsars are known
- PSR J1748-2446ad discovered in 2005 is until know the fastest one with P=1.39 ms (716 Hz)

Why Pulsars spin so fast ?

Conservation of the angular momentum and mass during the gravitational collapse of the iron core that will form the neutron star



If the initial iron core and the final neutron star are assumed to be rigid spheres with moment of inertia $I=(2/5)MR^2$

$$J_i = J_f \Rightarrow P_f = P_i\left(\frac{R_f}{R_i}\right)^2$$

Taking $P_i \sim 10^3$ s and $R_f/R_i \sim 10^{-2}$ one gets $P_f \sim 10^{-3}$ s

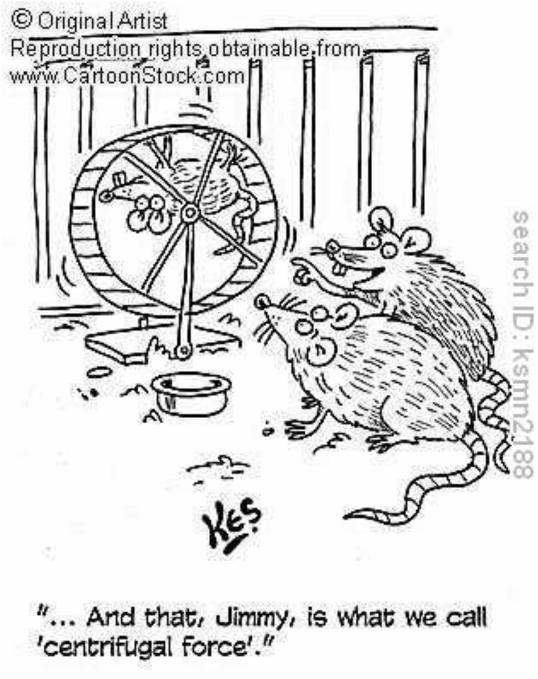
Minimum Rotational Period of a Neutron Star

Pulsar cannot spin arbitrarily fast. The absolute minimum rotational period is obtained when

Centrifugal Force = Gravitational Force

↓

Keplerian Frequency



In Newtonian Gravity

$$P_{\min} = 2\pi\sqrt{\frac{R^3}{GM}} \approx 0.55\left(\frac{M_{sun}}{M}\right)^{1/2}\left(\frac{R}{10km}\right)^{3/2} \quad ms$$

In General Relativity

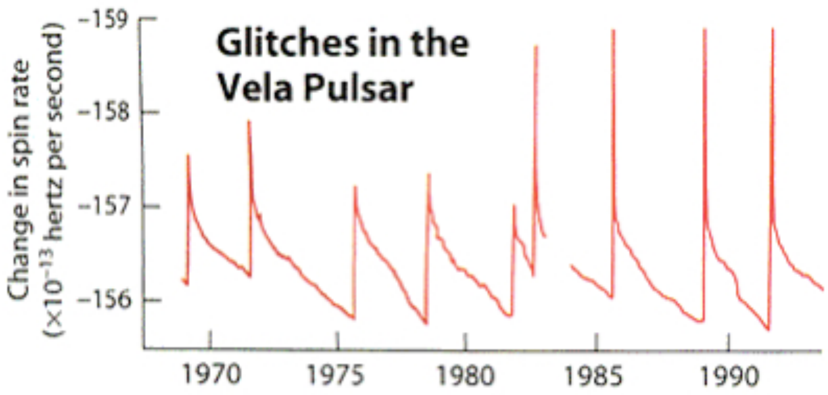
$$P_{\min} = 0.96\left(\frac{M_{sun}}{M}\right)^{1/2}\left(\frac{R}{10km}\right)^{3/2} \quad ms$$

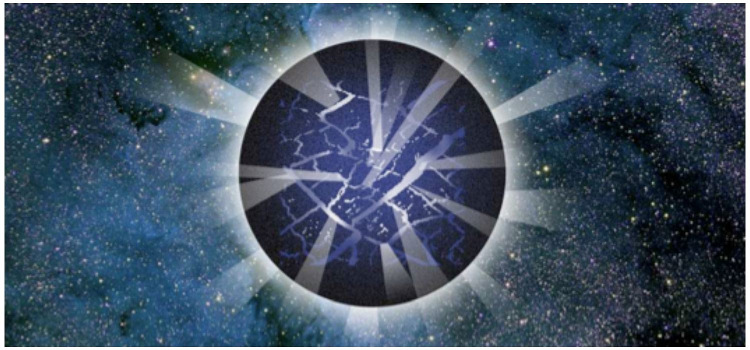
Actual record: PSR J1748-2446ad → P=1.39595482 ms

Pulsar glitches

Sometimes the period P of a pulsar decreases suddenly. These variations (glitches), although small, are observable

$$\frac{\Delta\Omega}{\Omega} \approx 10^{-9} - 10^{-5}$$



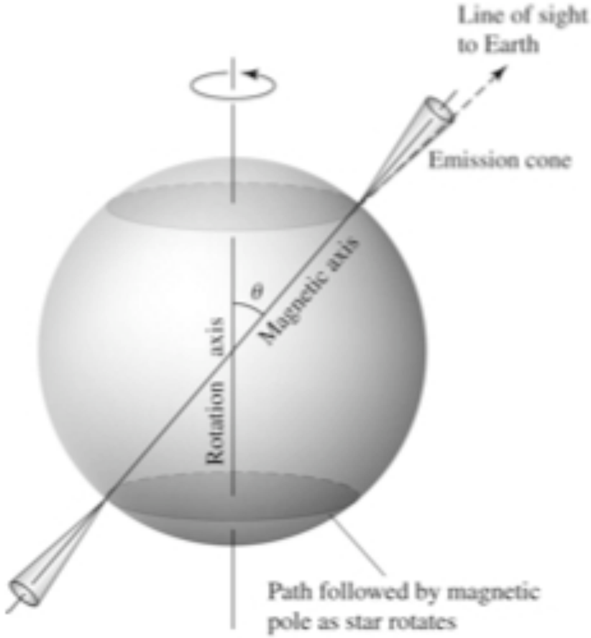


Small glitches are interpreted as starquakes ➔ indirect proof of the existence of NS crust

Big glitches and the observation of long relaxation times are a proof of the existence of superfluid matter in the NS interiors

Basic Model of a Pulsar: Magnetic Dipole

Pulsar are believed to be highly magnetized rotating neutron stars radiating at the expenses of their rotational energy



$$\dot{E}_{mag} = -\frac{2}{3c^3}\left|\ddot{\vec{\mu}}\right|^2 = \dot{E}_{rot}$$

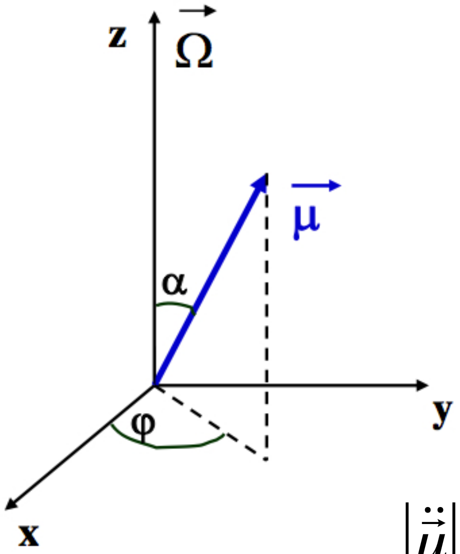
$\vec{\mu} \equiv$ Magnetic dipole moment

Pacini, Nature 216 (1967), 219 (1968)

Gold, Nature 218 (1968), 221 (1969)

Ostriker & Gunn, ApJ 157 (1969)

Basic Model of a Pulsar: Magnetic Dipole



$$\vec{\mu} = \mu \sin\alpha \cos\varphi \hat{e}_x + \mu \sin\alpha \sin\varphi \hat{e}_y + \mu \cos\alpha \hat{e}_z$$

supposing: α=const, μ=|$\vec{\mu}$|=const $\quad \Omega = \frac{d\varphi}{dt} = \dot{\varphi}$

$$\dot{\Omega}^2 << \Omega^4$$

$$\left|\ddot{\vec{\mu}}\right| = \mu^2 (\sin\alpha)^2 \left[\Omega^4 + \dot{\Omega}^2\right] \longrightarrow \left|\ddot{\vec{\mu}}\right| \approx \mu^2 (\sin\alpha)^2 \Omega^4$$

Therefore

$$\dot{E}_{mag} = -\frac{2}{3c^3} \mu^2 (\sin\alpha)^2 \Omega^4 = \dot{E}_{rot}$$

For a sphere with a pure dipole magnetic field

$$\mu = \frac{1}{2} B_P R^3$$

- ✓ B_p: magnetic field at the poles
- ✓ R: radius of the sphere

Basic Model of a Pulsar: Magnetic Dipole

Then $$\dot{E}_{mag} = -\frac{1}{6c^3} R^6 B_P^2 (\sin\alpha)^2 \Omega^4 = \dot{E}_{rot}$$

On the other hand $$E_{rot} = \frac{1}{2} I\Omega^2 \xrightarrow{\dot{I}=0} \dot{E}_{rot} = I\Omega\dot{\Omega}$$

One arrives to the PSR evolution differential equation

$$\dot{\Omega} = -K\Omega^3 \quad or \quad P\dot{P} = (2\pi)^2 K, \quad K = \frac{1}{6c^3}\frac{R^6}{I}(B_P \sin\alpha)^2$$

More generally, one can write the PSR evolution differential equation as

$$\dot{\Omega} = -K\Omega^n \quad or \quad P^{n-2}\dot{P} = (2\pi)^{n-1} K, \quad K = \frac{1}{6c^3}\frac{R^6}{I}\left(B_P \sin\alpha\right)^2$$

with solution

$$\Omega(t) = \frac{\Omega_0}{\left[(n-1)K\Omega_0^{n-1}t+1\right]^{1/(n-1)}}, \quad P(t) = P_0\left[(n-1)K\Omega_0^{n-1}t+1\right]^{1/(n-1)}$$

Differenciating it assuming K=const, one obtains

$$n = \frac{\Omega\ddot{\Omega}}{\dot{\Omega}^2} = 2 - \frac{P\ddot{P}}{\dot{P}^2} \qquad \text{braking index}$$

n=3 within the magnetic dipole model

The three quantities P, $\dot{P}$ & $\ddot{P}$ have been measured for few PSRs

The Pulsar Age

The solution of the PSR evolution differential equation can be rewritten as

$$t = -\frac{1}{n-1}\frac{\Omega(t)}{\dot{\Omega}(t)}\left[1-\left(\frac{\Omega(t)}{\Omega_0}\right)^{n-1}\right]$$

or

$$t = \tau - \left[(n-1)K\Omega_0^{n-1}\right]^{-1}$$

"True" Pulsar Age

(valid under the assumption K=const.)

with

$$\tau = -\frac{1}{n-1}\frac{\Omega(t)}{\dot{\Omega}(t)} = \frac{1}{n-1}\frac{P(t)}{\dot{P}(t)} \xrightarrow{n=3} \tau = -\frac{1}{2}\frac{\Omega(t)}{\dot{\Omega}(t)} = \frac{1}{2}\frac{P(t)}{\dot{P}(t)}$$

Pulsar Dipole Age

if $\Omega(t) << \Omega_0 \longrightarrow t \approx \tau$

(t=present time)

The measure of P and $\dot{P}$ gives the pulsar dipole age

Example: the age of the Crab Pulsar

SN explosion: 1054 AD

P=0.0330847 s, $\dot{P}$=4.22765x10^{-13} s/s

Braking index: n=2.515 +/- 0.005

t_{Crab}=(2014-1054) yr = 960 yr, τ=1238 yr (dipole age)

Assuming the validity of the pulsar dipole mode, using the previous equation for the true pulsar age we can infer the initial spin period of the Crab pulsar

$$P_0 = P\left(1 - \frac{t_{Crab}}{\tau}\right)^{1/2} \cong 0.016s$$

But $n \neq 3$

Measured value of the braking index n

PSR	n	P (s)	P dot (10^{-15} s/s)	Dipole age (yr)
PSR B0531+21 (Crab)	2.512 +/- 0.005	0.03308	422.765	1238
PSR B0833-45 (Vela)	1.4 +/- 0.2	0.08933	125.008	11000
PSR B0540-69	2.839 +/- 0.005	0.1506	1536.5	1554
PSR B0540-69	2.01 +/- 0.02	0.0505	478.924	1672
PSR J1119-6127	2.91 +/- 0.05	0.40077	4021.782	1580

Deviations of braking index n from 3 probably due to:

- ✓ Torque on the pulsar from outflow particles
- ✓ Change with t of "constant" K, i.e., I(t), B(t), α(t)

Pulsar evolutionary path on the P-$\dot{P}$ plane

Taking the logarithm of

$$P\dot{P} = (2\pi)^2 K, \quad K = \frac{1}{6c^3}\frac{R^6}{I}(B_P \sin\alpha)^2$$

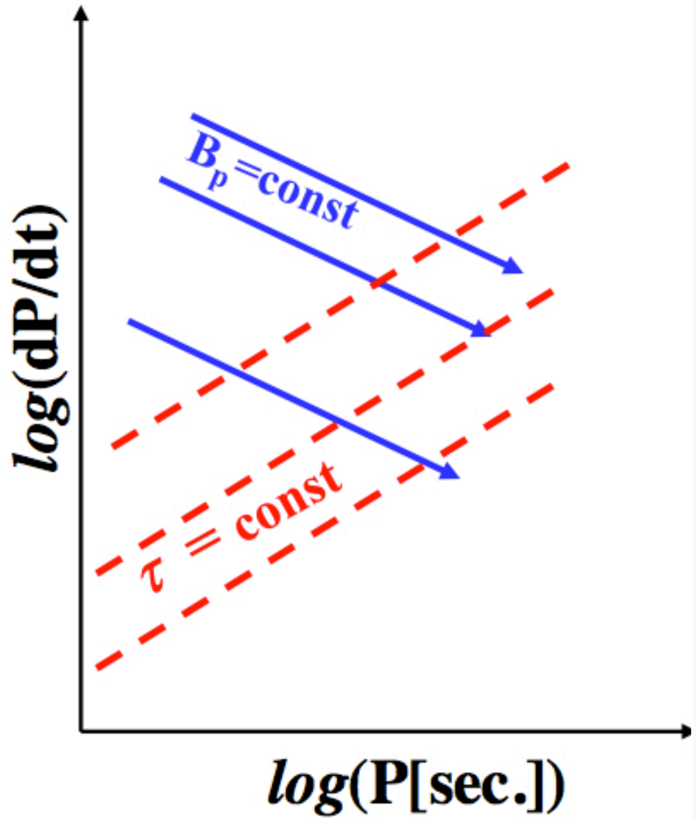
and

$$\tau = \frac{P}{2\dot{P}}$$

we get

$$\log \dot{P} = \log\left[\frac{(2\pi)^2 R^6}{6c^3 I} B_P^2 \sin^2\alpha\right] - \log P$$

$$\log \dot{P} = \log P - \log(2\tau)$$

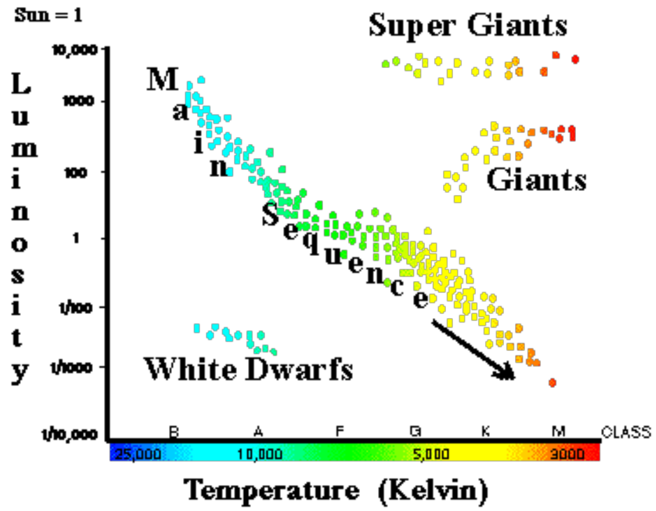


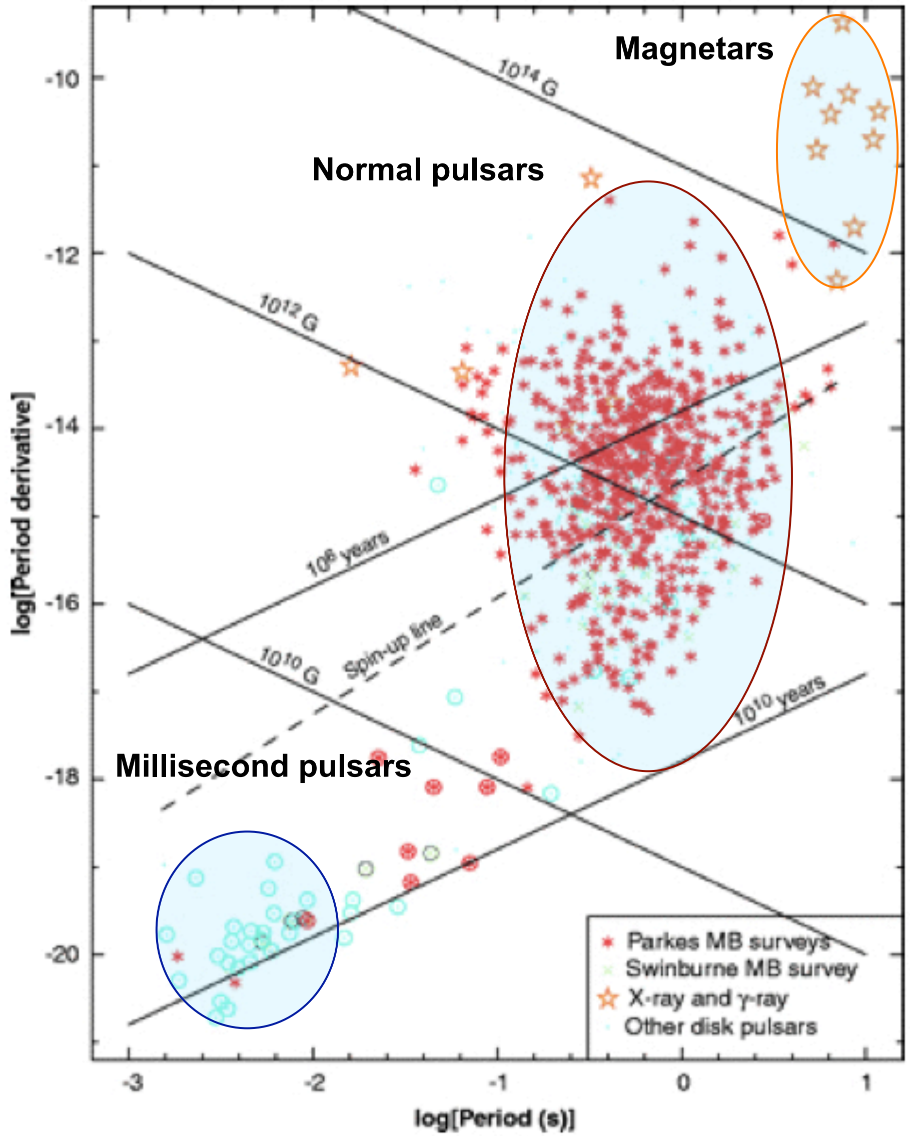
Pulsar distribution in the P-$\dot{P}$ plane

$$\log \dot{P} = \log\left[\frac{(2\pi)^2 R^6}{6c^3 I} B_P^2 \sin^2\alpha\right] - \log P$$

$$\log \dot{P} = \log P - \log(2\tau)$$

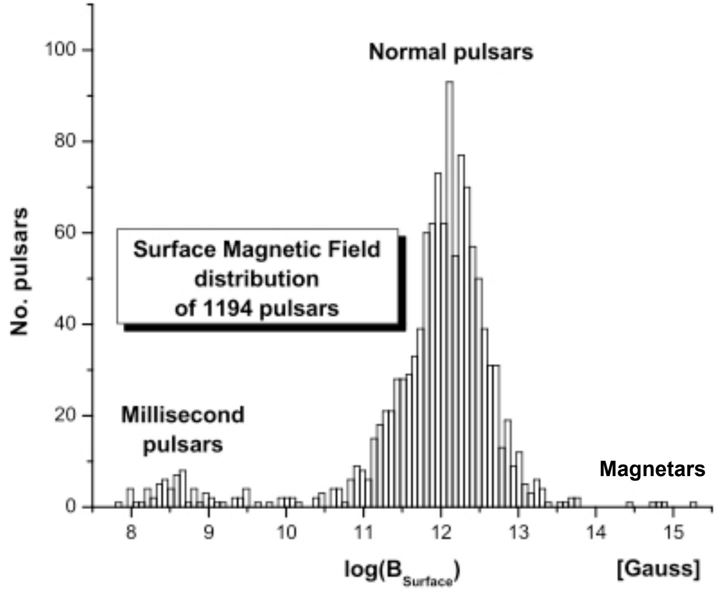
Pulsar equivalent of the Hertzprung-Russell diagram for ordinary stars





Magnetic Field of a Pulsar

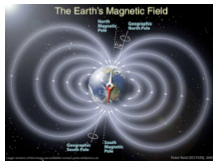
Type of Pulsar	Surface magnetic field
Millisecond	$10^8 – 10^9$ G
Normal	10^{12} G
Magnetar	$10^{14} – 10^{15}$ G



Extremely high compared to …

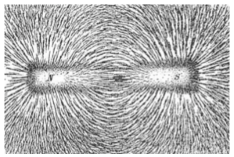
Earth

$0.3 - 0.5G$



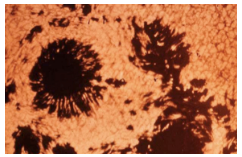
Magnet

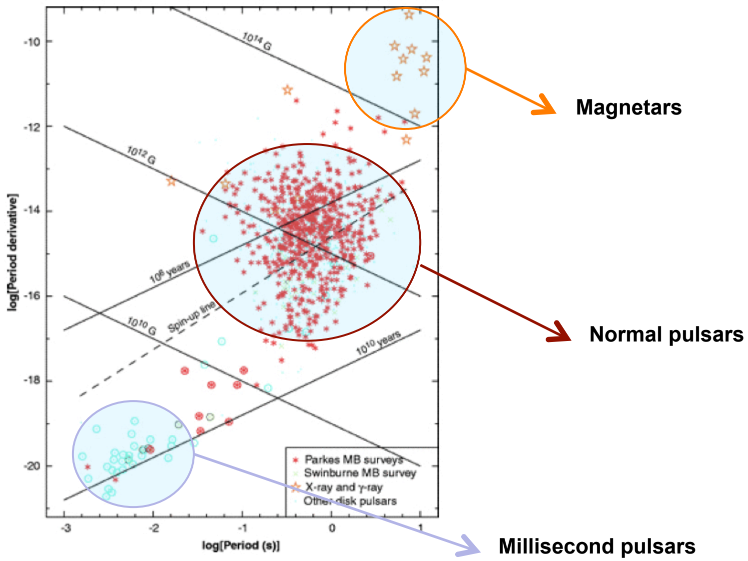
$10^3 - 10^4 G$



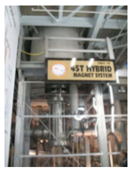
Sun spots

$10^5 G$





Largest continuous field in lab. (USA)



$4.5x10^5 G$

Largest magnetic pulse in lab. (Russia)

$2.8x10^7 G$

Where the NS magnetic field comes from ?

A satisfactory answer does not exist yet. Several possibilities have been considered:

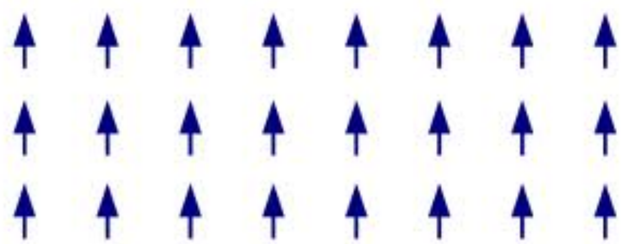
✧ Conservation of the magnetic flux during the gravitational collapse of the iron core

$$\phi_i = \phi_f \Rightarrow B_f = B_i \left(\frac{R_i}{R_f} \right)^2$$

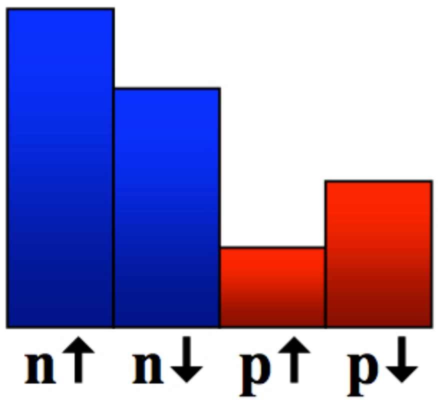
For a progenitor star with $B_i \sim 10^2$ G & $R_i \sim 10^6$ km we have $B_f \sim 10^{12}$ G

✧ Electric currents flowing in the highly conductive NS interior

✧ Spontaneous transition to a ferromagnetic state due to the nuclear interaction



Spin-polarized Isospin Asymmetric Nuclear Matter



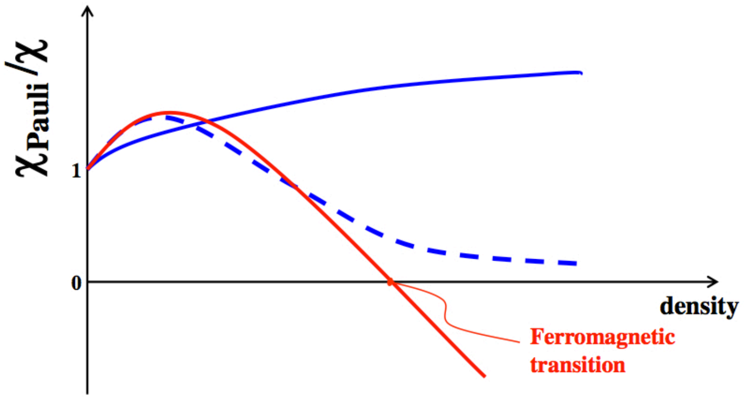
✧ Densities & Asymmetries

- ✓ $\rho_n = \rho_{n\uparrow} + \rho_{n\downarrow}, \quad \rho_p = \rho_{p\uparrow} + \rho_{p\downarrow}$
- ✓ $\rho = \rho_n + \rho_p, \quad \beta = \dfrac{\rho_n - \rho_p}{\rho}$
- ✓ $S_n = \dfrac{\rho_{n\uparrow} - \rho_{n\downarrow}}{\rho_n}, \quad S_p = \dfrac{\rho_{p\uparrow} - \rho_{p\downarrow}}{\rho_p}$

✧ Magnetic Susceptibility

$$\frac{1}{\chi_{ij}} = \frac{\rho}{\mu_i \rho_i \mu_j \rho_j} \frac{\partial^2 (E/A)}{\partial S_i \partial S_j}$$

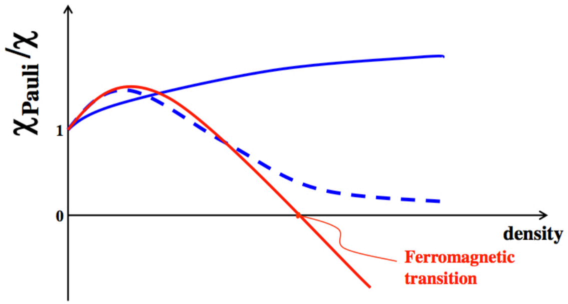
Stability againts spin fluctuations if $\chi > 0$



Ferromagnetic Transition

Considered by many authors with contradictory results:

Year	Autor/Model	Ferromagnetic Transition ?
1969	Brownell, Callaway, Rice (hard sphere gas)	Yes, k_F>2.3 fm^{-1}
1969	Clark & Chao	No
1970	Ostgard	Yes, k_F>4.1 fm^{-1}
1972	Pandharipande et al., (variational)	No
1975	Backman, Kallaman, Haensel (BHF)	No
1984	Vidaurre (Skyrme)	Yes, k_F>1.7-2.0 fm^{-1}
1991	S. Marcos et al., (DBHF)	No
2001	Fantoni et at. (AFDMC)	No
2002/2005	I.V., et al. (BHF)	No
2005/2006	I.V. et al., (Skyrme,Gogny)	Yes, k_F>2-3.4 fm^{-1}
2007-2011	F. Sammarruca (DBHF)	No



- ✧ Calculations based on phenomenological interactions (e.g., Skyrme, Gogny) predict the transition to occur at $(1\text{-}4)\rho_0$
- ✧ Calculations based on realistic NN & NNN forces (e.g., Monte Carlo, BHF, DBHF, LOCV) exclude such a transition

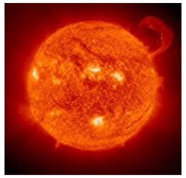
Neutron Star Structure: General Relativity or Newtonian Gravity ?

