Evolution Massive Stars and Core-collapse Supernova

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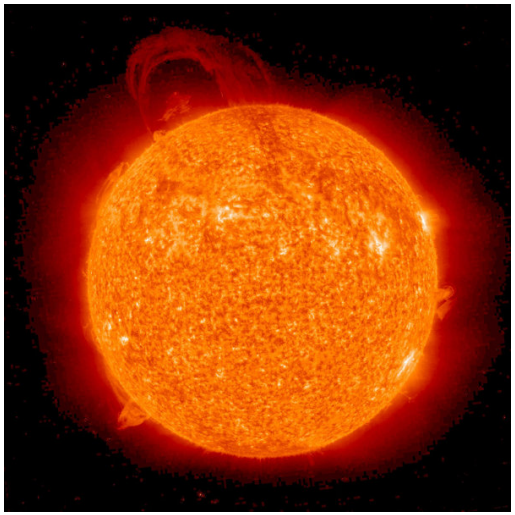
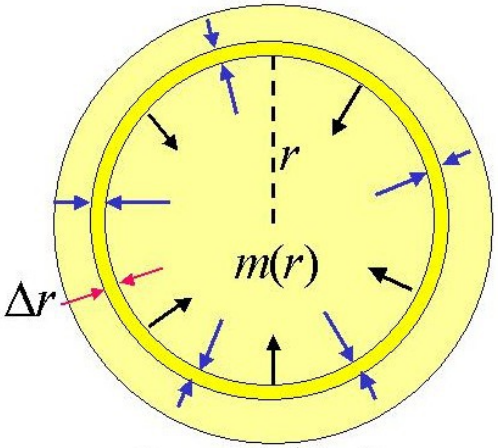
ELEMENTS

What is a star?





$\Delta m = (A\Delta r)\rho$

- A star is a self-luminous gaseous sphere.
- Stars produce energy by nuclear fusion reactions. A star is a self-regulated nuclear reactor.
- Gravitational collapse is balanced by pressure gradient: hydrostatic equilibrium.

$$dF_{\text{grav}} = -G\frac{mdm}{r^2} = [P(r+dr) - P(r)]dA = dF_{\text{pres}}$$

$$dm = 4\pi r^2 \rho dr$$

$$-G\frac{m\rho}{r^2} = \frac{dP}{dr}$$

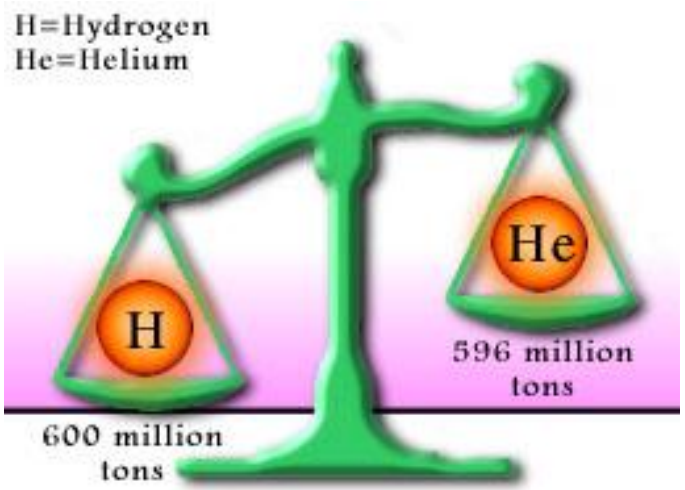
- Further equations needed to describe the transport of energy from the core to the surface, and the change of composition (nuclear reactions). Supplemented by an EoS: $P(\rho, T)$.

Where does the energy come from?

Energy comes from nuclear reactions in the core. Net effect is:

$$4\,^1\mathrm{H} \rightarrow {}^4\mathrm{He} + 2e^+ + 2\nu_e + 26.7\ \mathrm{MeV}$$

$$E = mc^2$$

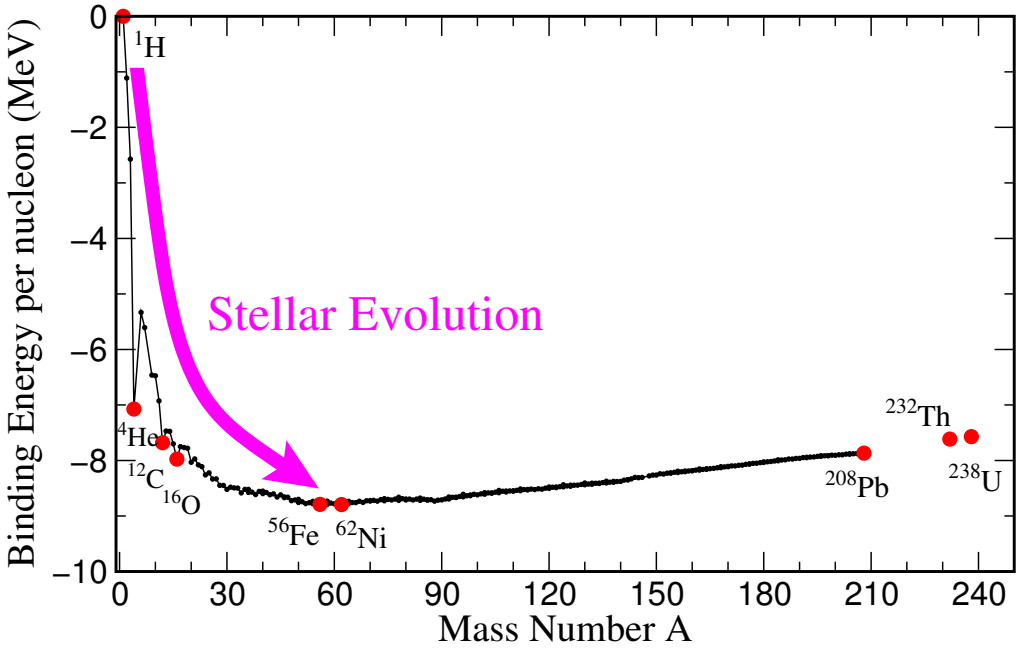


The Sun converts 600 million tons of hydrogen into 596 million tons of helium every second. The difference in mass is converted into energy. The Sun will continue burning hydrogen during 5 billions years.

Energy released by H-burning:
6.45×10^{18} erg g^{-1} = 6.7 MeV/nuc
Solar Luminosity: 3.85×10^{33} erg s^{-1}

Nuclear Binding Energy

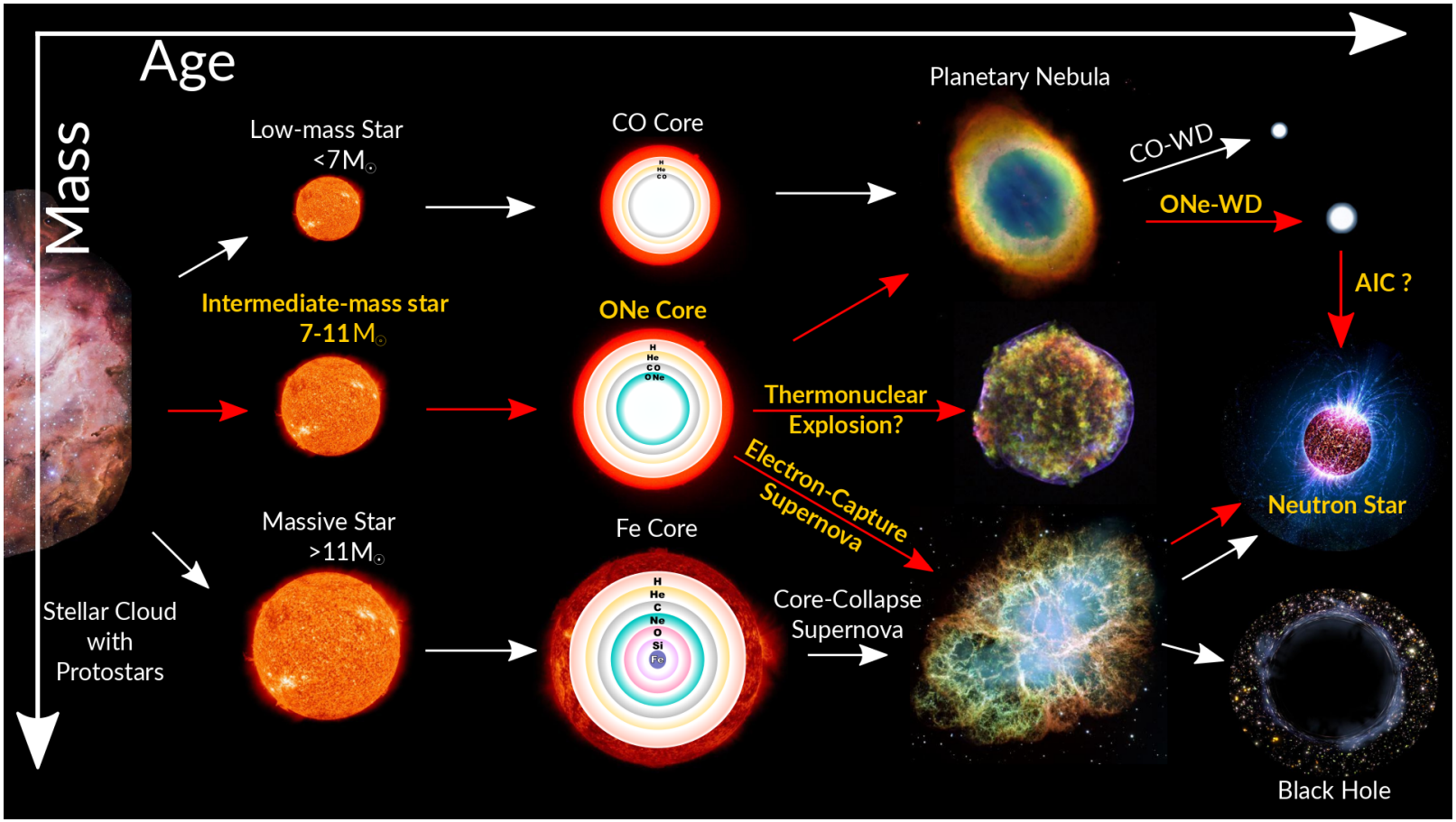
Liberated energy is due to the gain in nuclear binding energy.



- Rest mass energy is converted into kinetic energy of the particles. This translates into heat that raises the temperature unless energy transported away.
- During hydrostatic burning phases stars maintain an equilibrium: produced energy is transported and later emitted at the surface.

Stellar Evolution

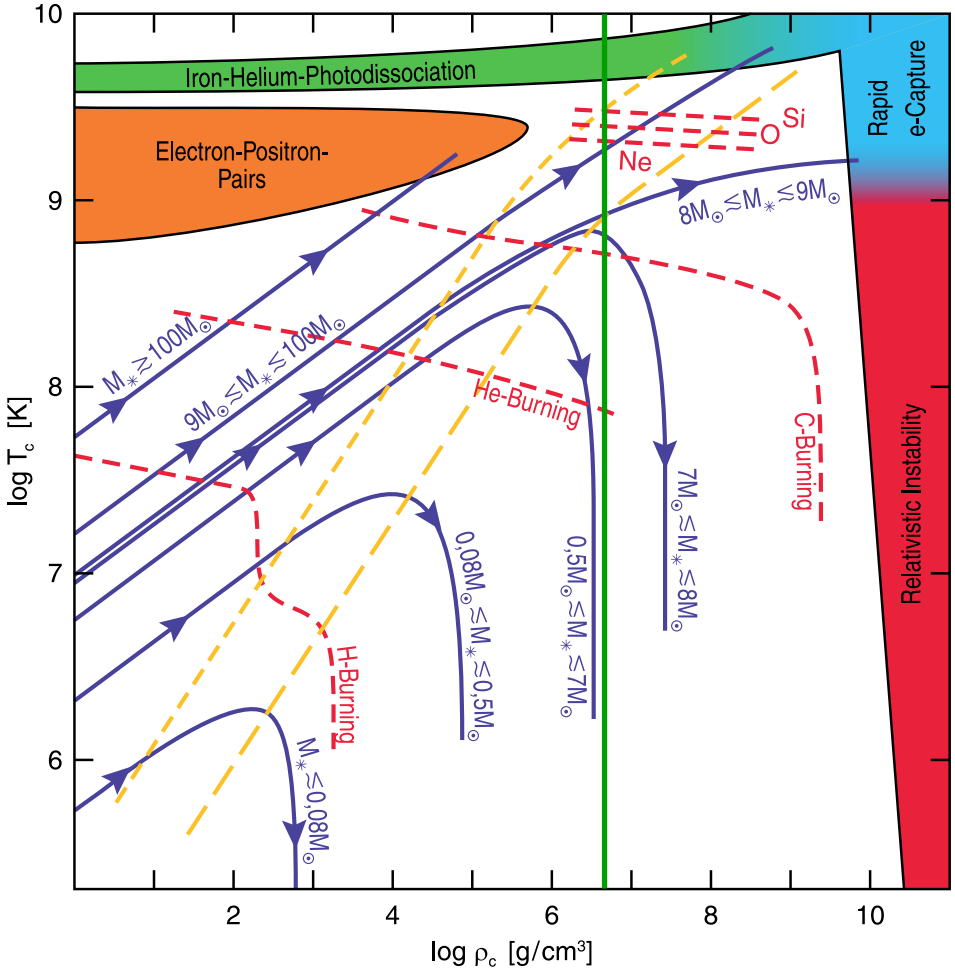
Star evolution, lifetime and death depends on mass.



Why such distinct outcomes?

Core evolution

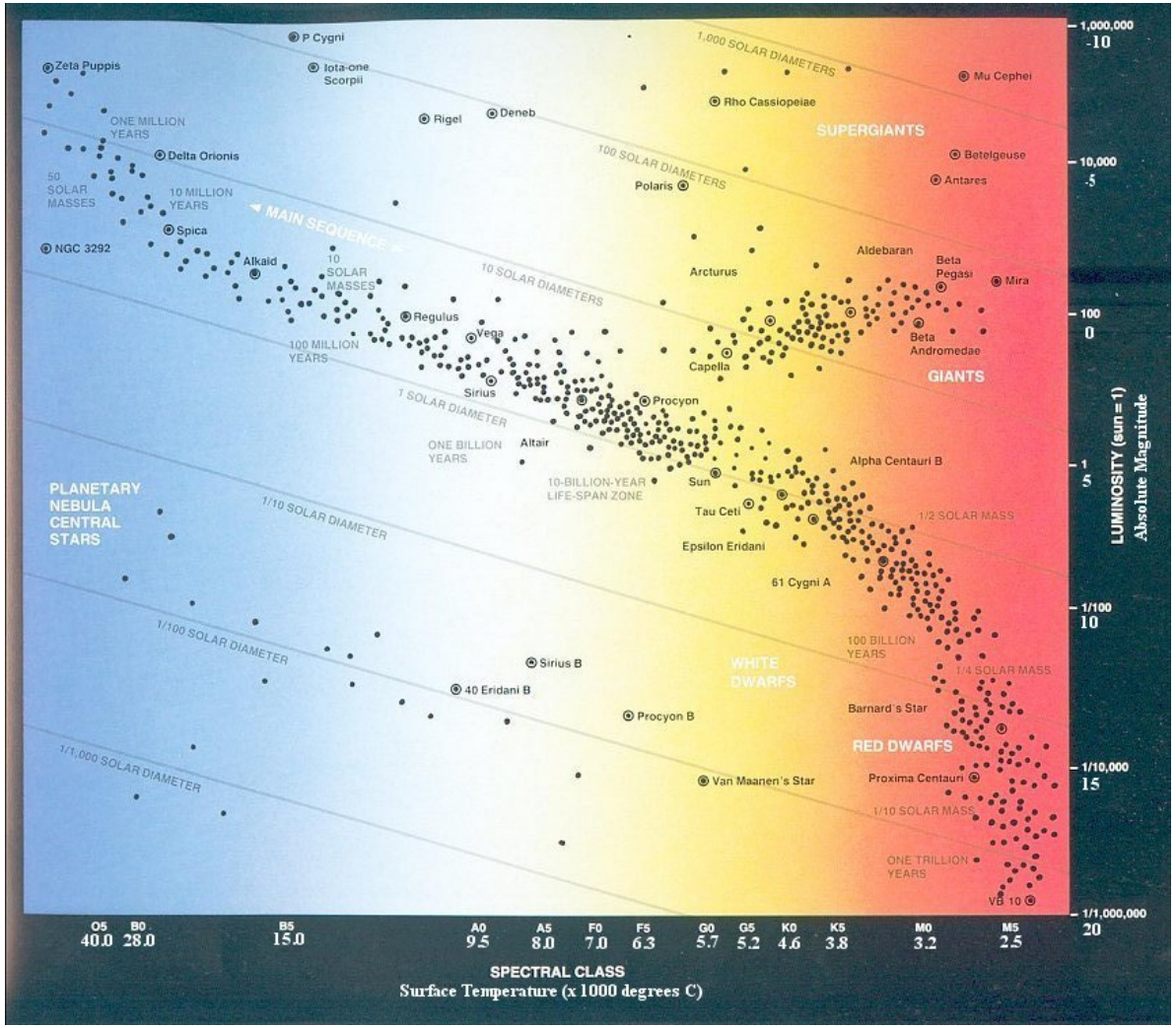
Hydrostatic equilibrium together with equation of state determines the evolution of the star core ($T_C \sim M^{2/3}\rho_c^{2/3}$)



Green: transition from relativistic to non-relativistic electrons.

from H.-T. Janka, Annu. Rev. Nucl. Part. Sci. **62**, 407 (2012).

Hertzspung-Russell diagram

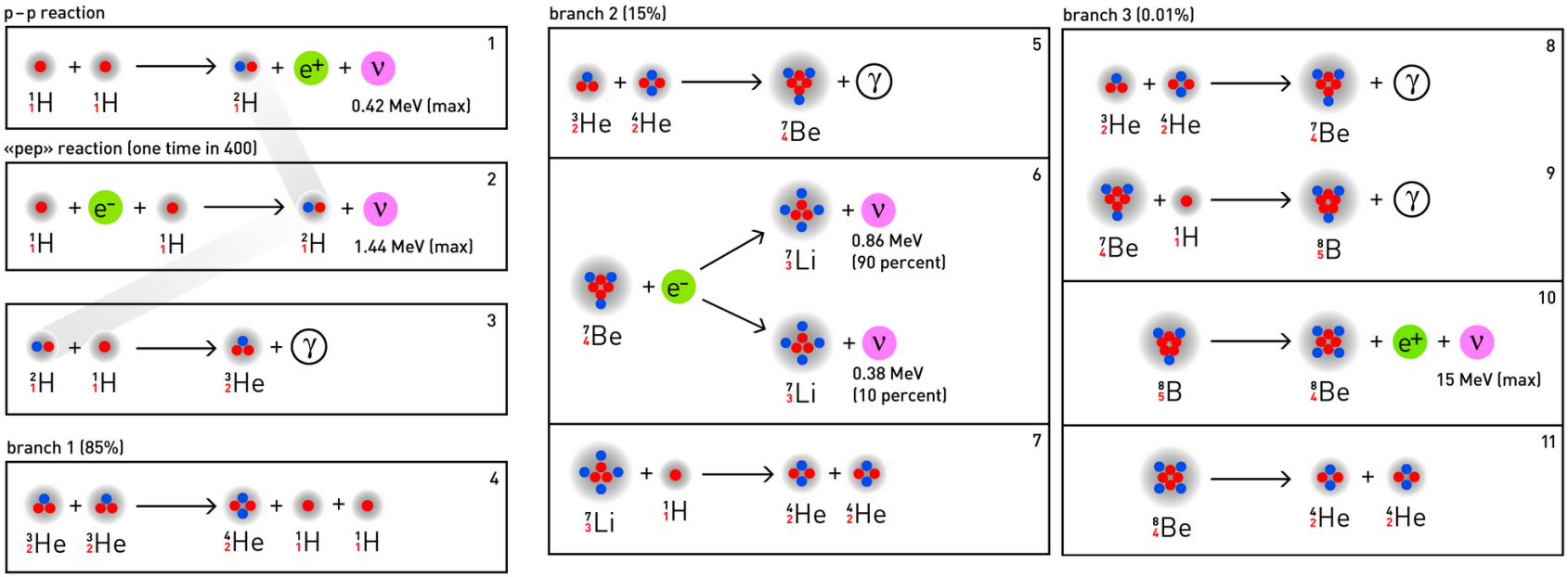


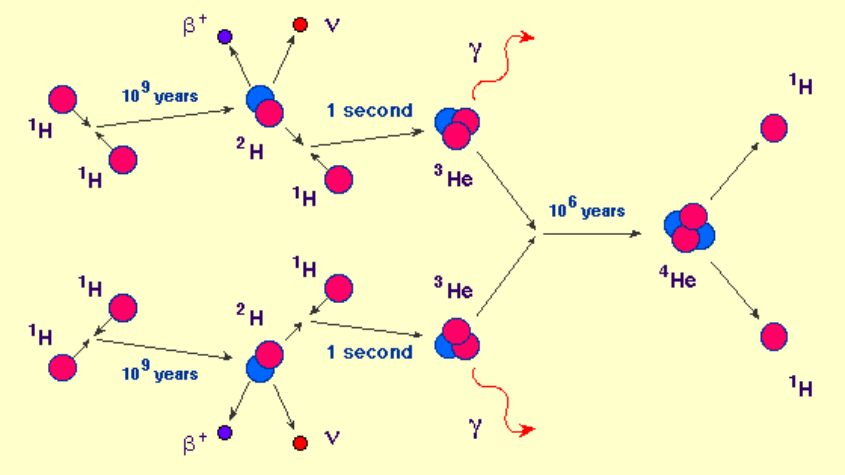
Assuming a star of radius R emitting like a blackbody with temperature T

$$L = 4\pi\sigma R^2 T^4$$

Hydrogen burning: pp chains

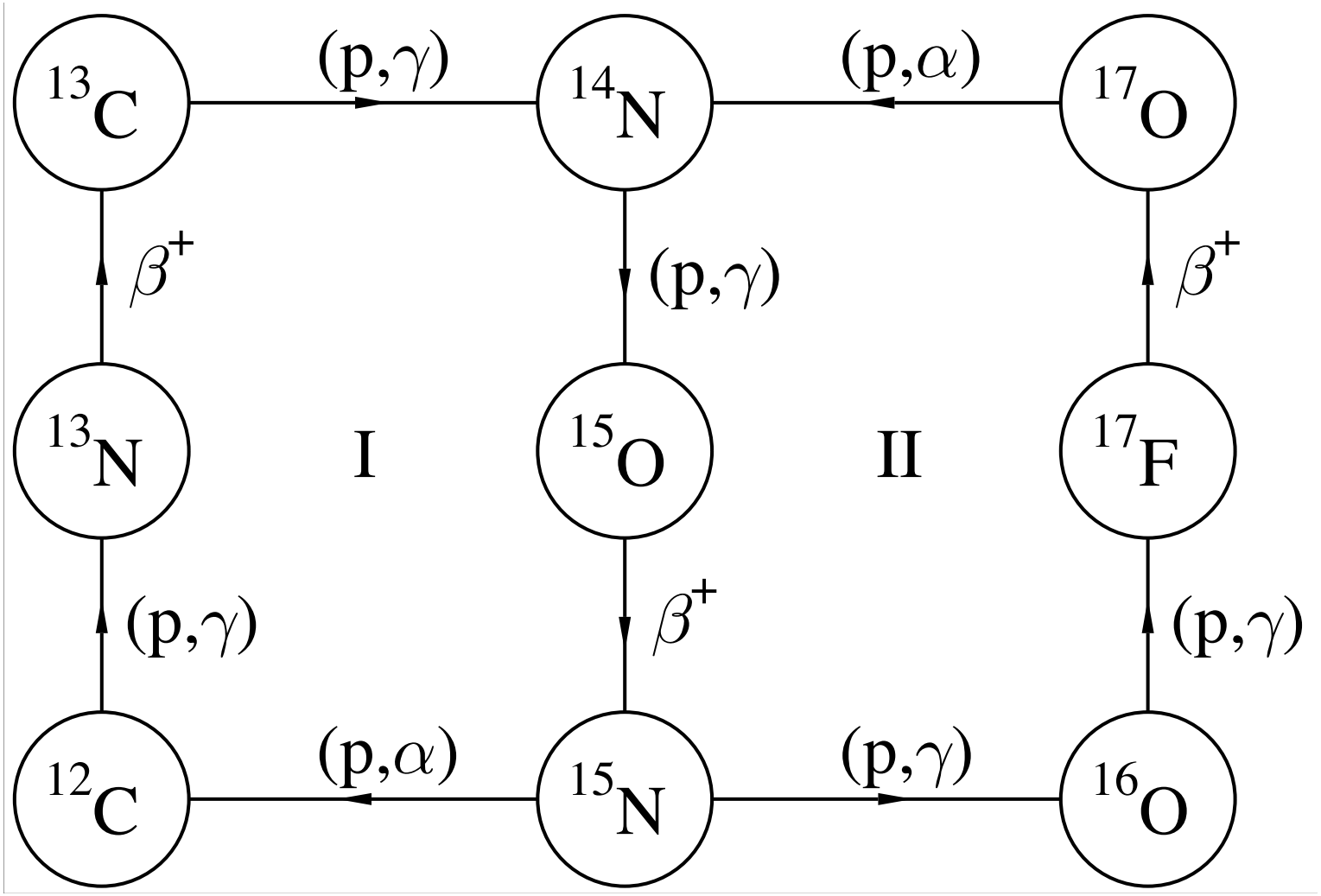
Once ^{4}He is produced can act as catalyst initializing the ppII and ppIII chains.



