Nuclear Astrophysics

III. Advanced burning and onset of collapse

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End of hydrogen core burning

When the hydrogen fuel in the core gets exhausted, an isothermal core of about 8% of the stellar mass can develop in the center. Continous hydrogen burning adds to the core mass which eventually rises over the Schönberg-Chandrasekhar mass limit. Then the core's temperature (and density) rise. Finally the central temperature is high enough ($T_c \approx 10^8$ K) to ignite helium core burning.

Hydrogen burning continues in a shell outside the helium core. This (hydrogen shell burning) occurs at higher temperatures than hydrogen core burning.

Which reaction can start helium burning?

Consider a supply of protons and ^{4}He.

We first note again that ^{5}Li is unbound. Although this nucleus is continuously formed by p+^{4}He reactions, the scattering is elastic and the formed ^{5}Li nuclei decay within 10^{-22} s.

As a consequence ^{4}He 'survives' in the core until sufficiently large temperatures are achieved to overcome the larger Coulomb barrier between ^{4}He nuclei. Unfortunately the ^{8}Be ground state, formed by elastic ^{4}He+^{4}He scattering, is a resonance too and decays within 10^{-16} s back to two ^{4}He nuclei.

The Salpeter-Hoyle suggestion

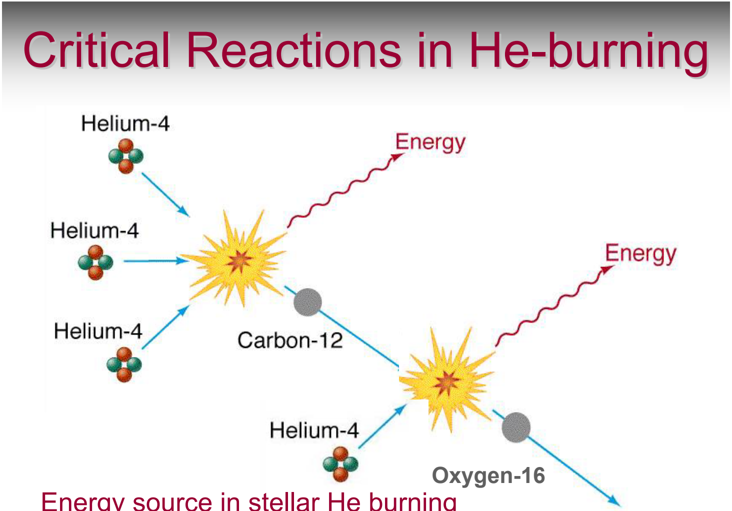
In 1952 Salpeter pointed out that the ^{8}Be lifetime might be sufficiently large that there is a chance that it captures another ^{4}He nucleus. This *triple-alpha reaction*

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

can then form ^{12}C and supply energy. However, the simultaneous collision of 3 ^{4}He (α-particles) is too rare to give the burning rate necessary in stellar models. So Hoyle predicted a resonance in ^{12}C to speed up the collision. And indeed, this *Hoyle state* was experimentally observed shortly after its prediction.
^{12}C can then react with another ^{4}He nucleus forming ^{16}O via

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

These two reactions make up helium burning.



Energy source in stellar He burning

Energy release determined by associated reaction rates

At the end of helium burning

Nucleosynthesis yields from stars may be divided into production by stars above or below $9M_\odot$.

- **stars with** $M \lesssim 9M_\odot$
  the stars are expected to shed their envelopes during helium burning and become white dwarfs. Most of the matter returned to the ISM is unprocessed.
- **stars with** $M > 9M_\odot$
  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 ISM.

Carbon Burning

Burning conditions:

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

T~ 600-700 Mio
ρ ~ 10^5-10^6 g/cm^3

Major reaction sequences:

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

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) 21

Neon Burning

Neon burning is very similar to carbon 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

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

this is followed by (using the liberated helium)

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

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

22

Oxygen Burning

Burning conditions:

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

Major reaction sequences:

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

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

27

Silicon Burning

Silicon burning is very similar to oxygen 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)

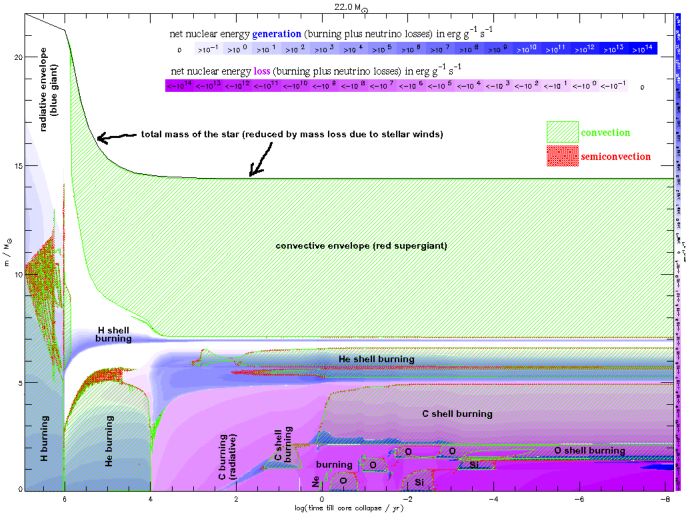
Nuclear burning stages

(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	$^{28}Si(\gamma,\alpha)$…

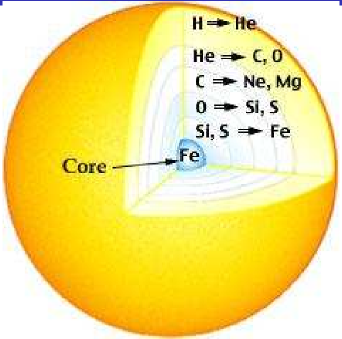
Kippenhahn diagram for a 22 $M_\odot$ star