Surface gravitational potential tell us how much compact an object is

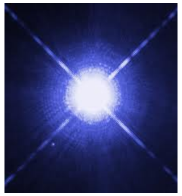
$$\frac{2GM}{c^2R}$$

➔ Relativistic effects are very important in Neutron Stars and General Relativity must be used to describe their structure

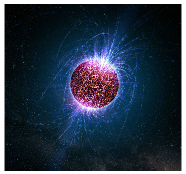
$\sim 10^{-10}$



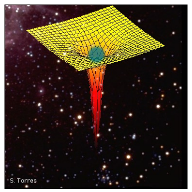
$\sim 10^{-5}$



$\sim 10^{-4} - 10^{-3}$



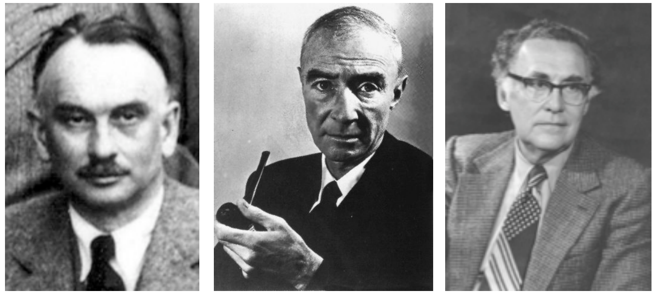
$\sim 0.2 - 0.4$



1

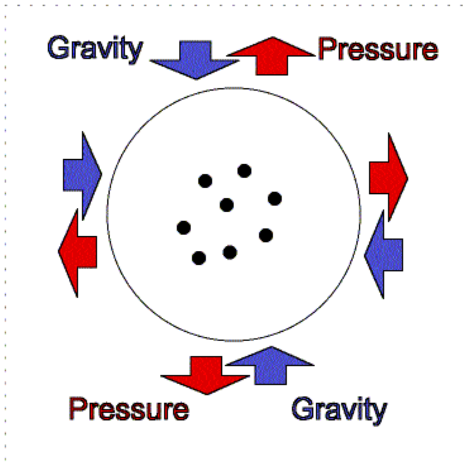
The Tolman-Oppenheimer-Volkoff Equations

In 1939 Tolman, Oppenheimer & Volkoff obtain the equations that describe the structure of a static star with spherical symmetry in General Relativity (Chandrasekhar & von Neumann obtained them in 1934 but they did not published their work)



Tolman, Phys. Rev. 55, 364 (1939)

Oppenheimer & Volkoff, Phys. Rev. 55, 374 (1939)

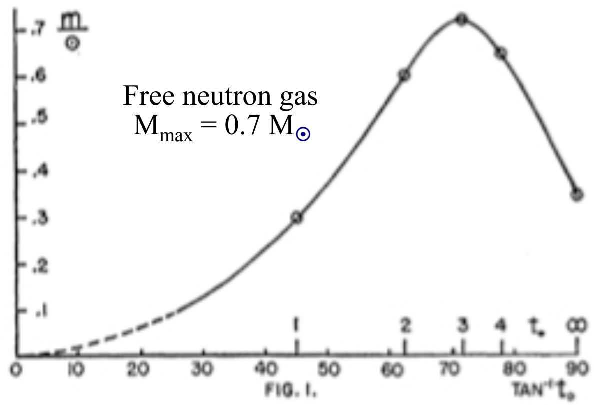


$$\frac{dP}{dr} = -G\frac{m(r)\varepsilon(r)}{r^2}\left(1+\frac{P(r)}{c^2\varepsilon(r)}\right)\left(1+\frac{4\pi r^3 P(r)m(r)}{c^2}\right)\left(1-\frac{2Gm(r)}{c^2 r}\right)^{-1}$$

$$\frac{dm}{dr} = 4\pi r^2 \varepsilon(r)$$

boundary conditions

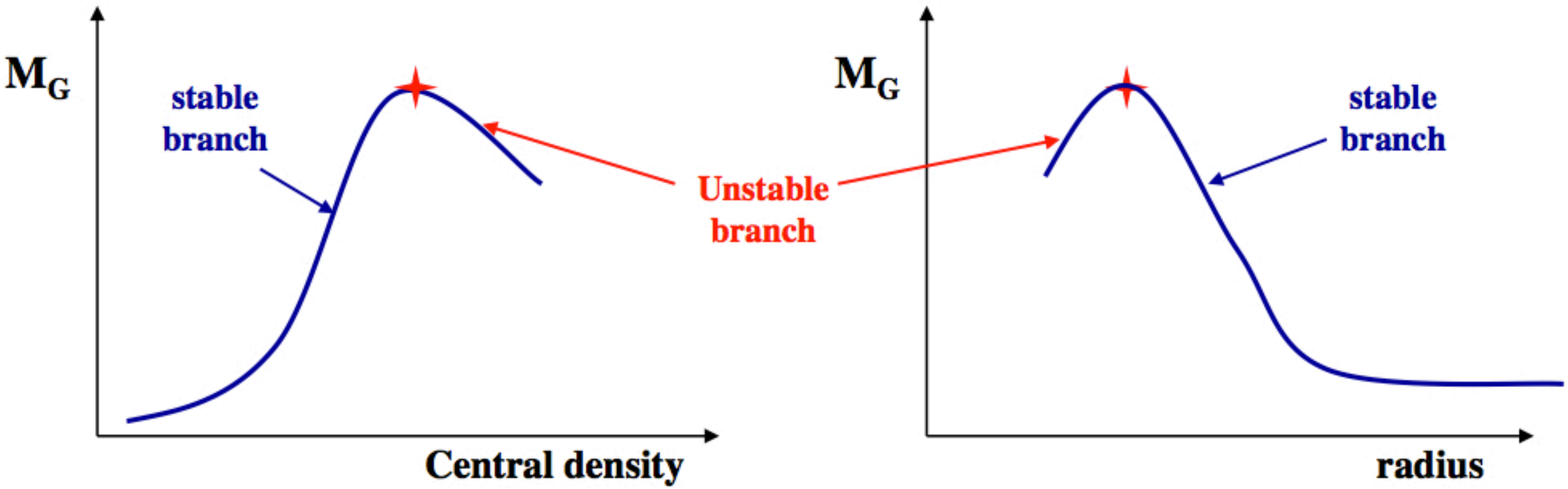
$$P(0) = P_o, \quad m(0) = 0$$
$$P(R) = 0, \quad m(R) = M$$



Stability solutions of the TOV equations

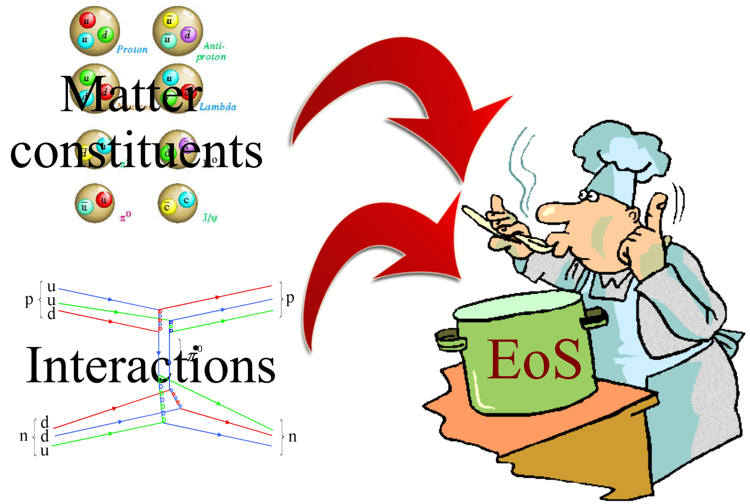
- The solutions of the TOV eqs. represent static equilibrium configurations
- Stability is required with respect to small perturbations

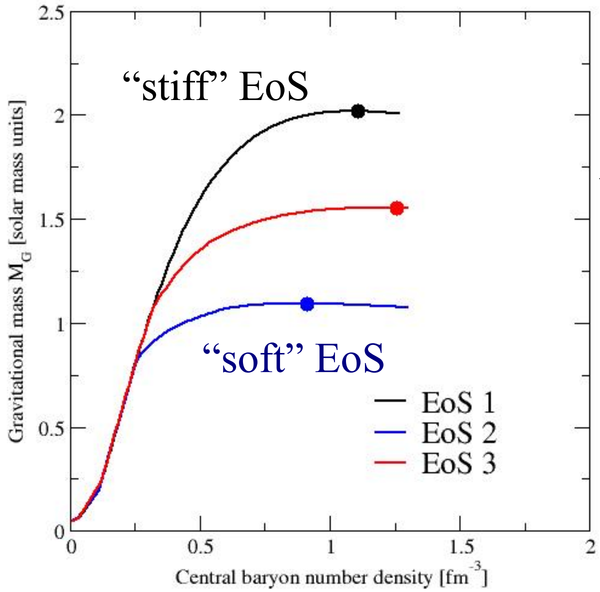
$$\frac{dM_G}{d\rho_c} > 0, \quad or \quad \frac{dM_G}{dr} < 0$$

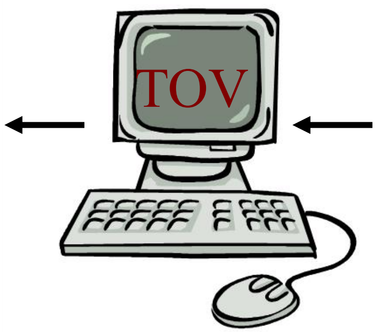


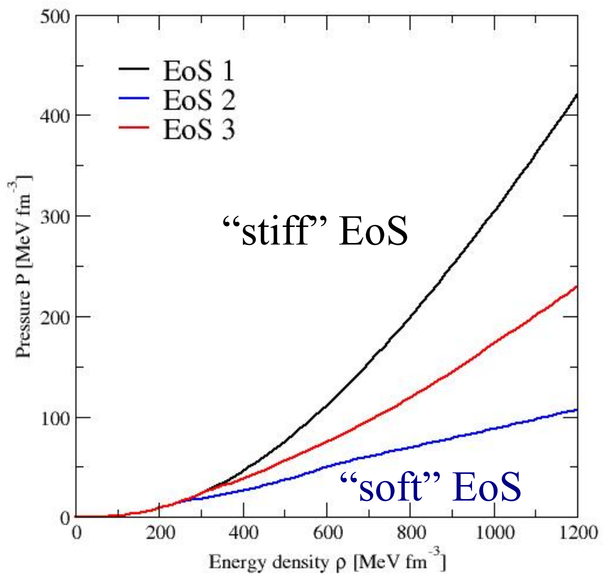
The role of the Equation of State

The only ingredient needed to solve the TOV equations is the (poorly known) EoS (i.e., p(ε)) of dense matter









General Features of a "realistic" neutron star matter EoS

Any "realistic" neutron star matter EoS must satisfy:

✧ Saturation Properties of Symmetric Matter

$$n_0 = 0.16 - 0.18\, fm^{-3}, \qquad \left(\frac{E}{A}\right)_0 = -16 \pm 1 \;\; MeV$$

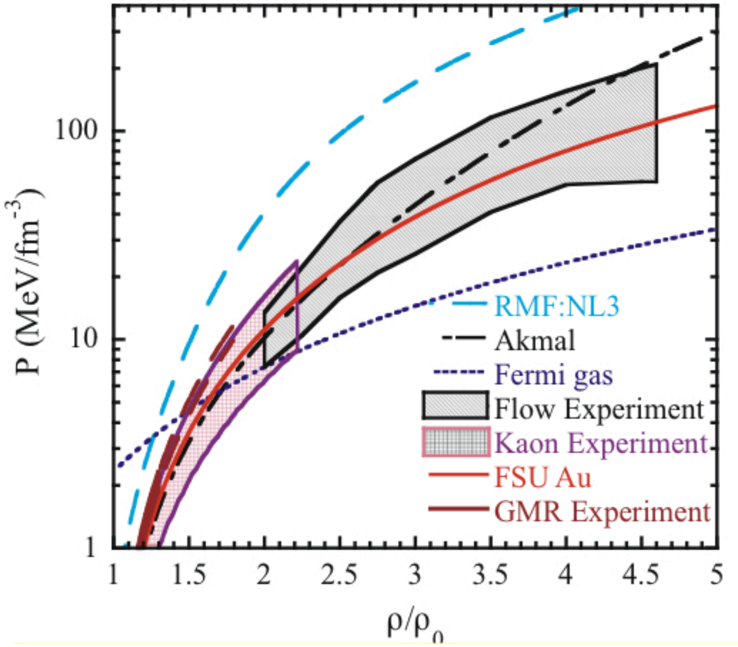
✧ Nuclear Symmetry Energy $E_{sym}(n_0) = 28 - 32 \;\; MeV$

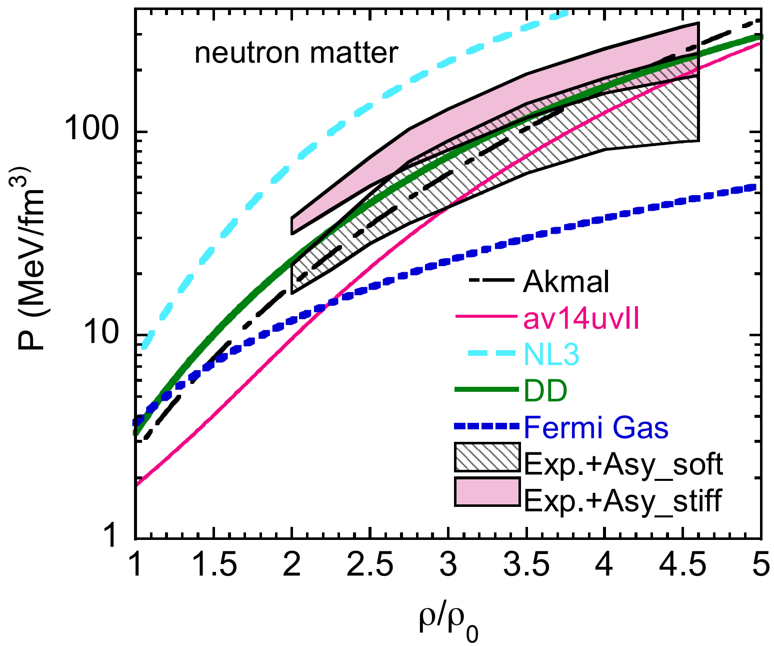
E_{sym} must be "well behaved" at high densities

✧ Nuclear Incompressibility $K_0 = 240 \pm 20 \;\; MeV$

✧ Causal Condition $c_S^2 = \dfrac{dP}{d\rho} \le c^2$

Constraints of the Nuclear EoS from HIC



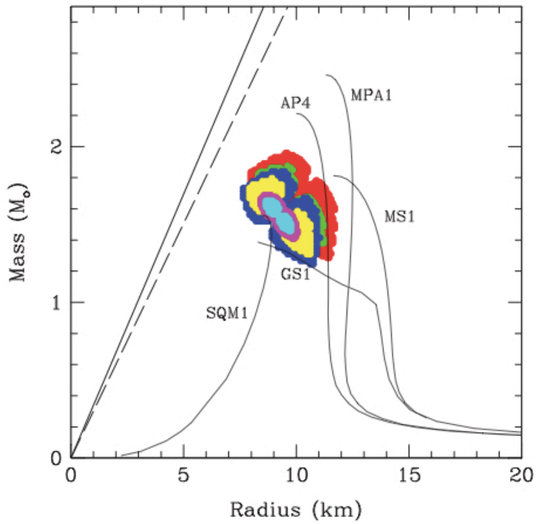


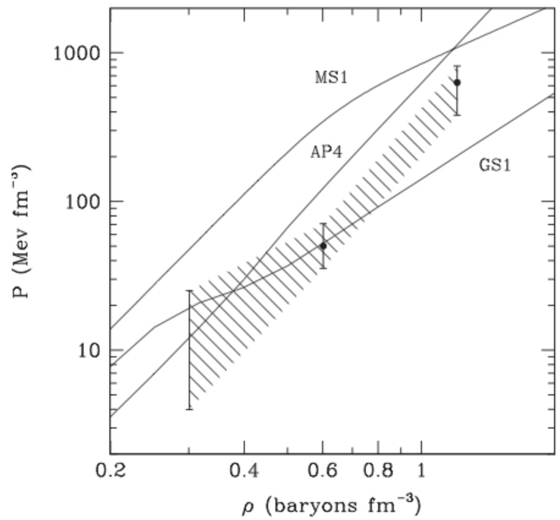
- Collective flow constraints confirms the softening of the EoS at high densities
- Constraints from kaon production are consistent with the flow constraints and bridge gap to GMR constraints
- Symmetry energy dominates the uncertainty in the neutron matter EoS

P. Danielewicz et al., Science 298, 1592 (2002)

Astrophysical determination of the Nuclear EoS

✧ Piecewise polytropic EoS above ρ_0 from mass-radius relation of 3 type-I X-ray bursts





❖ SLy below ρ_0

❖ Piecewise poytropic above ρ_0

$$\rho_{i-1} < \rho \leq \rho_i, \quad \varepsilon = \alpha_i \rho + \beta_i \rho^{\Gamma_i}, \quad P = \left(\Gamma_i - 1\right) \beta_i \rho^{\Gamma_i}$$

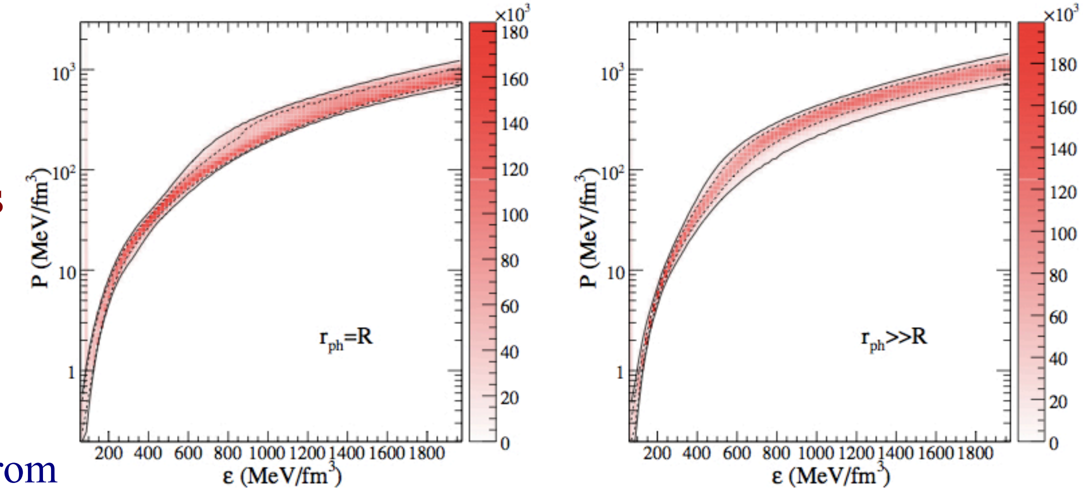
$\log P_0$ $(0.37\rho_{ns})$	$\log P_1$ $(1.85\rho_{ns})$	$\log P_2$ $(3.7\rho_{ns})$	$\log P_3$ $(7.4\rho_{ns})$
-0.64	[0.6–1.4]	$1.70^{+0.15}_{-0.15}$	$2.8^{+0.1}_{-0.2}$

F. Ozel & D. Psaltis, PRD 80, 103003 (2009)
F. Ozel, G. Baym & T. Guver, PRD 82, 101301(R) (2010)

Astrophysical determination of the Nuclear EoS

✧ Nuclear parameters determined in a Bayesian data analysis of:

- ❖ 3 type-I X-ray burst
- ❖ 3 transient low mass X-ray binaries
- ❖ Cooling of 1 isolated NS,
  RX J1856-3754



Parameters in the range expected from nuclear systematics & lab. experiments

Quantity	Lower Limit	Upper Limit
K (MeV)	180	280
K' (MeV)	−1000	−200
S_v (MeV)	28	38
γ	0.2	1.2
n_1 (fm^{-3})	0.2	1.5
n_2 (fm^{-3})	0.2	2.0
ε_1 (MeV fm^{-3})	150	600
ε_2 (MeV fm^{-3})	ε_1	1600

$$\varepsilon = n_B \left\{ m_B + B + \frac{K}{18}(u-1)^2 + \frac{K'}{162}(u-1)^3 + (1-2x)^2[S_k u^{2/3} + S_p u^{\gamma}] + \frac{3}{4}\hbar c x (3\pi^2 n_b x)^{1/3} \right\}$$

A. W. Steiner, J. M. Lattimer & E. F. Brown, ApJ 722, 33 (2010)

Theoretical Approaches to the Nuclear EoS

Phenomenological approaches

Based on effective density-dependent interactions with parameters adjusted to reproduce nuclear observables and compact star properties

- ❖ Liquid drop type: BPS, BBP, LS, OFN
- ❖ Thomas-Fermi: Shen
- ❖ HF: NV, Sk, BSk, PAL, RMF, RHF, QMC
- ❖ Statistical models: HWN, RG, HS

I apologize for all those approaches I have missed

Microscopic ab-initio approaches

Based on two- & three-body realistic interactions. The EoS is obtained by "solving" the complicated many-body problem

- ❖ Variational: APS, CBF, FHNC, LOVC
- ❖ Monte-Carlo: VMC, DMC, GFMC, AFDMC
- ❖ Diagrammatic: BBG (BHF), SCGF
- ❖ RG methods: $V_{low\ k}$ & SRG from χEFT potentials
- ❖ DBHF

Upper limit of the Maximum Mass

M_{max} depends mainly on the behaviour of EoS, P(ε), at high densities. Any realistic EoS must satisfy two conditions:

- Causality: $\frac{dP}{d\rho} \leq c^2$
- Stability: $\frac{dP}{d\rho} > 0$

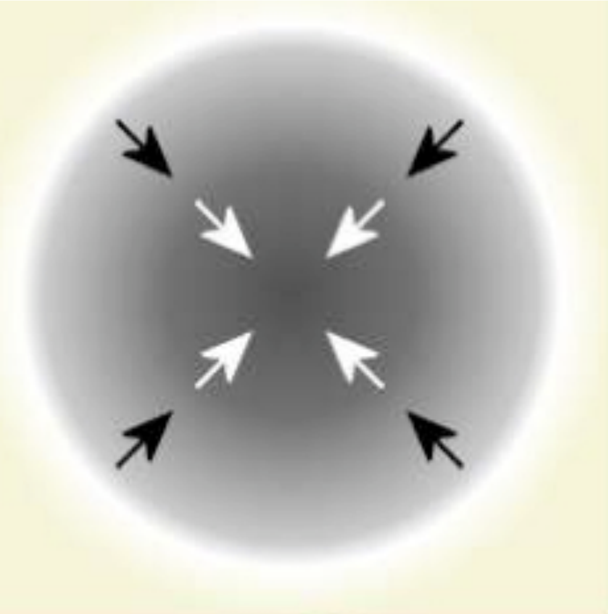
If the EoS is known up to ρ_r, these conditions imply:

$$M_{\max} \leq 3M_{\odot}\left(\frac{5x10^{14}\, g/cm^3}{\rho_r}\right)^{1/2}$$

If rotation is taken into account M_{max} can increase up to 20%:

$$M_{\max} \leq 3.89M_{\odot}\left(\frac{5x10^{14}\, g/cm^3}{\rho_r}\right)^{1/2}$$

Estimation of Neutron Star Mass & Radius



Imagine that a neutron star is:

- ✓ a sphere of uniform density
- ✓ made only of neutrons
- ✓ in addition to the nuclear force neutrons feel also the gravity

Idea: use Bethe-Weizsäker semi-empirical mass formula including the gravitational force

$$B(Z,A) = a_v A - a_s A^{2/3} - a_{coul}\frac{Z^2}{A^{1/3}} - a_{sim}\frac{(Z-N)^2}{A} + \delta a_p A^{-1/2}$$

Only Neutrons (Z=0) + Gravitational Energy (sphere with M=Nm$_n$ & R)

$$B(Z=0, A=N) = a_v N - a_s N^{2/3} - a_{sim} N + \delta a_p N^{-1/2} + \tfrac{3}{5}\frac{G(Nm_n)^2}{R}$$

Since N > N$^{2/3}$ & N$^{-1/2}$

$$B(Z=0, A=N) \approx \left(a_v - a_{sim}\right) N + \frac{3}{5}\frac{Gm_n^2}{r_0} N^{5/3}$$

$$R = r_0 N^{1/3} = 1.15 \times 10^{-15} N^{1/3}\, m$$

The minimum number of neutrons needed to bound gravitationally is obtained imposing

B > 0

The B > 0 tell us that:

$$(a_v - a_{sim})N + \frac{3}{5}\frac{Gm_n^2}{r_0}N^{5/3} > 0 \Rightarrow N > \left(\frac{5}{3}\frac{(a_{sim} - a_v)r_0}{Gm_n^2}\right)^{3/2}$$

Using the values:

$$a_v = 16\,MeV, \quad a_{sim} = 30\,MeV, \quad G = 6.707\times 10^{-39}\hbar c\left(\frac{c^4}{GeV^2}\right), \; m_n = 0.939\frac{GeV}{c^2}$$

We finally arrive to:

$$N \sim 10^{56} - 10^{57} \qquad M \sim 1 \; M_\odot \qquad R \sim 10\,km$$

Which gives an average density of:

$$\rho \sim 10^{14} - 10^{15}\, g/cm^3$$

$$N \sim 10^{56} - 10^{57}$$

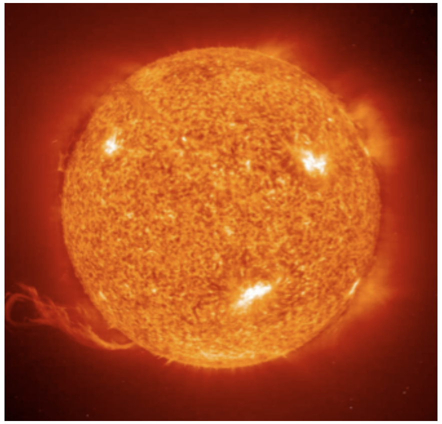
$$M \sim 1 \ \ M_{\odot}$$

A neutron star
is a kind of GIANT
ATOMIC NUCLEUS
in which particles are
gravitationally bound

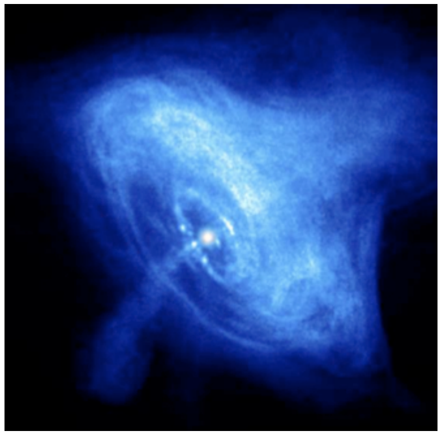
$$R \sim 10\, km$$

$$\rho \sim 10^{14} - 10^{15}\ g/cm^3$$

A neutron star
has a mass similar to that
of the Sun, but with a
radius about
70.000 smaller !!!