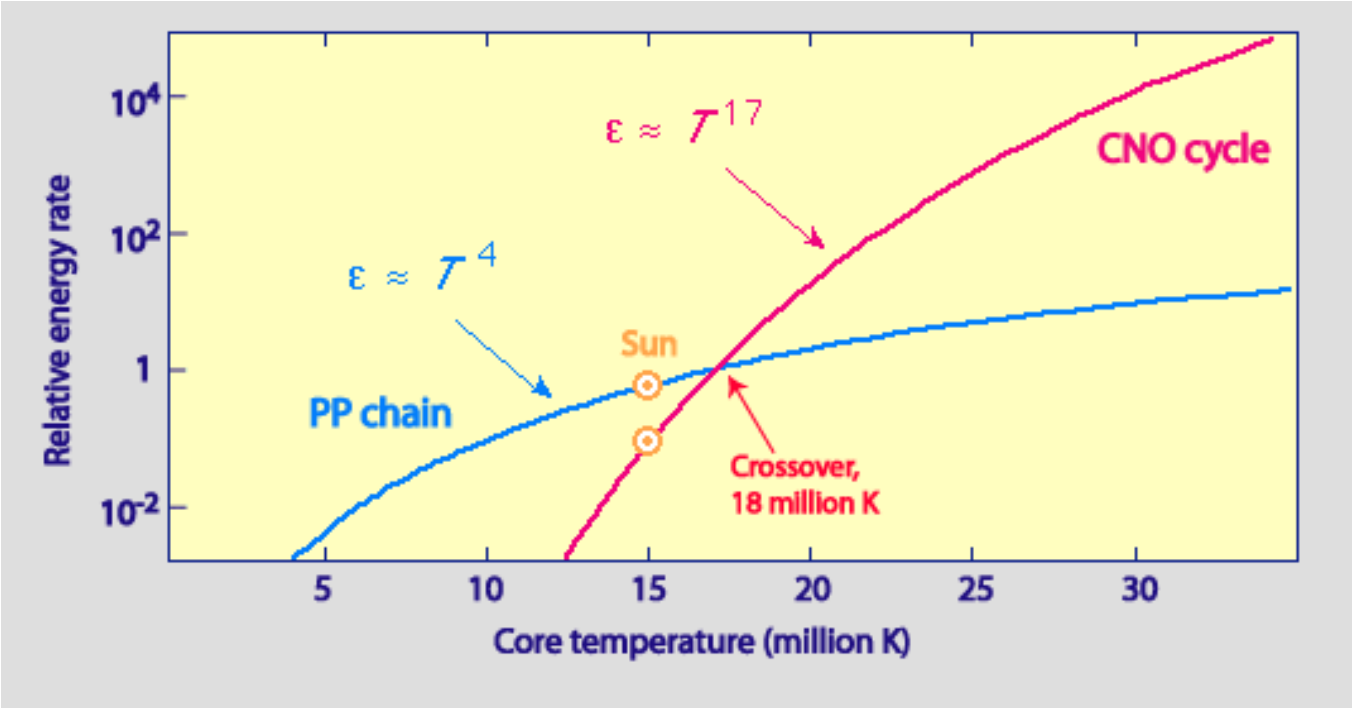


Hydrogen burning: CNO bicycle

Hydrogen can also be burning by the CNO cycle provided C, N, O previously available in the star.



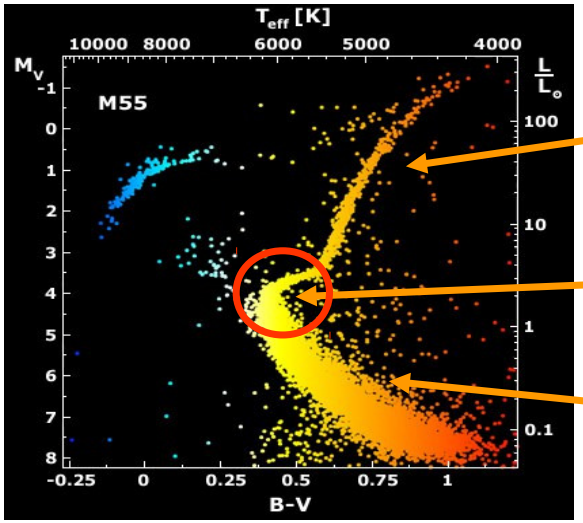
Energy generation: CNO cycle vs pp-chains



Consequences

- Stars slightly heavier than the Sun burn hydrogen via CNO cycle.
- CNO cycle goes significantly faster. Such stars have much shorter lifetimes

Mass ($M_\odot$)	lifetime (yr)
0.8	1.4×10^{10}
1.0	1×10^{10}
1.7	2.7×10^{9}
3.0	2.2×10^{8}
5.0	6×10^{7}
9.0	2×10^{7}
16.0	1×10^{7}
25.0	7×10^{6}
40.0	1×10^{6}



High-mass stars evolved onto the giant branch

Turn-off point

Low-mass stars still on the main sequence

Turn-off sensitive to $^{14}N(p,\gamma)$ rate.

Helium Burning

- Once hydrogen is exhausted the stellar core is made mainly of helium. Hydrogen burning continues in a shell surrounding the core.
- $^4\mathrm{He} + p$ produces $^5\mathrm{Li}$ that decays in 10^{-22} s.
- Helium survives in the core till the temperature become large enough ($T \approx 10^8$ K) to overcome the coulomb barrier for $^4\mathrm{He} + {}^4\mathrm{He}$. The produced $^8\mathrm{Be}$ decays in 10^{-16} s. However, the lifetime is large enough to allow the capture of another $^4\mathrm{He}$:

$$3\,{}^4\mathrm{He} \rightarrow {}^{12}\mathrm{C} + \gamma$$

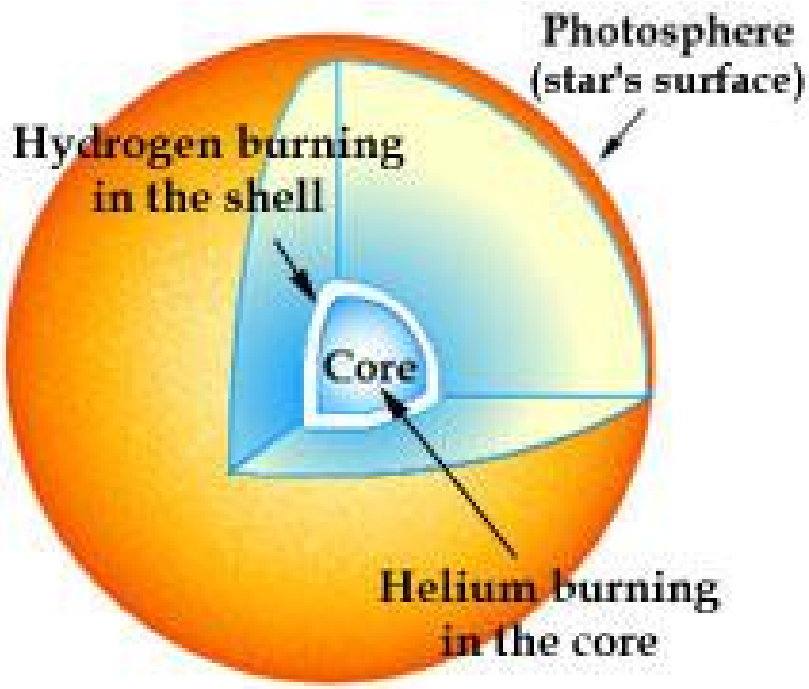
- Hoyle suggested that in order to account for the large abundance of Carbon and Oxygen, there should be a resonance in $^{12}\mathrm{C}$ that speeds up the production.
- $^{12}\mathrm{C}$ can react with another $^4\mathrm{He}$ producing $^{16}\mathrm{O}$

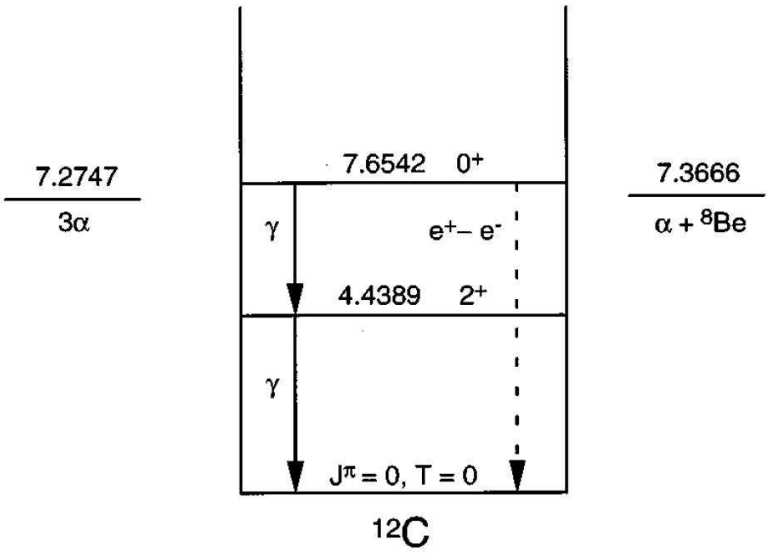
$$^{12}\mathrm{C} + \alpha \rightarrow {}^{16}\mathrm{O} + \gamma$$

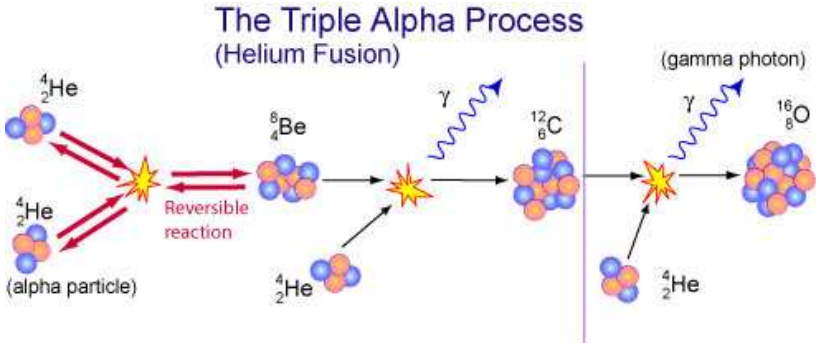
- These are the two main reactions during helium burning.
- Beginning with helium burning whenever a new burning process starts in the core previous burning processes continue in the surrounding shells

Hoyle State and tripple α reaction

Red giant structure

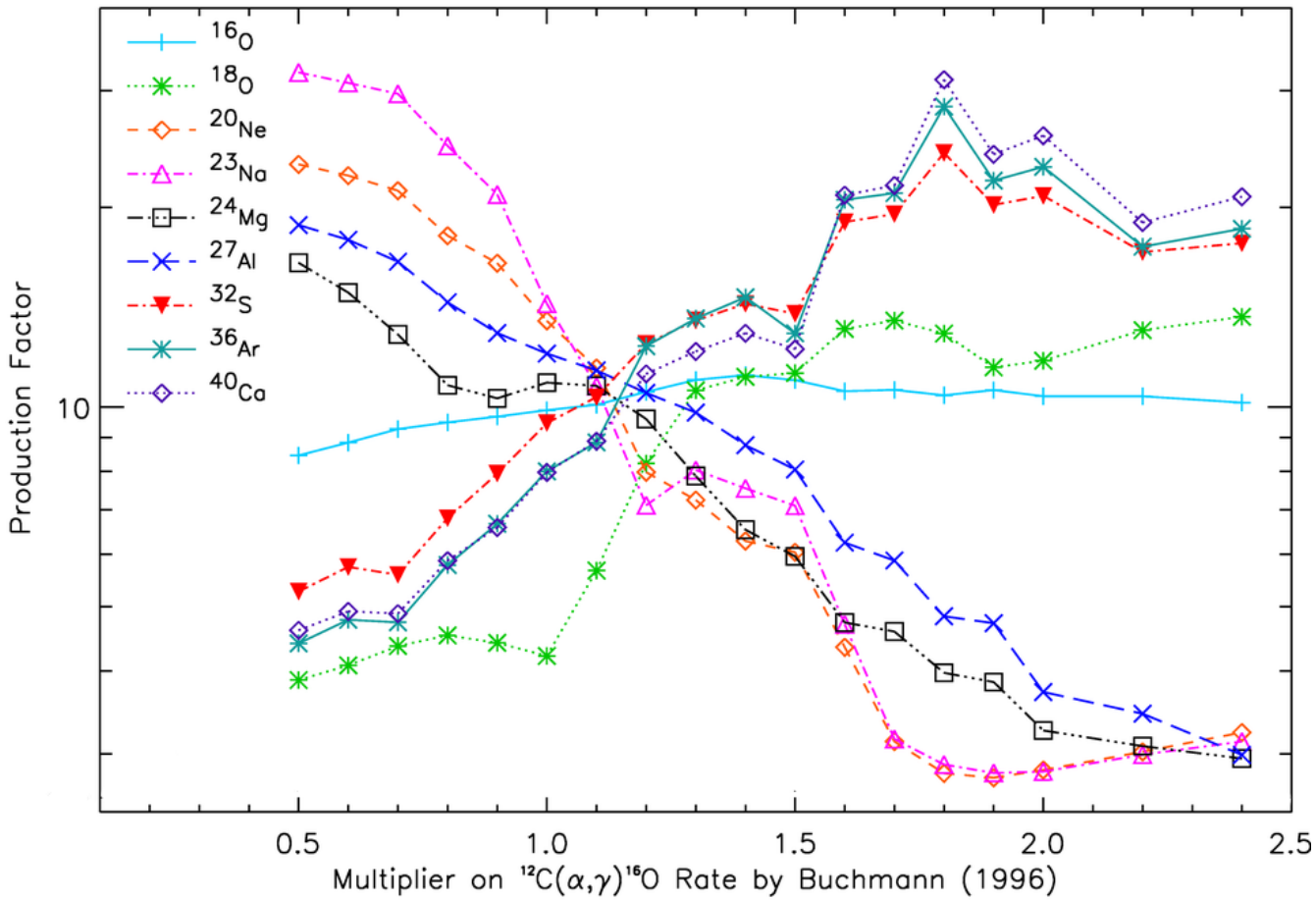






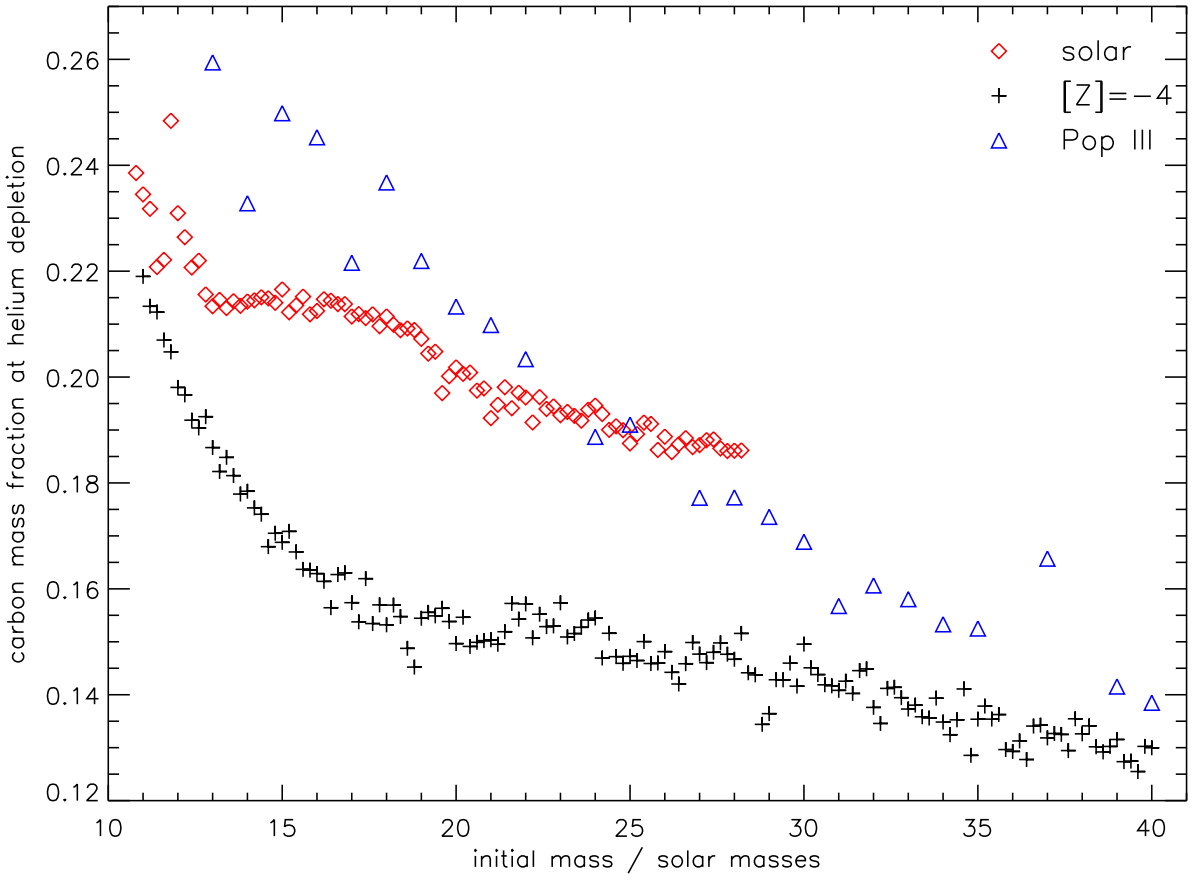
Influence $^{12}C(\alpha,\gamma)$

The $^{12}C(\alpha,\gamma)$ reaction plays a fundamental role as it determines the C/O ratio of the core, the posterior evolution and the nucleosynthesis



Influence $^{12}C(\alpha,\gamma)$

Central carbon abundance at the end of helium burning.



Woosley, Heger & Weaver, RMP **74**, 1015 (2002)

End of helium burning

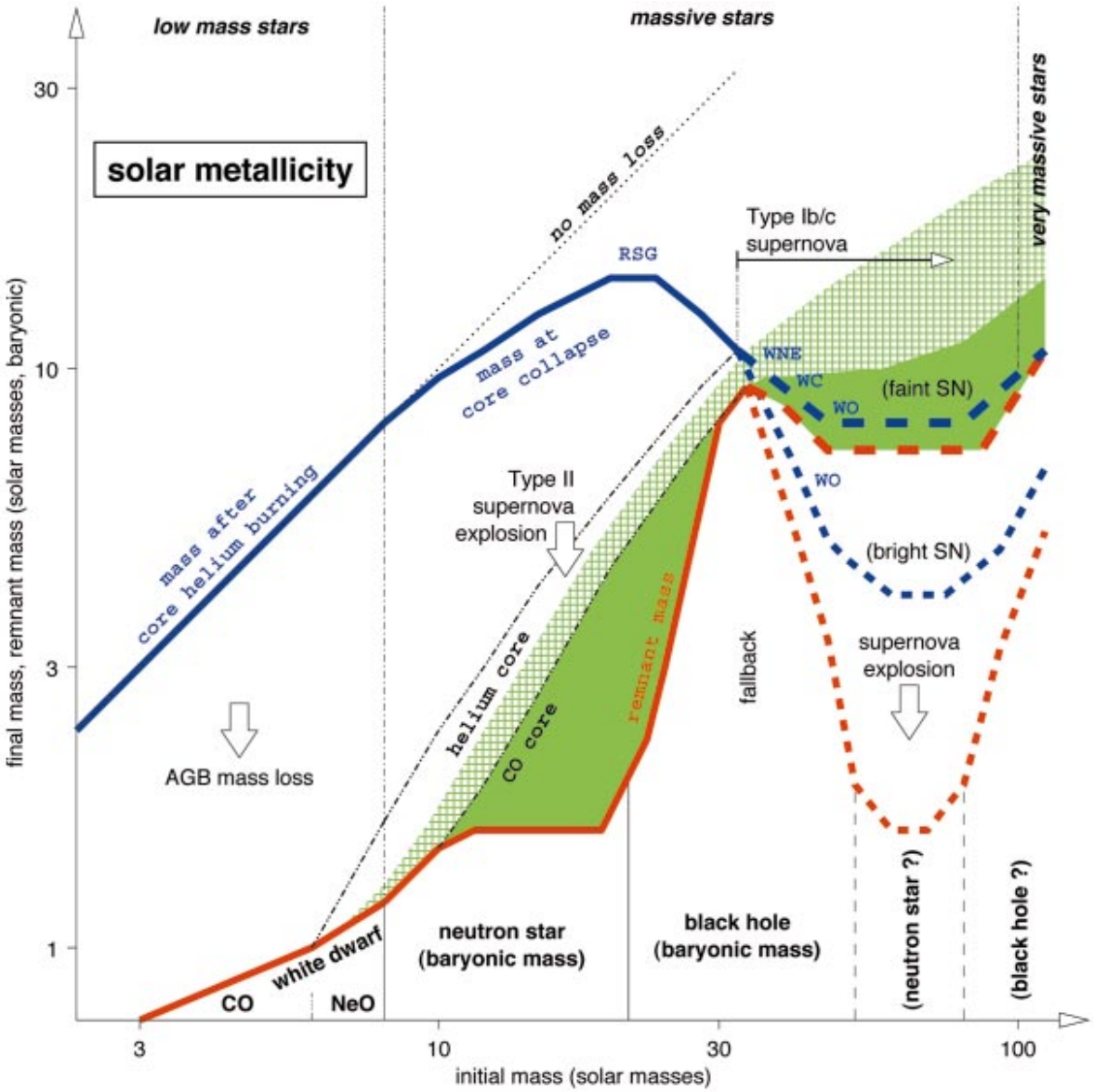
Nucleosynthesis yields from stars depend on mass:

stars with $M \lesssim 8\,\mathrm{M}_\odot$ These stars eject their envelopes during helium shell burning producing planetary nebula and white dwarfs. Constitute the site for the s process.

stars with $8\,\mathrm{M}_\odot \lesssim M \lesssim 12\,\mathrm{M}_\odot$ The end products of these stars is still under discussion and may depend on metallicity. Potentially they can produce ONe white-dwarfs (planetary nebula), thermonuclear supernova, and/or electron capture supernova.