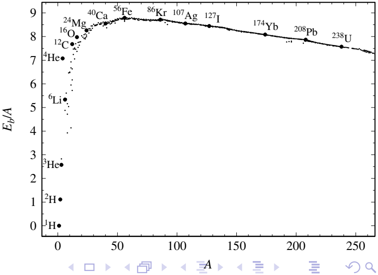
(A. Heger and S. Woosley)



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 once it grows above around 1.4 solar masses.





The early iron core

- The core is made of heavy nuclei (iron-mass range $A \sim 45 - 65$) and electrons. There are Y_e electrons per nucleon.
- The mass of the core M_c is determined by the nucleons.
- There is no nuclear energy source which adds to the pressure. Thus, the pressure is mainly due to the degenerate electrons, with a small correction from the electrostatic interaction between electrons and nuclei.
- As long as $M_c < M_{ch} = 1.44(2Y_e)^2 M_\odot$ (plus slight corrections for finite temperature), the core can be stabilized by the degeneracy pressure of the electrons. M_{ch} is the famous Chandrasekhar mass which sets the limit mass of White Dwarfs - dead stars which are stabilized by the degeneracy pressure of electrons.

Initial collapse conditions

If we approximate the pressure by the one of a relativistic degenerate electron gas ($P \sim \rho^{4/3}$) one has

$$P/\rho \approx \frac{1}{4} Y_e \mu_e$$

$$\mu_e \approx 1.1 (\rho_7 Y_e)^{1/3} \ \mathrm{MeV}$$

Hence P/ρ is given in MeV, with

$$1 \text{ MeV} = 0.96 \times 10^{18} \text{ erg/g}$$

Note that the electron chemical potential μ_e is nearly 1 MeV at $\rho_7 = 1$ and hence reaches the nuclear energy scale. Hence it might be energetically favorable to capture high-energy electrons by nuclei.

Onset of collapse

However, there are two processes which make the situation unstable.

1. Silicon burning is continuing in a shell around the iron core. This adds mass to the iron core, thus M_c grows.
2. Electrons can be captured by nuclei.

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

   This reduces the pressure and cools the core, as the neutrinos leave. In other words, Y_e and hence the Chandrasekhar mass M_{ch} is reduced.

The core finally collapses.

Electron capture

The cross section for electron capture on *free protons* at rest is

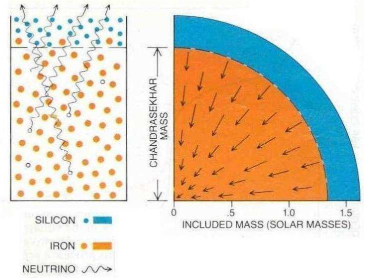
$$\sigma_p = 4.5 \times 10^{-44} E_\nu^2 \ \mathrm{cm}^2$$

where E_ν is the energy of the emitted neutrino in MeV. The rate of electron capture on free protons then is

$$r = \sigma_p N_A Y_p = 0.016 \rho_7 E_\nu^2 Y_p [s^{-1}]$$

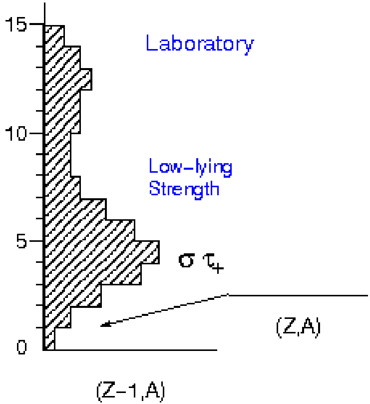
Although the capture cross section for those nuclei present in the core is usually smaller than the one of free protons (due to the larger energy threshold between parent and daughter nucleus), the abundance of free protons is quite low, so that the total electron capture rate is dominated by nuclei. This is an interesting nuclear structure problem which was first tackled within the Independent Particle Model (IPM) and then within the interacting shell model.