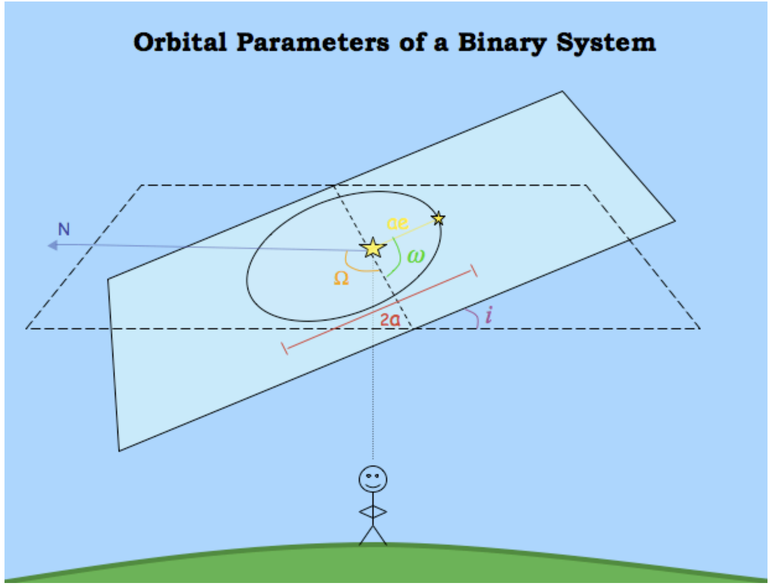
Radius: ~ 700.000 km
Mass: 1.989×10^{30} kg



Radius ~ 10 km
Mass ~ 1.989×10^{30} kg

How to Measure Neutron Star Masses

Use Doppler variations in spin period to measure orbital velocity changes along the line-of-sight



- 5 Keplerian parameters can normally be determined:

  P, a sin i, ε, T_0 & ω

- 3 unknowns: M_1, M_2, i

Kepler's 3rd law

$$\frac{G(M_1+M_2)}{a^3}=\left(\frac{2\pi}{P}\right)^2 \rightarrow f(M_1,M_2,i)\equiv\frac{(M_2\sin i)^3}{(M_1+M_2)^2}=\frac{Pv^3}{2\pi G}$$

mass function

In few cases small deviations from Keplerian orbit due to GR effects can be detected

Measure of at least 2 post-Keplerian parameters ➔ High precision NS mass determination

$$\dot{\omega} = 3T_{\otimes}^{2/3}\left(\frac{P_b}{2\pi}\right)^{-5/3}\frac{1}{1-\varepsilon}\left(M_p + M_c\right)^{2/3}$$ → Periastron precession

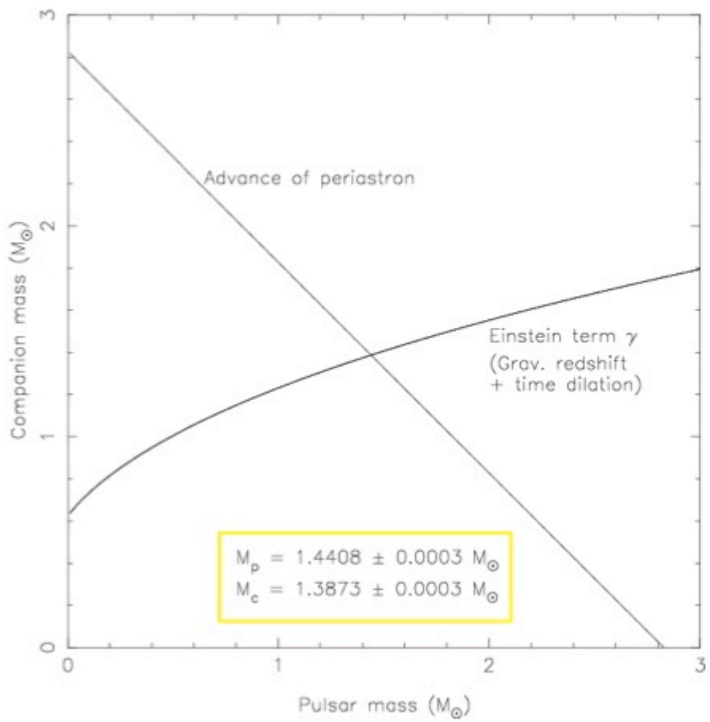
$$\gamma = T_{\otimes}^{2/3}\left(\frac{P_b}{2\pi}\right)^{1/3}\varepsilon\frac{M_c\left(M_p + 2M_c\right)}{\left(M_p + M_c\right)^{4/3}}$$ → Time dilation and grav. redshift

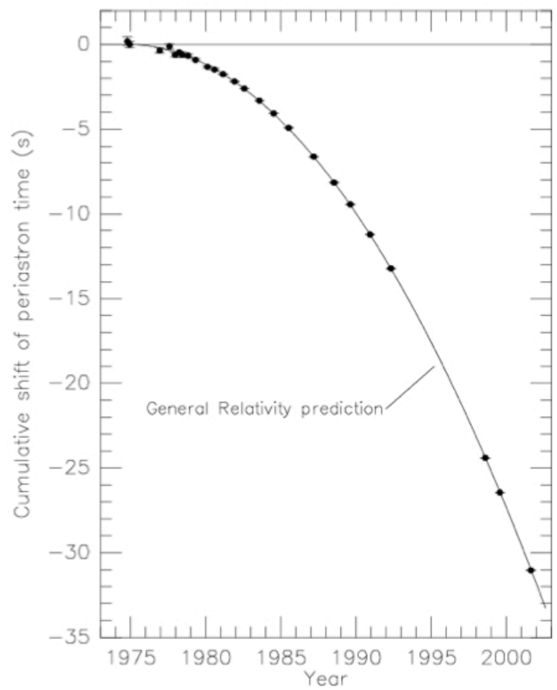
$$r = T_{\otimes}M_c$$ → Shapiro delay “range”

$$s = \sin i = T_{\otimes}^{-1/3}\left(\frac{P_b}{2\pi}\right)^{-2/3}x\frac{\left(M_p + M_c\right)^{2/3}}{M_c}$$ → Shapiro delay “shape”

$$\dot{P}_b = -\frac{192\pi}{5}T_{\otimes}^{5/3}\left(\frac{P_b}{2\pi}\right)^{-5/3}f(\varepsilon)\frac{M_pM_c}{\left(M_p + M_c\right)^{1/3}}$$ → Orbit decay due to GW emission

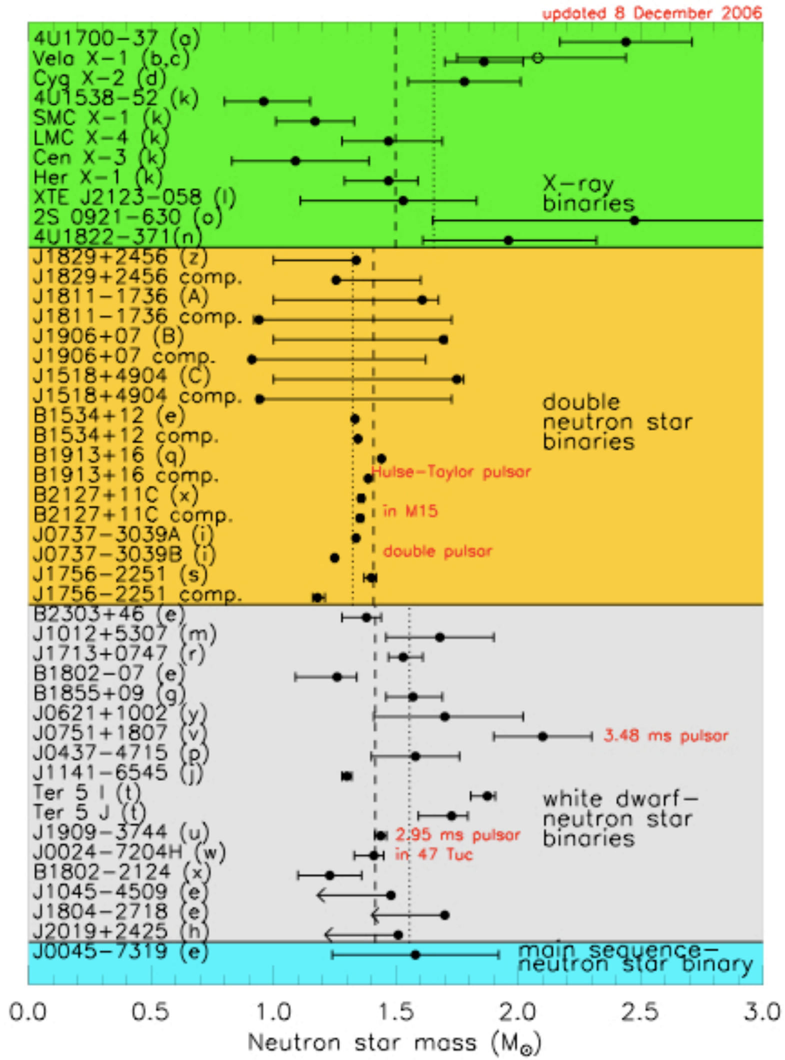
An example: the mass of the Hulse-Taylor pulsar (PSR J1913+16)



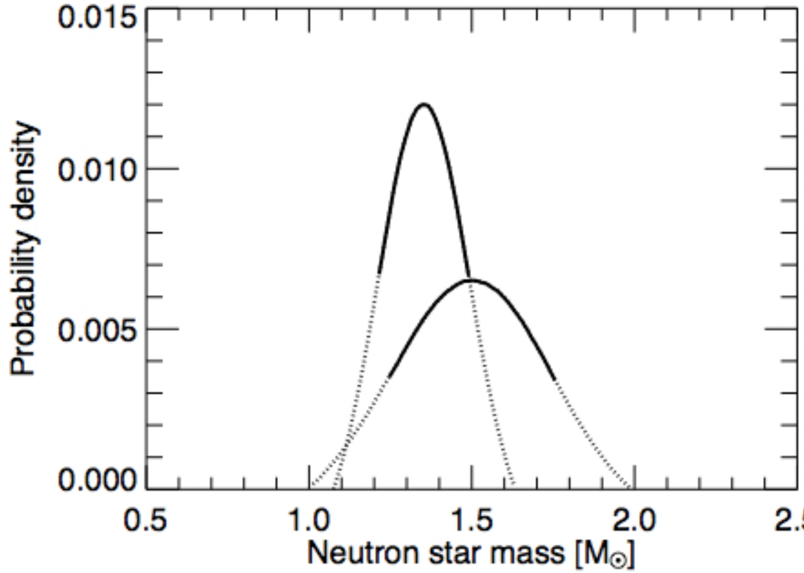


Parameter	Value
Orbital period P_b (d)	0.322997462727(5)
Projected semi-major axis x (s)	2.341774(1)
Eccentricity e	0.6171338(4)
Longitude of periastron ω (deg)	226.57518(4)
Epoch of periastron T_0 (MJD)	46443.99588317(3)
Advance of periastron $\dot{\omega}$ (deg yr^{-1})	4.226607(7)
Gravitational redshift γ (ms)	4.294(1)
Orbital period derivative $(\dot{P}_b)^{obs}$ (10^{-12})	−2.4211(14)

Measured Neutron Star Masses (up to ~ 2006-2008)



(Lattimer & Prakash 2007)



up to ~ 2006-2008 any valid EoS should predict

$$M_{max}\left[EoS\right] > 1.4 - 1.5 M_{\odot}$$

N.B. I will comment on more recent measurements latter when talking about the “hyperon problem”

Limits on the Neutron Star Radius

The radius of a neutron star with mass M cannot be arbitrarily small

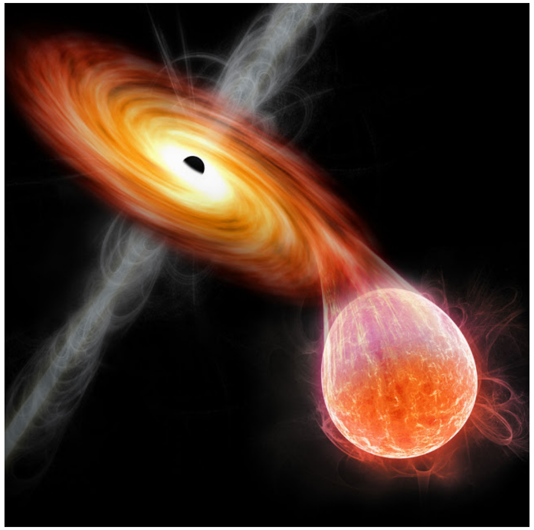
General Relativity: a Neutron Star is not a Black Hole	$R > \frac{2GM}{c^2}$
Finite Pressure: Neutron Star matter cannot be arbitrarily compressed	$R > \frac{9}{4}\frac{GM}{c^2}$
Causality: speed of sound must be smaller than c	$R > 2.9\frac{GM}{c^2}$

How to measure Neutron Star Radii

Radii are very difficult to measure because NS:

- ✧ are very small (~ 10 km)
- ✧ are far from us (e.g., the closest NS, RX J1856.5-3754, is at ~ 400 ly)

A possible way to measure it is to use the thermal emission of low mass X-ray binaries:



NS radius can be obtained from

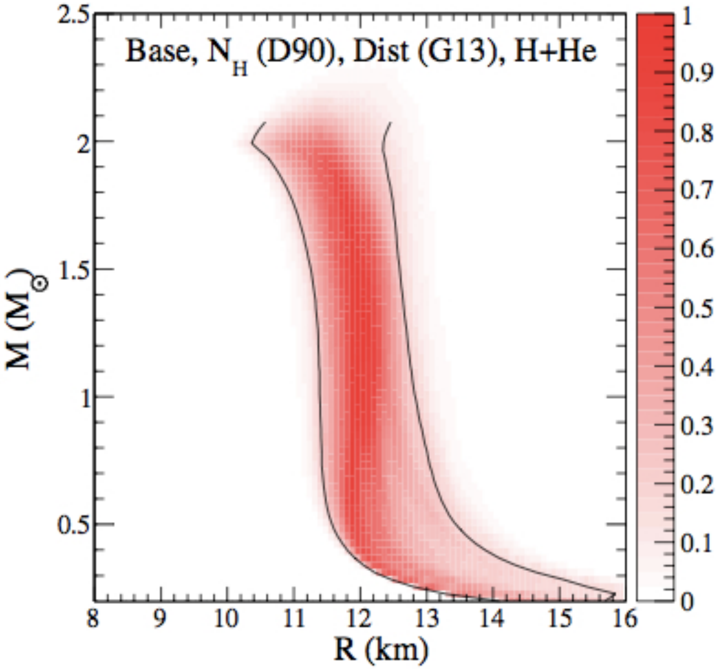
- ✧ Flux measurement +Stefan-Boltzmann's law
- ✧ Temperature (Black body fit+atmosphere model)
- ✧ Distance estimation (difficult)
- ✧ Gravitational redshift z (detection of absorption lines)

$$R_\infty = \sqrt{\frac{FD^2}{\sigma_{SB} T^4}} \rightarrow R_{NS} = \frac{R_\infty}{1+z} = R_\infty \sqrt{1 - \frac{2GM}{R_{NS} c^2}}$$

Recent Estimations of Neutron Star Radii

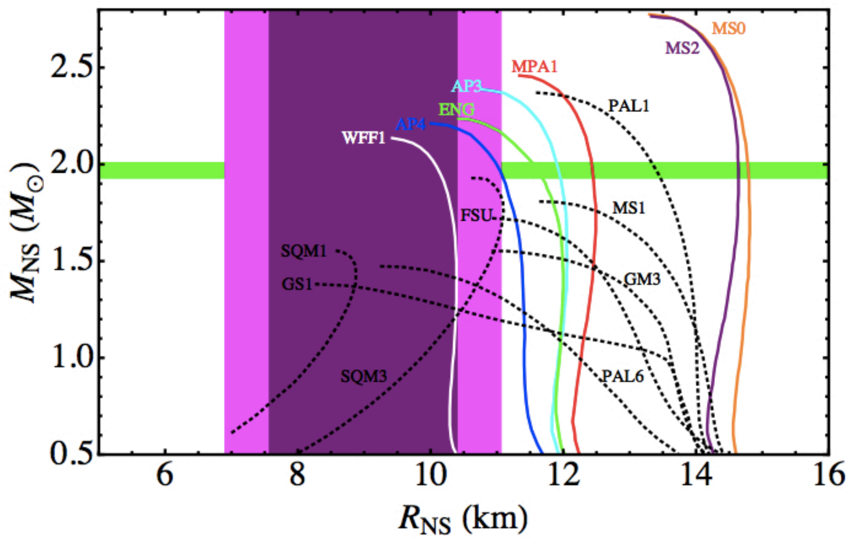
The recent analysis of the thermal spectrum from 5 quiescent LMXB in globular clusters is still controversial

Steiner et al. (2013, 2014)



$$R = 12.0 \pm 1.4 km$$

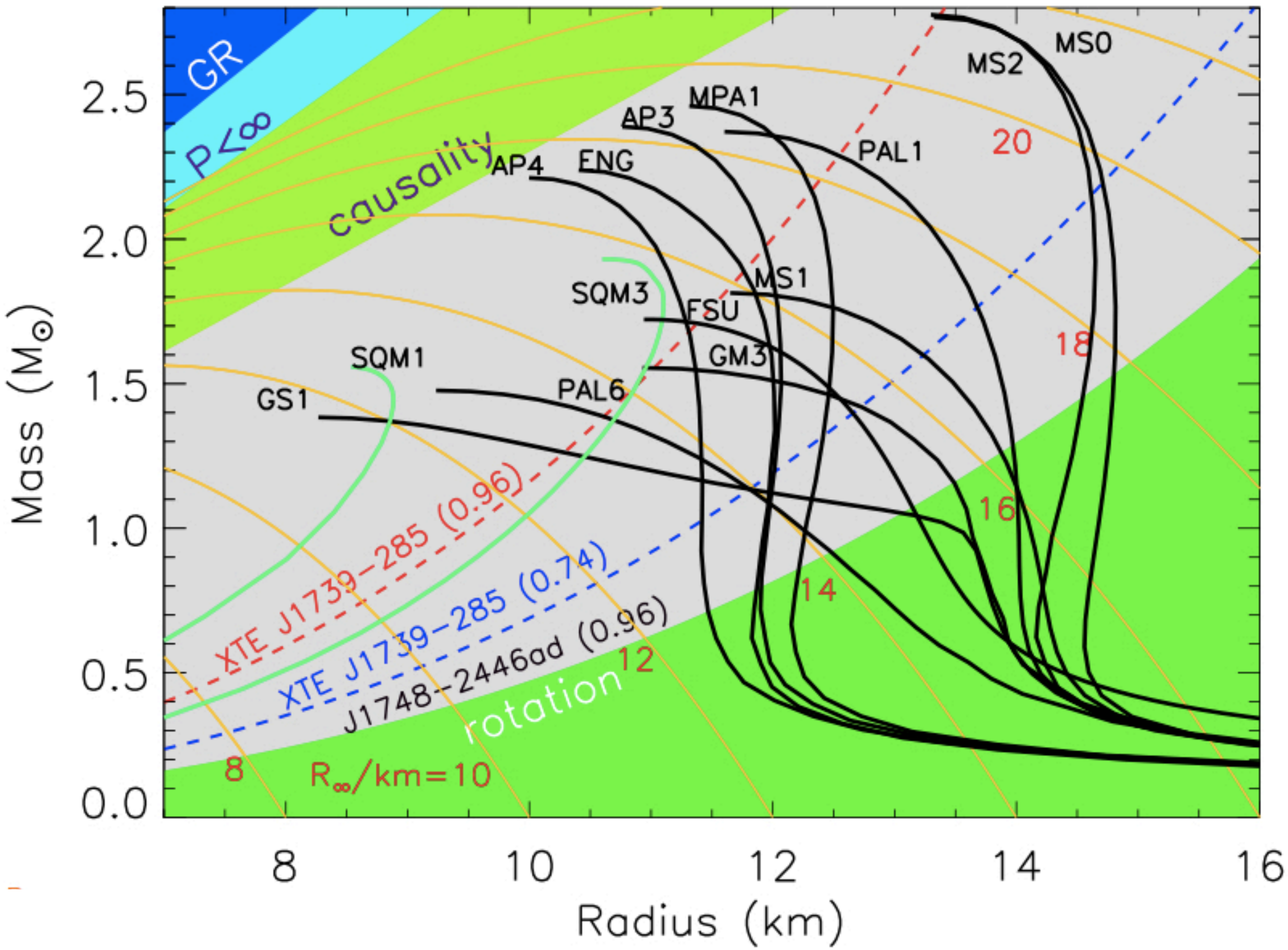
Guillot et al. (2013, 2014)



$R = 9.1^{+1.3}_{-1.5} km$ 2013 analysis

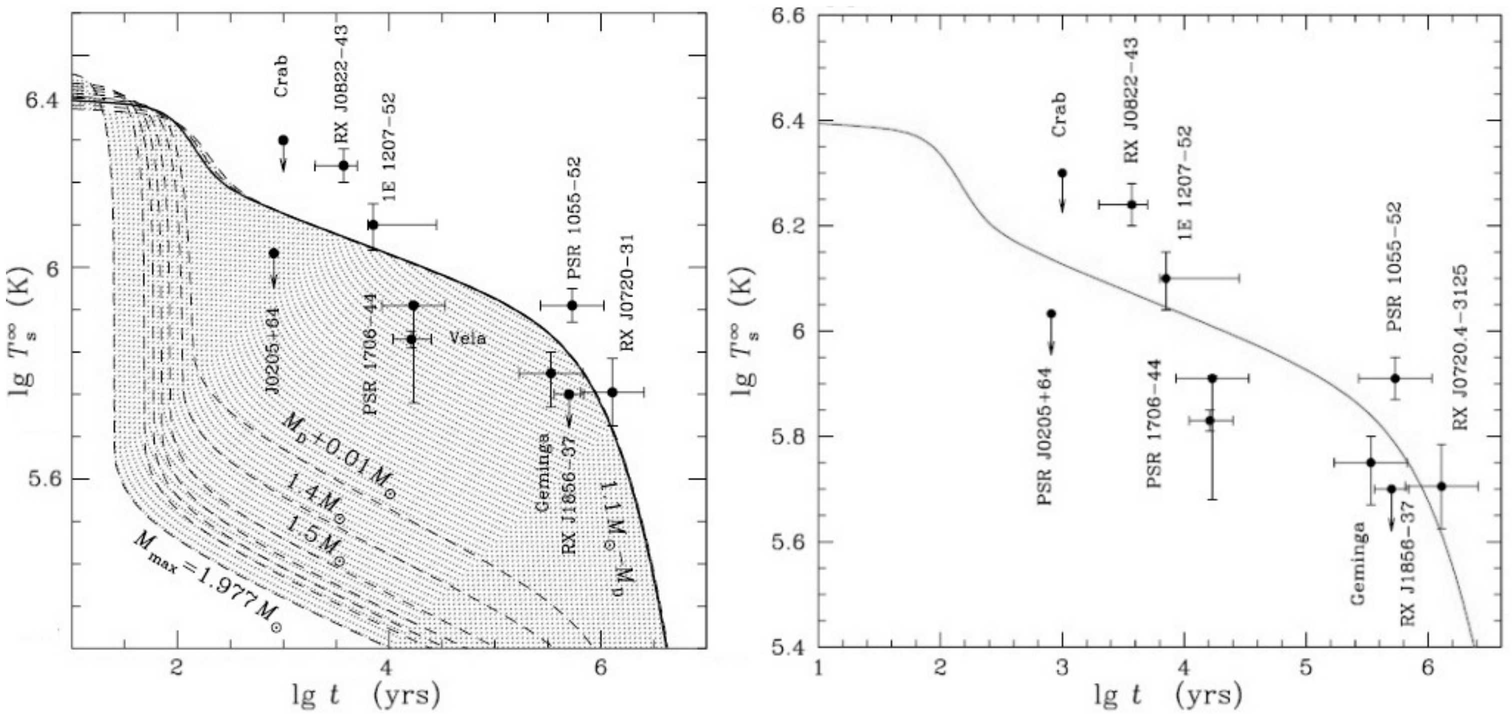
$R = 9.4 \pm 1.2 km$ 2014 analysis

Limits of the Mass & Radius of a Neutron Star



Thermal Evolution of Neutron Stars

Information, complementary to that from mass & radius, can be also obtained from the measurement of the temperature (luminosity) of neutron stars

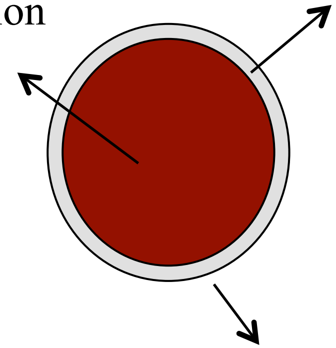


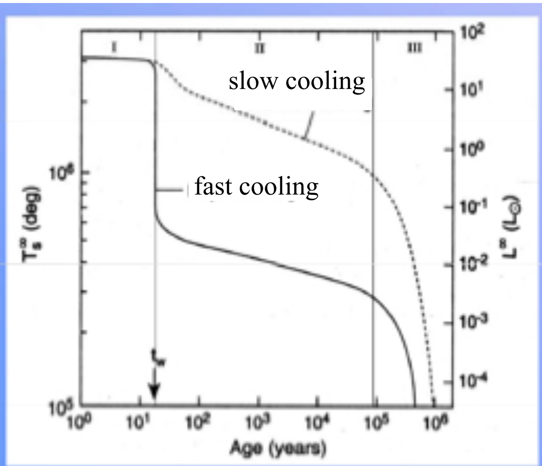
D. G. Yakovlev & C. J. Pethick, A&A 42, 169 (2004)