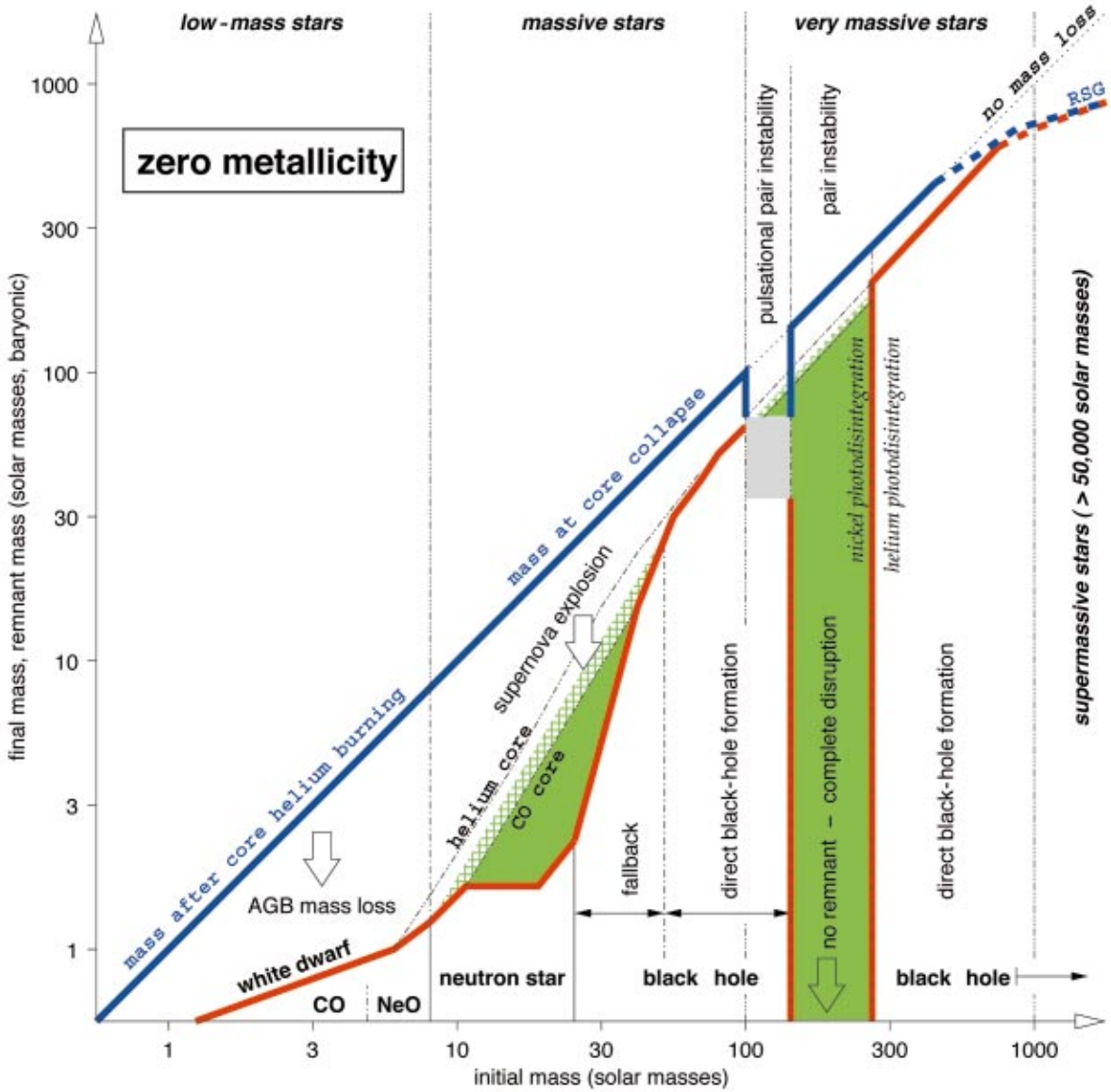
stars with $M \gtrsim 12\,\mathrm{M}_\odot$ These stars will ignite carbon burning under non-degenerate conditions. The subsequent evolution proceeds in most cases to core collapse. These stars make the bulk of newly processed matter that is returned to the interstellar medium. The end product is a neutron star or black hole depending on the mass and metalicity of the star.

End products stellar evolution (solar metallicity)



Woosley, Heger & Weaver, RMP 76, 1015 (2002)

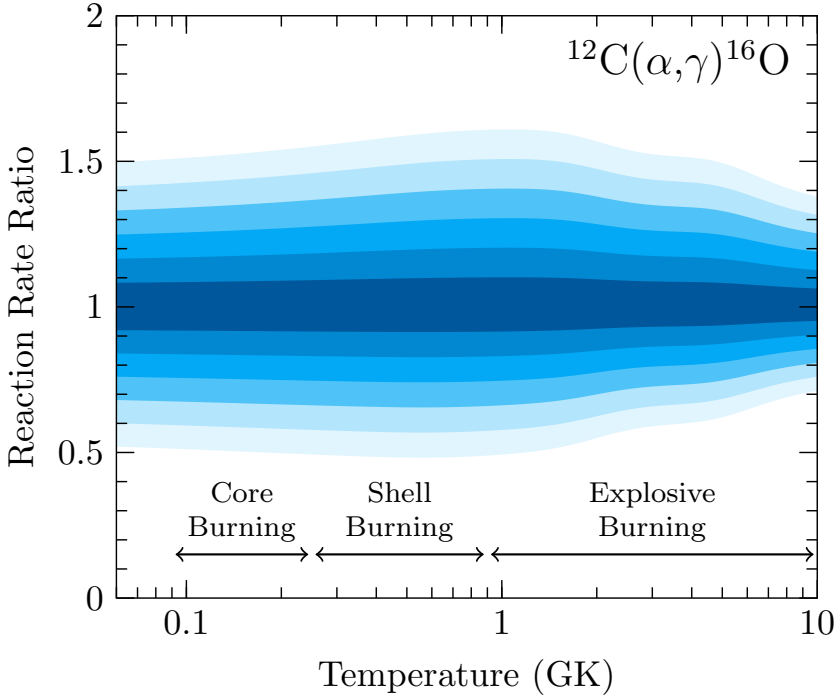
End products stellar evolution (zero metallicity)

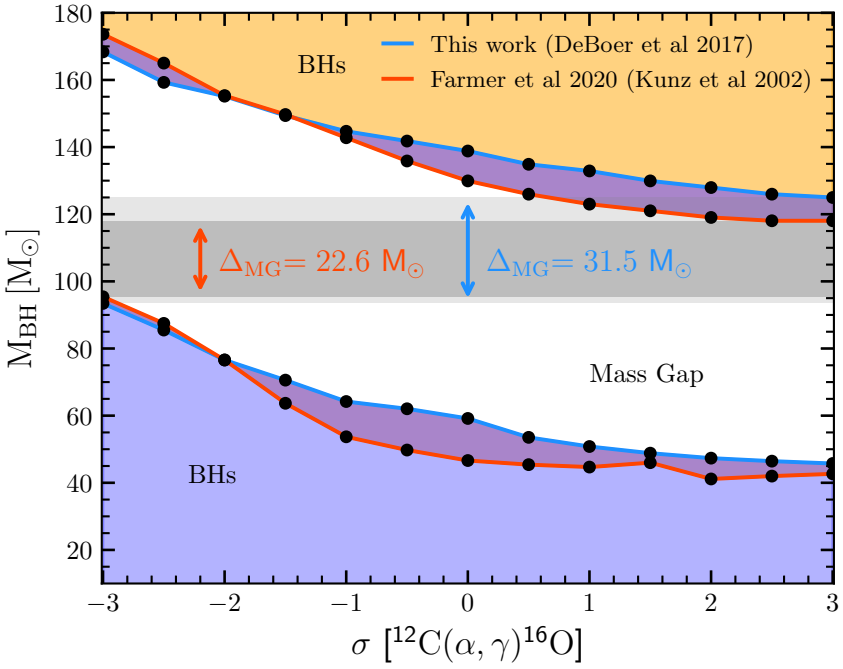


Woosley, Heger & Weaver, RMP **74**, 1015 (2002)

Mass gap dependence on $^{12}C(\alpha,\gamma)$

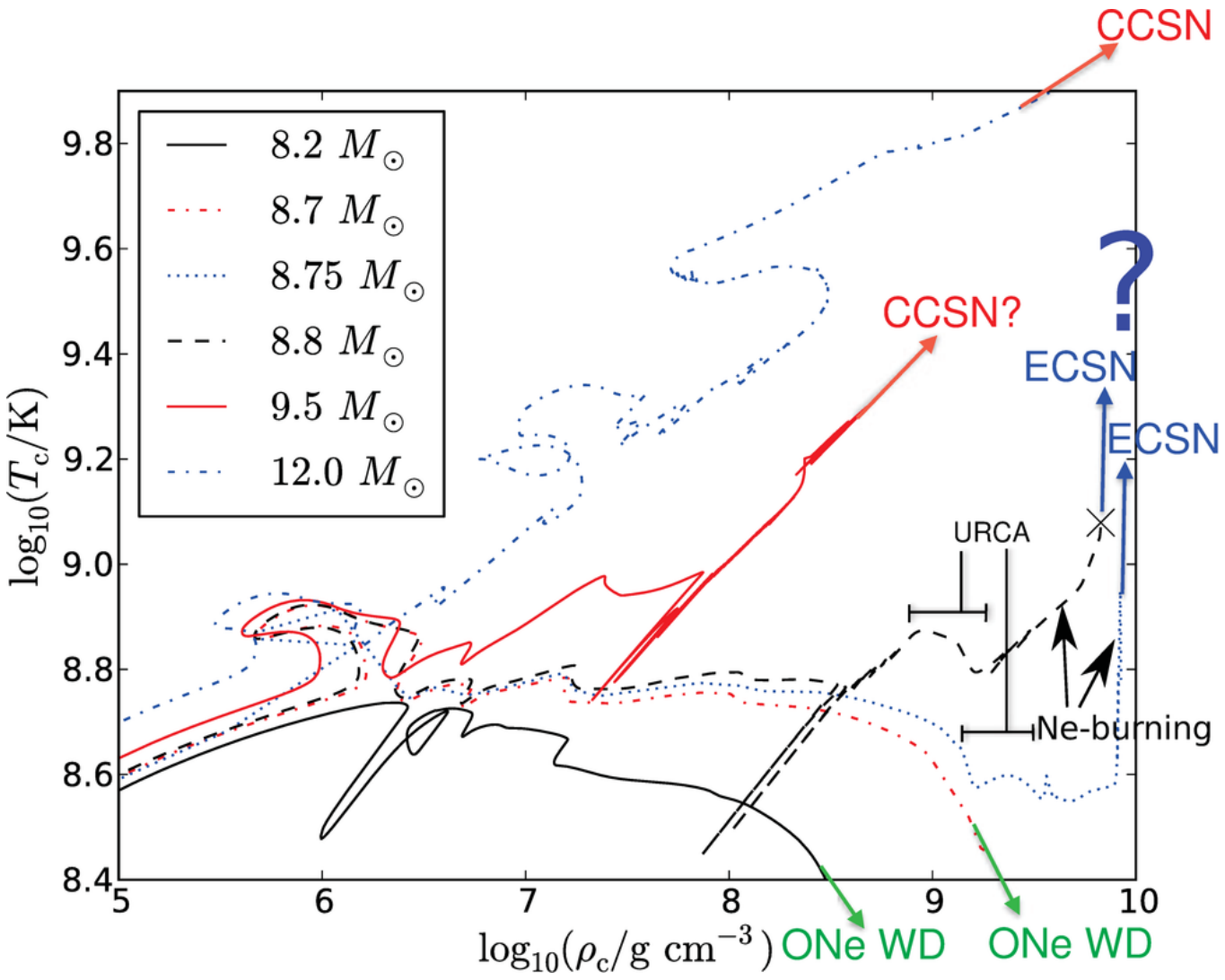
The location of the mass gap vs uncertainty in $^{12}C(\alpha,\gamma)$





Mehta *et al*, arXiv:2105.06366

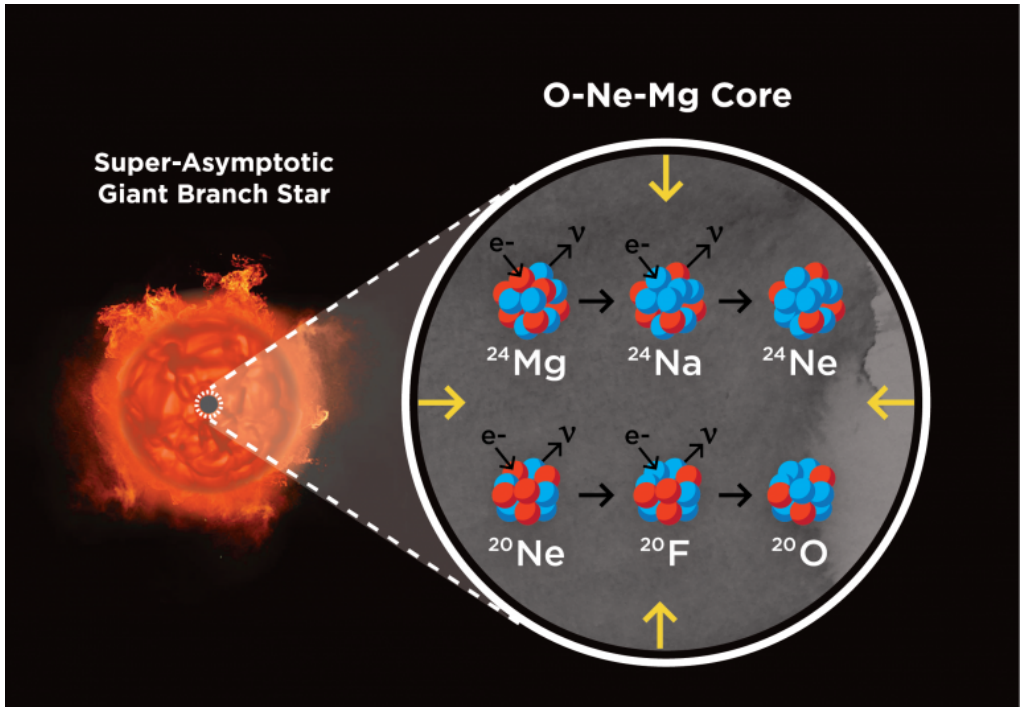
Core evolution (intermediate stars)



Jones *et al.*, ApJ **772**, 150 (2013)

Evolution intermediate mass stars

- After the formation of an ONe core, shell burning keeps adding mass to the core. As the density increases electron capture processes become possible.
- At some point Oxygen burning will start.
- There is a critical density such as for lower values burning leads to a thermonuclear supernova while for higher values the end product is a neutron star and a core-collapse supernova.



Urca processes

Consider a nucleus (Z, A) which β decays:

$$(Z, A) \rightarrow (Z + 1, A) + e^- + \bar{\nu}_e$$

In the stellar environment electrons get degenerate with increasing density. The electron energy, $\epsilon_F \propto \rho^{1/3}$, can be high enough to enforce:

$$(Z + 1, A) + e^- \rightarrow (Z, A) + \nu_e$$

An Urca cycle can take place producing a ν_e and $\bar{\nu}_e$ that carry away energy. The cycle operates without changing composition and contributing very effectively to the cooling.

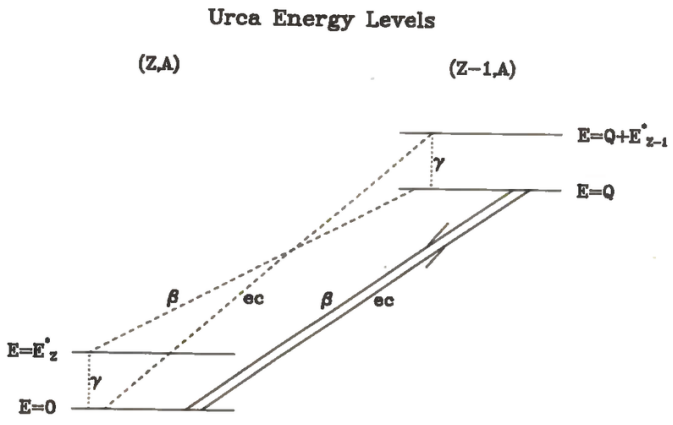
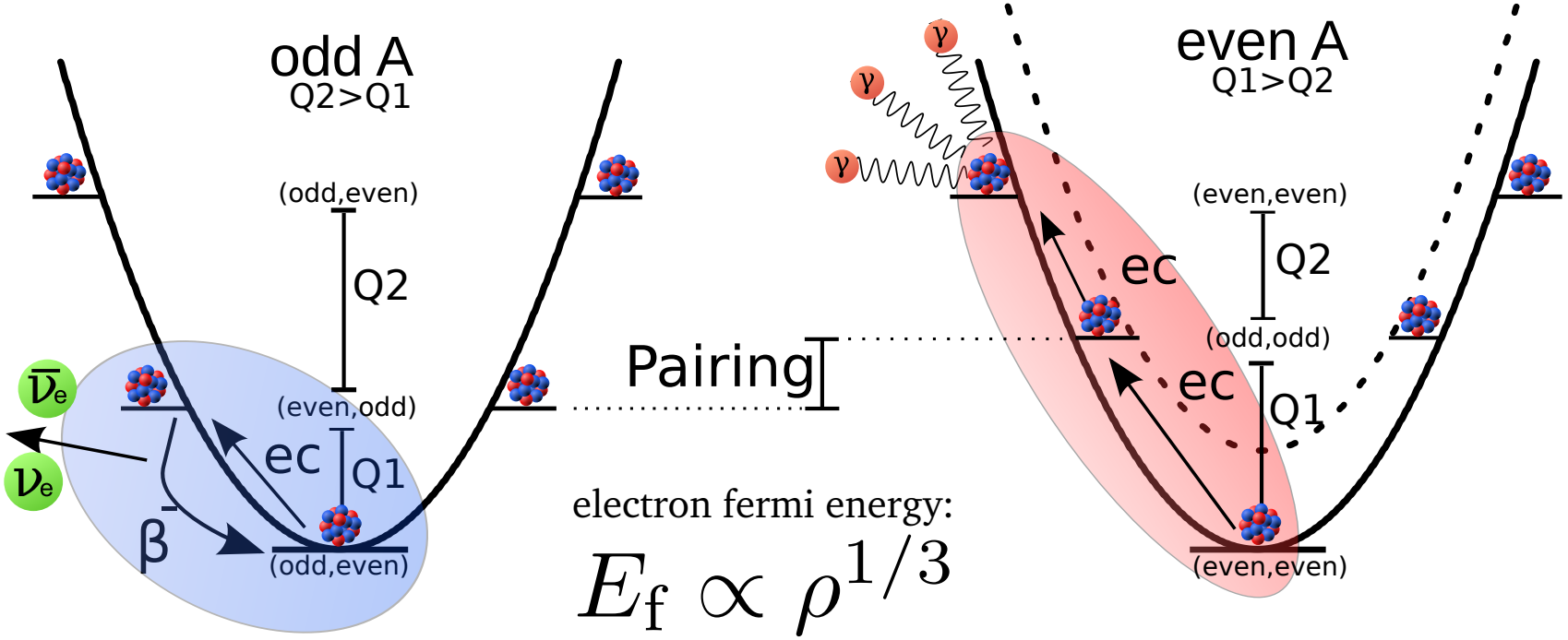