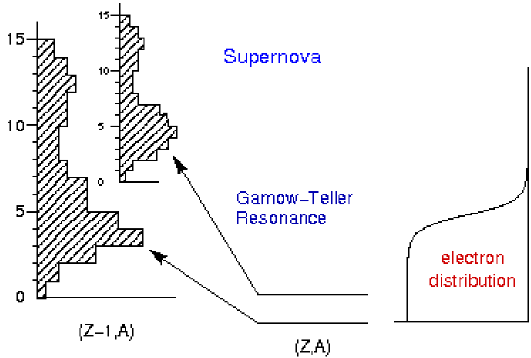
Presupernova evolution



- $T = 0.1\text{–}0.8\ \mathrm{MeV}$, $\rho = 10^7\text{–}10^{10}\ \mathrm{g\ cm^{-3}}$. Composition of iron group nuclei (A = 45–65)
- 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)
- Calculated within large-scale shell model, constrained by data from charge-exchange (d,^{2}He), (n,p) experiments

Laboratory vs stellar electron capture



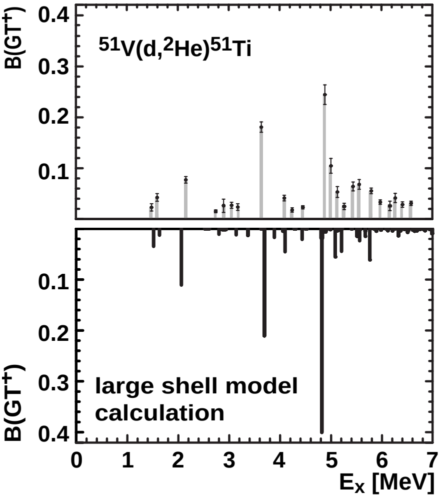


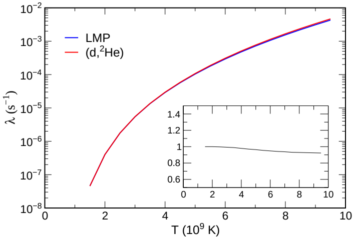
capture of K-shell electrons to tail of GT strength distribution; parent nucleus in ground state

capture of electrons from high-energy tail of FD distribution; capture of strong GT transitions possible; thermal ensemble of initial states

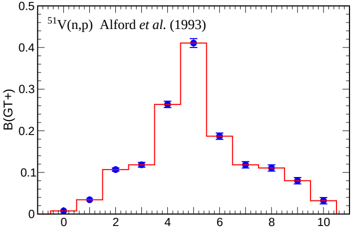
Shell model and (d,^{2}He) GT strengths

C. Bäumer *et al.* PRC **68**, 031303 (2003)





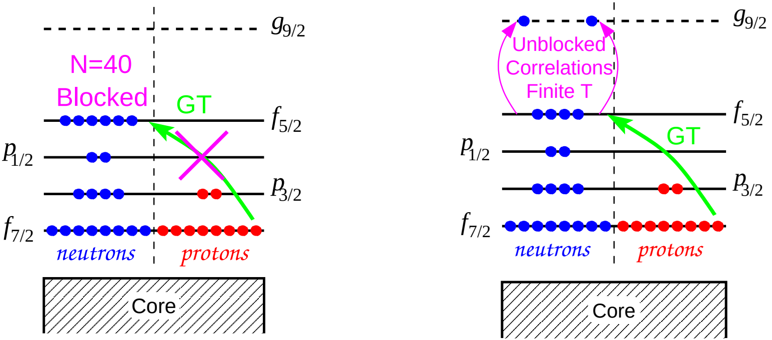
Old (n, p) data



Consequences of electron capture

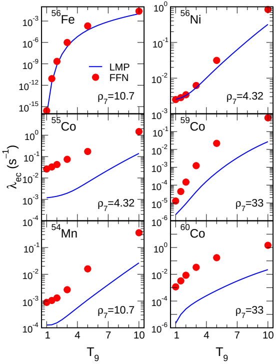
- Electron captures reduce the number of electrons and hence the pressure force against the gravitational collapse.
- Neutrinos carry energy out of the core. The core is kept cool and at a low entropy.
- Low entropy implies high order. Nuclei survive during the collapse phase.
- Electron captures also reduce the number of protons, but keeps the number of nucleons conserved.
- Nuclei become more neutron rich. Large neutron excess favors heavy nuclei. The nuclear composition is driven to more neutron-rich and heavy nuclei.
- Why do the neutron-rich nuclei not decay? β-decays are strongly suppressed by Pauli blocking as the electron chemical is high and effectively closes the phase space.

Pauli blocking of Gamow-Teller transition

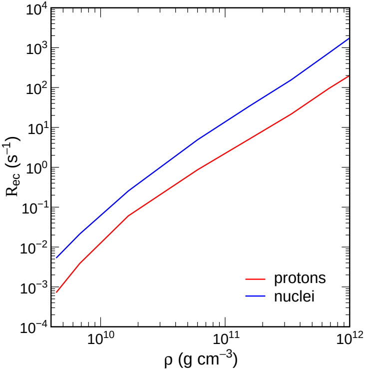


- Unblocking mechanism: correlations and finite temperature
- calculation of rate in SMMC + RPA model

How do shell-model rates compare to previous rates?



$A < 65$: SM rates smaller



$A > 65$: SM rates larger

Thus at 'low' densities electron capture is slower than previously believed, but at high densities previous simulations had switched off the main cooling mechanism (capture on nuclei).