Neutron Star Cooling in a Nutshell

Two cooling regimes

Slow	Fast
Low NS mass	High NS mass



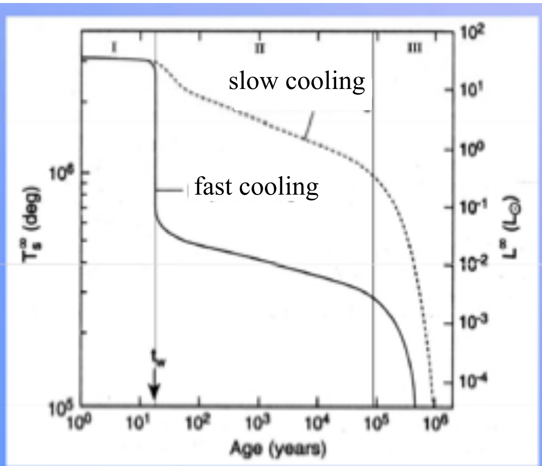


I. Core relaxation epoch
II. Neutrino cooling epoch
III. Photon cooling epoch

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

- ✓ C_v: specific heat
- ✓ L_γ: photon luminosity
- ✓ L_ν: neutrino luminosity
- ✓ H: "heating"

Neutron Star Cooling & Symmetry Energy



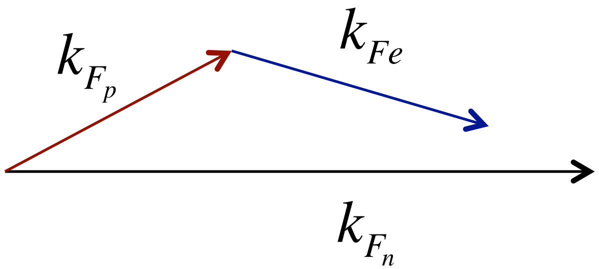
I. Core relaxation epoch
II. Neutrino cooling epoch
III. Photon cooling epoch

- Fast: e.g., Direct URCA

$$n \rightarrow p + l + \bar{\nu}_l$$
$$l + p \rightarrow n + \nu_l$$

- Slow: e.g., Modified URCA

$$N + n \rightarrow N + p + l + \bar{\nu}_l$$
$$N + l + p \rightarrow N + n + \nu_l$$



Direct URCA cannot occur unless x_p> 11%-15%

Larger Symmetry Energy ➔ Larger x_p ➔ Earlier onset of Direct URCA

$$\mu_n - \mu_p = 4(1-2x_p)S_2(\rho) = \mu_l - \mu_{\upsilon_l} \Rightarrow \frac{x_p}{1-2x_p} = \frac{4S_2(\rho)}{\hbar c(3\pi^2\rho)^{1/3}}$$

Neutrino Emission

Name	Process	Emissivity (erg cm^{-3} s^{-1})	
Modified Urca cycle (neutron branch)	$n + n \to n + p + e^- + \bar{\nu}_e$ $n + p + e^- \to n + n + \nu_e$	$\sim 2{\times}10^{21}\, R\, T_9^8$	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	$n + n \to [nn] + \nu + \bar{\nu}$ $p + p \to [pp] + \nu + \bar{\nu}$	$\sim 5{\times}10^{21}\, R\, T_9^7$ $\sim 5{\times}10^{19}\, R\, T_9^7$	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

Anything beyond just neutrons & protons results in an enhancement of the neutrino emission

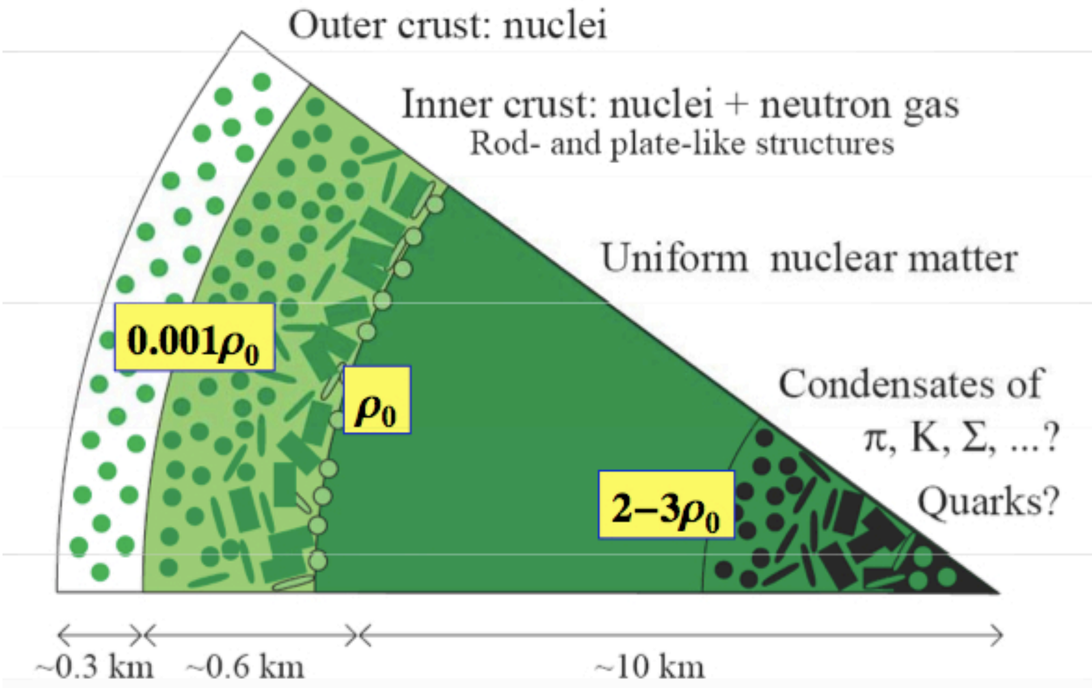
Anatomy of a Neutron Star

Equilibrium composition determined by

✓ Charge neutrality

$$\sum_i q_i \rho_i = 0$$

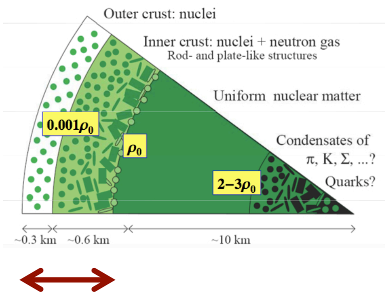
✓ Equilibrium with respect to weak interacting processes

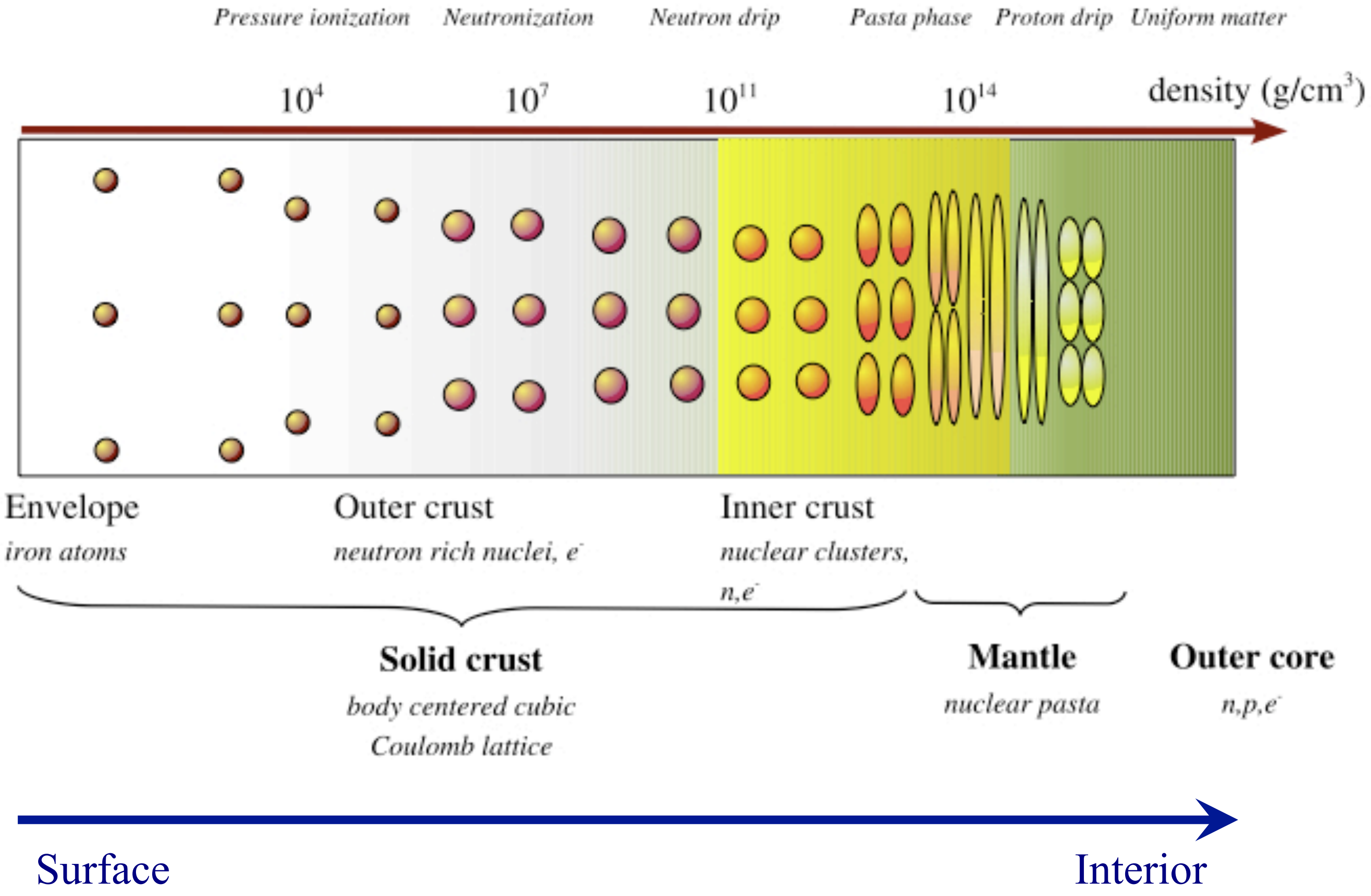


$$b_1 \rightarrow b_2 + l + \bar{\nu}_l$$
$$b_2 + l \rightarrow b_1 + \nu_l$$

$$\longrightarrow \quad \mu_i = b_i \mu_n - q_i \left(\mu_e - \mu_{\nu_e} \right), \quad \mu_i = \frac{\partial \varepsilon}{\partial \rho_i}$$

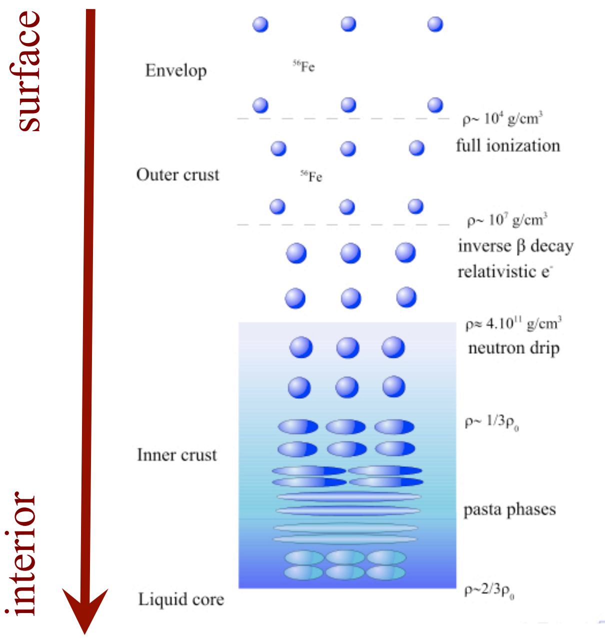
Crust of a Neutron Star

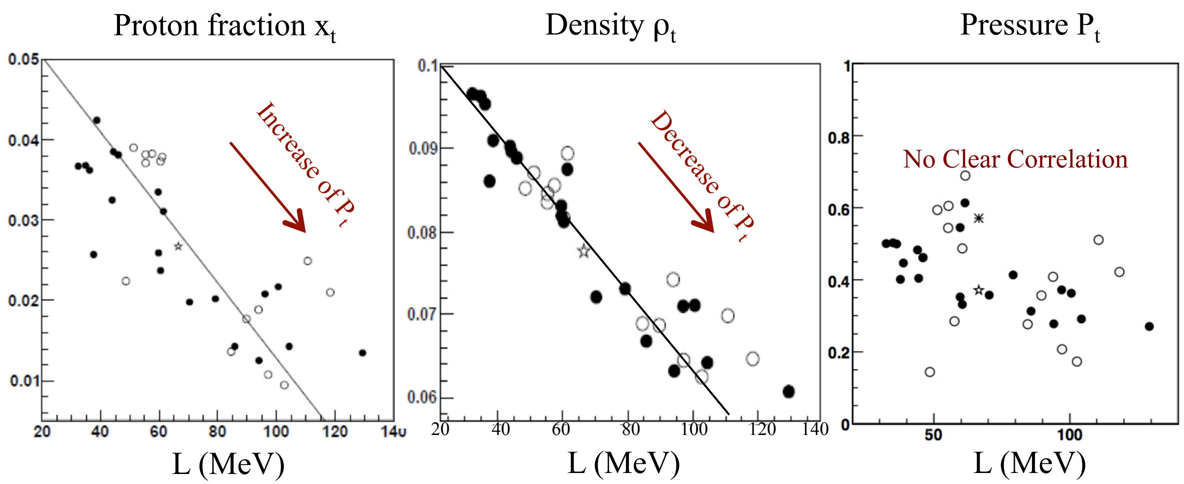


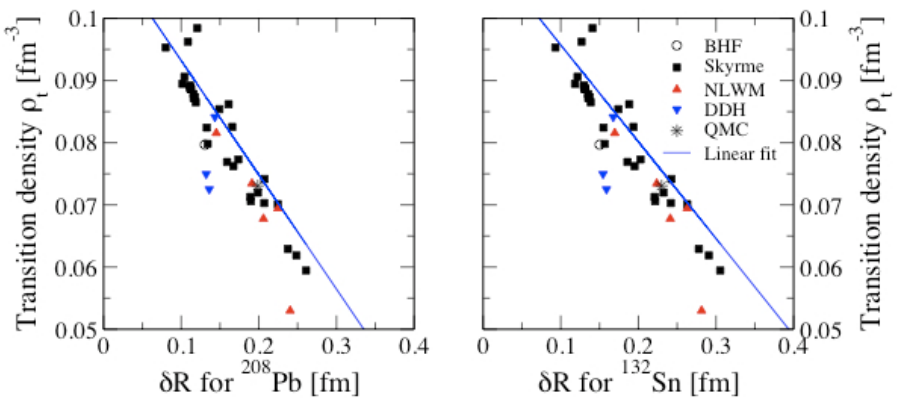


Crust-core Transition & Symmetry Energy

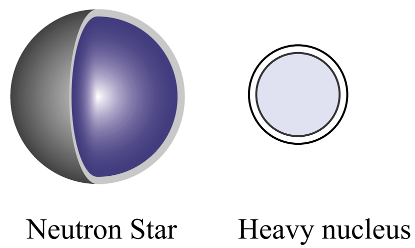
$$P(\rho,\beta)=\frac{\rho^2}{3\rho_0}\left(L\beta^2+\left(K_0+K_{sym}\beta^2\right)\frac{\rho-\rho_0}{3\rho_0}+\cdots\right)$$







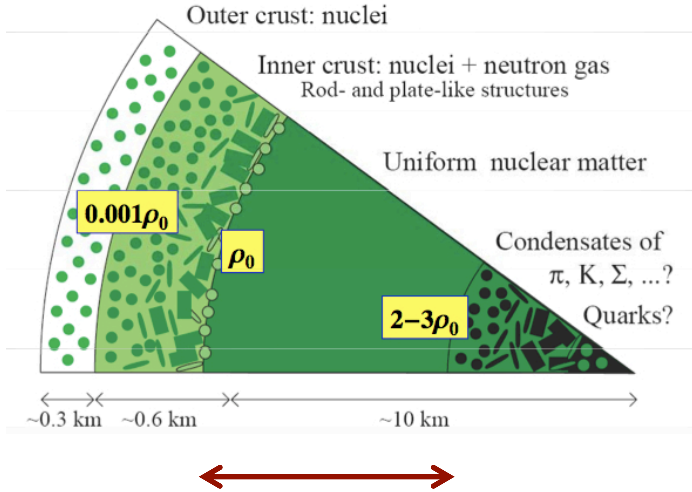
Inverse correlation between dR and ρ_t
(Horowitz & Piekarewicz)

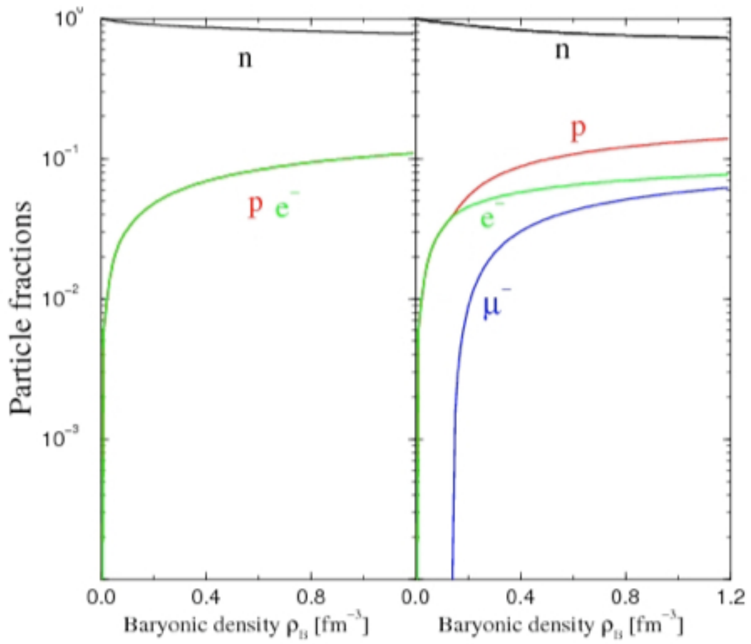


Crust & Neutron Skin made out of neutron rich matter at similar densities ➔ Both governed by EoS at $\rho < r_0$ (particularly by $S_2(\rho)$ & its derivatives)

External Core of a Neutron Star

The external core of a neutron star is mainly a fluid of neutron-rich matter in equilibrium with respect to weak interaction processes (β-stable matter)





$$n \rightarrow p + e^- + \bar{\nu}_{e^-}$$

$$p + e^- \rightarrow n + \nu_{e^-}$$

$$\Rightarrow \mu_p = \mu_n - \mu_{e^-} + \mu_{\nu_{e^-}}$$

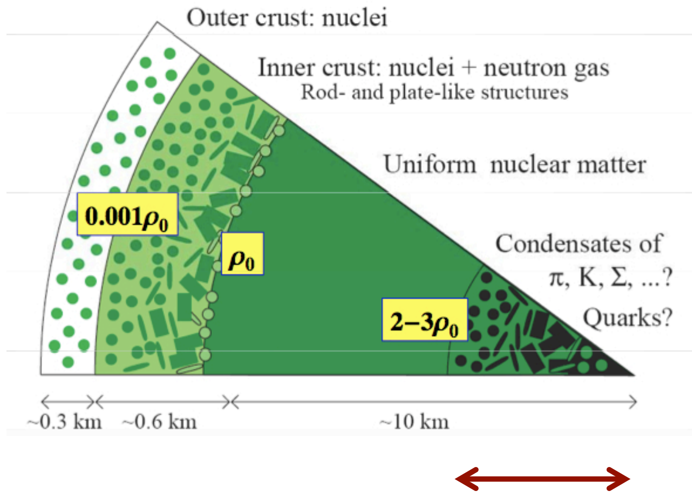
Internal Core of a Neutron Star

Since:

✧ The value of the central density is very high: $\rho_c \sim (4\text{-}8)\rho_0$

$(\rho_0 = 0.17\ \mathrm{fm}^{-3} = 2.8 \times 10^{14}\ \mathrm{g/cm}^3)$

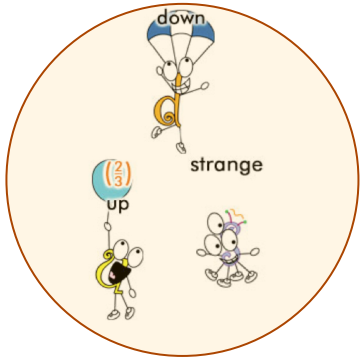
✧ Nucleon chemical potential increases rapidly with the density ρ



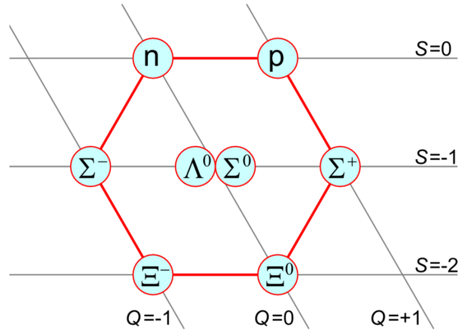
The presence of exotic degrees of freedom is expected in the Neutron Star interior (π, K^- condensates, hyperons, quarks,…)

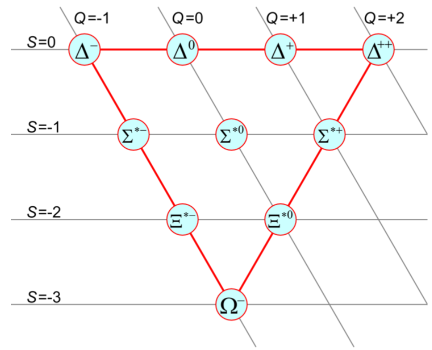
What is a hyperon ?

- A hyperon is a baryon made of one , two or three strange quarks



Hyperon	Quarks	$I(J^P)$	Mass (MeV)
Λ	uds	$0(1/2^+)$	1115
Σ^+	uus	$1(1/2^+)$	1189
Σ^0	uds	$1(1/2^+)$	1193
Σ^-	dds	$1(1/2^+)$	1197
Ξ^0	uss	$1/2(1/2^+)$	1315
Ξ^-	dss	$1/2(1/2^+)$	1321
Ω^-	sss	$0(3/2^+)$	1672





Hyperons in Neutron Stars

Hyperons in NS considered by many authors since the pioneering work of Ambartsumyan & Saakyan (1960)

Phenomenological approaches

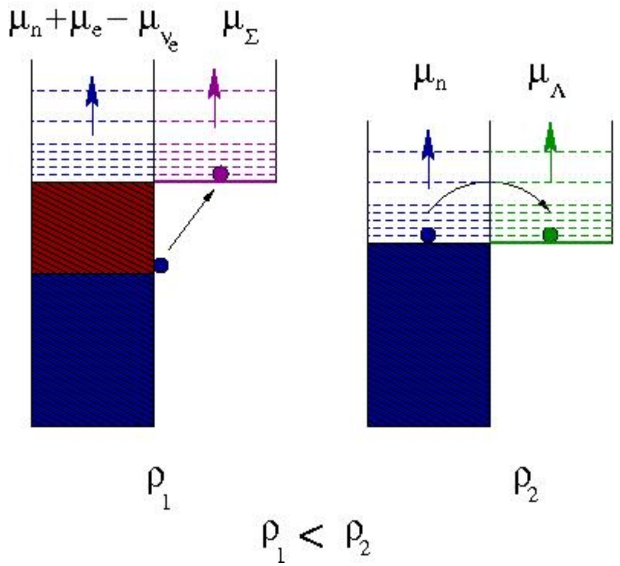
- ✧ Relativistic Mean Field Models: Glendenning 1985; Knorren et al. 1995; Shaffner-Bielich & Mishustin 1996, Bonano & Sedrakian 2012, …
- ✧ Non-realtivistic potential model: Balberg & Gal 1997
- ✧ Quark-meson coupling model: Pal et al. 1999, …
- ✧ Chiral Effective Lagrangians: Hanauske et al., 2000
- ✧ Density dependent hadron field models: Hofmann, Keil & Lenske 2001

Microscopic approaches

- ✧ Brueckner-Hartree-Fock theory: Baldo et al. 2000; I. V. et al. 2000, Schulze et al. 2006, I.V. et al. 2011, Burgio et al. 2011, Schulze & Rijken 2011
- ✧ DBHF: Sammarruca (2009), Katayama & Saito (2014)
- ✧ $V_{low\ k}$: Djapo, Schaefer & Wambach, 2010
- ✧ Quantum Monte Carlo: Lonardoni et al., (2014)

Sorry if I missed somebody

Hyperons are expected to appear in the core of neutron stars at $\rho \sim (2\text{-}3)\rho_0$ when μ_N is large enough to make the conversion of N into Y energetically favorable.



$$n + n \rightarrow n + \Lambda$$

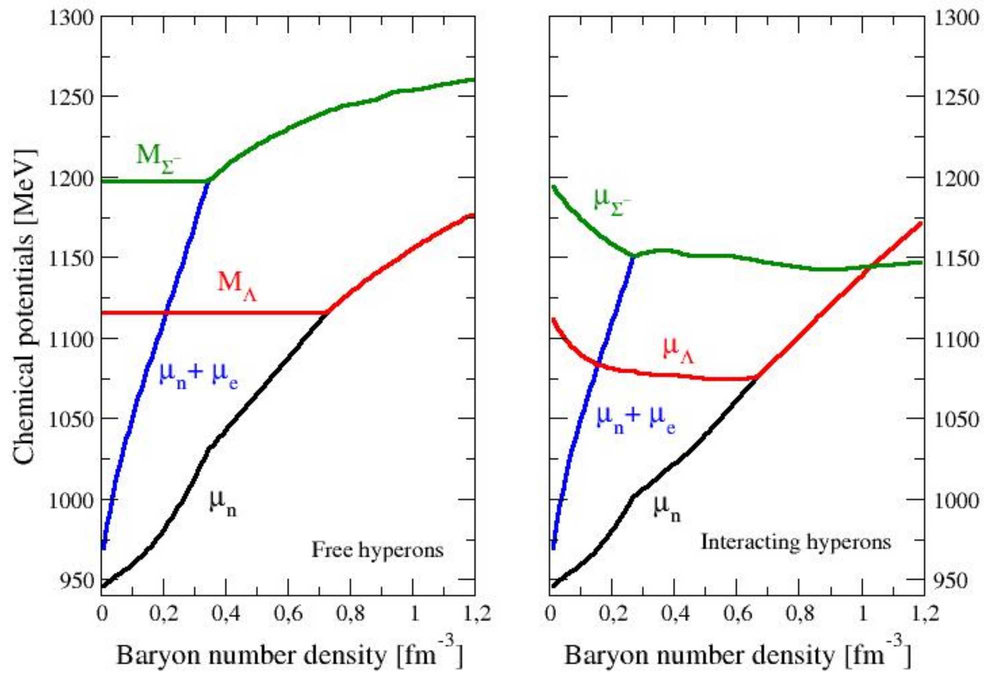
$$p + e^- \rightarrow \Lambda + \nu_{e^-}$$

$$n + n \rightarrow p + \Sigma^-$$

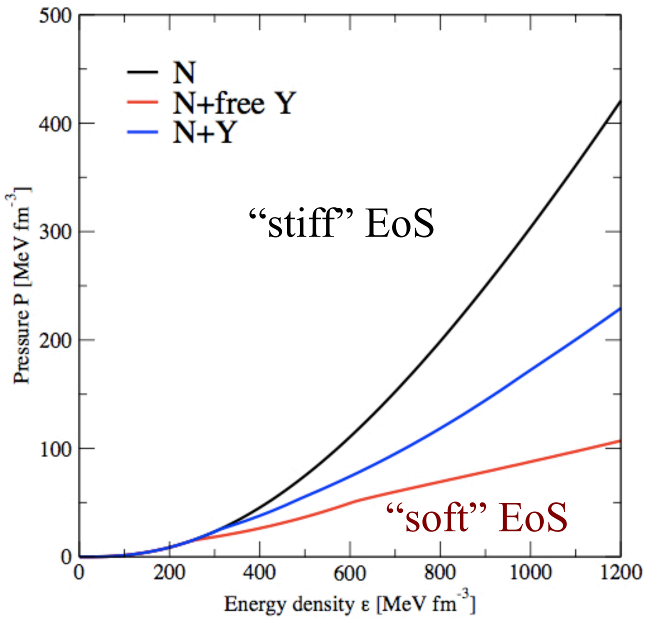
$$n + e^- \rightarrow \Sigma^- + \nu_{e^-}$$