Fig. 11.3. Urca Energy Levels

Urca pairs: cooling vs heating

mass parabola for isobaric chain



electron fermi energy:

$$E_{\mathrm{f}} \propto \rho^{1/3}$$

Urca cooling

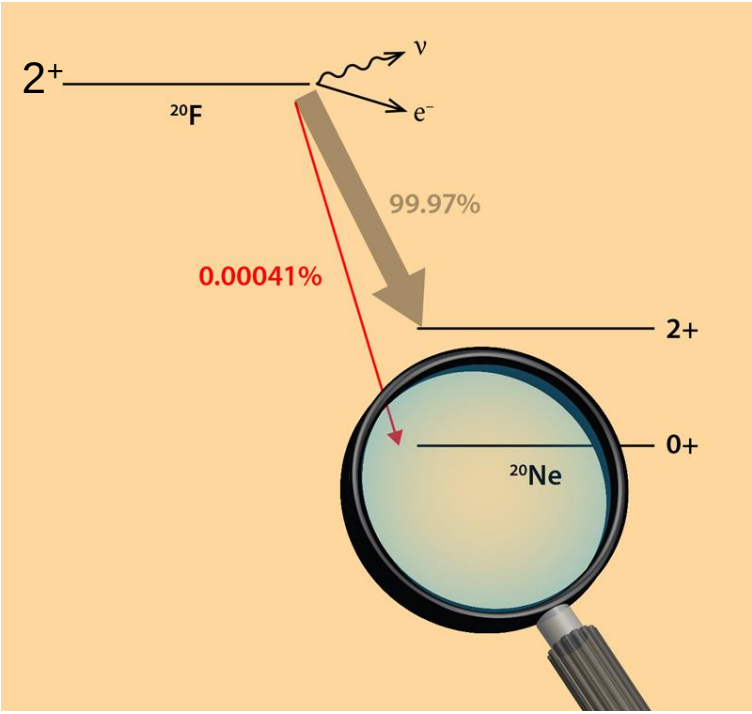
$$(\mathrm{A,Z}) + e^- = (\mathrm{A,Z\text{-}1}) + \nu_e$$
$$(\mathrm{A,Z\text{-}1}) = (\mathrm{A,Z}) + e^- + \bar{\nu}_e$$

heating

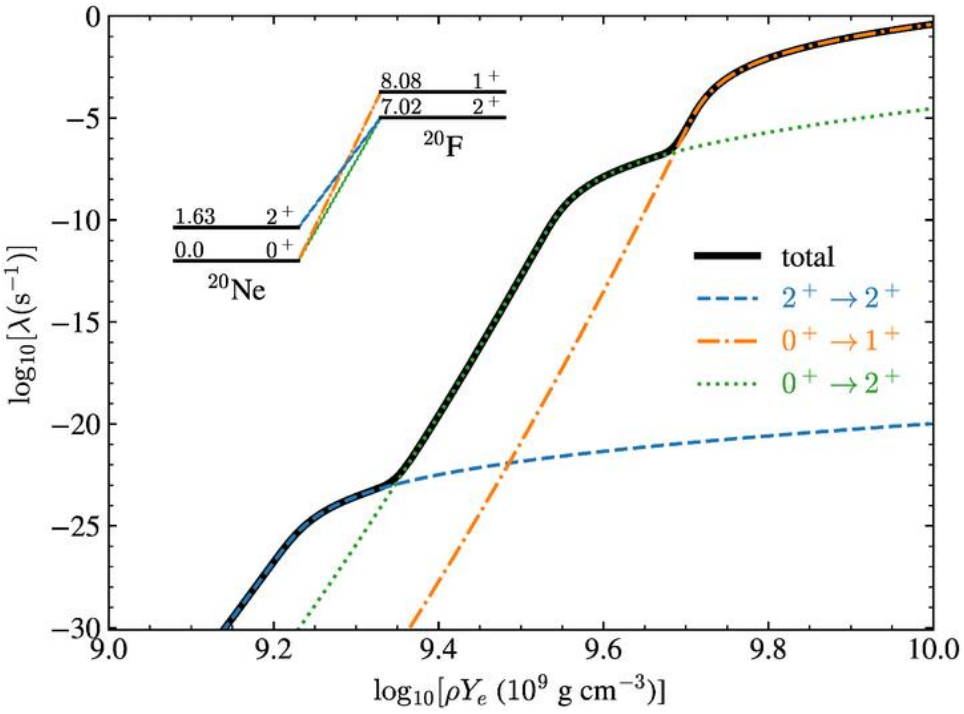
$$(\mathrm{A,Z}) + e^- = (\mathrm{A,Z\text{-}1}) + \nu_e$$
$$(\mathrm{A,Z\text{-}1}) + e^- = (\mathrm{A,Z\text{-}2})^* + \nu_e + \gamma$$

Intermediate mass stars

- Core consists mainly of ^{16}O and ^{20}Ne. Accretion of mass into the core increases the density and the energy of the electrons $\varepsilon \sim \rho^{1/3}$ reaching several MeV. Electrons are captured by ^{20}Ne.
- Capture rate determined by a transition between ground states of ^{20}F and ^{20}Ne



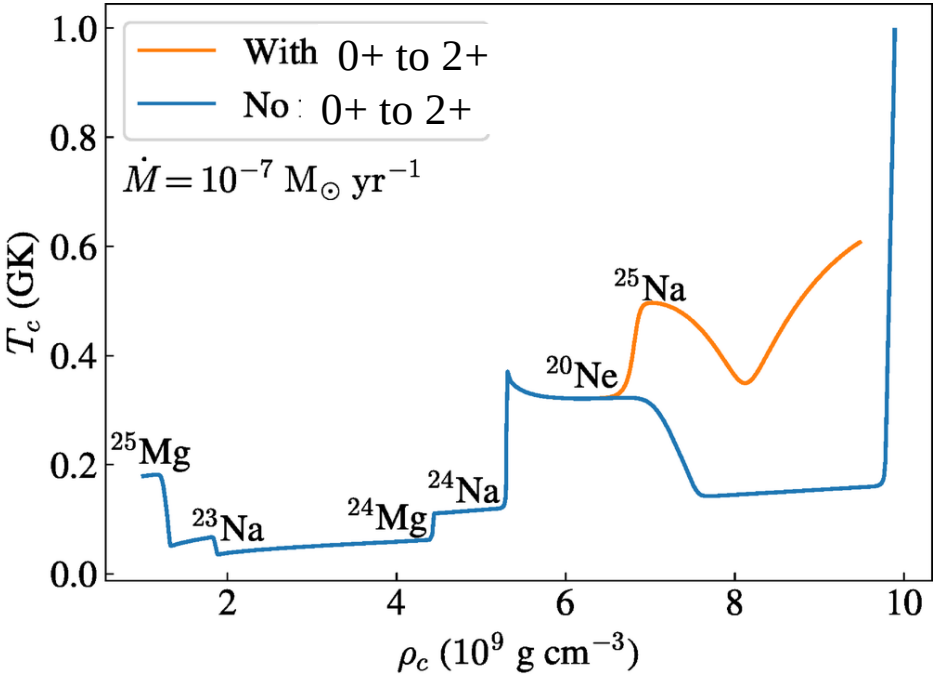
https://physics.aps.org/articles/v12/151

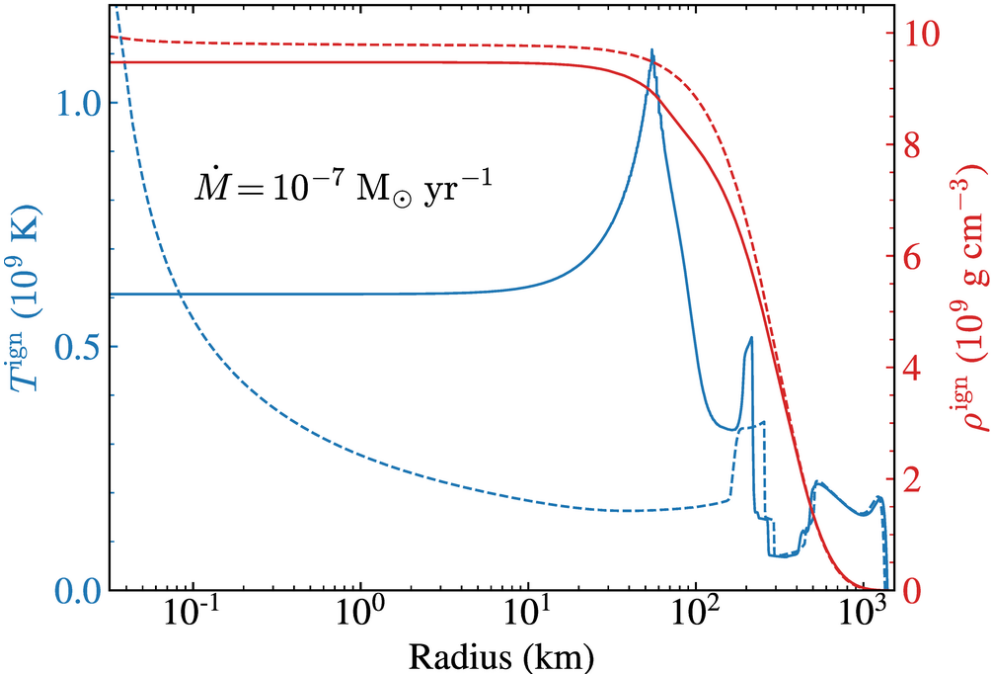


Kirsebom, et al, PRL123, 262701 (2019)

Impact

- With new rate Oxygen ignition occurs at lower densities and liberates enough energy to unbound part of the star producing a thermonuclear supernova and a low mass ONeFe white-dwarft instead of a low mass neutron star.





Kirsebom, et al, PRL 123, 262701 (2019), PRC 100, 065805 (2019)
See Hiramatsu et al, Nature Astronomy 5, 903 (2021) for a ECSN candidate

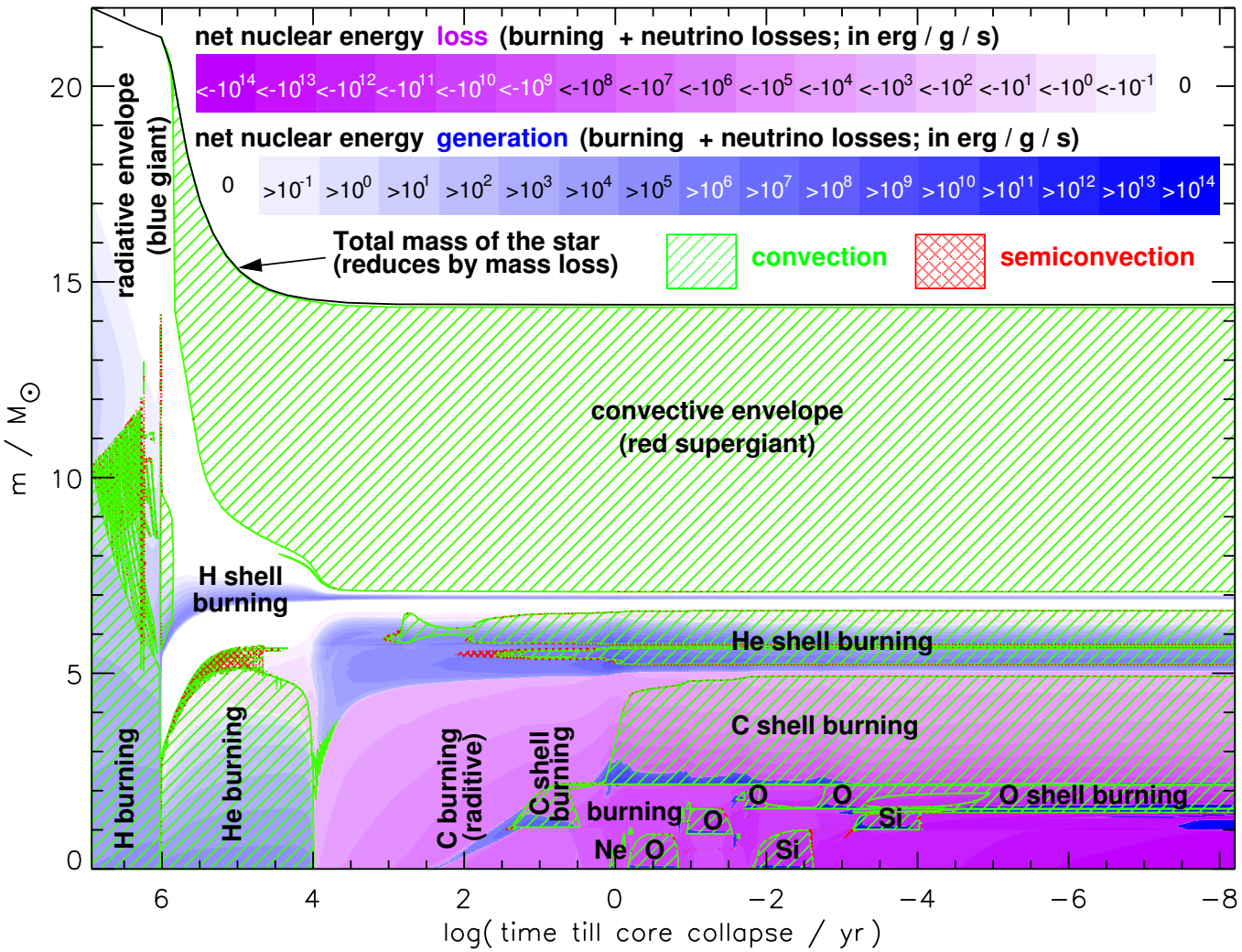
Stellar life (Massive star)

Nuclear burning stages

(e.g., 20 solar mass star)

Fuel	Main Product	Secondary Product	T (10^9 K)	Time (yr)	Main Reaction
H	He	^{14}N	0.02	10^7	$4\,H \xrightarrow{CNO} {}^4He$
He	O, C	^{18}O, ^{22}Ne s-process	0.2	10^6	$3\,He^4 \rightarrow {}^{12}C$; ${}^{12}C(\alpha,\gamma){}^{16}O$
C	Ne, Mg	Na	0.8	10^3	${}^{12}C + {}^{12}C$
Ne	O, Mg	Al, P	1.5	3	${}^{20}Ne(\gamma,\alpha){}^{16}O$; ${}^{20}Ne(\alpha,\gamma){}^{24}Mg$
O	Si, S	Cl, Ar, K, Ca	2.0	0.8	${}^{16}O + {}^{16}O$
Si	Fe	Ti, V, Cr, Mn, Co, Ni	3.5	0.02	${}^{28}Si(\gamma,\alpha)\ldots$

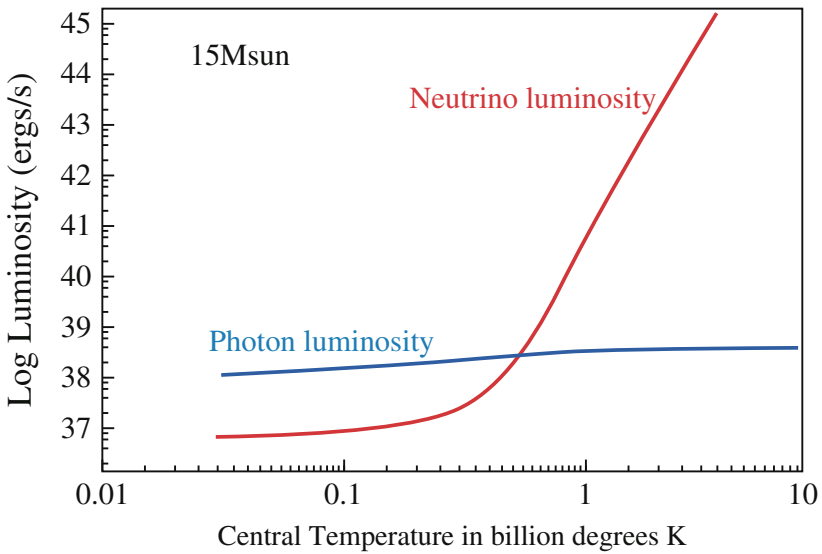
Energy generation vs energy loss



22 $M_\odot$ evolution, from Heger and Woosley 2001.

Photon vs neutrino luminosities

Neutrino luminosity is larger than photon luminosity. Weak interaction processes start to determine the evolution.



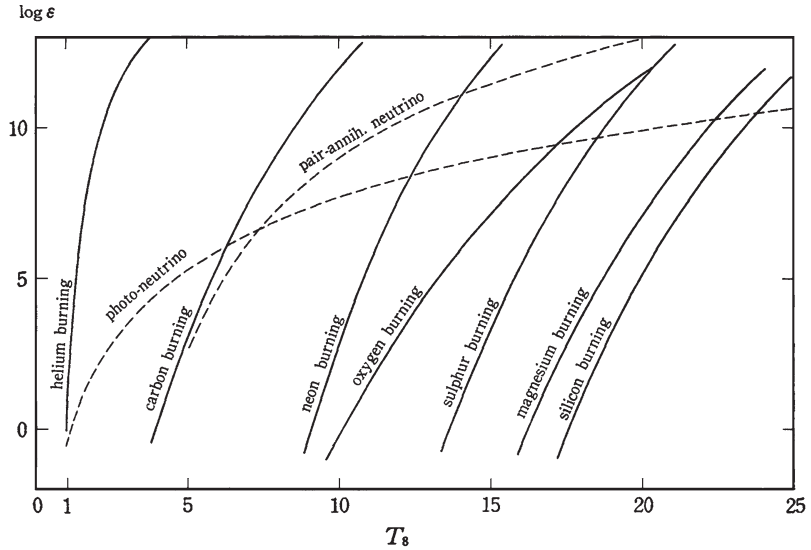


Fig. 9.2 A comparison between the neutrino losses and various nuclear energy generation reactions according to Hayashi et al. (1962). The rates are given in erg/g/s as a function of temperature in units of 10^8 K. The density is 10^4 g/cm^3. Note that several reactions shown in the figure are no longer considered relevant to massive stars. The neutrino rates have been updated since 1962, but the general picture of neutrino losses taking over energy generation remains today

From G. Shaviv, *The Synthesis of the Elements*

Carbon Burning

Burning conditions:

for stars > 8 M_o (solar masses) (ZAMS)

T~ 600-700 Mio
$\rho \sim 10^5$-10^6 g/cm^3

Major reaction sequences:

$$^{12}C + {}^{12}C \rightarrow {}^{24}Mg^* \rightarrow {}^{23}Mg + n - 2.62\ \text{MeV}$$
$$\rightarrow {}^{20}Ne + \alpha + 4.62\ \text{MeV}$$
$$\rightarrow {}^{23}Na + p + 2.24\ \text{MeV}.$$

dominates by far

of course p's, n's, and a's are recaptured … ^{23}Mg can b-decay into ^{23}Na

Composition at the end of burning:

mainly ^{20}Ne, ^{24}Mg, with some ^{21,22}Ne, ^{23}Na, ^{24,25,26}Mg, ^{26,27}Al

of course ^{16}O is still present in quantities comparable with ^{20}Ne (not burning … yet)

Neon Burning

Burning conditions:

for stars > 12 M_o (solar masses) (ZAMS)

T~ 1.3-1.7 Bio K
$\rho \sim 10^6$ g/cm^3

Why would neon burn before oxygen ???

Answer:

Temperatures are sufficiently high to initiate **photodisintegration** of ^{20}Ne

$$\left.\begin{array}{l} {}^{20}\mathrm{Ne}+\gamma \rightarrow {}^{16}\mathrm{O} + \alpha \\ {}^{16}\mathrm{O}+\alpha \rightarrow {}^{20}\mathrm{Ne} + \gamma \end{array}\right\}$$ equilibrium is established

this is followed by (using the liberated helium)

$$^{20}\mathrm{Ne}+\alpha \rightarrow {}^{24}\mathrm{Mg} + \gamma$$

so net effect: $$\boxed{2\ {}^{20}\mathrm{Ne} \rightarrow {}^{16}\mathrm{O} + {}^{24}\mathrm{Mg} + 4.59\ \mathrm{MeV}.}$$

Oxygen Burning

Burning conditions:

T~ 2 Bio
$\rho \sim 10^7$ g/cm^3

Major reaction sequences:

$$^{16}\mathrm{O} + {}^{16}\mathrm{O} \rightarrow {}^{32}\mathrm{S}^* \rightarrow {}^{31}\mathrm{S} + n + 1.45\ \mathrm{MeV} \quad (5\%)$$
$$\rightarrow {}^{31}\mathrm{P} + p + 7.68\ \mathrm{MeV} \quad (56\%)$$
$$\rightarrow {}^{30}\mathrm{P} + d - 2.41\ \mathrm{MeV} \quad (5\%)$$
$$\rightarrow {}^{28}\mathrm{Si} + \alpha + 9.59\ \mathrm{MeV}. \quad (34\%)$$

plus recapture of n,p,d,α

Main products:

^{28}Si,^{32}S (90%) and some ^{33,34}S,^{35,37}Cl,$^{36.38}$Ar, ^{39,41}K, ^{40,42}Ca

Silicon Burning

Burning conditions:

T~ 3-4 Bio
$\rho \sim 10^9$ g/cm^3

Reaction sequences:

- Silicon burning is fundamentally different to all other burning stages.
- **Complex network of fast (γ,n), (γ,p), (γ,a), (n,γ), (p,γ), and (a,γ) reactions**
- The net effect of Si burning is: 2 ^{28}Si --> ^{56}Ni,

need new concept to describe burning:

Nuclear Statistical Equilibrium (NSE)

Quasi Statistical Equilibrium (QSE)

Nuclear Statistical Equilibrium

Processes mediated by the strong and electromagnetic interactions proceed in a time scale much sorter than any other evolutionary time scale of the system. Implies that there is an equilibrium between the processes responsible for the creation and destruction of nuclei:

$$A(Z, N) \rightleftarrows Zp + Nn + \gamma$$

that corresponds to a relation between chemical potentials

$$\mu(Z, A) = (A - Z)\mu_n + Z\mu_p$$

Nuclear abundances in NSE

Nuclei follow Boltzmann statistics:

$$\mu(Z,A) = m(Z,A)c^2 + kT \ln\left[\frac{n(Z,A)}{G_{Z,A}(T)}\left(\frac{2\pi\hbar^2}{m(Z,A)kT}\right)^{3/2}\right]$$

with $G_{Z,A}(T)$ the partition function:

$$G_{Z,A}(T) = \sum_i (2J_i+1)e^{-E_i(Z,A)/kT} \approx \frac{\pi}{6akT}\exp(akT)\ (a \sim A/9\ \mathrm{MeV})$$

Results in Saha equation for abundances:

$$Y(Z,A) = \frac{G_{Z,A}(T)A^{3/2}}{2^A}\left(\frac{\rho}{m_u}\right)^{A-1} Y_p^Z Y_n^{A-Z}\left(\frac{2\pi\hbar^2}{m_u kT}\right)^{3(A-1)/2} e^{B(Z,A)/kT}$$

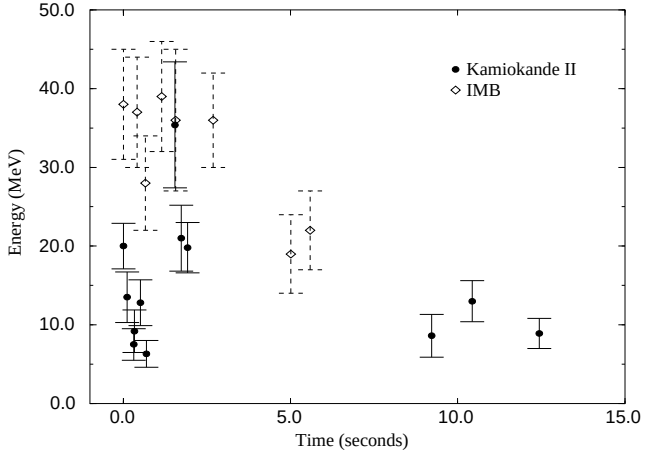
Composition depends on two parameters: Y_p, Y_n, that are fully determined knowing ρ, T, and Y_e.
Large partition functions warrant that composition remains on nuclei during collapse.

SN1987A

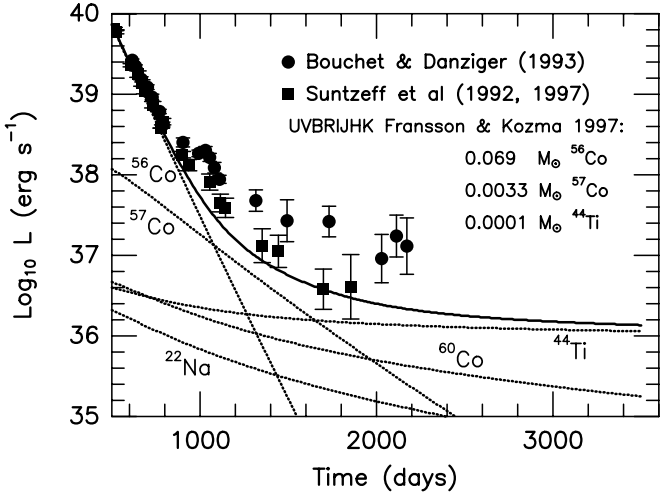
Type II supernova in LMC
(~ 55 kpc)

- $E_{\text{grav}} \approx 10^{53}$ erg
- $E_{\text{rad}} \approx 8 \times 10^{49}$ erg
- $E_{\text{kin}} \approx 10^{51}$ erg = 1 Bethe

neutrinos $E_\nu \approx 2.7 \times 10^{53}$ erg

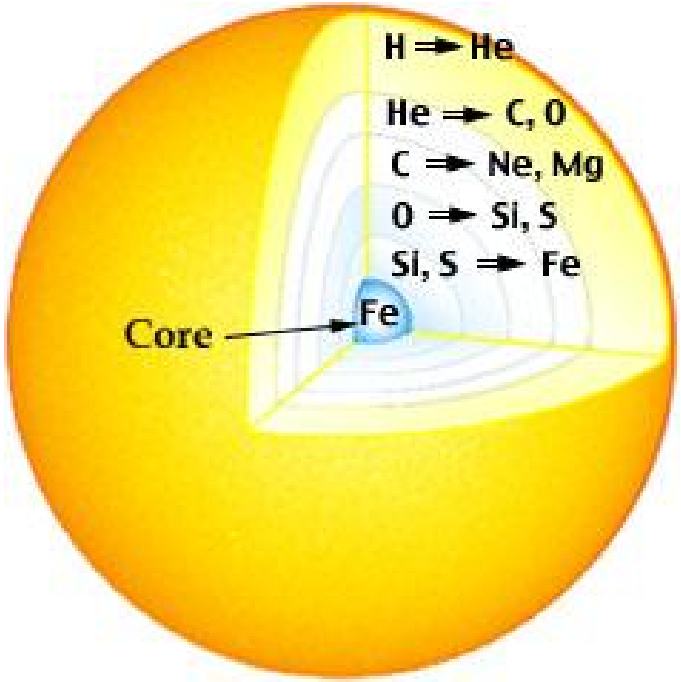


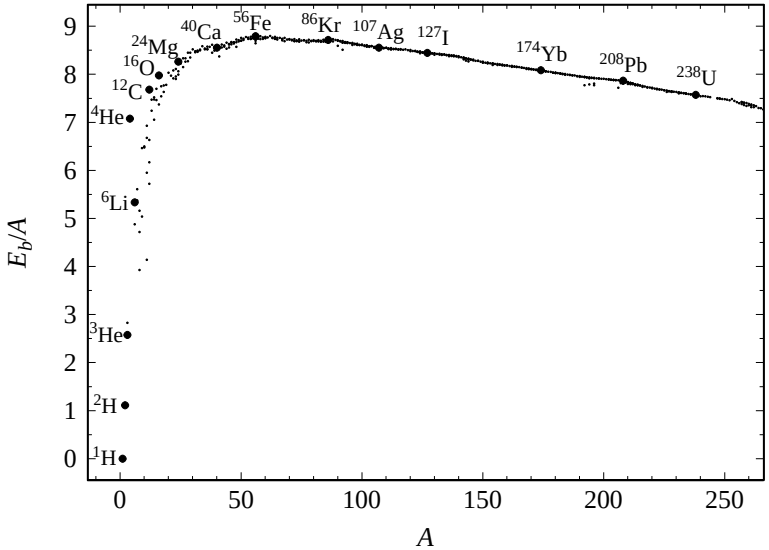
light curve



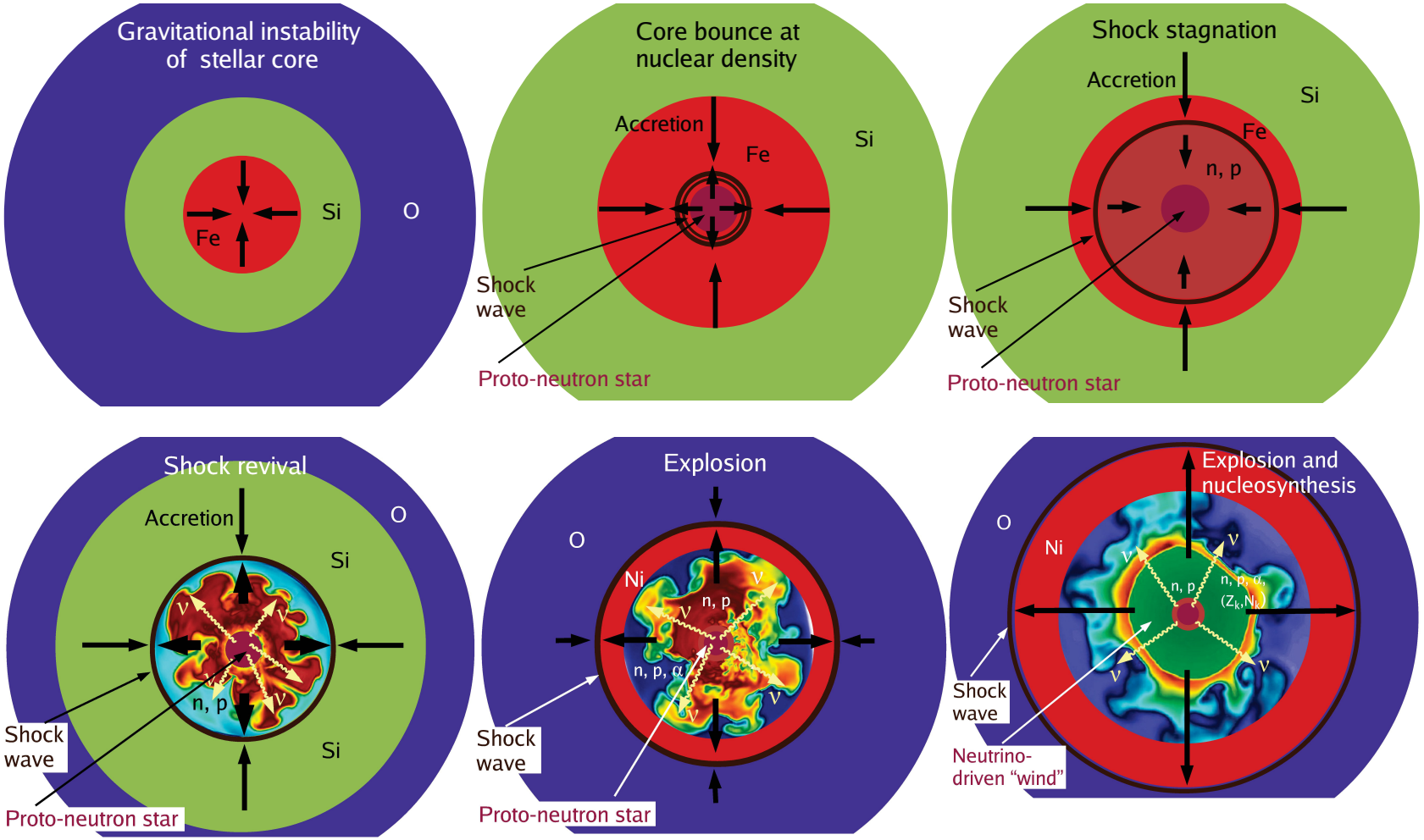
Presupernova Star

- Star has an onion like structure.
- Iron is the final product of the different burning processes.
- As the mass of the iron core grows it becomes unstable and collapses when it reaches around 1.4 $M_\odot$.



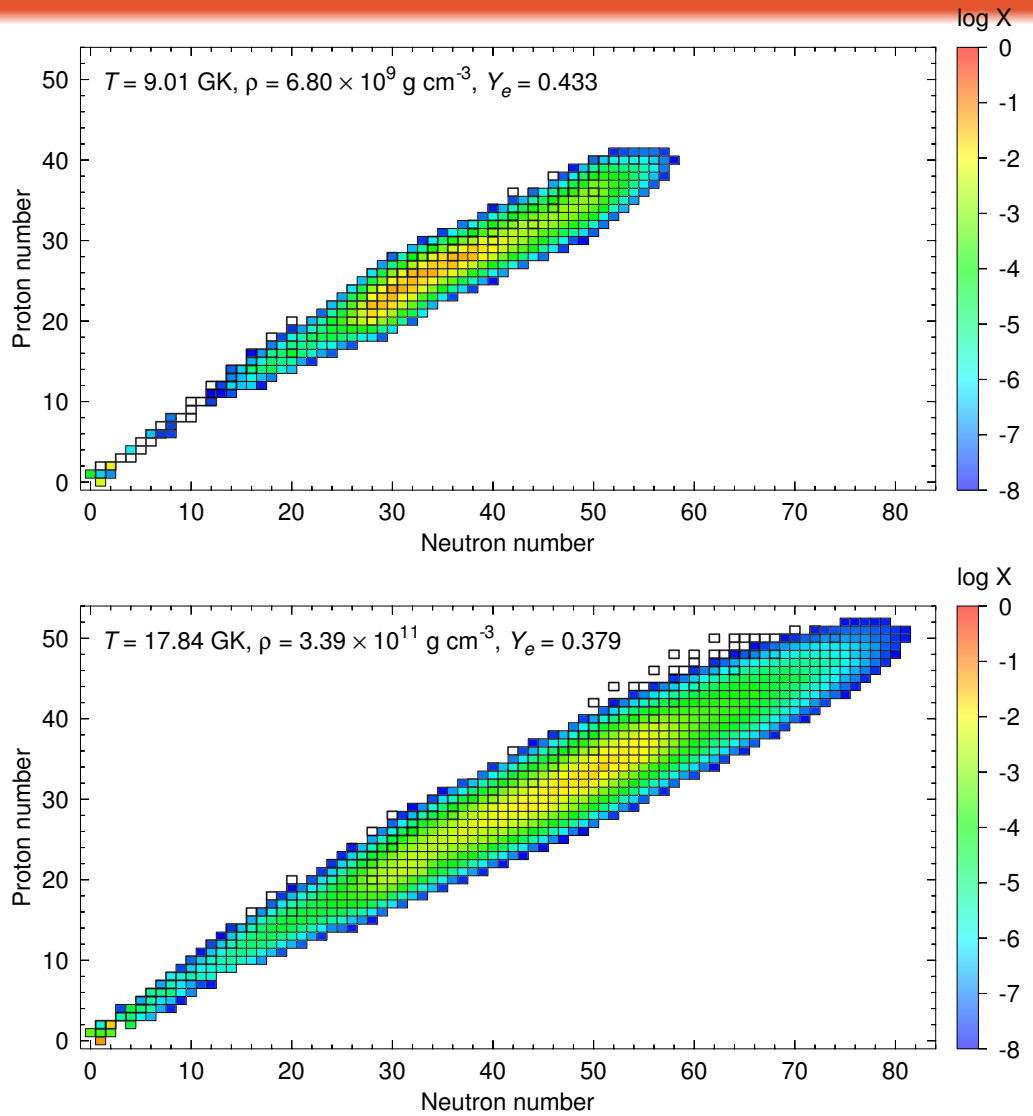


Schematical evolution



H.-Th. Janka, *et al*, PTEP 01A309 (2012)

Composition in NSE



Composition is given by heavy nuclei till bounce. Transition to “nuclear matter” occurs after bounce.

Onset of collapse

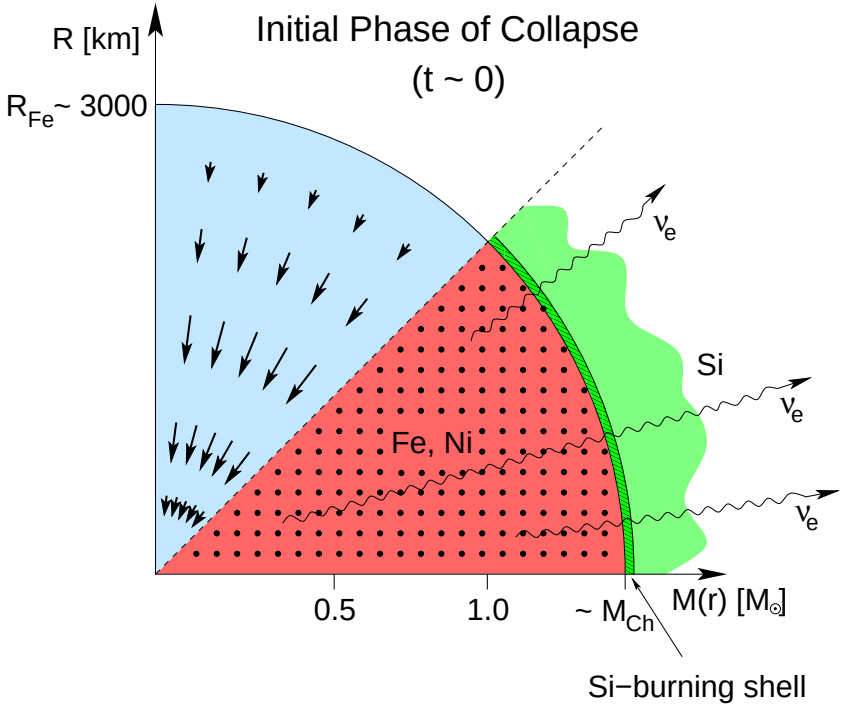
There are two processes that make the situation unstable:

1. Silicon burning is continuing in a shell around the iron core. This adds mass to the iron core increasing M_c.
2. Electrons can be captured by protons (free or in nuclei):

$$e^- + A(Z, N) \rightarrow A(Z-1, N+1) + \nu_e.$$

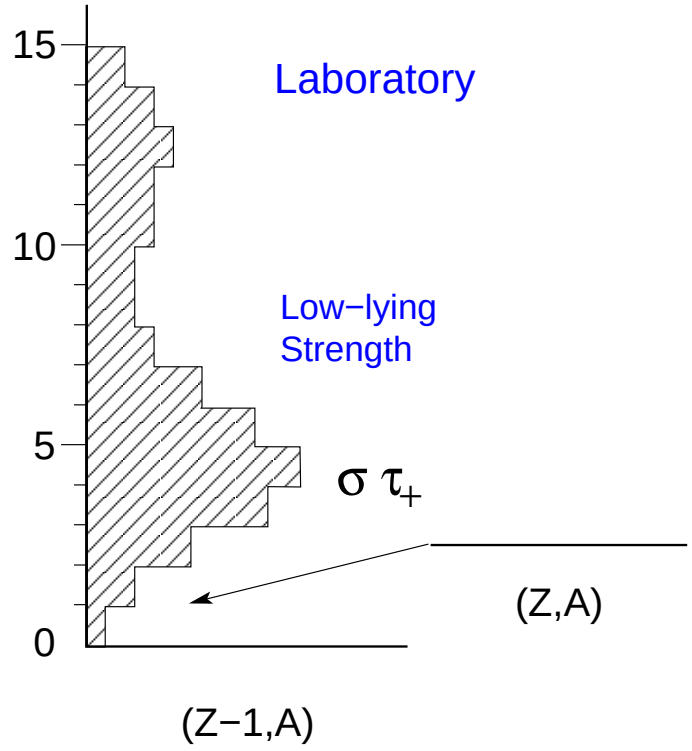
This reduces the pressure and keep the core cold, as the neutrinos leave. The net effect is a reduction of Y_e and consequently of the Chandrasekhar mass (M_{ch})

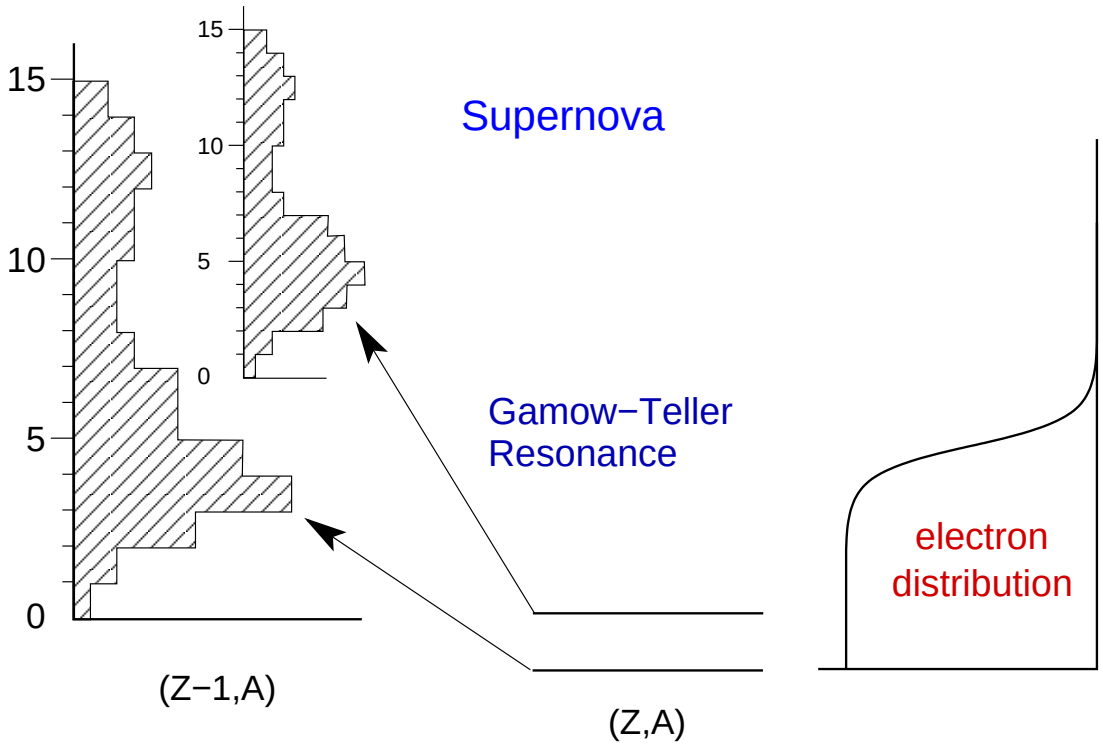
Presupernova evolution



- $T = 0.1$–0.8 MeV, $\rho = 10^7$–10^{10} g cm^{-3}. Composition of iron group nuclei.
- Important processes:
  - electron capture: $e^- + (N,Z) \rightarrow (N+1, Z-1) + \nu_e$
  - β^- decay: $(N,Z) \rightarrow (N-1, Z+1) + e^- + \bar{\nu}_e$
- Dominated by allowed transitions (Fermi and Gamow-Teller)
- Evolution decreases number of electrons (Y_e) and Chandrasekar mass ($M_{\rm ch} \approx 1.4(2Y_e)^2 M_\odot$)

Laboratory vs. stellar electron capture



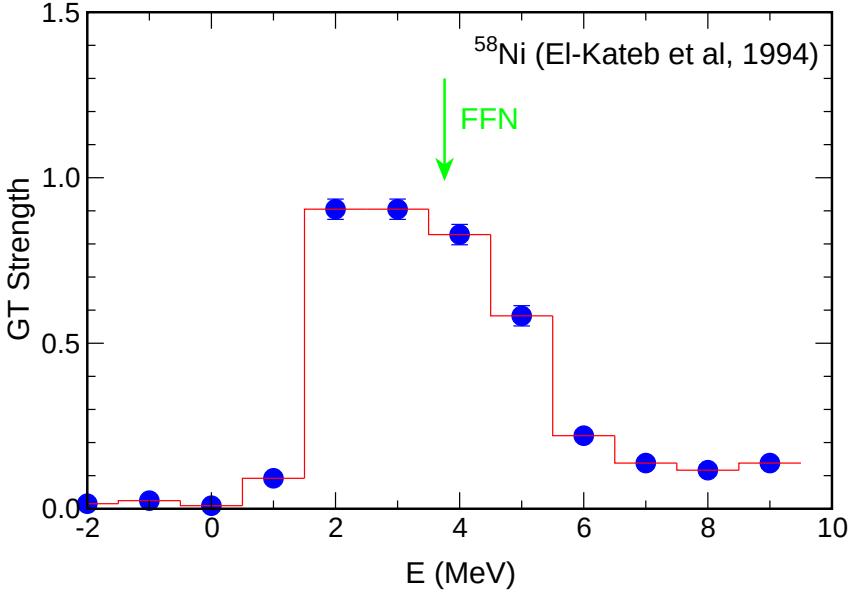


Capture of K-shell electrons to tail of GT strength distribution. Parent nucleus in the ground state

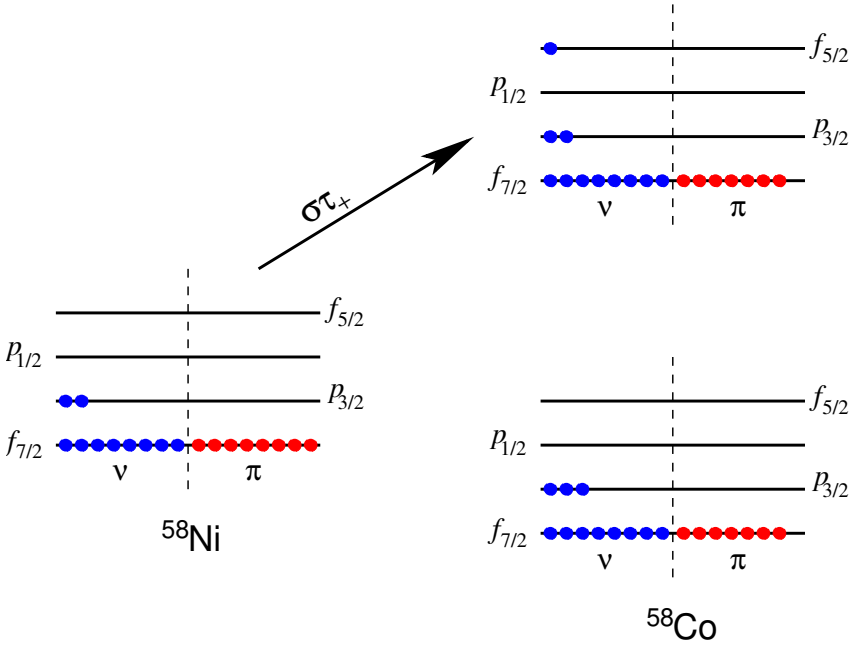
Capture of electrons from the high energy tail of the FD distribution. Capture to states with large GT matrix elements (GT resonance). Thermal ensemble of initial states.

Independent Particle Model

GT$_+$ strength in ^{58}Ni measured in (n, p).