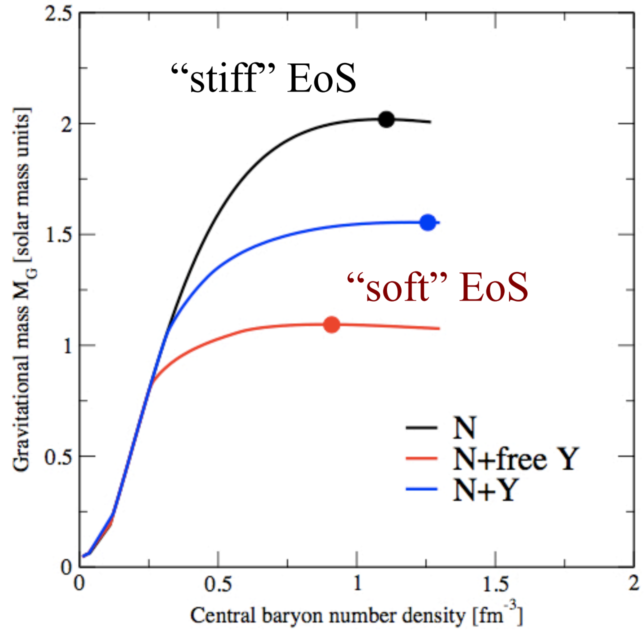
$$\mu_{\Sigma^-} = \mu_n + \mu_{e^-} - \mu_{\nu_{e^-}}$$

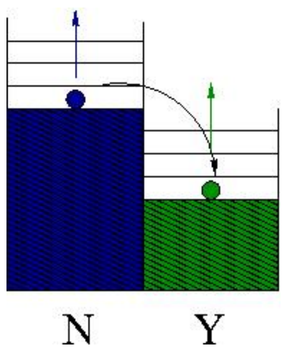
$$\mu_\Lambda = \mu_n$$



Effect of Hyperons in the EoS and Mass of Neutron Stars



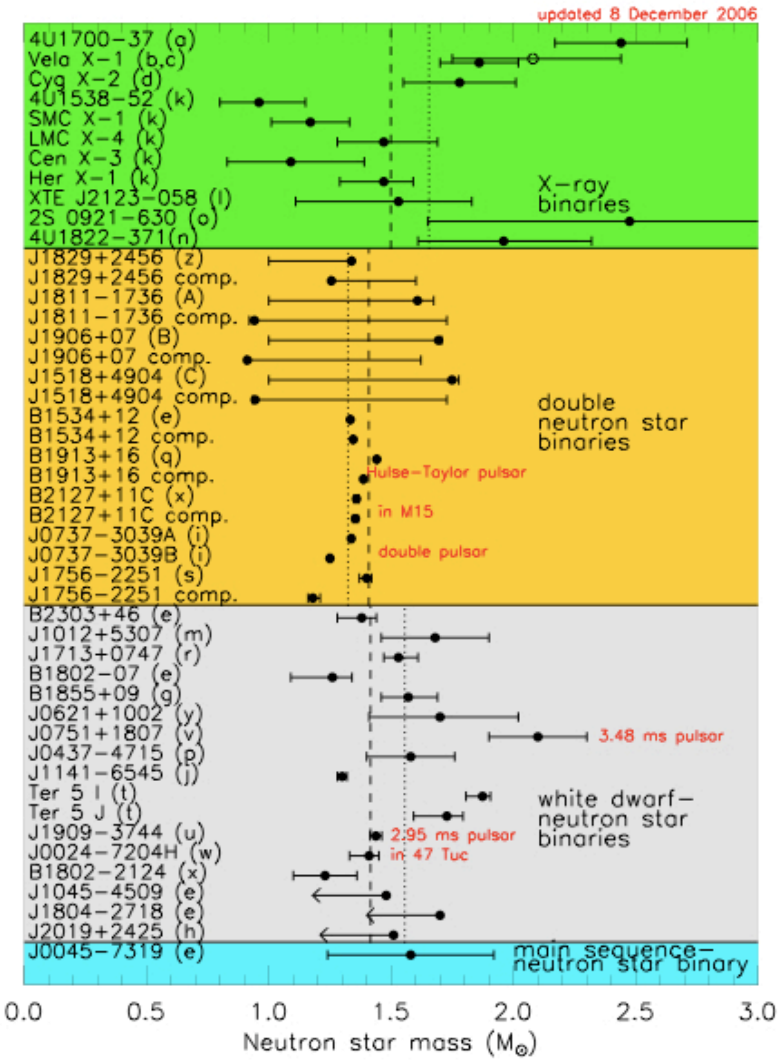




Relieve of Fermi pressure due to the appearance of hyperons ➔

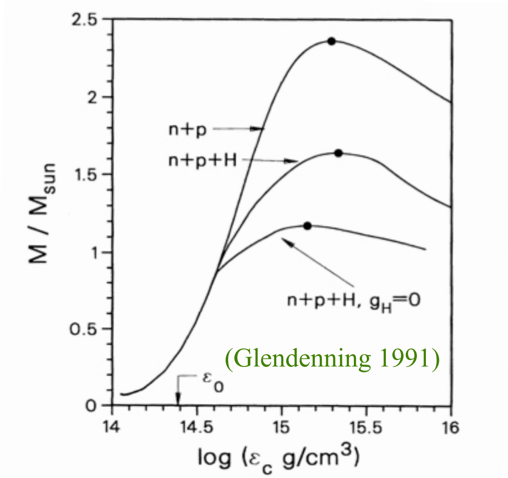
EoS softer ➔ reduction of the mass

Hyperons in NS
(up to ~ 2006-2008)



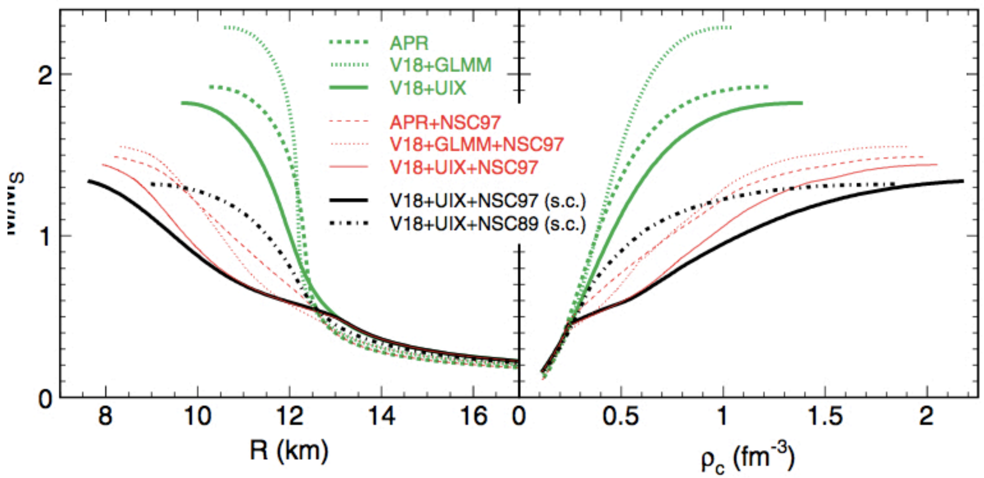
(Lattimer & Prakash 2007)

Phenomenological:
M_{max} compatible with 1.4-1.5 $M_\odot$



(Glendenning 1991)

Microscopic : M_{max} < 1.4-1.5 $M_\odot$



(Schulze, Polls, Ramos & IV 2006)

Recent measurements of high masses → life of hyperons (and theoretitians) more difficult

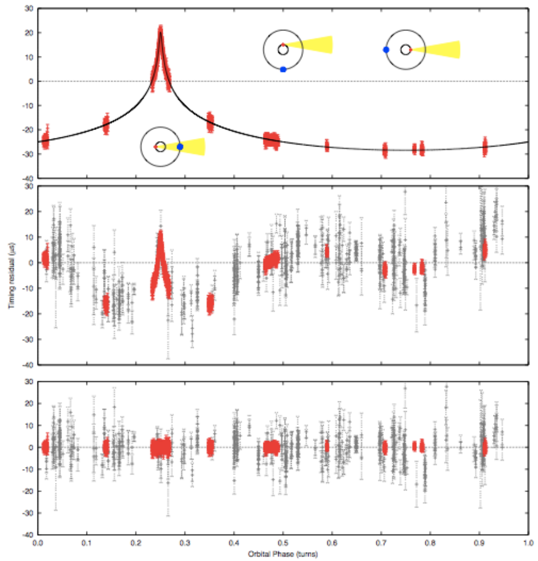
- PSR J164-2230 (Demorest et al. 2010)

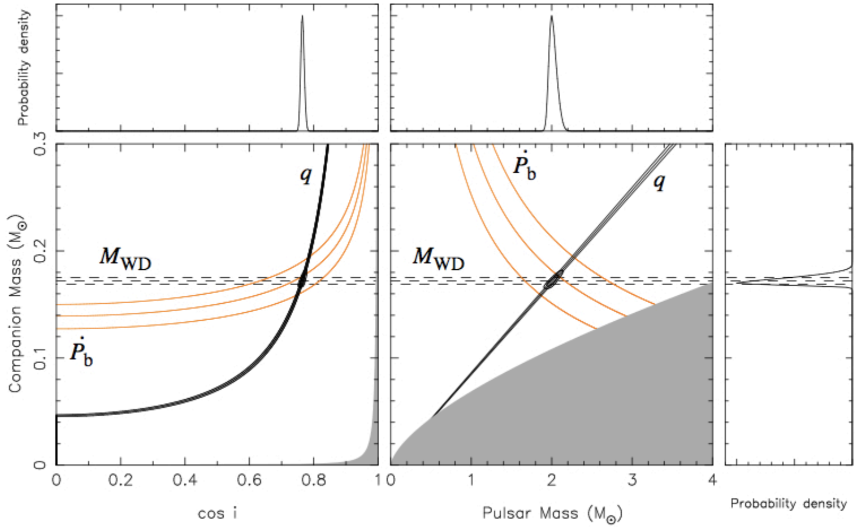
✓ binary system (P=8.68 d)

✓ low eccentricity ($\varepsilon=1.3 \times 10^{-6}$)

✓ companion mass: $\sim 0.5 M_\odot$

✓ pulsar mass: $M = 1.928 \pm 0.017 M_\odot$





- PSR J0348+0432 (Antoniadis et al. 2013)

✓ binary system (P=2.46 h)

✓ very low eccentricity

✓ companion mass: $0.172 \pm 0.003 M_\odot$

✓ pulsar mass: $M = 2.01 \pm 0.04 M_\odot$

Formation of Binary Systems

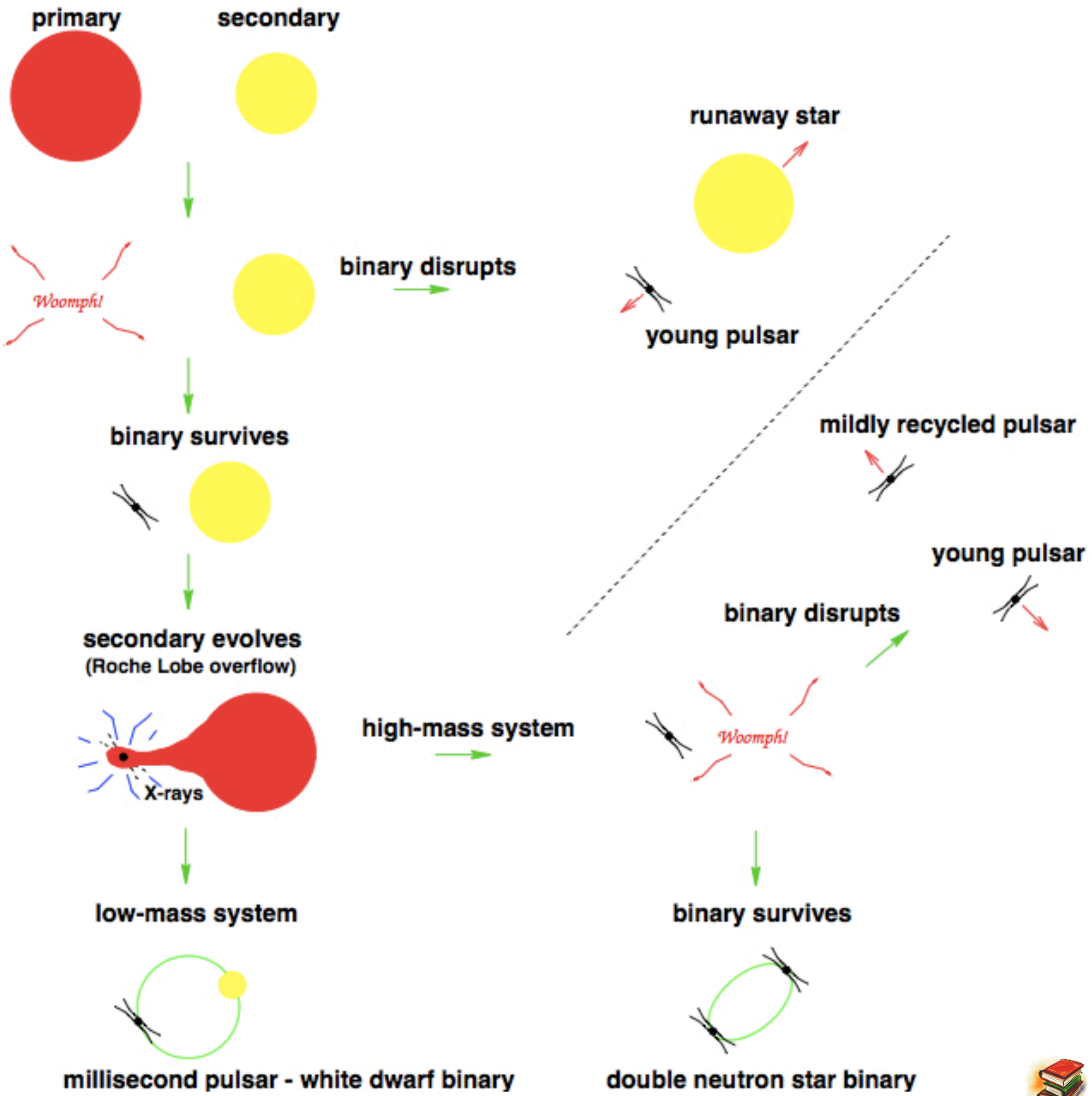
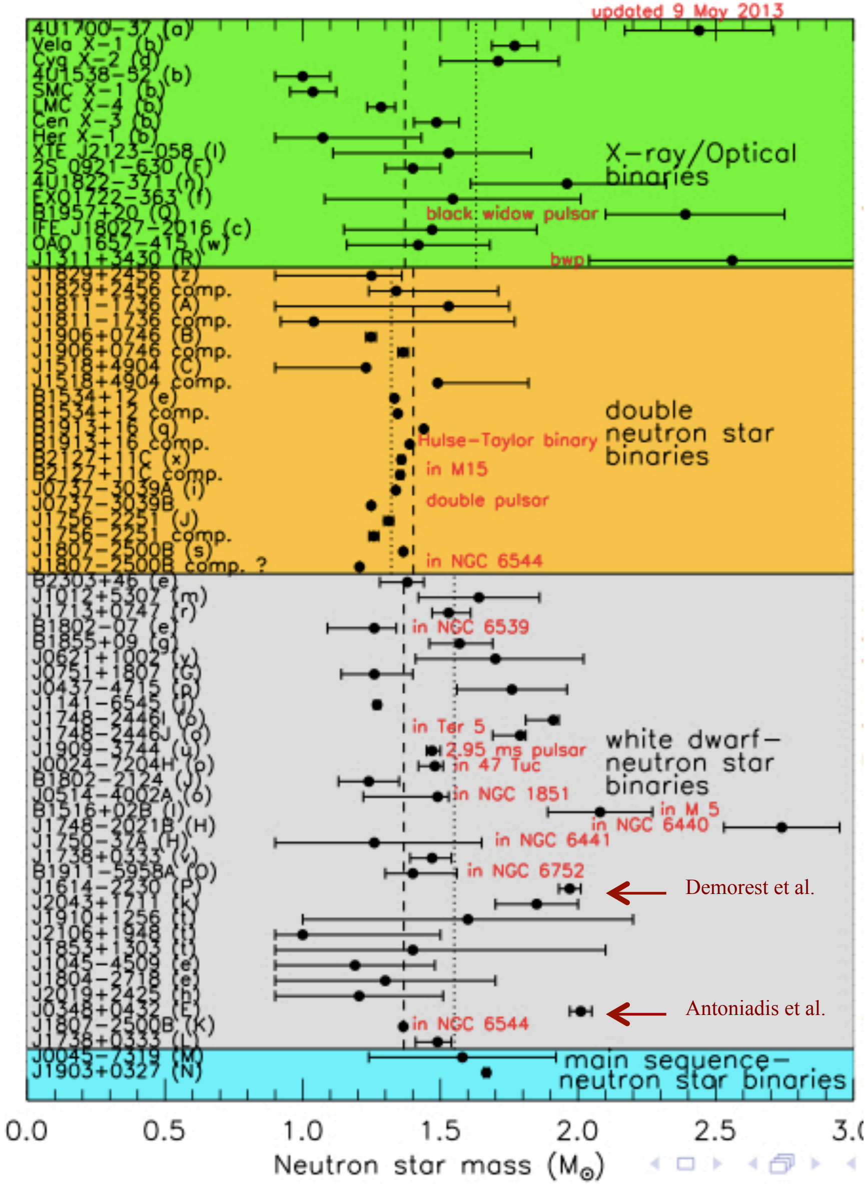


Figure by P.C.C. Freire

Measured Neutron Star Masses (2017)



updated from Lattimer 2013

Observation of ~ 2 $M_\odot$ neutron stars

↓

Dense matter EoS stiff enough is required such that

$$M_{max}\left[EoS\right] > 2M_\odot$$

A natural question arises:

Can hyperons, or strangeness in general, still be present in the interior of neutron stars in view of this constraint ?

The Hyperon Puzzle

"Hyperons ➔ "soft (or too soft) EoS" not compatible (mainly in microscopic approaches) with measured (high) masses. However, the presence of hyperons in the NS interior seems to be unavoidable."

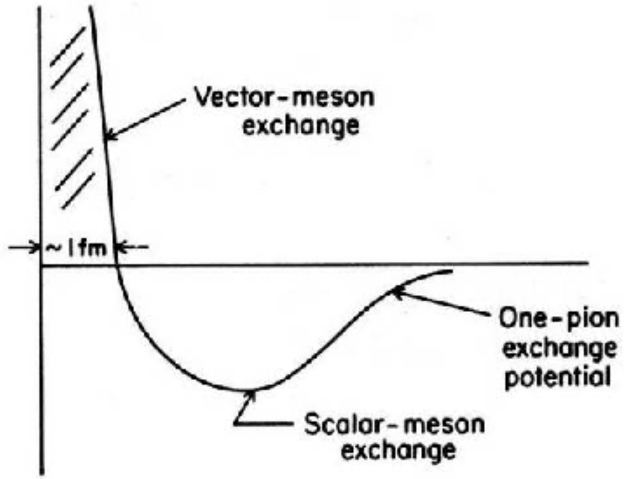
- ✓ can YN & YY interactions still solve it ?
- ✓ or perhaps hyperonic three-body forces ?
- ✓ what about quark matter ?

Solution I: YY vector meson repulsion

(explored in the context of RMF models)

General Feature:

Exchange of scalar mesons generates attraction (softening), but the exchange of vector mesons generates repulsion (stiffening)



Add vector mesons with hidden strangeness (ϕ) coupled to hyperons yielding a strong repulsive contribution at high densities

Dexhamer & Schramm (2008), Bednarek et al, (2012), Weissenborn et al., (2012)
Oertel et al. (2014), Maslov et al. (2015)

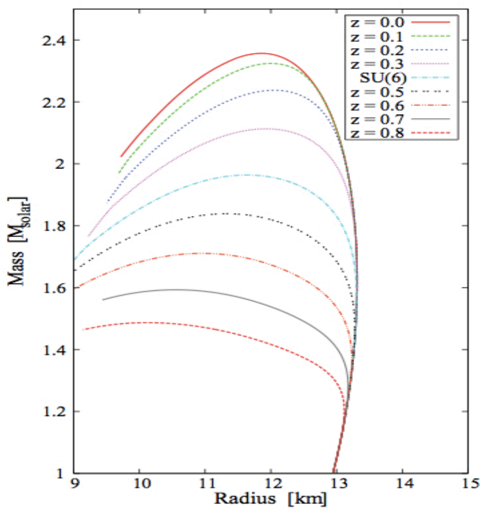
Weissenborn et al. (2012)

- ✓ σ^2, σ^3, σ^4 terms
- ✓ ρ^2, ω^2, ω^4 terms
- ✓ "hidden strangeness" mesons: σ^*, ϕ

  (σ^{*2}, ϕ^2)
- ✓ g_{YV} couplings: from SU(6) to SU(3)

  vary $z=g_8/g_1$ & $\alpha=F/(F+D)$
- ✓ g_{YS} couplings adjusted by fitting $U_B^{(N)}$

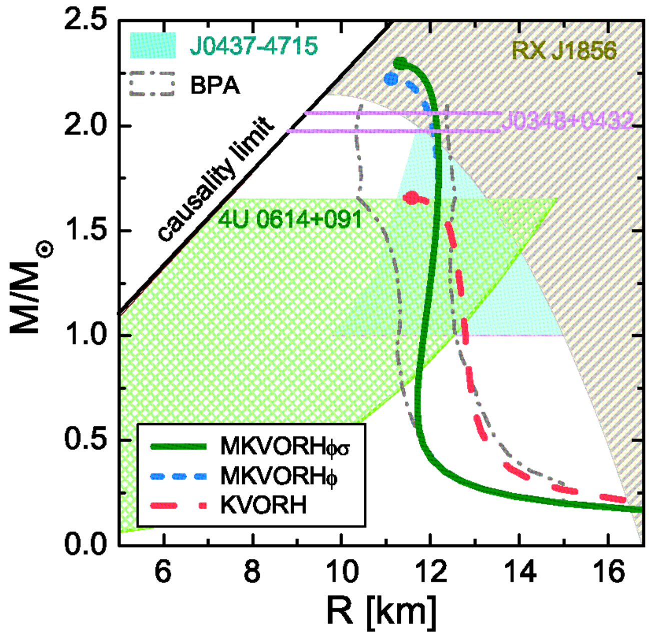
  ($U_\Lambda^{(N)}$=-30, $U_\Sigma^{(N)}$=+30, $U_\Xi^{(N)}$=-28 MeV)



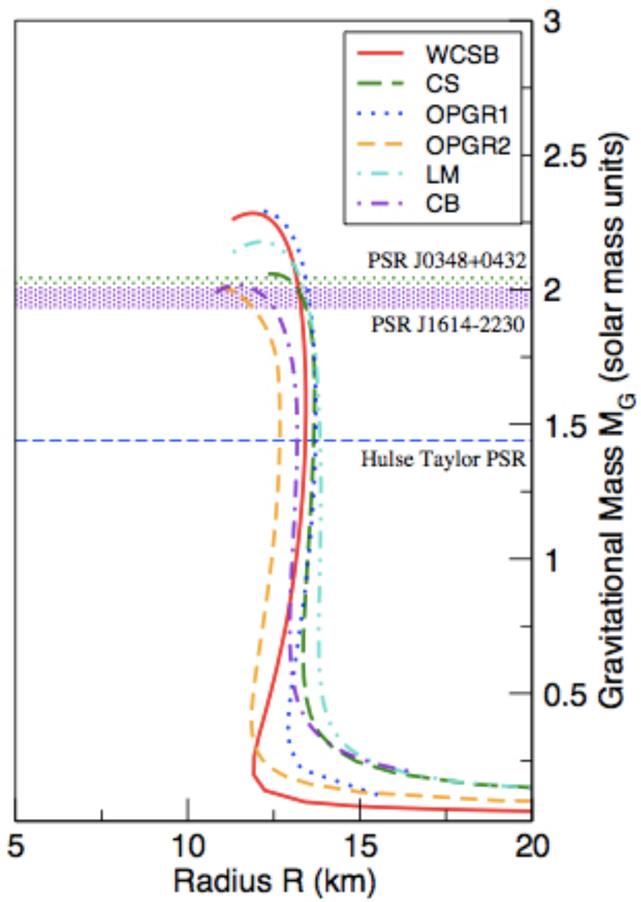
M_{max} compatible with $2M_\odot$

Maslov et al. (2015)

- ✓ RMF with scaled hadron masses (universal) & coupling constants (not universal)
- ✓ Model flexible enough to satisfy constraints from HIC & astrophysical data
- ✓ Hyperon puzzle <u>partially solved if a reduction of ϕ meson mass is included</u>



Although these and other similar models are able to reconcile the presence of hyperons in the NS interior with the existence of $2M_{\odot}$ NS, <u>one must be cautious</u> !!



D. Chatterjee & I. V. (2015)

✧ These models contain several free parameters which most of the times are arbitrarily chosen being the only jutification our still "scarce" knowledge of the YY interaction.

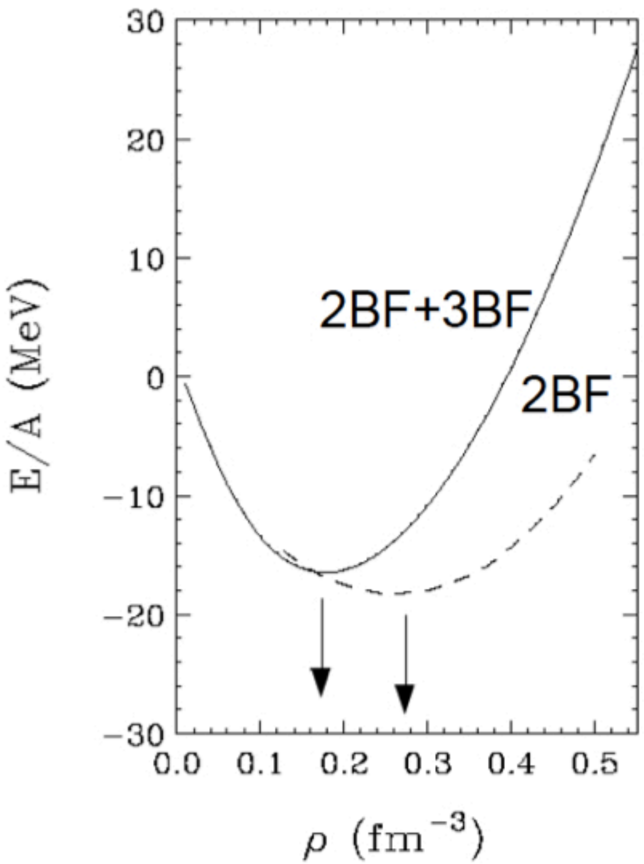
Hence:

In absence of sufficient experimental data on multi-strange hypernuclei and YY scattering the validity of these models is still questionable.