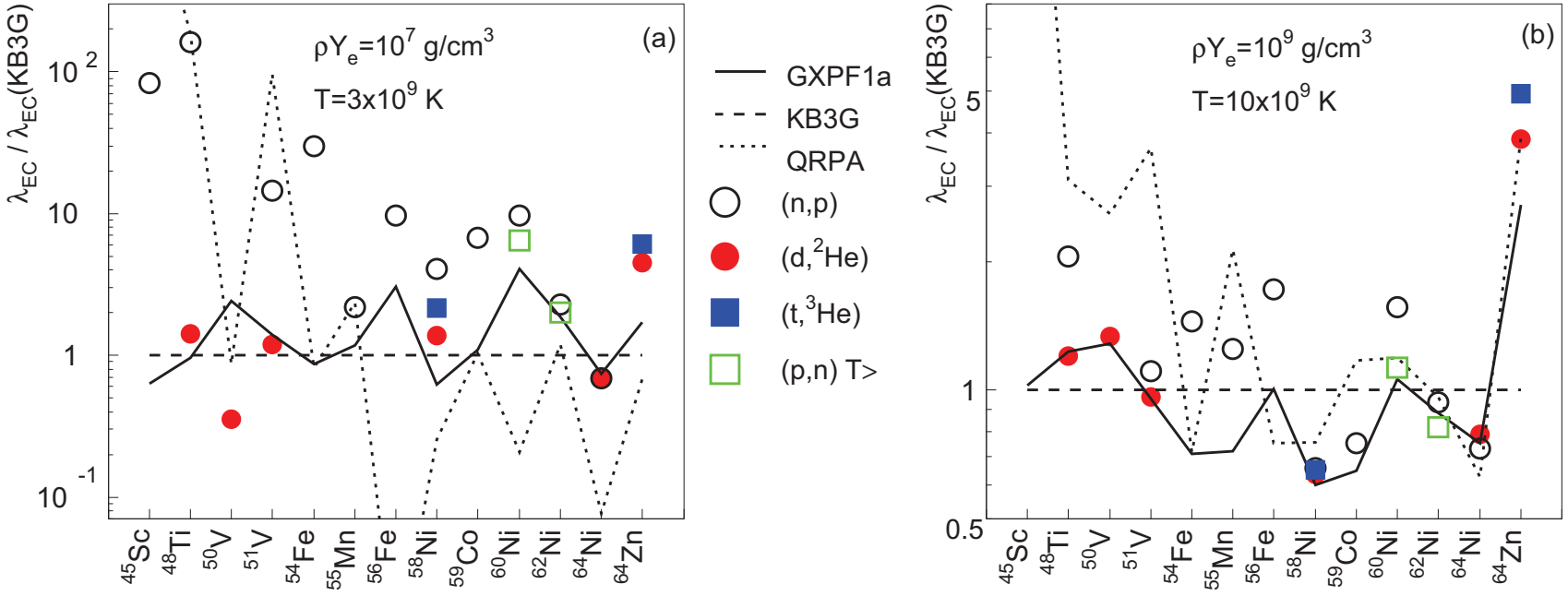
Independent particle model.



- The IPM allows for a single transition ($f_{7/2} \rightarrow f_{5/2}$). It does not correctly reproduce the fragmentation of GT strength (correlations).
- To account for correlations, it is necessary to explicitly consider the "residual" interaction between nucleons.
- Large experimental and theoretical effort dedicated to constrain the rates.

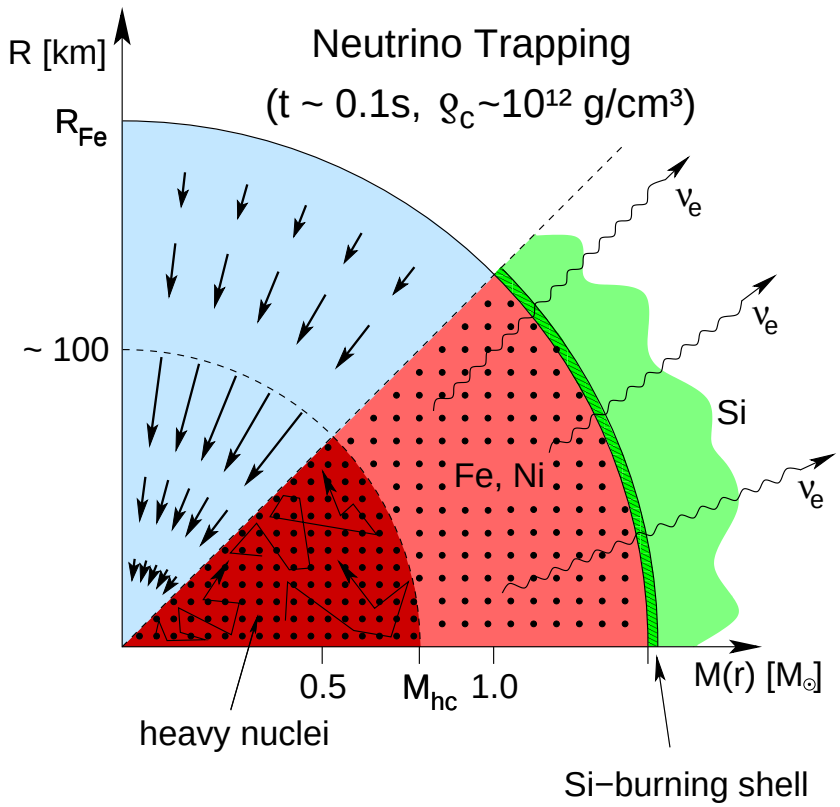
Systematic comparison theory vs experiment

A. L. Cole *et al.*, PRC **86**, 015809 (2012)



- Rates for iron-group nuclei are under control
- With increasing density, less sophisticated models like QRPA may suffice.

Collapse phase



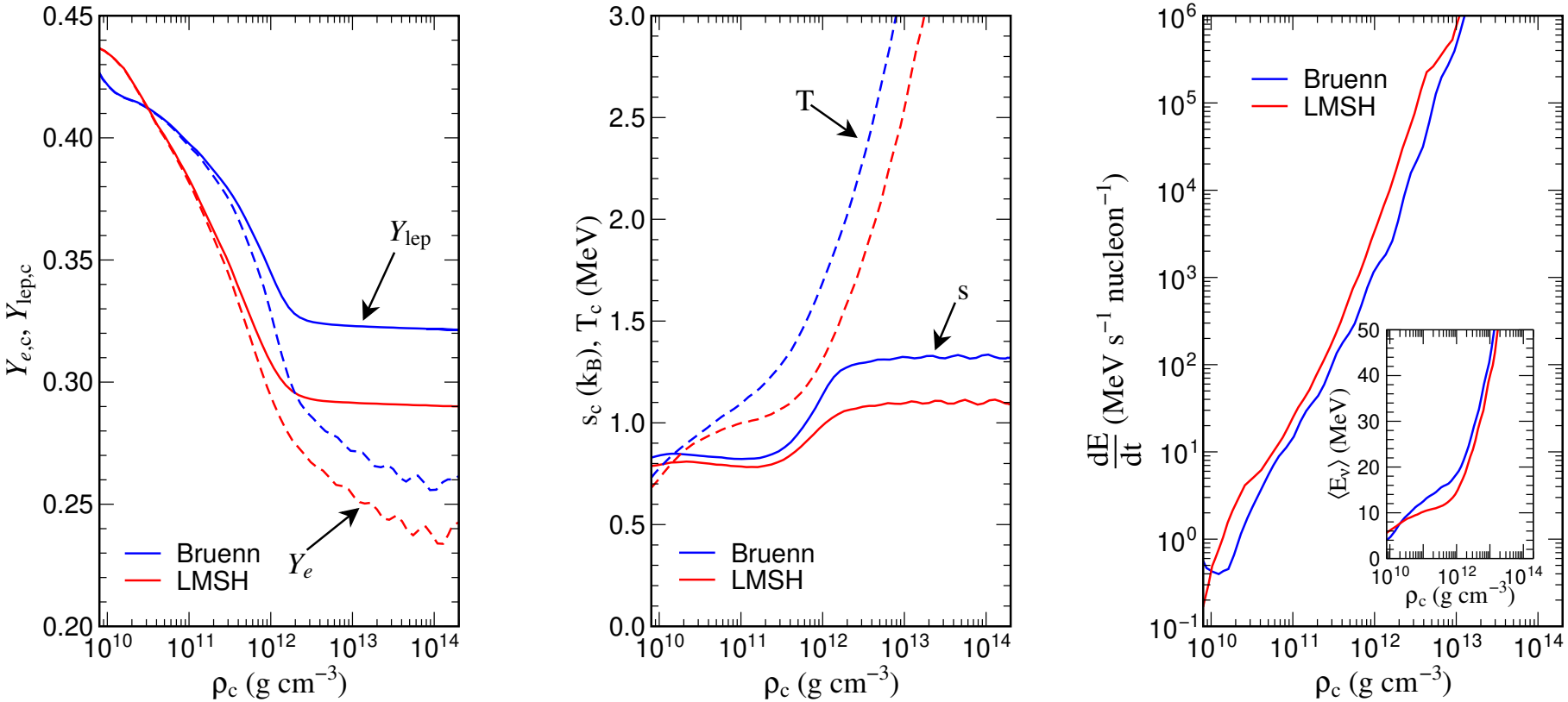
Important processes:

- Neutrino transport (Boltzmann equation):
  $\nu + A \rightleftarrows \nu + A$ (trapping)
  $\nu + e^- \rightleftarrows \nu + e^-$ (thermalization)

  cross sections $\sim E_\nu^2$
- electron capture on protons:
  $e^- + p \rightleftarrows n + \nu_e$
- electron capture on nuclei:
  $e^- + A(Z,N) \rightleftarrows A(Z-1,N+1) + \nu_e$

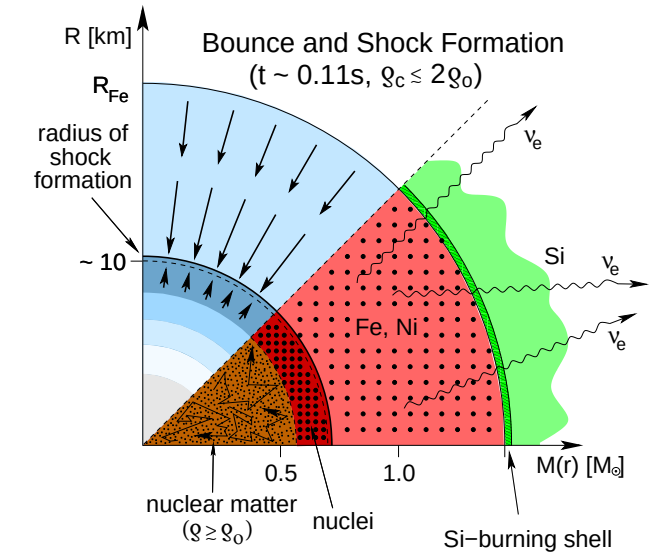
Evolution relevant quantities

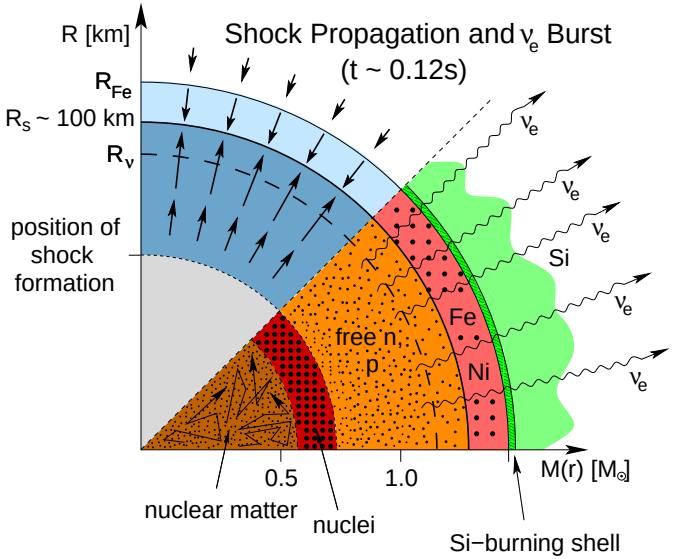
Marek, Rampp, Janka & Buras (Approx. General Relativistic model)
15 $M_\odot$ presupernova model from A. Heger & S. Woosley



$$Y_{\rm lep} = Y_e + Y_\nu$$

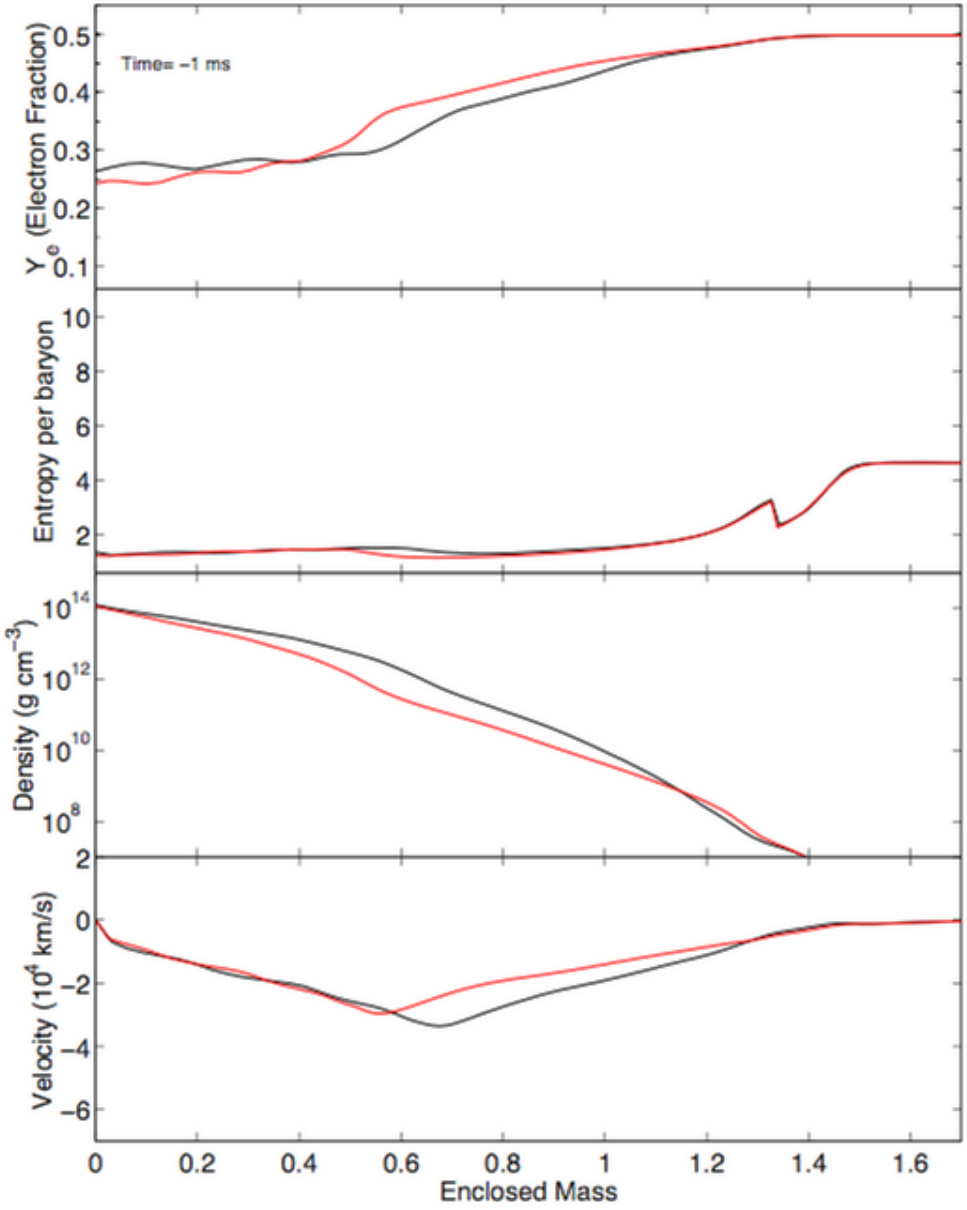
Bounce and ν_e burst



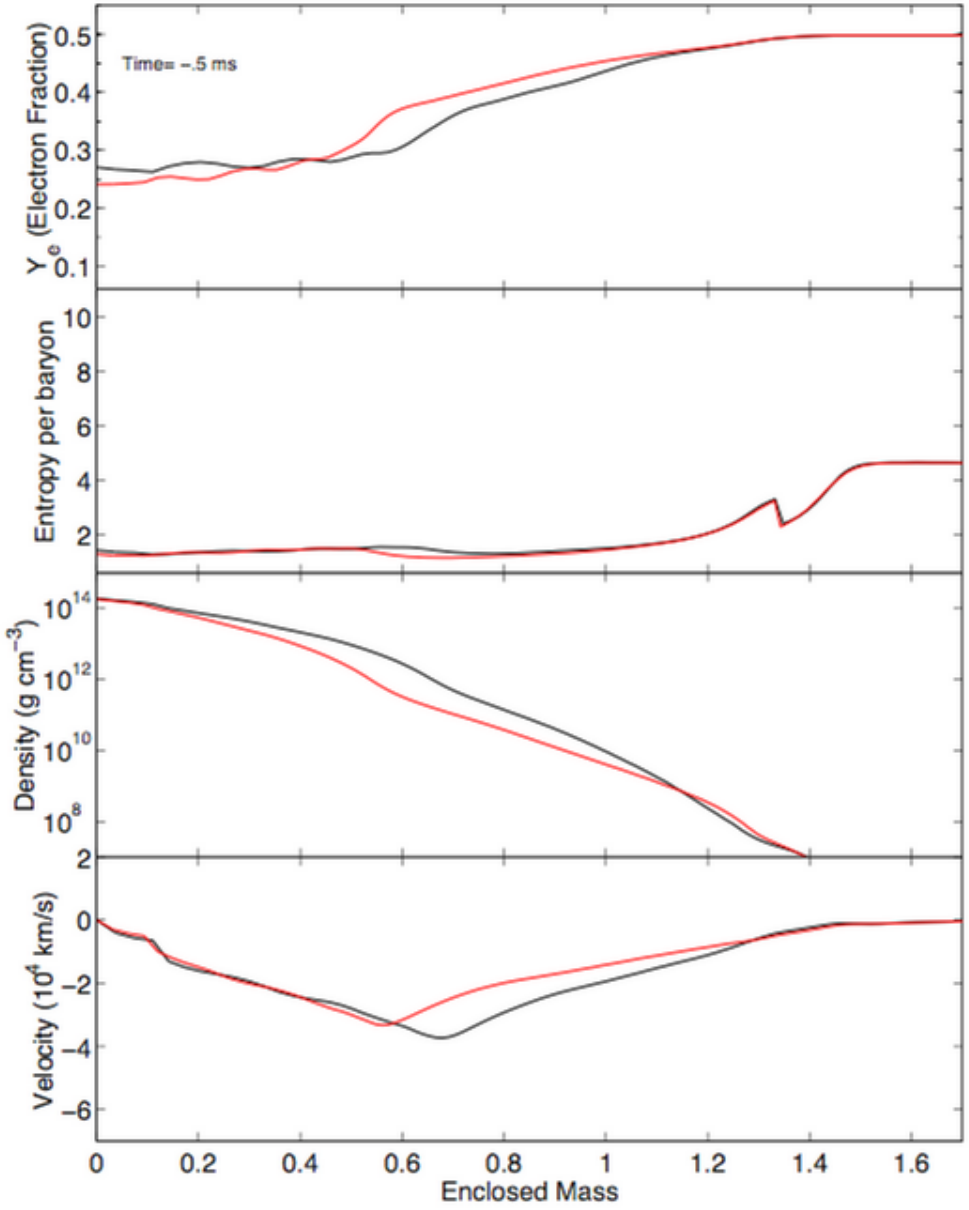


- Collapse continues until central density becomes around twice nuclear matter density.
- Sudden increase in nuclear pressure stops the collapse and a shock wave is launched at the sonic point. The energy of the shock depends on the Equation of State.
- The passage of the shock dissociates nuclei into free nucleons which costs $\sim 8\ \mathrm{MeV}$/nucleon. Additional energy is lost by neutrino emission produced by electron capture (ν_e burst).
- Shock stalls at a distance of several 100 km.

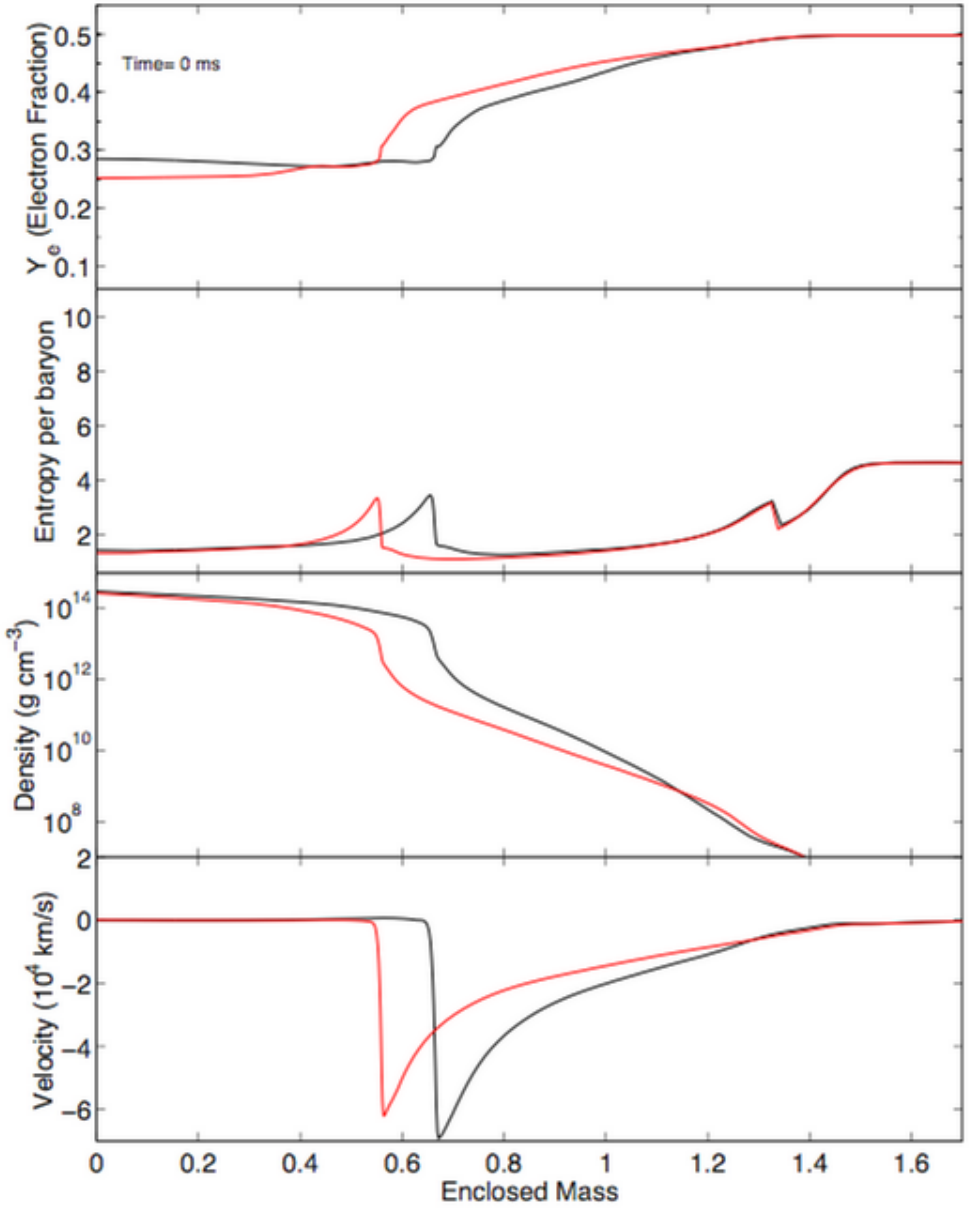
Spherical simulation



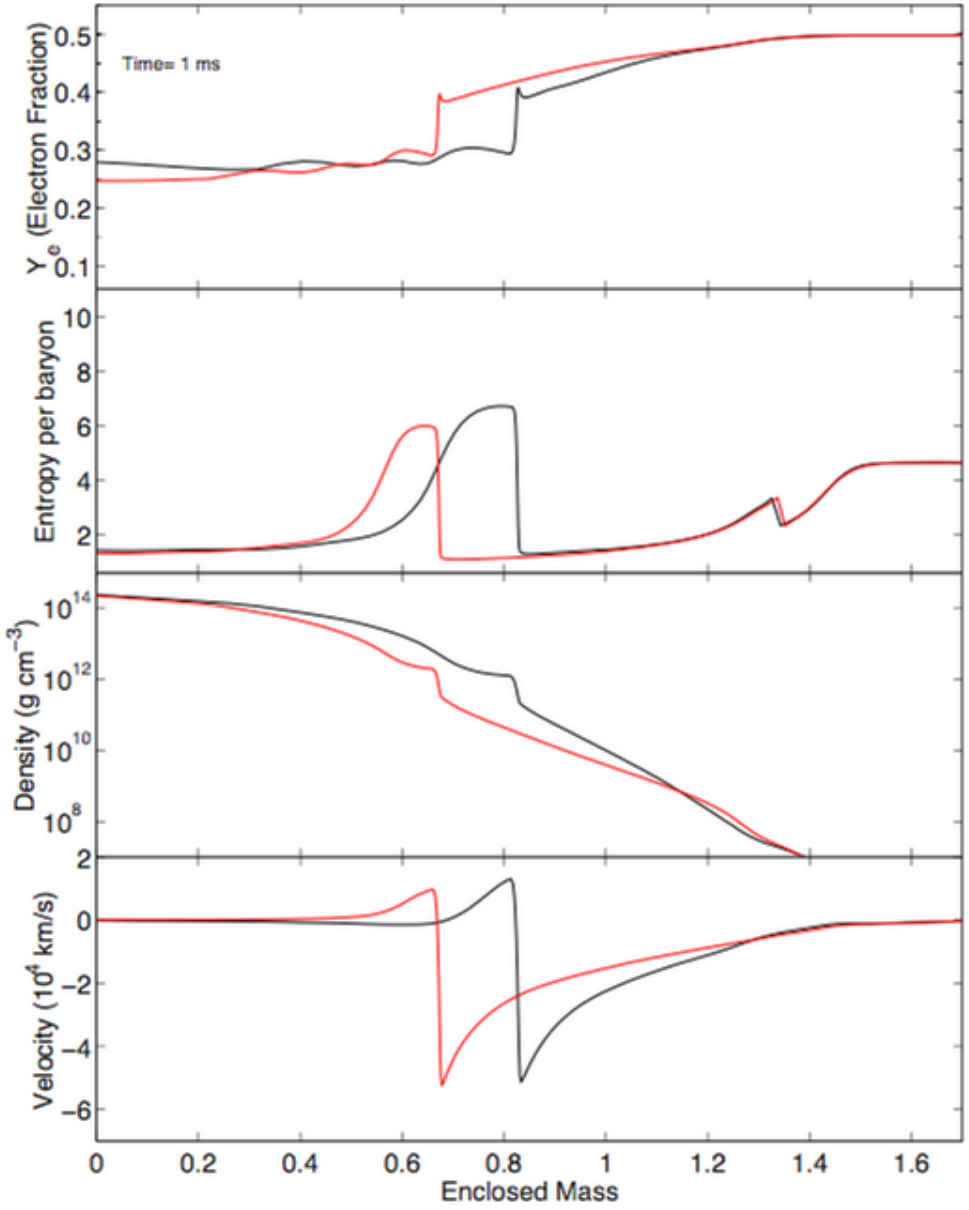
Spherical simulation



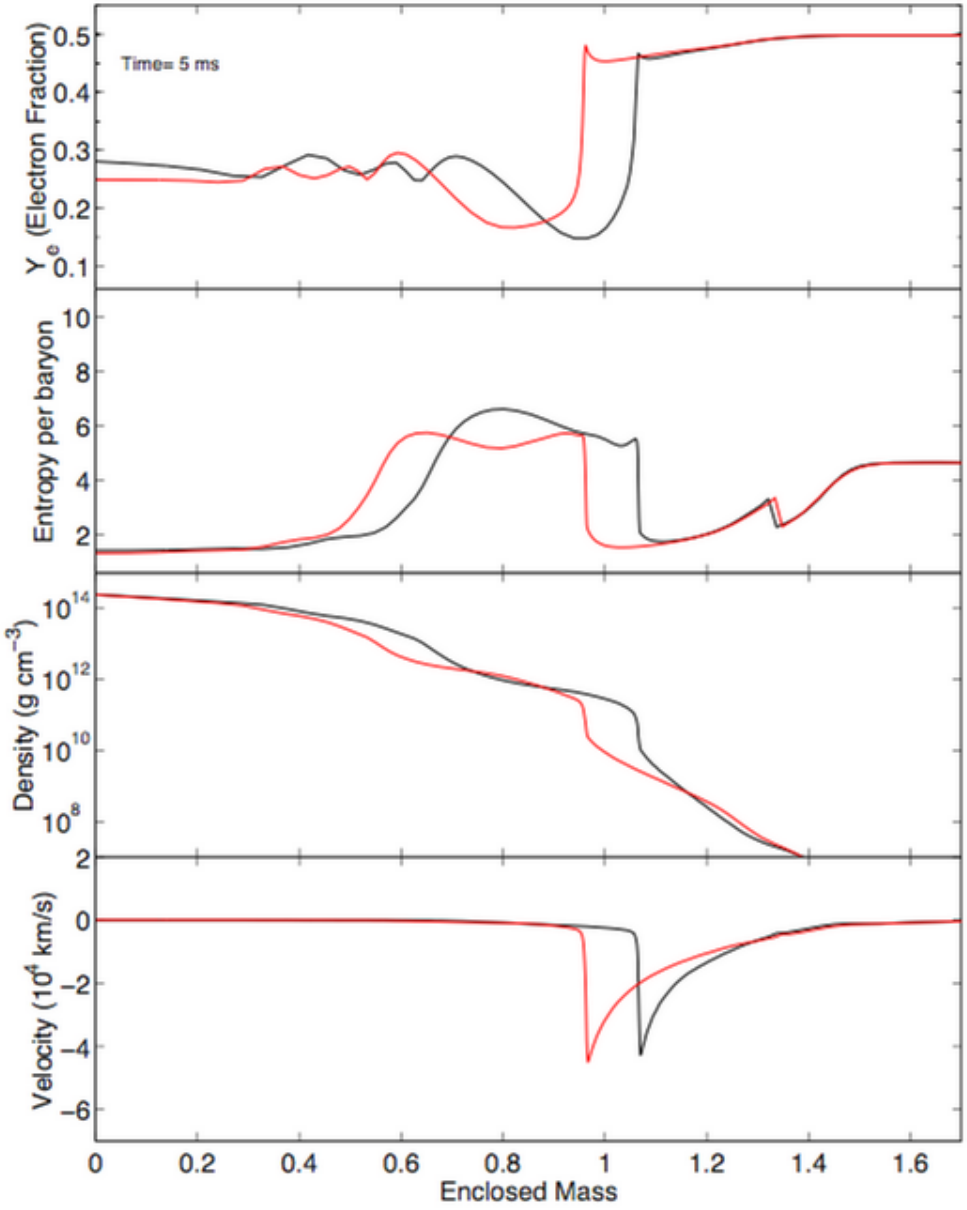
Spherical simulation



Spherical simulation



Spherical simulation



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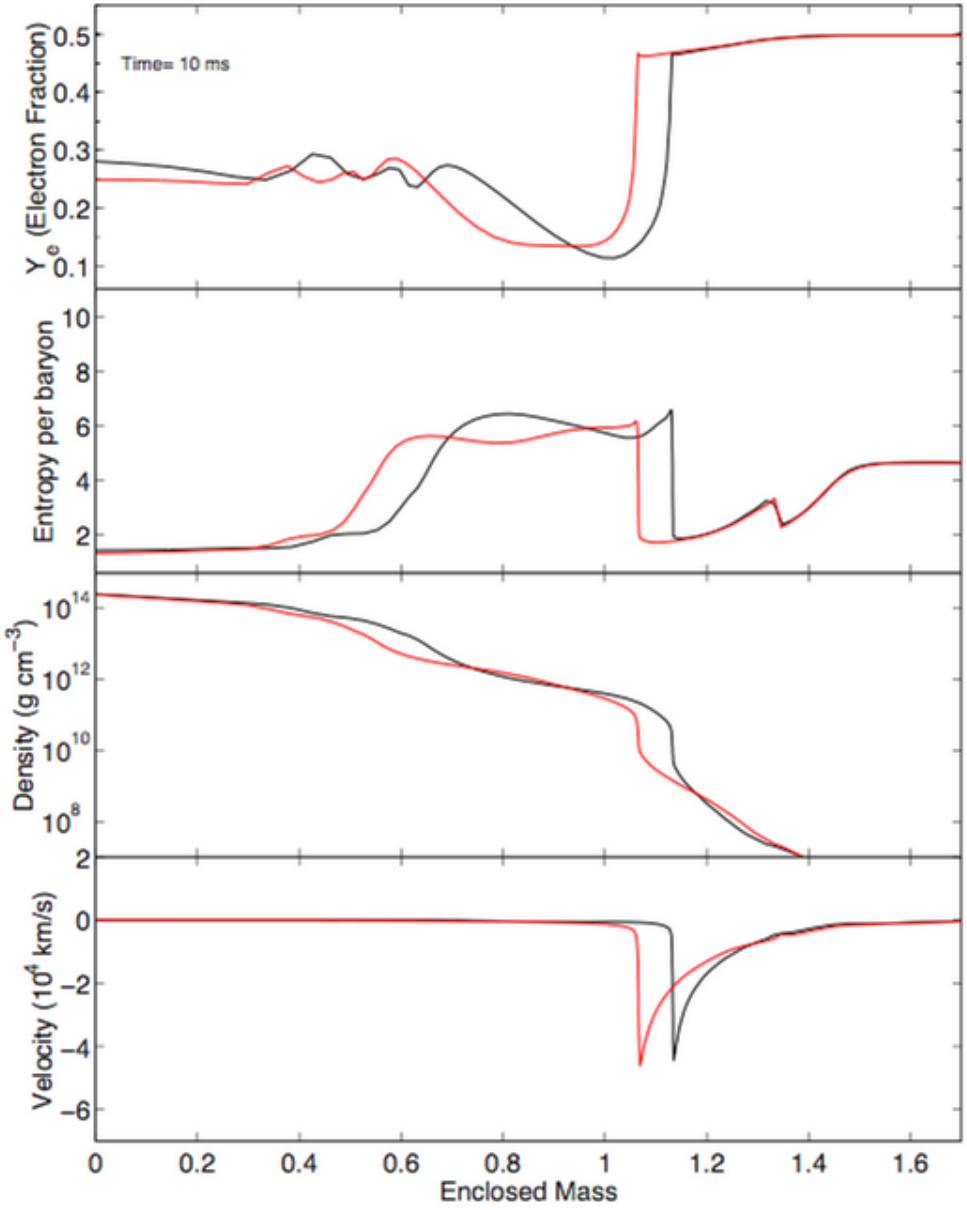
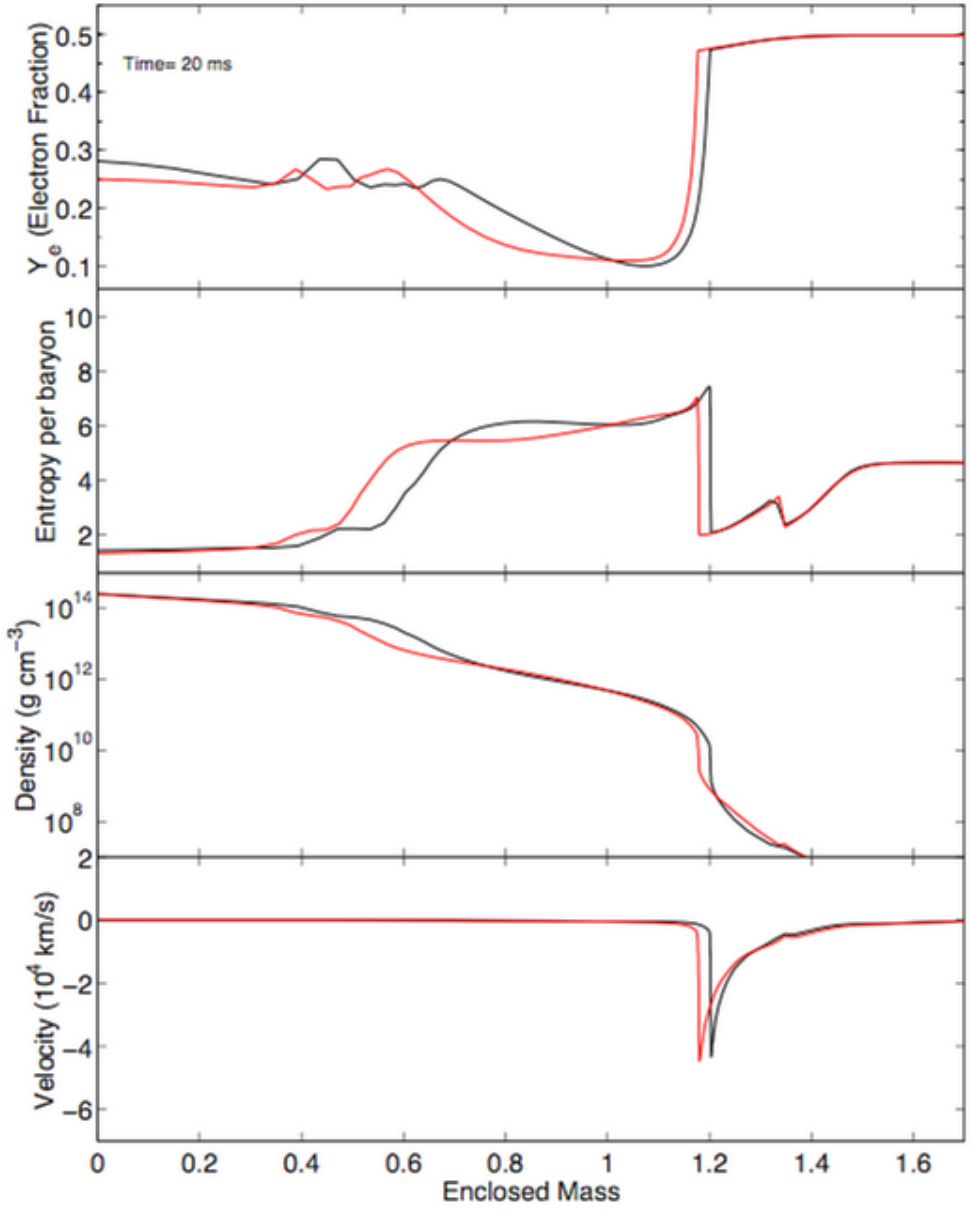
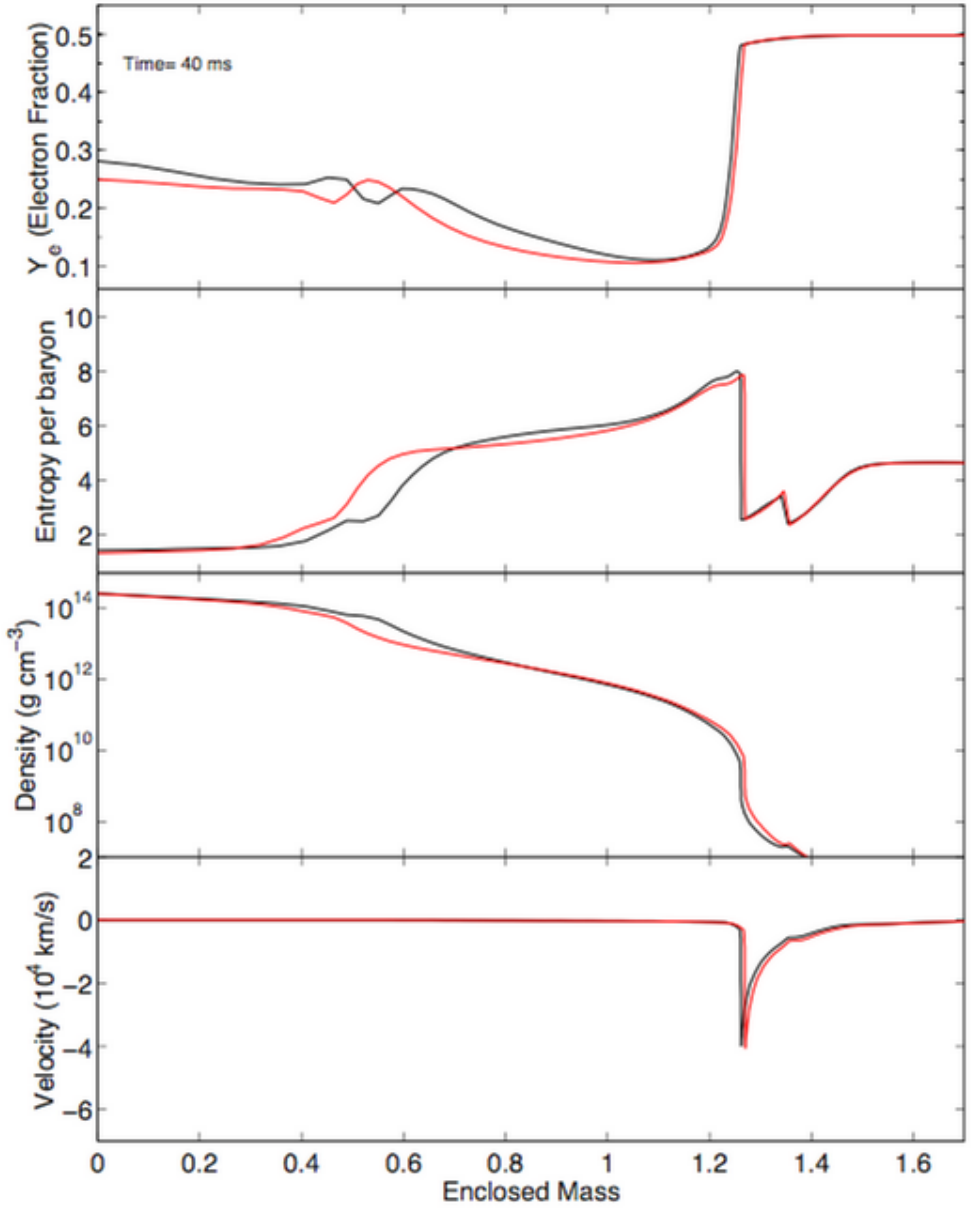


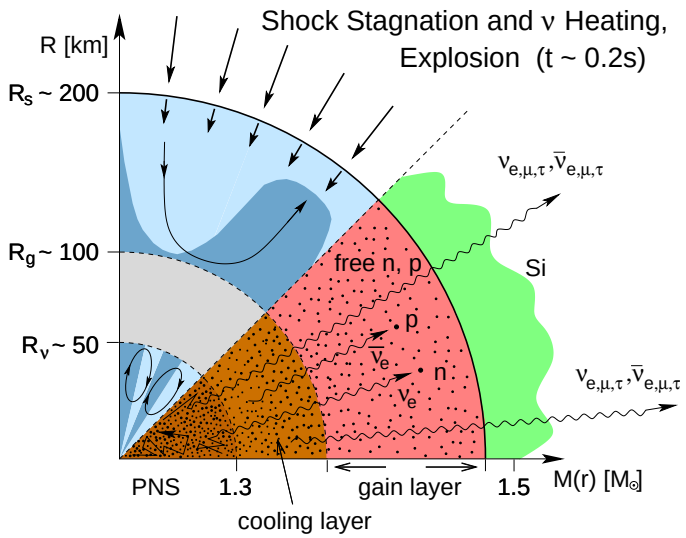
Spherical simulation

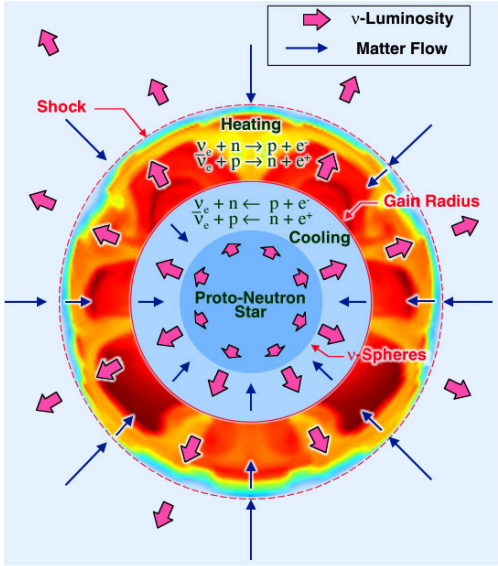


Spherical simulation



Delayed explosion mechanism: neutrino heating





Main processes:

$$\nu_e + n \rightleftarrows p + e^-$$
$$\bar{\nu}_e + p \rightleftarrows n + e^+$$

Concept of gain radius due to Bethe. Corresponds to the region where cooling (electron positron capture) and heating (neutrino antineutrino absorption) are equal.