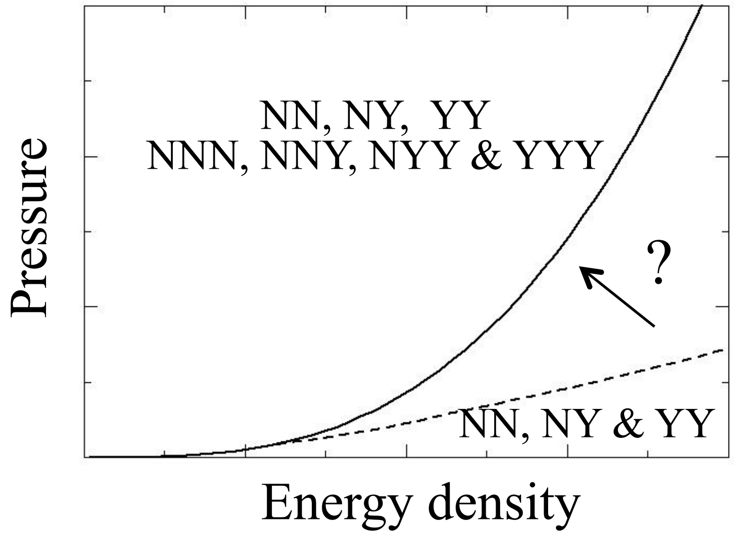
Solution II: can Hyperonic TBF solve this puzzle ?

Natural solution based on: Importance of NNN force in Nuclear Physics
(Considered by several authors: Chalk, Gal, Usmani, Bodmer, Takatsuka, Loiseau, Nogami, Bahaduri, IV)

NNN Force

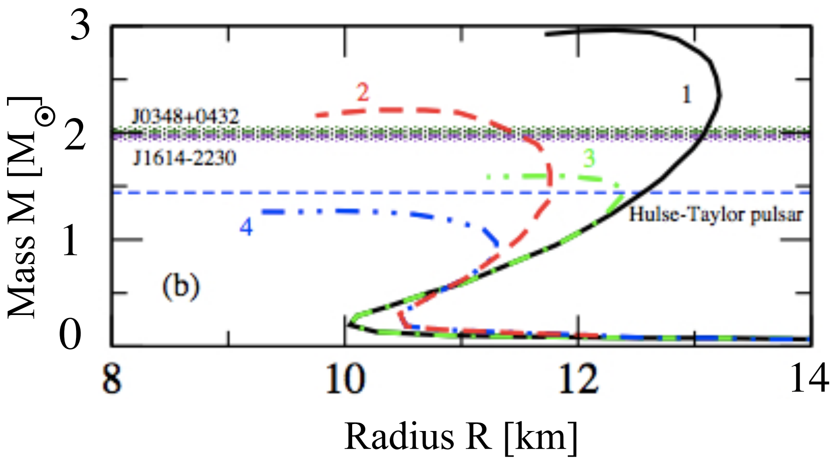


NNY, NYY & YYY Forces



Can hyperonic TBF provide enough repulsion at high densities to reach $2M_\odot$?

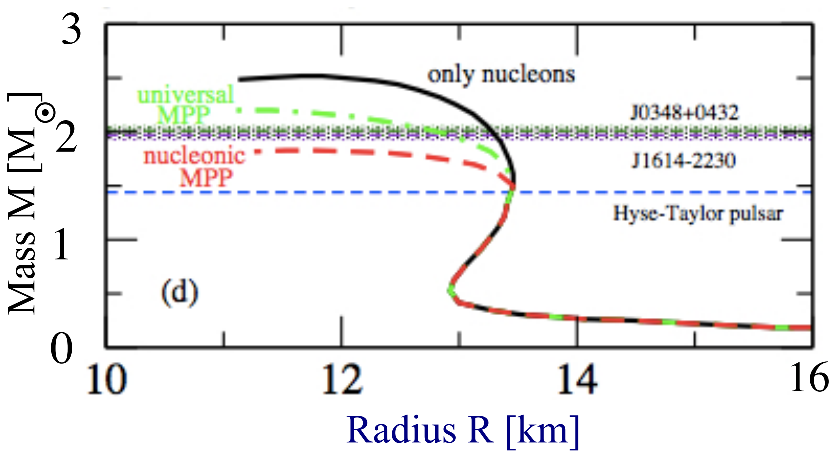
The results are contradictory



I. V. et al. (2011)

BHF with NN+YN+phenomenological YTBF. Different strength of YTBF including the case of universal TBF

$$1.27 < M_{max} < 1.6 M_\odot$$

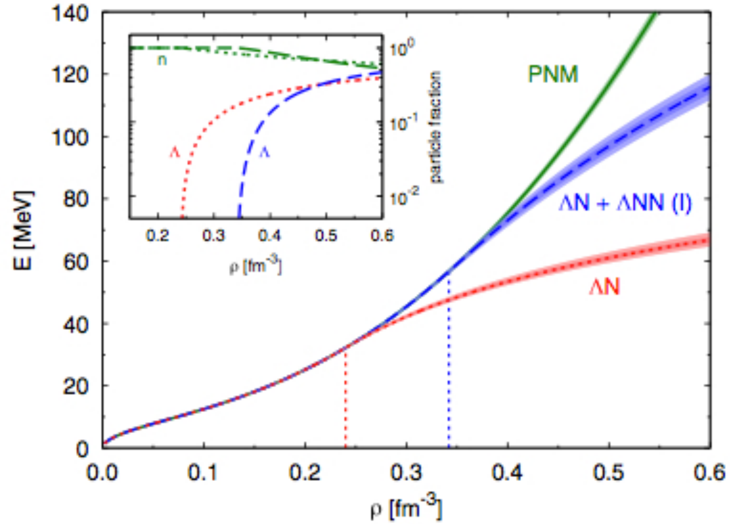


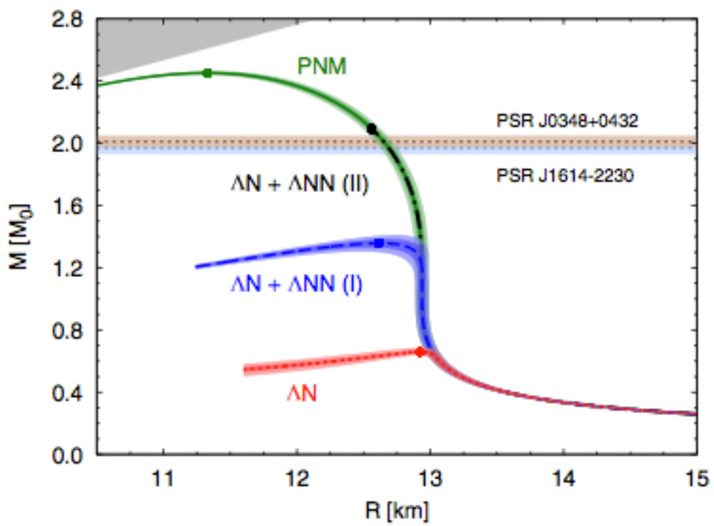
Yamamoto et al. (2015)

BHF with NN+YN+universal repulsive TBF (multipomeron exchange mecanism)

$$M_{max} > 2 M_\odot$$

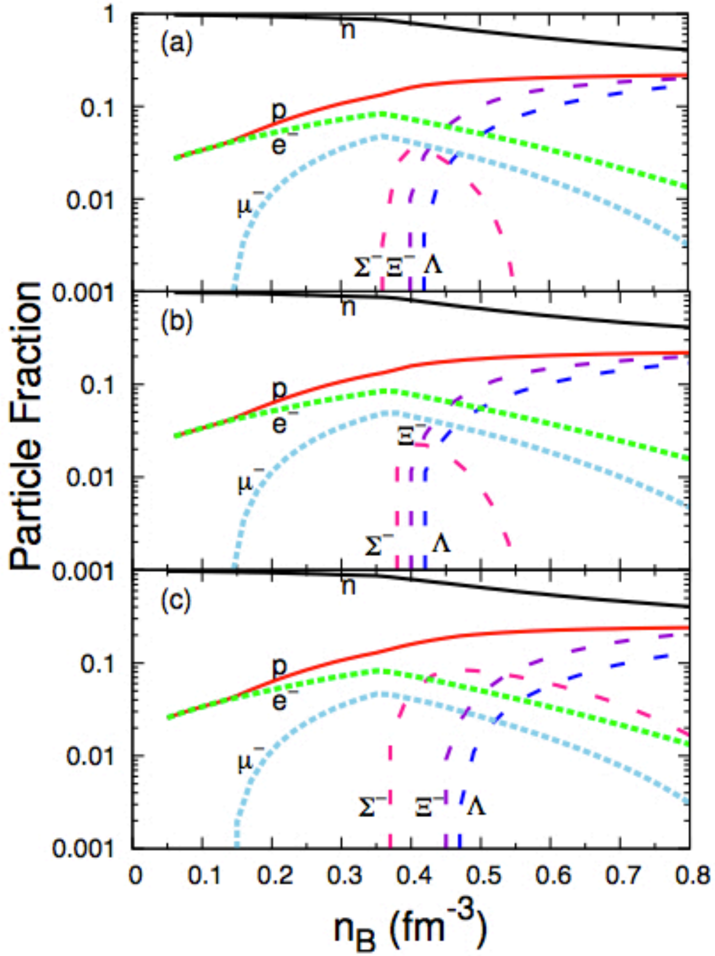
It should be mentioned also the recent Quantum Monte Carlo calculation by Lonardoni et al. (2015)

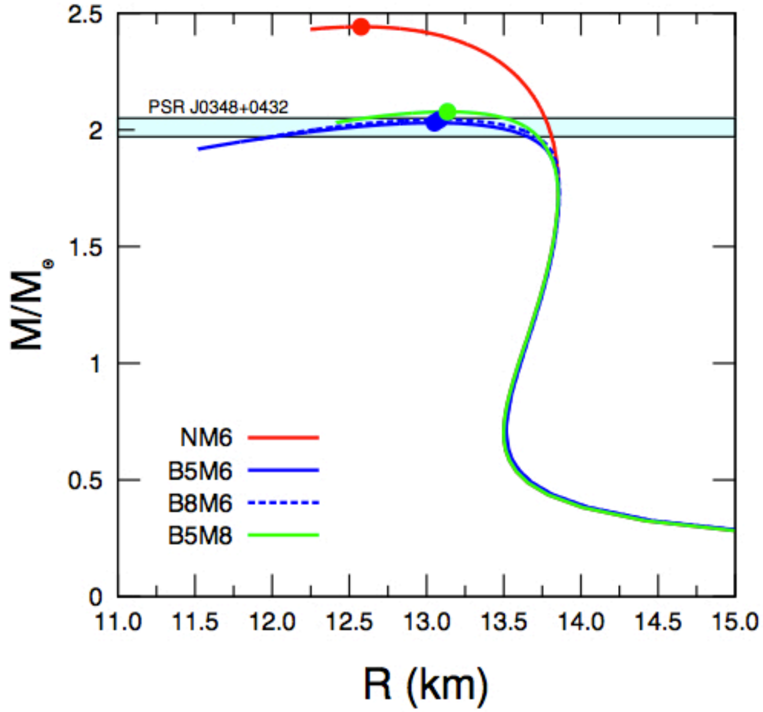




- First Quantum Monte Carlo calculation on neutron+Λ matter
- Strong dependence of Λ onset on Λnn force
- Some of the parametrizations of the Λnn force give maximum masses compatible with $2M_\odot$ but the onset of Λ is above the maximum density considered (~0.56 fm^{-3}). So in fact, no Λs are present in NS interior

and the recent DBHF calculation of hyperonic matter by Katayama & Saito (2014)





- DBHF includes some TBF effects in a natural way
- M_{max} compatible with $2M_{\odot}$
- But the construction of YN is a bit obscure in this work

Take Away Message

- ✧ It is still an open question whether hyperonic TBFs can, by themselves, solve completely the hyperon puzzle or not.

- ✧ It seems, however, that even if they are not the full solution, most probably they can contribute to it in an important way.

Solution III: Quark Matter Core

General Feature:

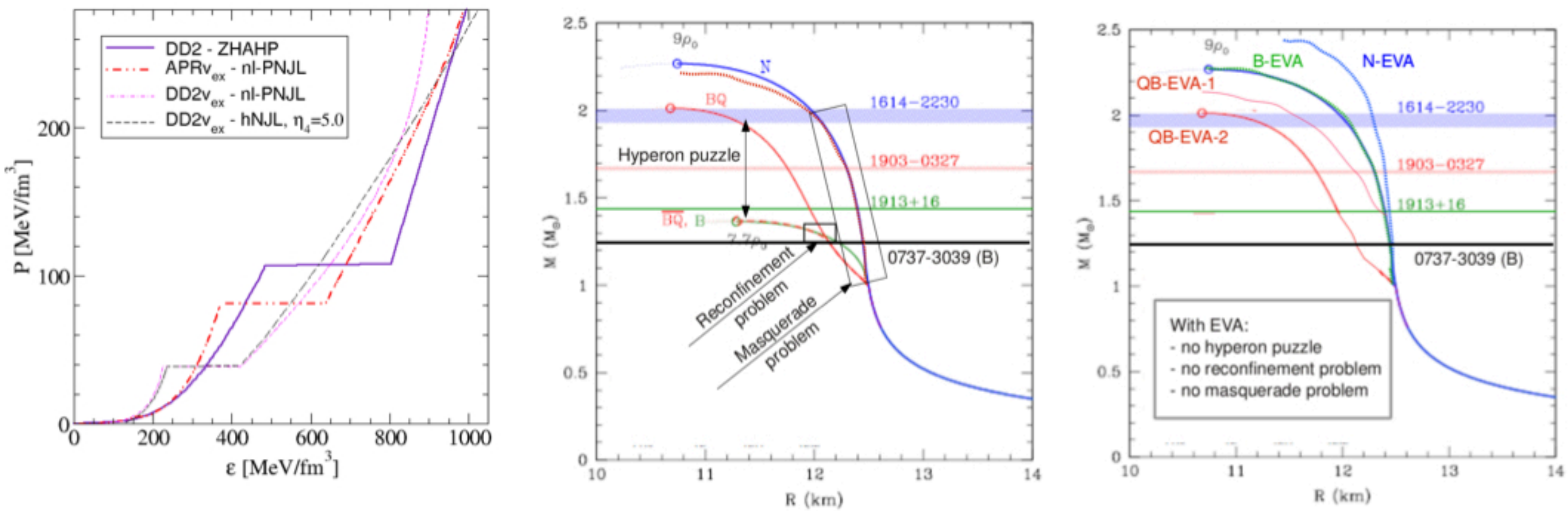
Some authors have suggested an early phase transition to deconfined quark matter as solution to the hyperon puzzle. Massive neutron stars could actually be hybrid stars with a stiff quark matter core.

To yield $M_{\max} > 2M_{\odot}$ Quark Matter should have:

- significant overall quark repulsion $\longrightarrow$ stiff EoS
- strong attraction in a channel $\longrightarrow$ strong color superconductivity

Ozel et al., (2010), Weissenborn et al., (2011), Klaehn et al., (2011), Bonano & Sedrakian (2012), Lastowiecki et al., (2012), Zdunik & Haensel (2012)

A recent work by D. Blaschke & D. Alvarez-Castillo (2015)

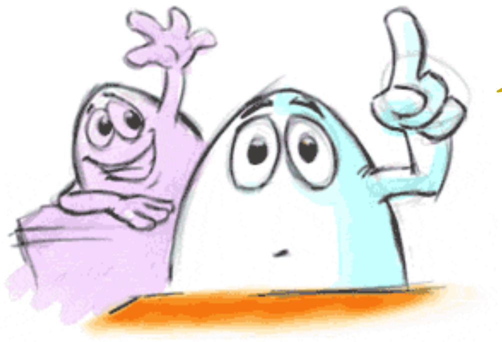


Compositeness of baryons (by excluded volume and/or quark Pauli blocking) on the hadronic side + confinement and stiffening effects on the quark matter:

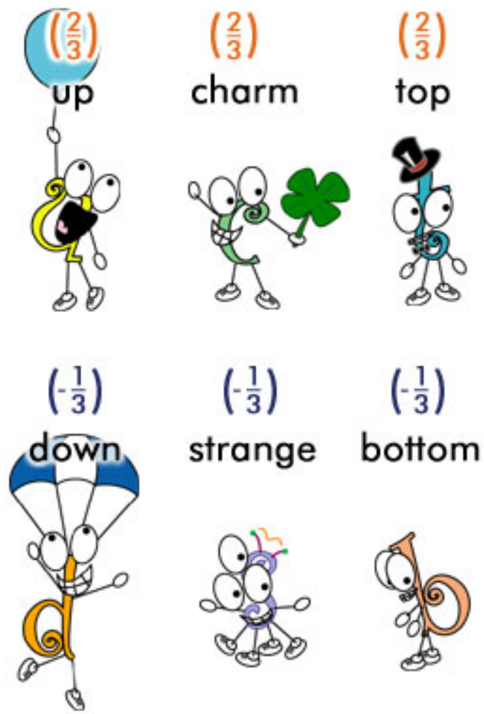
→

Earlier phase transition to QM with sufficient stiffening at high densities to solve: hyperon puzzle, masquerade problem & reconfinement puzzle

What quark flavors are expected in a Neutron Stars ?



Flavor	Mass	Charge [e]
u	~ 5 MeV	2/3
d	~ 10 MeV	-1/3
s	~ 200 MeV	-1/3
c	~ 1.3 GeV	2/3
b	~ 4.3 GeV	-1/3
t	~ 175 GeV	2/3



Suppose: ✓ u, d, s non-interacting
✓ $m_{u=}m_d=m_s=0$ ⟶ i.e., ideal ultra-relativistic Fermi gas (*)

❖ Threshold density for the c quark (similar for b & t)

$$s \rightarrow c + e^{-} + \bar{\nu}_e \Rightarrow \mu_s = \mu_c + \mu_e + \mu_{\bar{\nu}_e}$$

but ✓ u, d, s in β-equilibrium
✓ $Q_{tot=}=0$ ⟶

$$n_B = n_u = n_d = n_s$$
$$n_e = n_{\bar{\nu}_e} = 0$$

then

$$\mu_s = E_{F_s} = \hbar c\left(\pi^2 n_s\right)^{1/3} = \hbar c\left(\pi^2 n_B\right)^{1/3} \geq m_c = 1.3 \quad GeV$$

$$\Rightarrow n_B \geq 29 \quad fm^{-3} \sim 180 n_0$$

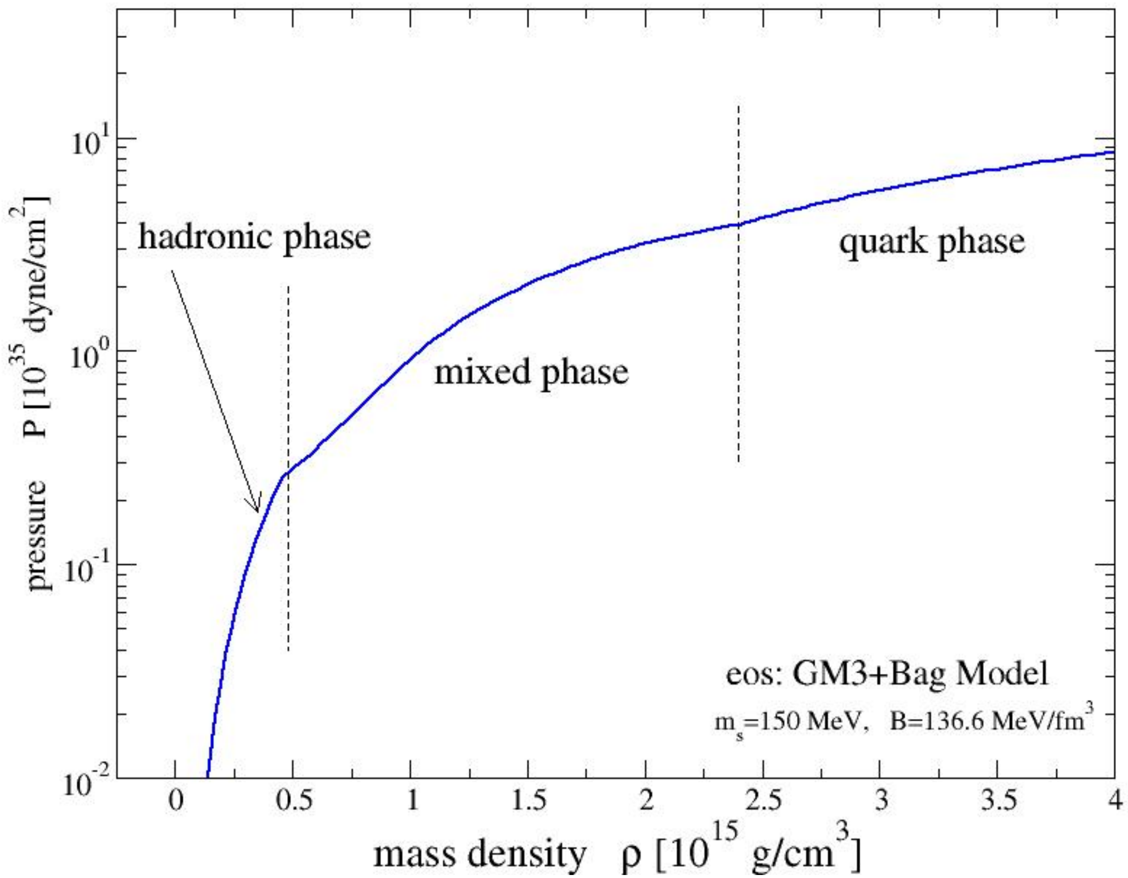
Only u,d,s quarks are expected in Neutron Stars

The Equation of State for Hybrid Stars

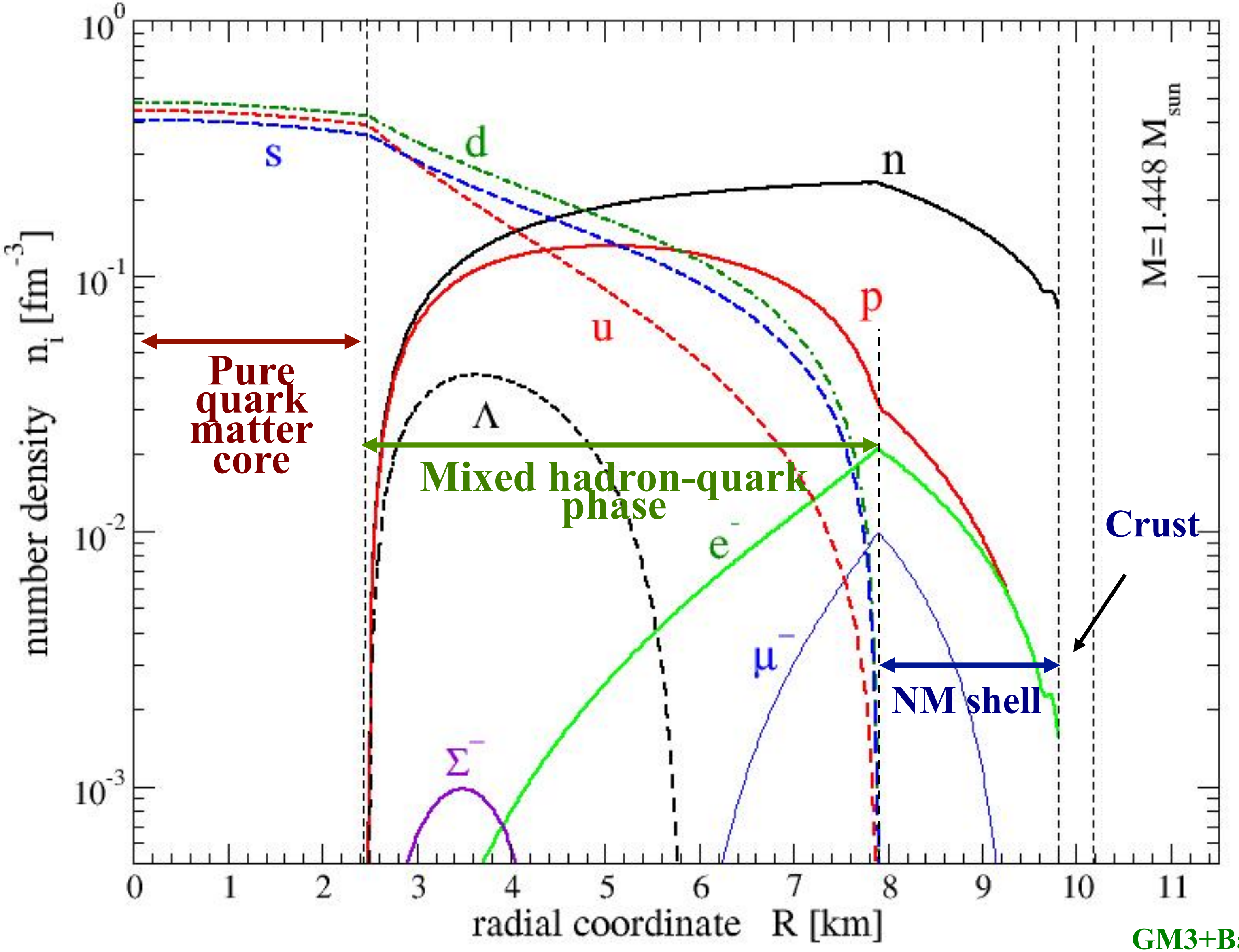
- **Hadronic phase :**
  RMF Models
  Microscopic BHF

- **Quark phase :**
  EOS based on the MIT bag model for hadrons.
  [Farhi, Jaffe, Phys. Rev. D46(1992)]

- **Mixed phase :**
  Gibbs construction for a multicomponent system with two conserved "charges". [Glendenning, Phys. Rev. D46 (1992)]



Hybrid Star Composition



GM3+Bag model
m_s=150 MeV, B=13.6.6Mev/fm^3

The Strange Matter Hypothesis

Bodmer (1971), Terezawa (1979) & Witten (1984)

Three-flavour u,d,s quark matter in equilibrium with respect to the weak interactions, could be the true ground state of strongly interacting mater, rather than ^{56}Fe

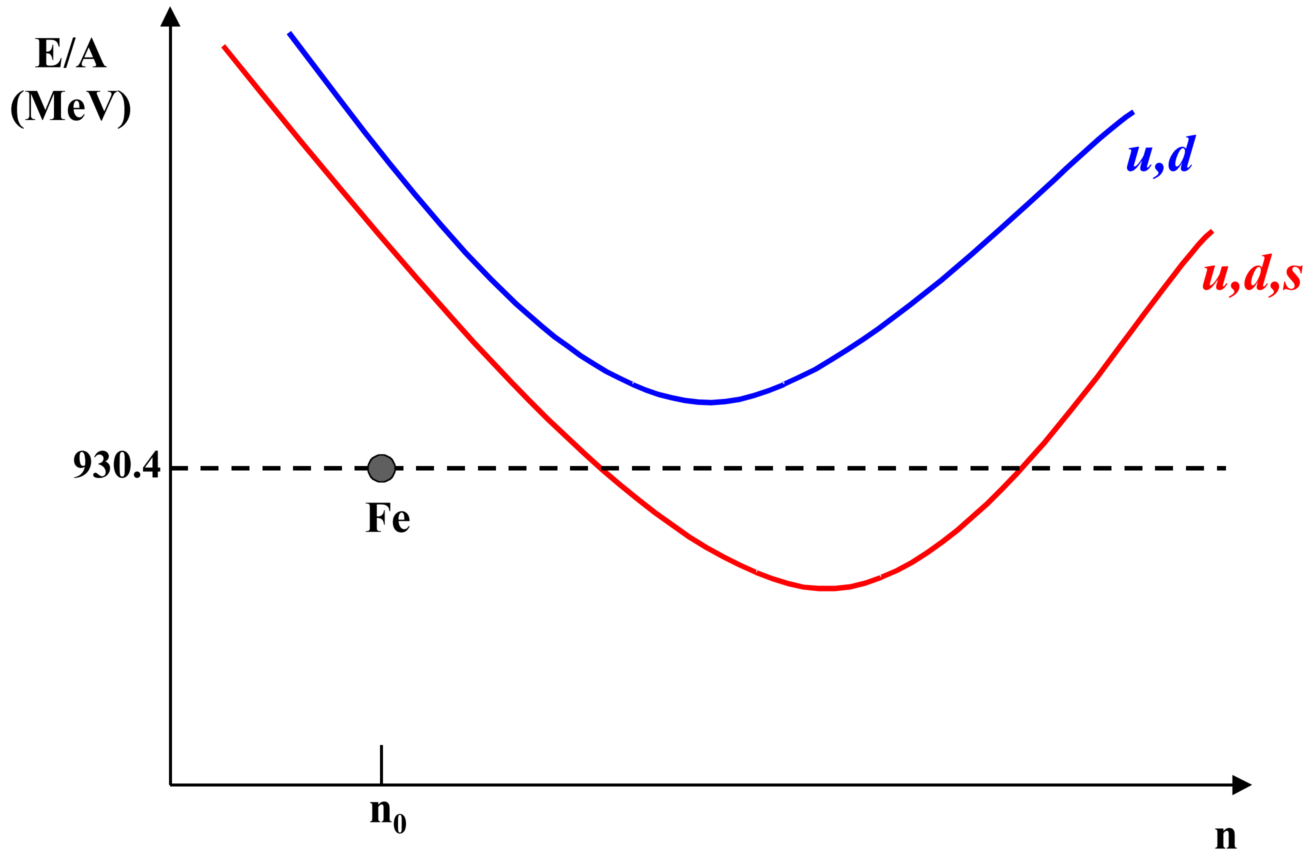
$$E/A|_{SQM} < E(^{56}Fe)/56 \sim 930 \text{ MeV}$$

Stability of nuclei with respect to u,d quark matter

The success of traditional nuclear physics provides a clear indication that quarks in the atomic nuclei are confined within neutrons and protons

$$E/A|_{ud} > E(^{56}Fe)/56 \sim 930 \text{ MeV}$$

Schematically



- If the SQM hypothesis is true, why nuclei do not decay into SQM droplets (strangelets) ?