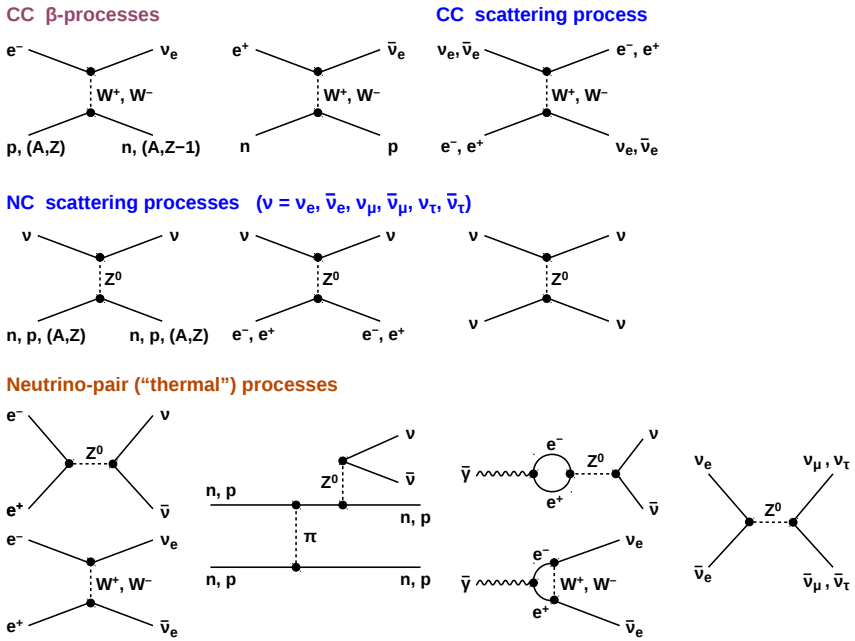
$$\text{Cooling: } 143 \left(\frac{kT}{2\,\text{MeV}} \right)^6 \text{MeV/s}$$

$$\text{Heating: } 110 \left(\frac{L_{\nu_e,52}\epsilon^2_{\nu_e}}{r_7^2} Y_n + \frac{L_{\bar{\nu}_e,52}\epsilon^2_{\bar{\nu}_e}}{r_7^2} Y_p \right) \text{MeV/s}$$

Gravitational energy of a nucleon at 100 km: 14 MeV
Energy transfer induces convection and requires multidimensional simulations.

Relevant neutrino processes

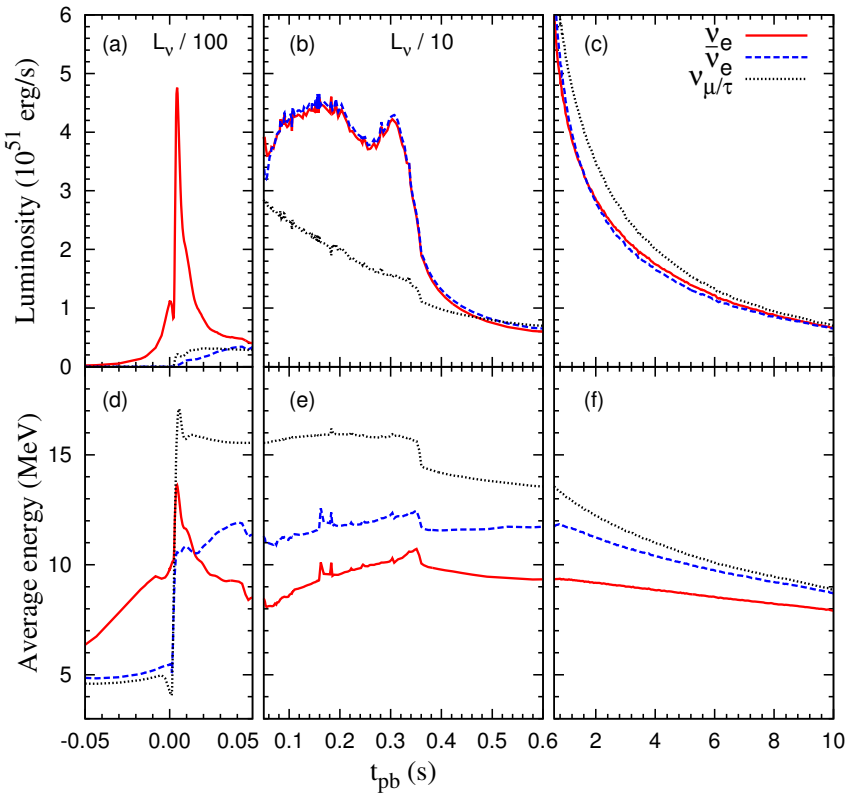
Supernova simulations are very sensitive to the treatment of neutrino processes in high density matter. Many different neutrino processes need to be modelled within Boltzmann radiation transport.



- Neutral current processes determine the diffusion of neutrinos out of the neutron star.
- Charged current processes determine energy differences between flavors (nucleosynthesis) and energy transfer to the shock.

H.-Th. Janka, arXiv:1702.08713 [astro-ph.HE]

Neutrino signatures

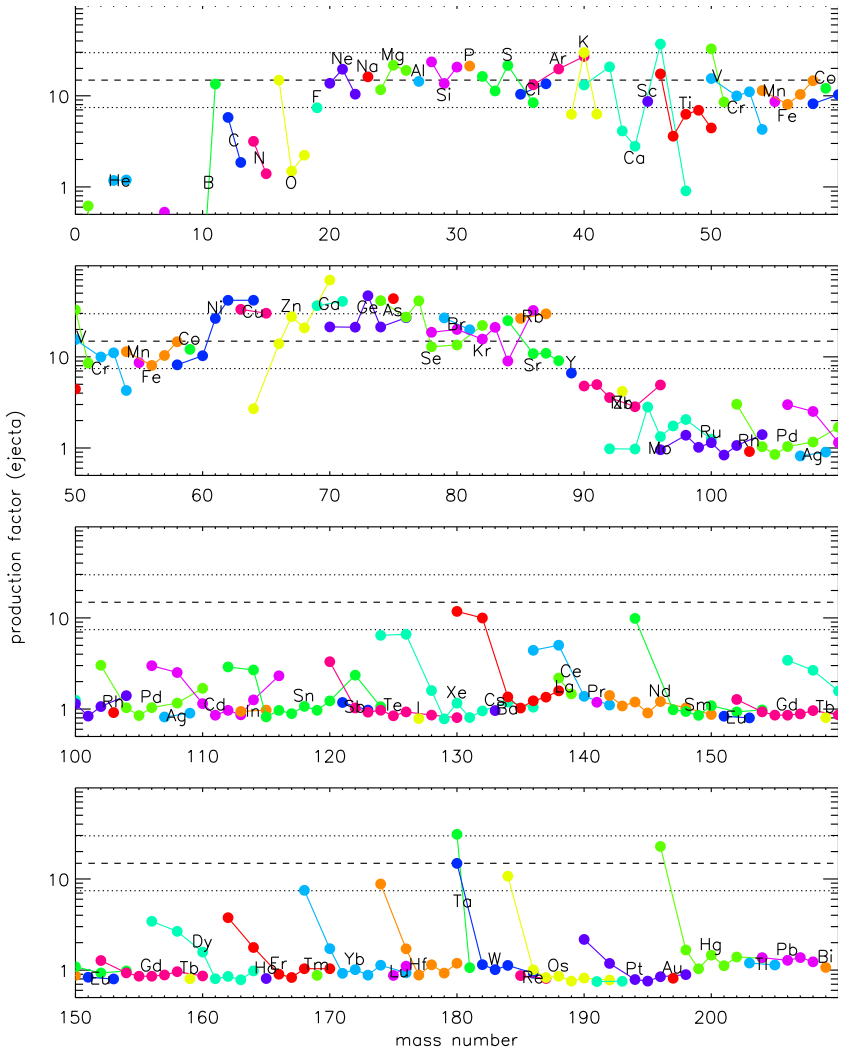


Neutrino emission provides signatures of different dynamical phases

- Collapse
- ν_e burst: shock break-out protoneutron star
- Accretion phase before onset explosion
- Protoneutron-star cooling phase

Nucleosynthesis

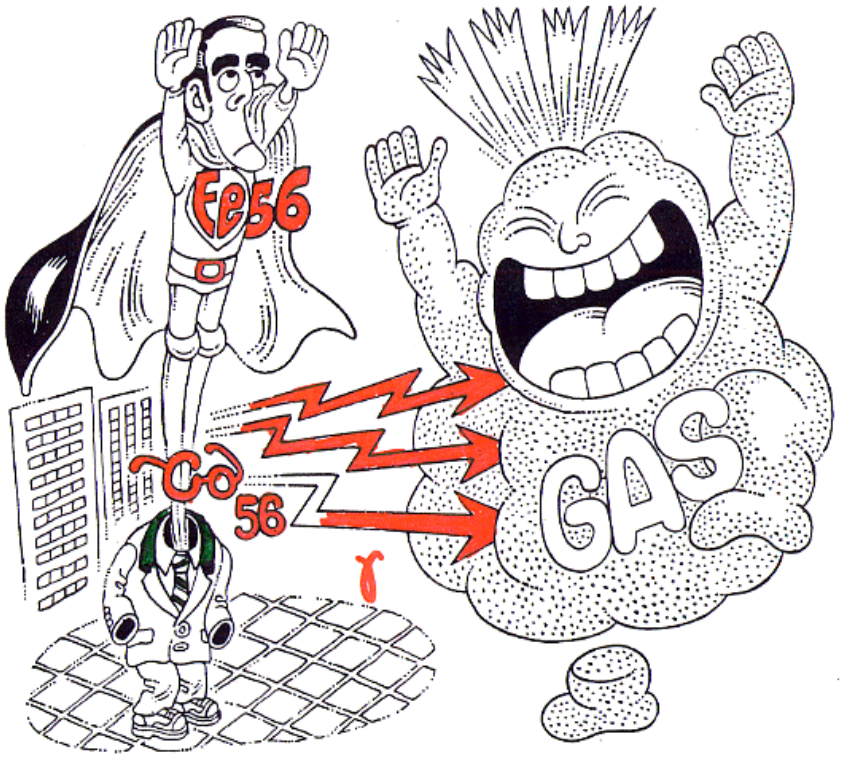
Main products are O, C and Iron group nuclei. producion factor: $(M_i/M_{\mathrm{ej}})/(X_{i,\odot})$



Electromagnetic Signatures of nucleosynthesis

Light curve follows the decay of Nickel.

$$^{56}\mathrm{Ni} \xrightarrow{6\ \mathrm{days}} {}^{56}\mathrm{Co} \xrightarrow{77\ \mathrm{days}} {}^{56}\mathrm{Fe}$$



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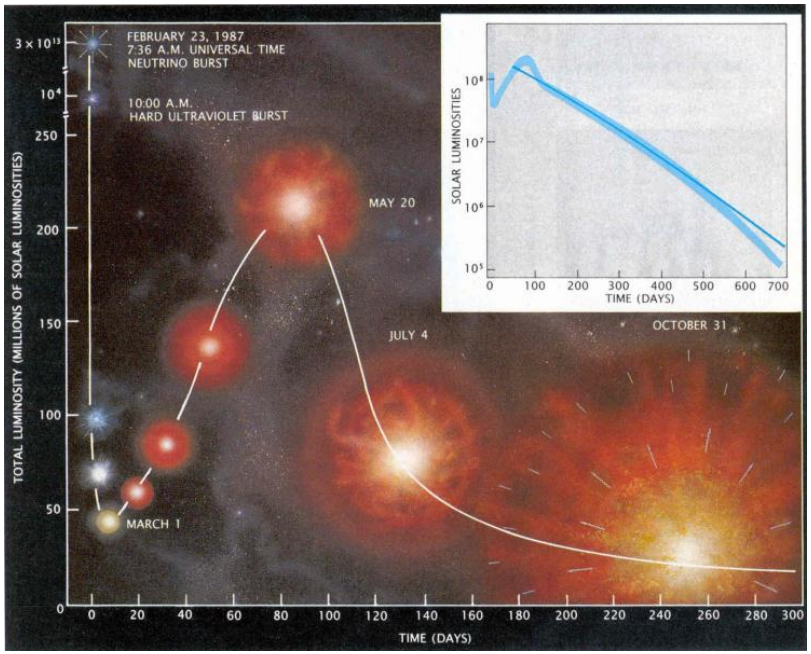
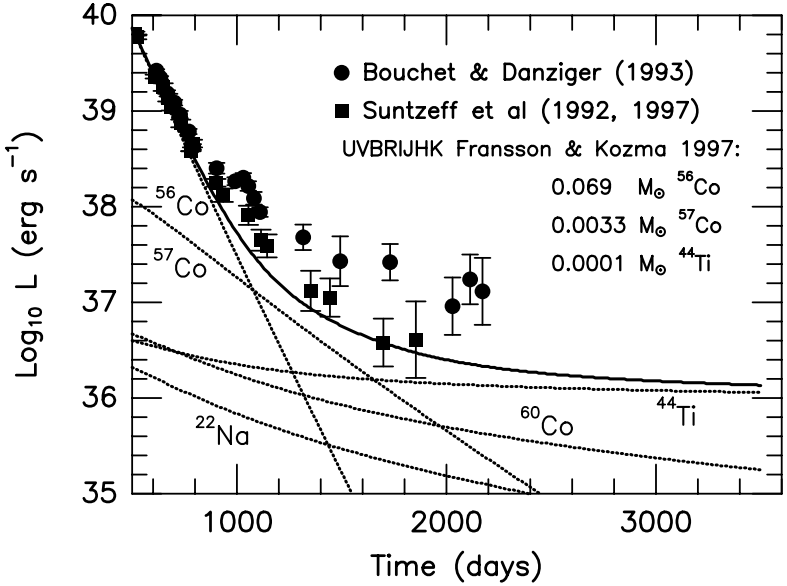
Diehl & Timmes, PASP 110, 637 (1998)



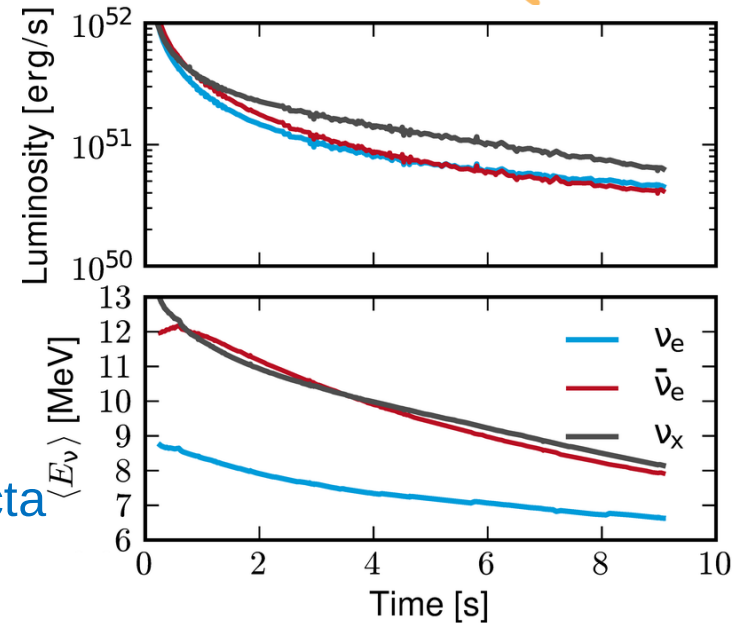


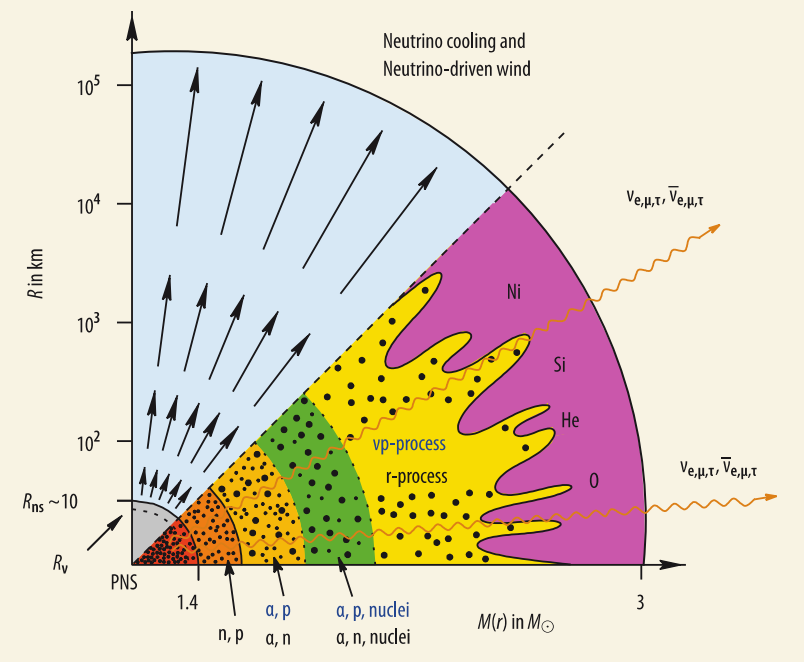
Heavy elements in supernova?

Heavy elements produced in neutrino winds from protoneutron star cooling. Neutrino interactions determine proton-to-nucleon ratio, Y_e

$$\nu_e + n \rightleftarrows p + e^-$$
$$\bar{\nu}_e + p \rightleftarrows n + e^+$$

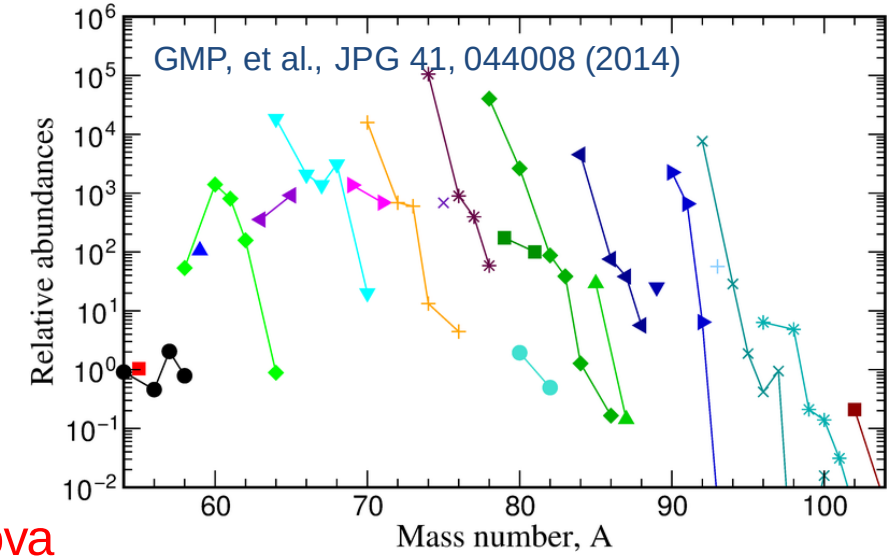
Very similar spectra ν_e and $\bar{\nu}_e$ → proton rich ejecta





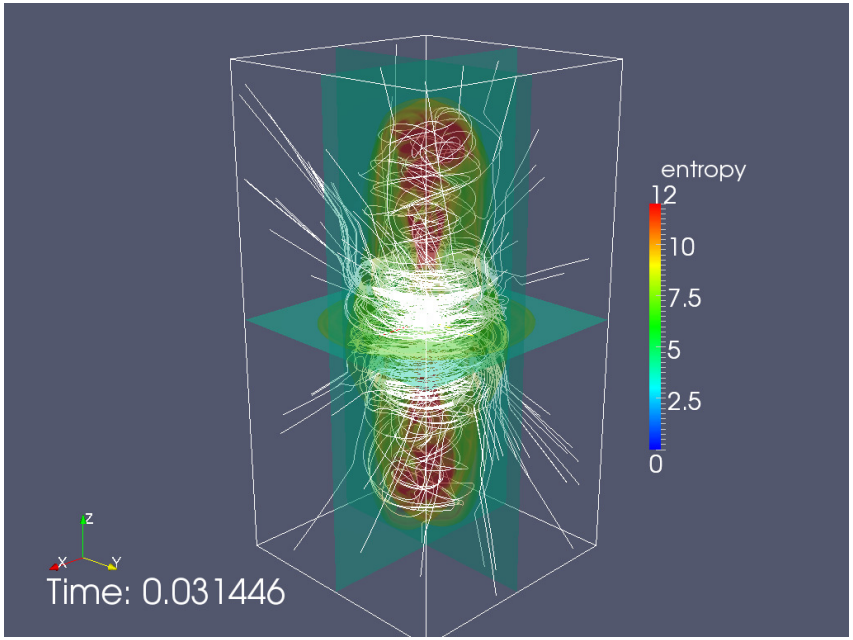
No r-process in typical core-collapse supernova

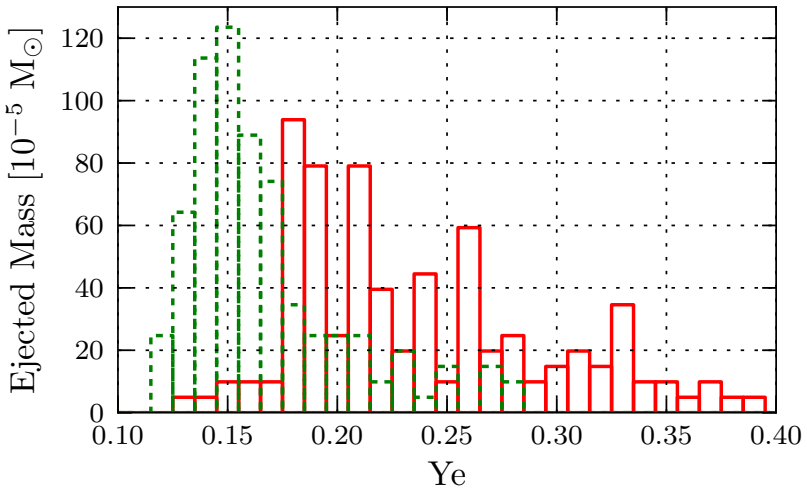
Supernova produce only medium mass nuclei

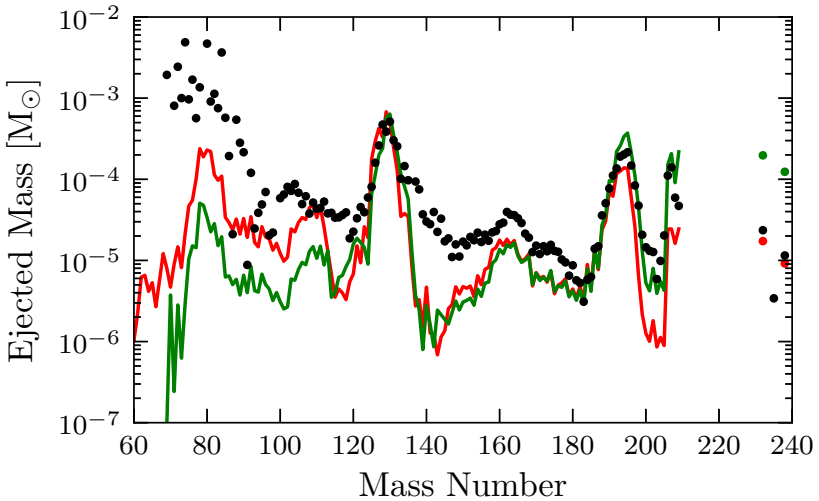


Magnetorotationally Driven Supernovae

Winteller *et al*, ApJ **750**, L22 (2012)







- Rotation and magnetic fields may play an important role in some supernova explosions (relation GRBs?).
- Origin highly magnetized neutron stars (magnetars)
- Site r-process nucleosynthesis.

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