- One should explain the existence of atomic nuclei in Nature

Stability of Nuclei with respect to SQM

- Direct decay of ^{56}Fe to a SQM droplet

$$^{56}Fe \rightarrow {}^{56}(SQM) \quad \Longrightarrow$$

~ 56 simultaneous strangeness changing weak process

$$u \rightarrow s + e^{+} + \nu_e$$
$$d + u \rightarrow s + u$$

The probability for the direct decay is $P \sim (G_F^2)^{56} \sim 0$ and the mean-life time of ^{56}Fe with respect to the direct decay to a drop of SQM is

$\tau >>$ age of the Universe

- Step by step decay of ^{56}Fe to a SQM droplet

$$^{56}Fe \rightarrow X_{\Lambda}^{56} \rightarrow Y_{\Lambda\Lambda}^{56} \rightarrow \ldots \rightarrow {}^{56}(SQM)$$

$$^{56}Fe \rightarrow Fe_{\Lambda}^{56}$$

$$^{56}Fe \rightarrow Mn_{\Lambda}^{56}$$

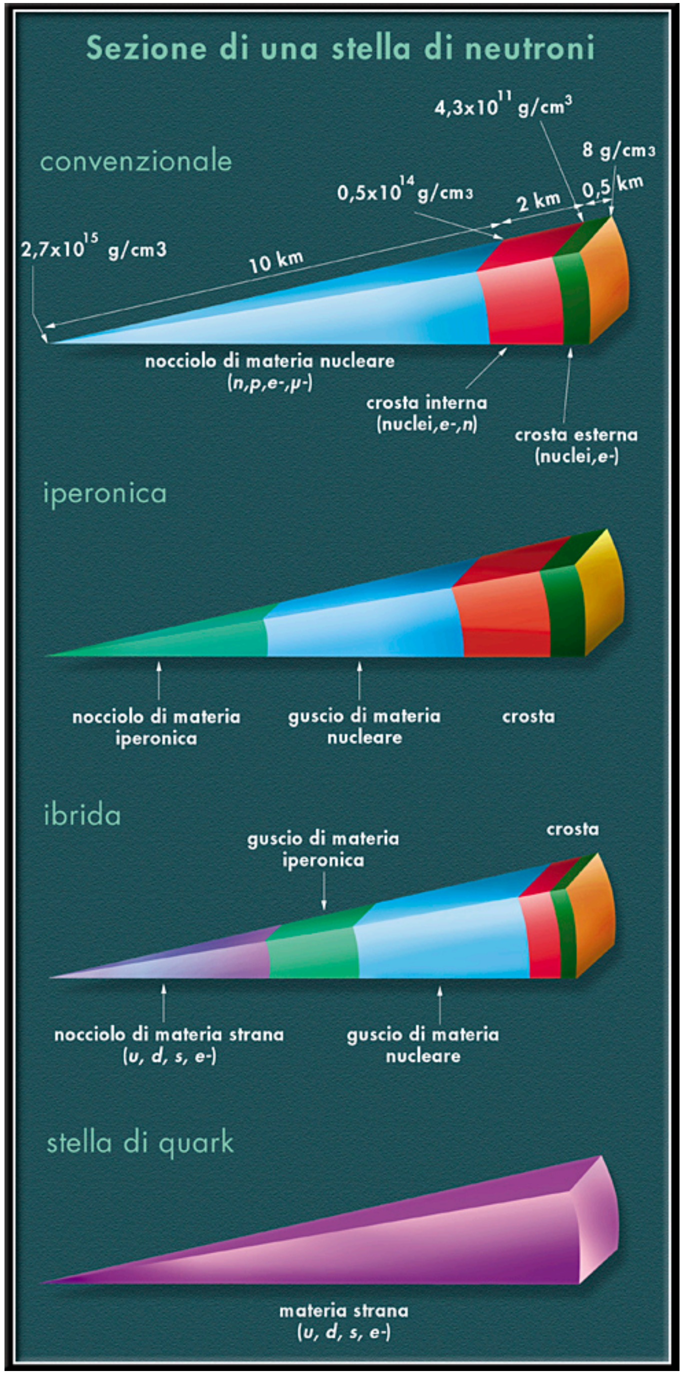
These processes are not energetically possible since

$$Q = M(^{56}Fe) - M(X_{\Lambda}^{56}) < 0$$

Thus, according with the Bodmer-Terezawa-Witten hypothesis, nuclei are metastable states of strong interacting matter with a mean-life time

τ >> age of the Universe

Two families of Neutron Stars



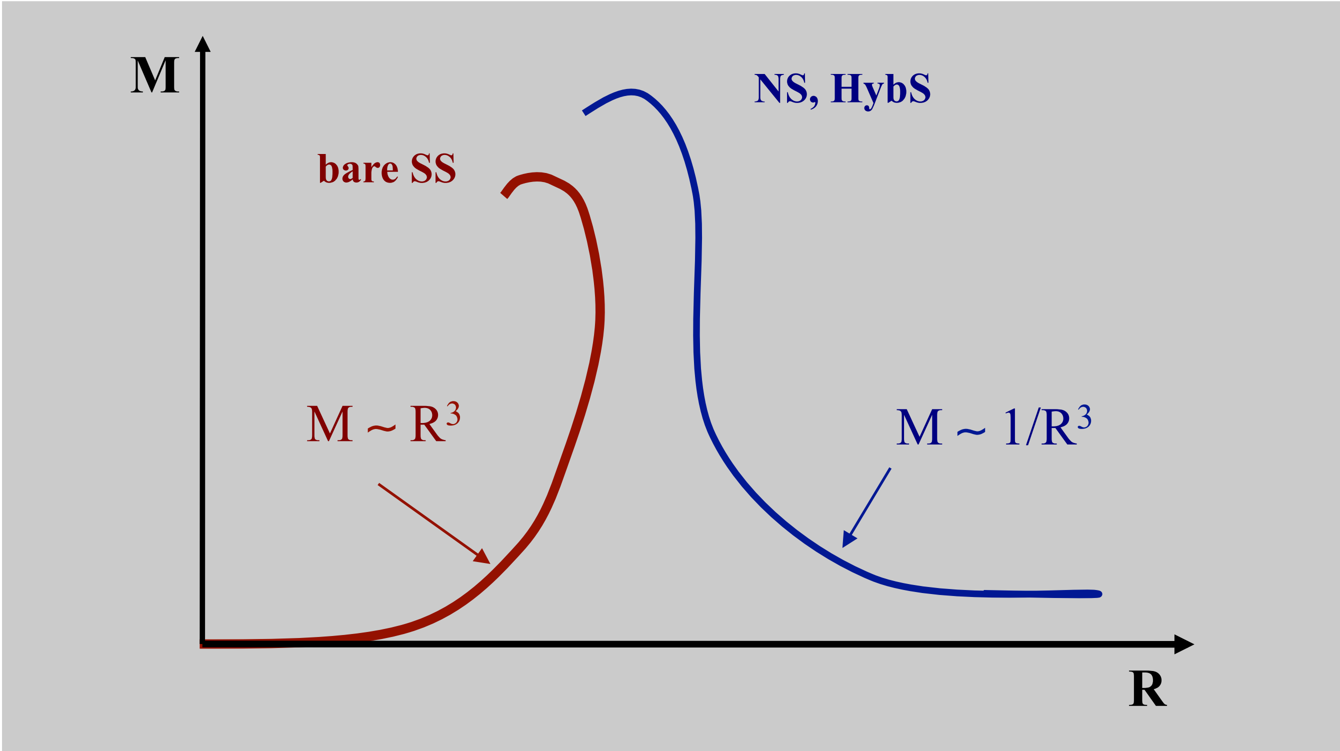
Hadron Stars (HS)

- Nucleonic Stars
- Hyperonic Stars

Quark Stars (QS)

- Hybrid Stars
- Strange Stars

Mass-radius relation



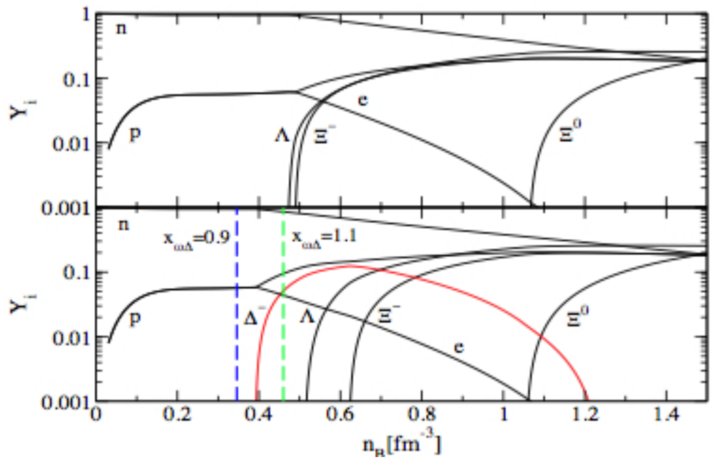
- ✧ Strange Stars are self-bound bodies i.e., bound by the strong interactions
- ✧ Hadronic or Hybrid Stars are bound by gravity.

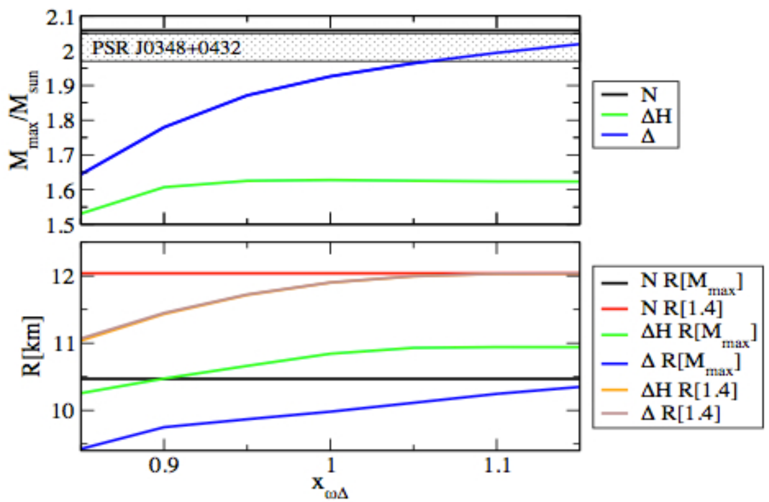
But also in this case we must pay attention

Currently theoretical descriptions of quark matter at high density rely on phenomenological models which are constrained using the few available experimental information on high density baryonic matter from heavy-ion collisions.

Is there also a Δ isobar puzzle ?

The recent work by Drago et al. (2014) calculation have studied the role of the Δ isobar in neutron star matter



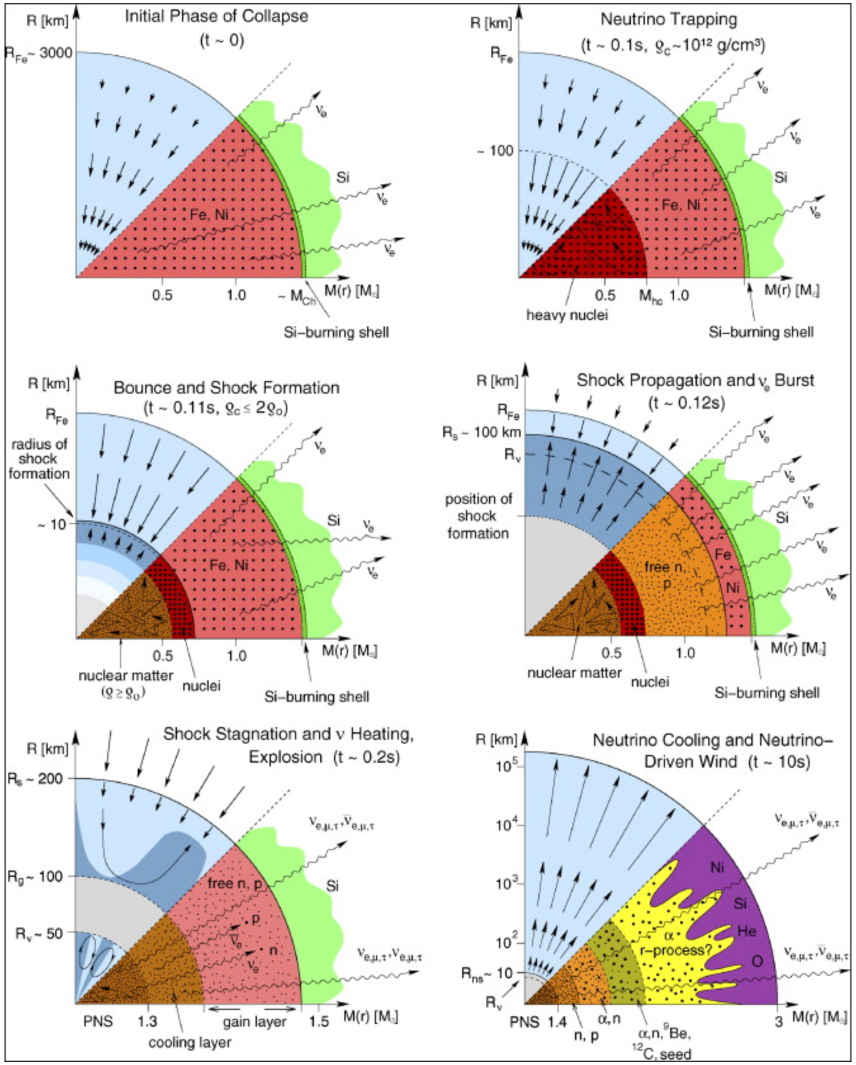


- Constraints from L indicate an early appearance of Δ isobars in neutron stars matter at ~ 2-3 ρ_0 (same range as hyperons)
- Appearance of Δ isobars modify the composition & structure of hadronic stars
- M_{max} is dramatically afected by the presence of Δ isobars

If Δ potential is close to that indicated by π−, e-nucleus or photoabsortion nuclear reactions then EoS is too soft ⟶ Δ puzzle similar to the hyperon one

Hyperon Stars at Birth

Proto-Neutron Stars



(Janka, Langanke, Marek, Martinez-Pinedo & Muller 2006)

New effects on PNS matter:

- Thermal effects

$$T \cong 30-40 \quad MeV$$
$$S/A \cong 1-2$$

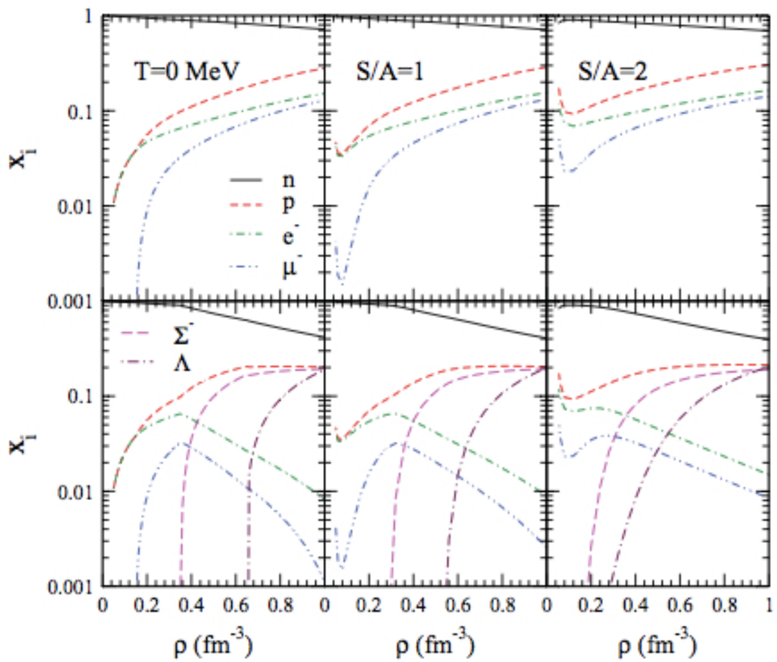
- Neutrino trapping

$$\mu_\nu \neq 0$$
$$Y_e = \frac{\rho_e + \rho_{\nu_e}}{\rho_B} \approx 0.4$$
$$Y_\mu = \frac{\rho_\mu + \rho_{\nu_\mu}}{\rho_B} \approx 0$$

Proto-Neutron Stars: Composition

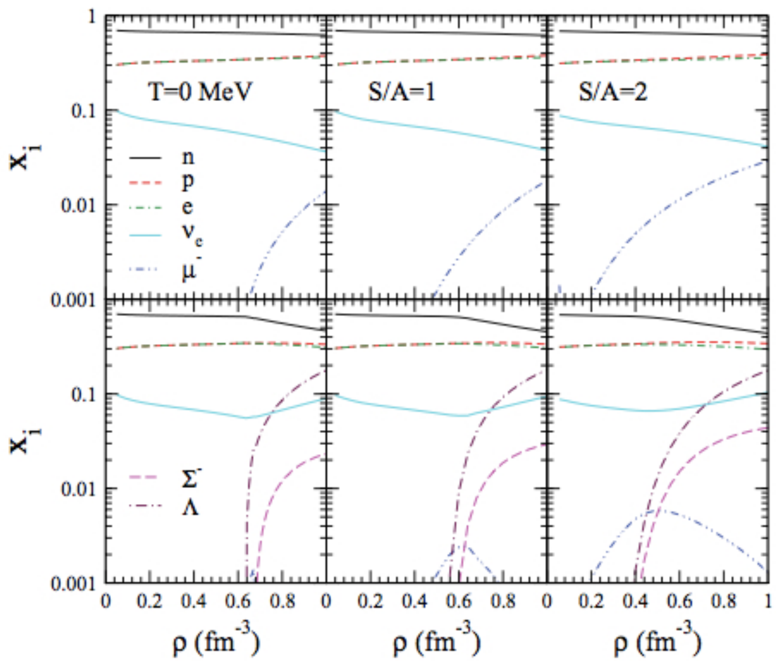
- Neutrino free $\mu_\nu = 0$

(Burgio & Schulze 2011)



- Neutrino trapped $\mu_\nu \neq 0$

(Burgio & Schulze 2011)

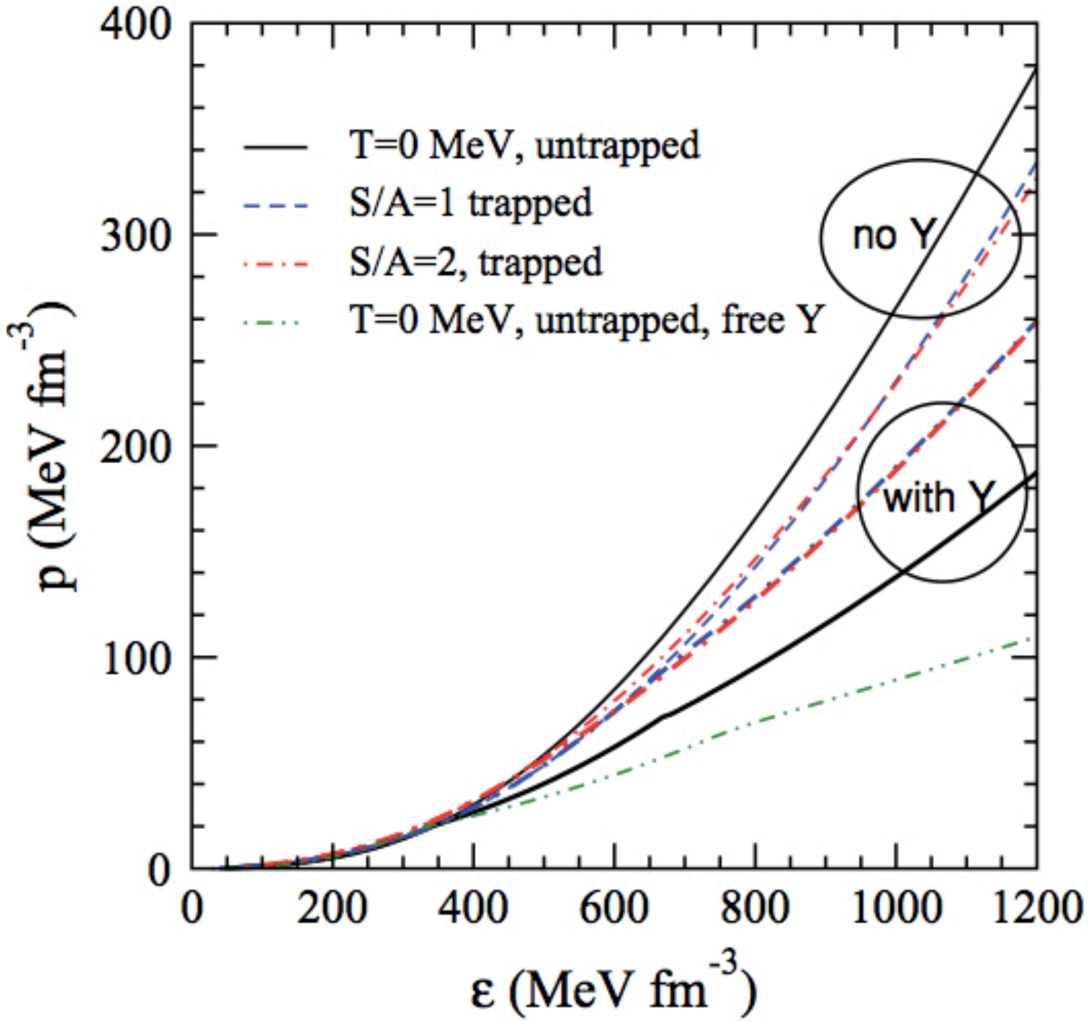


<u>Neutrino trapped</u>

- ✓ Large proton fraction
- ✓ Small number of muons
- ✓ Onset of $\Sigma^{-}(\Lambda)$ shifted to higher (lower) density
- ✓ Hyperon fraction lower in ν-trapped matter

Proto-Neutron Stars: EoS

(Burgio & Schulze 2011)



- Nucleonic matter

✧ ν-trapping + temperature → softer EoS

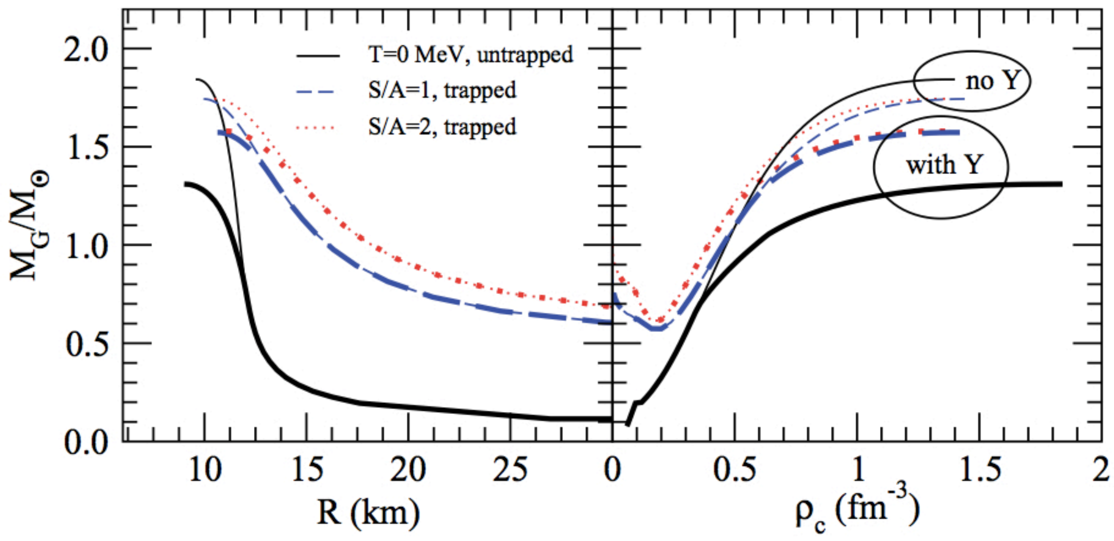
- Hyperonic matter

✧ ν-trapping + temperature → stiffer EoS

✧ More hyperon softening in ν-untrapped matter (larger hyperon fraction)

Proto-Neutron Stars: Structure

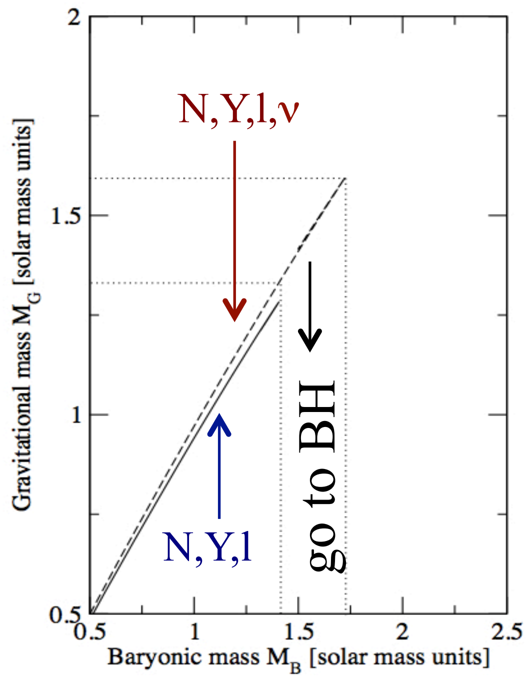
(Burgio & Schulze 2011)



- Nucleonic matter

  ν-trapping + T: reduction of M_{max}

(IV et al. 2003)



- Hyperonic matter

  ν-trapping + T: increase of M_{max} → delayed formation of a low mass BH

Hyperons & Neutron Star Cooling

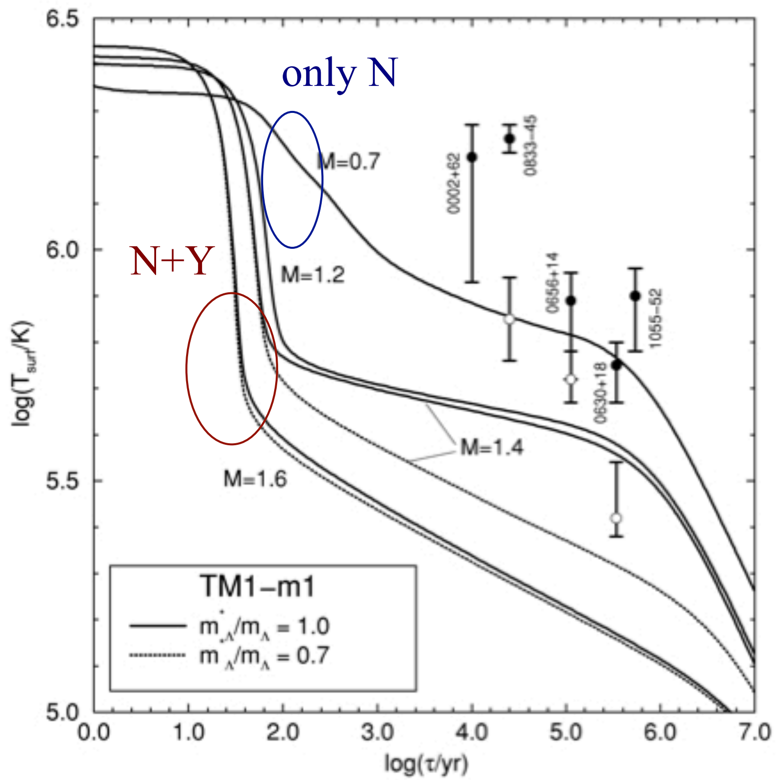
Hyperonic DURCA processes possible
as soon as hyperons appear
(nucleonic DURCA requires $x_p > 11\text{-}15\ \%$)

→ **Additional Fast Cooling Processes**

Process	R
$\Lambda \to p + l + \bar{\nu}_l$	0.0394
$\Sigma^- \to n + l + \bar{\nu}_l$	0.0125
$\Sigma^- \to \Lambda + l + \bar{\nu}_l$	0.2055
$\Sigma^- \to \Sigma^0 + l + \bar{\nu}_l$	0.6052
$\Xi^- \to \Lambda + l + \bar{\nu}_l$	0.0175
$\Xi^- \to \Sigma^0 + l + \bar{\nu}_l$	0.0282
$\Xi^0 \to \Sigma^+ + l + \bar{\nu}_l$	0.0564
$\Xi^- \to \Xi^0 + l + \bar{\nu}_l$	0.2218

+ partner reactions generating neutrinos,
Hyperonic MURCA, …

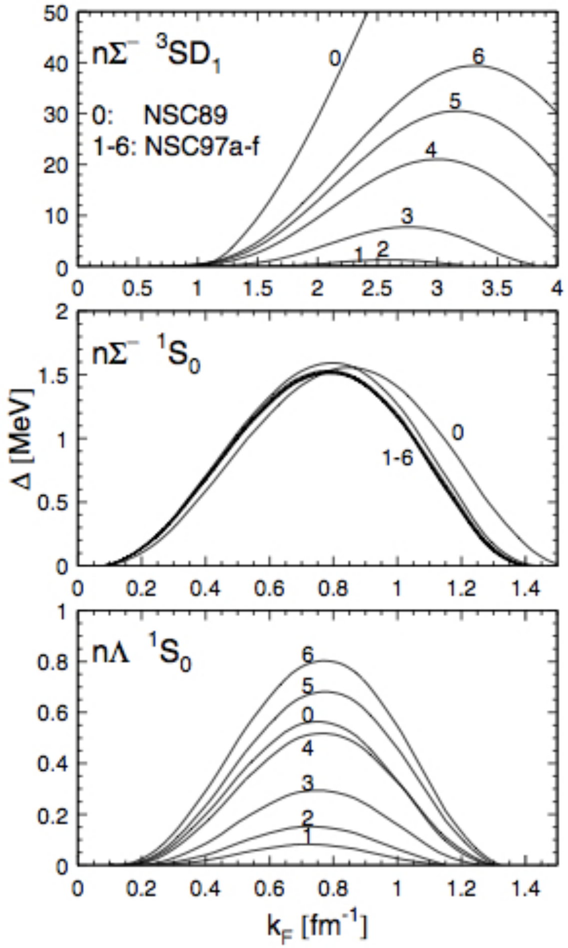
(Schaab, Shaffner-Bielich & Balberg 1998)



R: relative emissitivy w.r.t. nucleonic DURCA

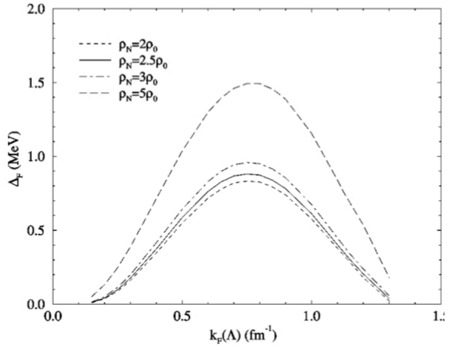
Pairing Gap ⟶ suppression of C_v & ε by $\sim e^{(-\Delta/k_B T)}$

- $^1S_0$, $^3SD_1$ ΣN & $^1S_0$ ΛN gap

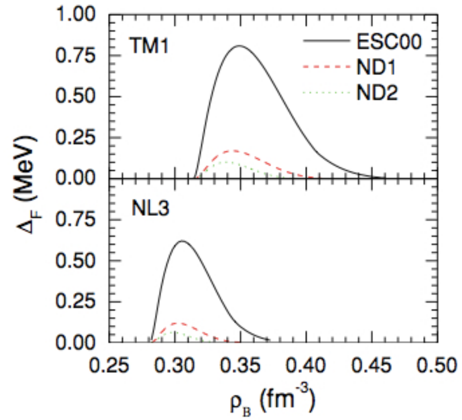


(Zhou, Schulze, Pan & Draayer 2005)

- $^1S_0$ ΛΛ gap

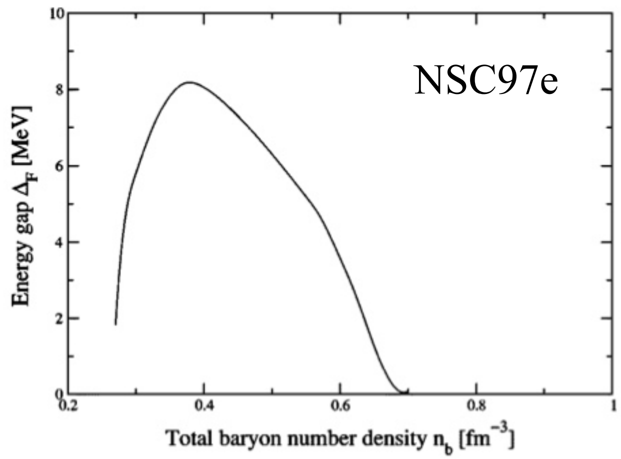


(Balberg & Barnea 1998)



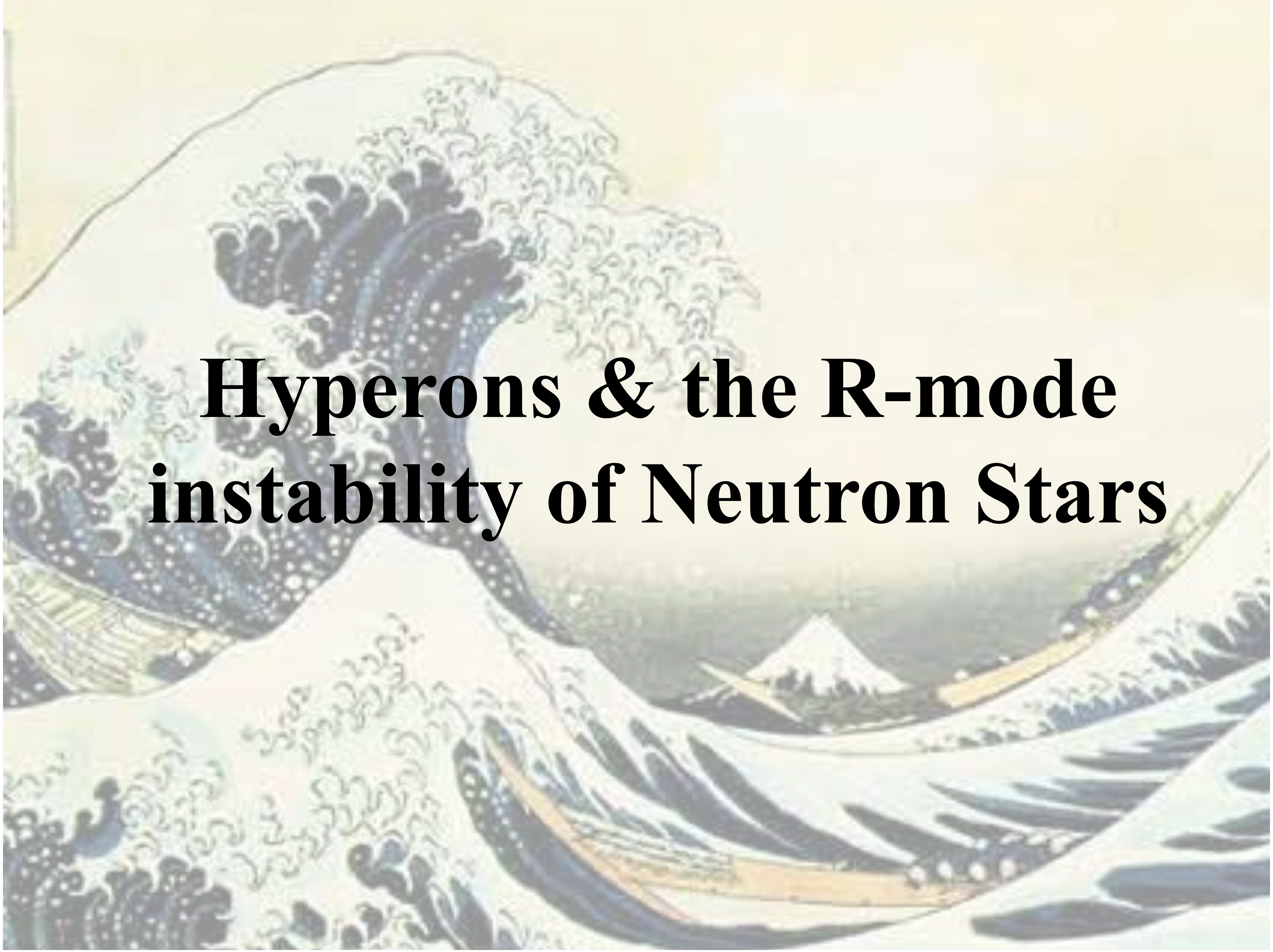
(Wang & Shen 2010)

- $^1S_0$ ΣΣ gap



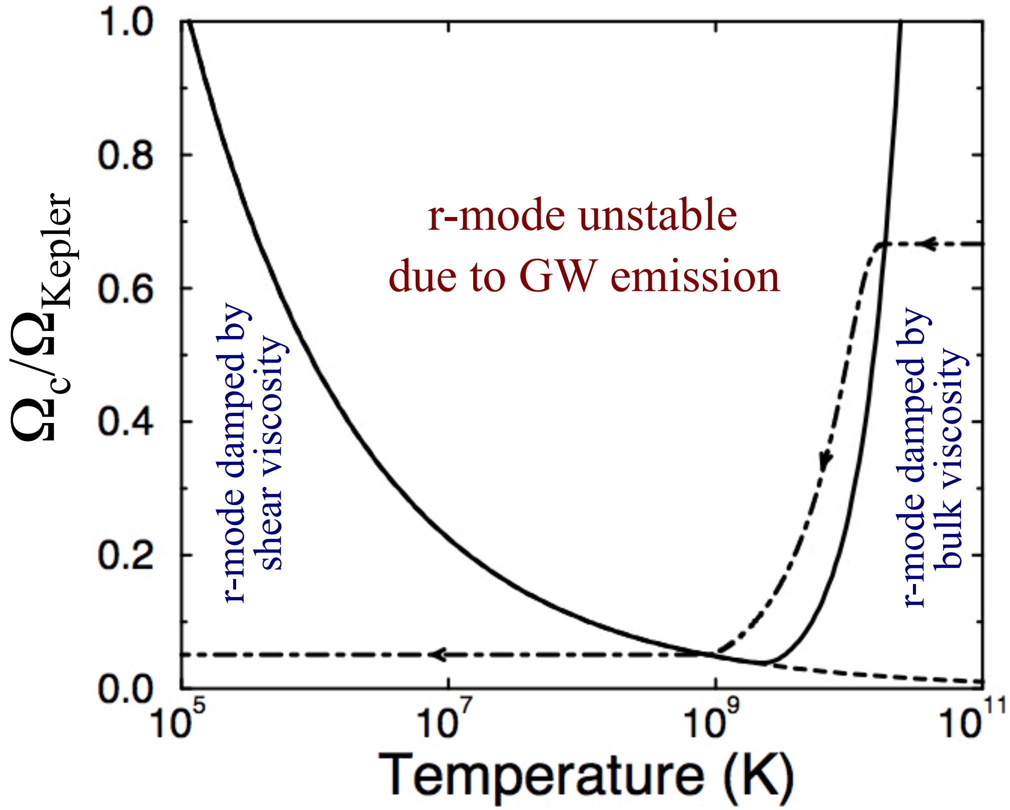
(IV & Tolós 2004)

Hyperons & the R-mode instability of Neutron Stars



The r-mode Instability

Ω_{Kepler} : Absolute Upper Limit of Rot. Freq.



Instabilities prevent NS to reach Ω_{Kepler}

r-mode Instability : toroidal mode of oscillation

✓ restoring force: Coriolis

✓ emission of GW in hot & rapidly rotating NS (CFS mechanism)

- GW makes the mode unstable
- Viscosity stabilizes the mode

$$A \propto A_0 e^{-i\omega(\Omega) - t/\tau(\Omega,T)}$$

$$\frac{1}{\tau(\Omega,T)} = -\frac{1}{\tau_{GW}(\Omega)} + \frac{1}{\tau_{Viscosity}(\Omega,T)}$$

Hyperon Bulk Viscosity ξ_Y

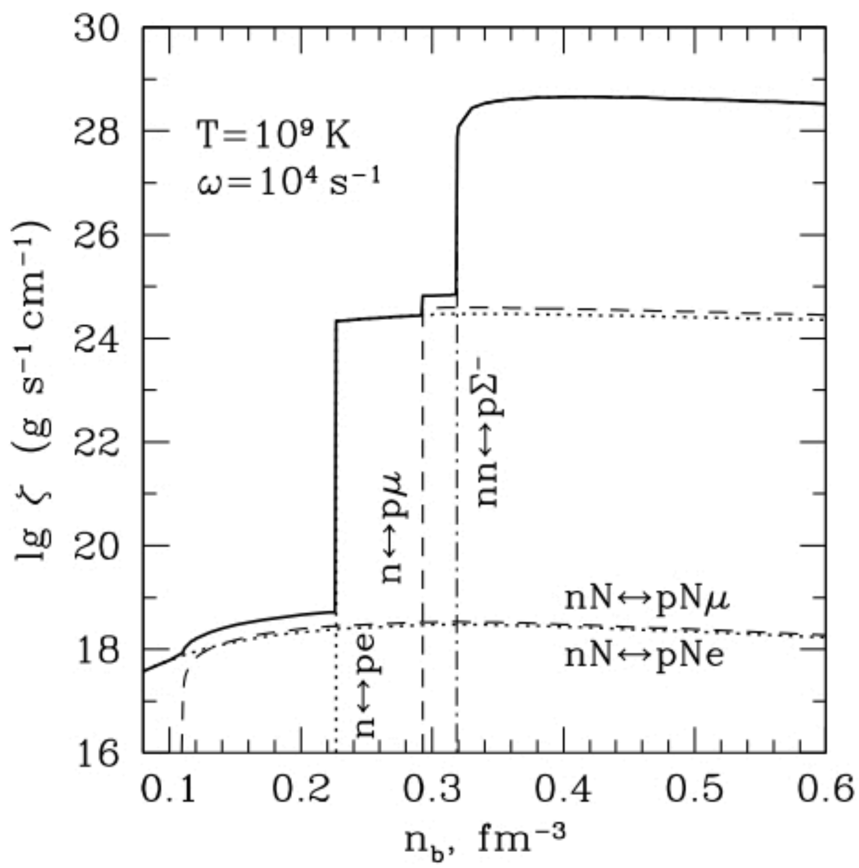
(Lindblom et al. 2002, Haensel et al 2002, van Dalen et al. 2002, Chatterjee et al. 2008, Gusakov et al. 2008, Shina et al. 2009, Jha et al. 2010,…)

Sources of ξ_Y:

non-leptonic weak reactions	$N + N \leftrightarrow N + Y$ $N + Y \leftrightarrow Y + Y$
Direct & Modified URCA	$Y \rightarrow B + l + \bar{\nu}_l$ $B' + Y \rightarrow B' + B + l + \bar{\nu}_l$
strong reactions	$N + Y \leftrightarrow N + Y$ $N + \Xi \leftrightarrow Y + Y$ $Y + Y \leftrightarrow Y + Y$

Reaction Rates & ξ_Y reduced by Hyperon Superfluidity

(Haensel, Levenfish & Yakovlev 2002)



Critical Angular Velocity of Neutron Stars

- r-mode amplitude: $A \propto A_o e^{-i\omega(\Omega)t - t/\tau(\Omega)}$

$$\frac{1}{\tau(\Omega,T)} = -\frac{1}{\tau_{GW}(\Omega)} + \frac{1}{\tau_{\xi}(\Omega,T)} + \frac{1}{\tau_{\eta}(T)}$$

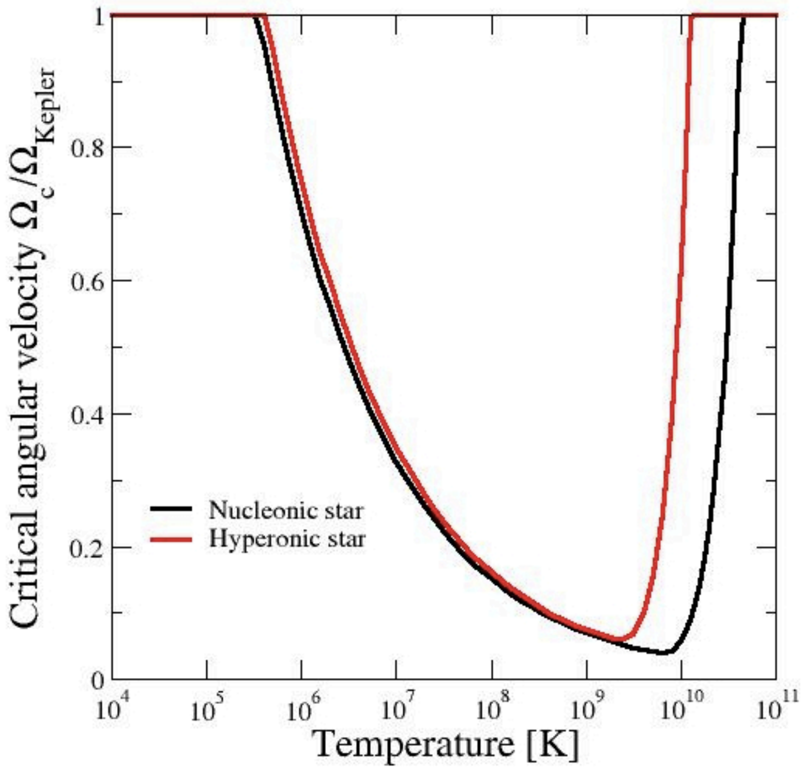
➔ $\frac{1}{\tau(\Omega_c,T)} = 0$ r-mode instability region

$\Omega < \Omega_c$ stable

$\Omega > \Omega_c$ unstable

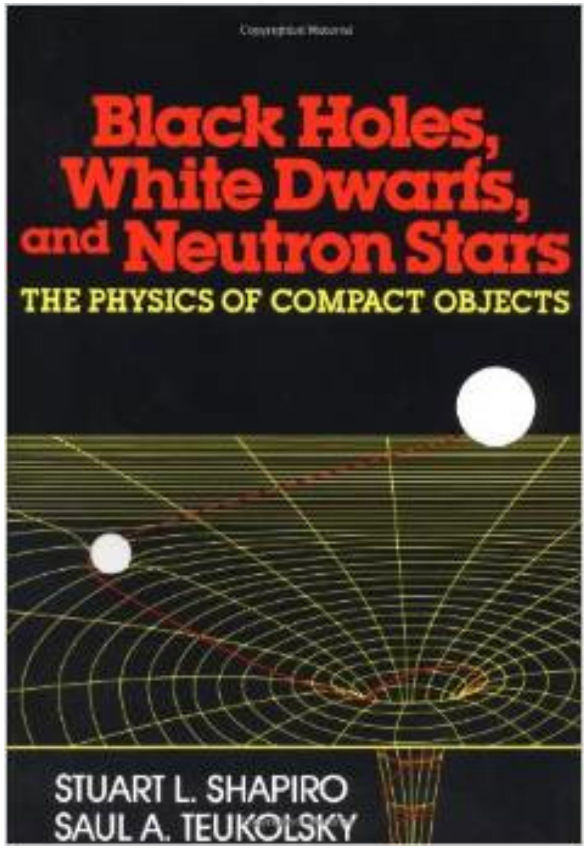
As expected:
smaller r-mode instability region
due to hyperons

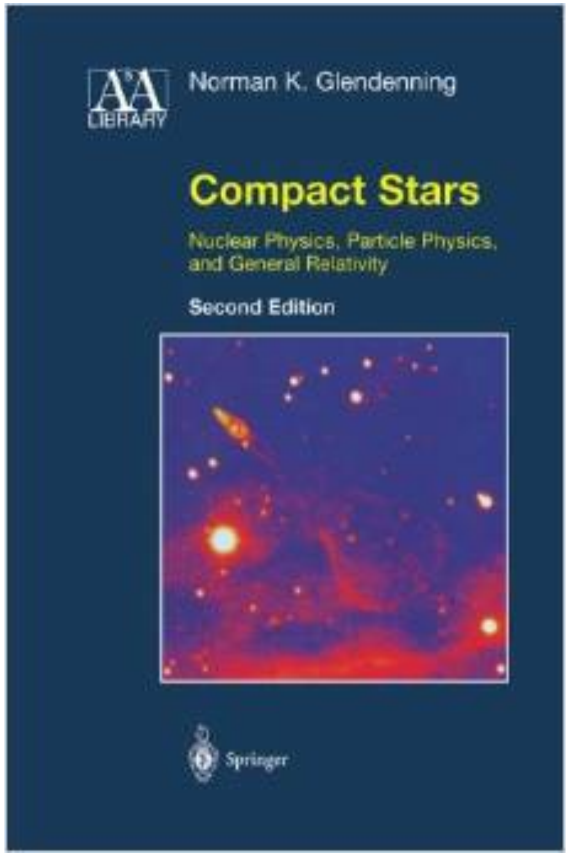
(I.V. & C. Albertus in preparation)

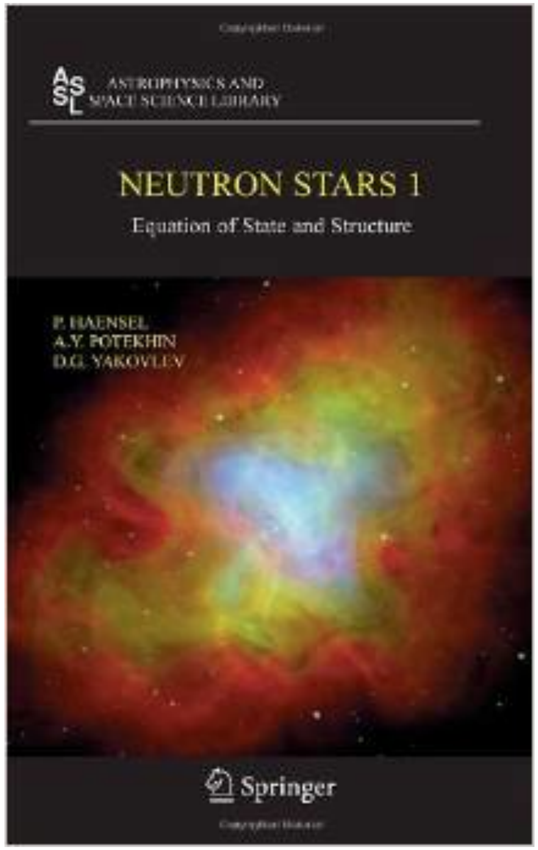


BHF: NN (Av18)+NY (NSC89)
(M=1.27$M_\odot$)

This short talk is just a brush-stroke on the physics of neutron stars. Three excellent monographs on this topic for interested readers are:







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- …

The final message of this talk

Neutron stars are excellent observatories to test fundamental properties of matter under extreme conditions and offer an interesting interplay between nuclear processes and astrophysical observables

- You for your time & attention
- The organizers for their invitation