

Lecture 12

Advanced Stages of Stellar Evolution – II Silicon Burning and NSE

The solution of reaction networks.

Consider the reaction pair:

$$\frac{dY_1}{dt} = -Y_1\lambda_1 + Y_2\lambda_2$$

$$\frac{dY_2}{dt} = Y_1\lambda_1 - Y_2\lambda_2$$

An explicit solution of the linearized equations would be

$Y_{new} = Y_{old} + \delta Y$ where

$$\delta Y_1 = (-Y_1\lambda_1 + Y_2\lambda_2) \Delta t$$

$$\delta Y_2 = (Y_1\lambda_1 - Y_2\lambda_2) \Delta t$$

For large Δt , the answer could oscillate.

In the usual case that the two species were not in equilibrium $Y_1\lambda_1 \neq Y_2\lambda_2$, a large time step, $\Delta t \rightarrow \infty$, would lead to a divergent value for the change in Y, including negative values.

On the other hand, forward or "implicit" differencing would give

$$\frac{\delta Y_1}{\Delta t} = -(Y_1 + \delta Y_1)\lambda_1 + (Y_2 + \delta Y_2)\lambda_2$$

$$\frac{\delta Y_2}{\Delta t} = -(Y_2 + \delta Y_2)\lambda_2 + (Y_1 + \delta Y_1)\lambda_1$$

$$\delta Y_1 \left(\frac{1}{\Delta t} + \lambda_1 \right) + \delta Y_2 (-\lambda_2) = -Y_2\lambda_2 + Y_1\lambda_1$$

$$\delta Y_1 (-\lambda_1) + \delta Y_2 \left(\frac{1}{\Delta t} + \lambda_2 \right) = Y_2\lambda_2 - Y_1\lambda_1$$

In general an
n x n matrix
n = 2 here

Add equations $\Rightarrow \delta Y_1 = -\delta Y_2$; substituting, one also has :

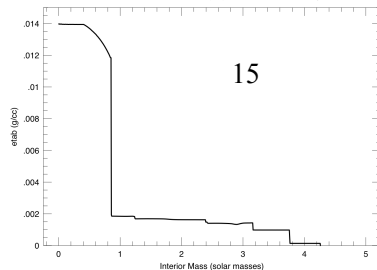
$$\delta Y_1 = \left(\frac{Y_1\lambda_1 - Y_2\lambda_2}{1/\Delta t + \lambda_1 + \lambda_2} \right) \quad (\text{if } 1/\Delta t \gg \lambda, \text{ same as the explicit solution})$$

Even if $\Delta t \rightarrow \infty$ the change in Y is finite and converges on the equilibrium value $Y_1\lambda_1 = Y_2\lambda_2$

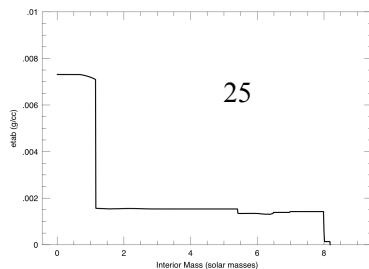
Silicon Burning

Silicon burning proceeds in a way different from any nuclear process discussed so far. It is analogous, in ways, to neon burning in that it proceeds by photodisintegration and rearrangement, but it involves many more nuclei and is quite complex.

The reaction $^{28}\text{Si} + ^{28}\text{Si} \rightarrow (^{56}\text{Ni})^*$ does not occur owing to the large Coulomb inhibition. Rather a portion of the silicon (and sulfur, argon, etc.) “melt” by photodisintegration reactions into a sea of neutrons, protons, and alpha-particles. These lighter constituents add onto the remaining silicon and heavier elements, gradually increasing the mean atomic weight until species in the iron group are most abundant.



Neutron excess after oxygen core depletion in 15 and 25 solar mass stars.



The inner core is becoming increasingly “neutronized”, especially the lower mass stars.

Initial Composition for Silicon Burning

The initial composition is mostly Si and S but which isotopes depends on whether one is discussing the inner core or locations farther out in the star. It is quite different for silicon core burning in a presupernova star and the explosive variety of silicon burning we will discuss later.

In the *center of the star*, one typically has, after oxygen burning, and a phase of electron capture that goes on between oxygen depletion and silicon ignition:

^{30}Si , ^{34}S , ^{38}Ar and a lot of other less abundant nuclei

Farther outside of the core where silicon might burn explosively, one has species more characteristically with $Z = N$

^{28}Si , ^{32}S , ^{36}Ar , ^{40}Ca , etc.

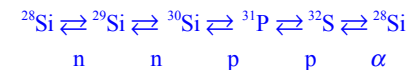
Historically, Si burning was discussed for a ^{28}Si rich composition

Quasi-equilibrium

This term is used to describe a situation where groups of adjacent isotopes, but not all isotopes globally, have come into equilibrium with respect to the exchange of n, p, a, and g.

It begins in neon burning with $^{20}\text{Ne} + \gamma \rightleftharpoons ^{16}\text{O} + \alpha$ and continues to characterize an increasing number of nuclei during oxygen burning. In silicon burning, it becomes the rule rather than the exception.

A typical "quasiequilibrium cluster" might include the equilibrated reactions:



By which one means $Y(^{28}\text{Si}) Y_n \rho \lambda_{n\gamma} (^{28}\text{Si}) \approx Y(^{29}\text{Si}) \lambda_{\gamma n} (^{29}\text{Si})$

$$Y(^{30}\text{Si}) Y_n \rho \lambda_{n\gamma} (^{29}\text{Si}) \approx Y(^{30}\text{Si}) \lambda_{\gamma n} (^{30}\text{Si})$$

etc.

Late during oxygen burning, many isolated clusters grow and merge until, at silicon ignition, there exist only two large QE groups

$$24 \leq A \leq 45 \quad 46 \leq A \leq 60 \quad (\text{at least})$$

Reactions below ^{24}Mg , e.g., $^{20}\text{Ne}(\alpha, \gamma)^{24}\text{Mg}$ and $^{12}\text{C}(\alpha, \gamma)^{16}\text{O}$ are, in general, *not* in equilibrium with their inverses (exception, $^{16}\text{O}(\alpha, \gamma)^{20}\text{Ne}$ which has been in equilibrium since neon burning).

Within the groups heavier than $A = 24$, except at the boundaries of the clusters, the abundance of any species is related to that of another by successive application of the Saha equation.

$$\begin{aligned} \text{e.g., } \frac{Y(^{40}\text{Ca})}{Y(^{28}\text{Si})} &= \left(\frac{Y(^{32}\text{S})}{Y(^{28}\text{Si})} \right) \left(\frac{Y(^{36}\text{Ar})}{Y(^{32}\text{S})} \right) \left(\frac{Y(^{40}\text{Ca})}{Y(^{36}\text{Ar})} \right) \\ &= \left(\frac{\rho Y_\alpha \lambda_{\alpha\gamma}(^{28}\text{Si})}{\lambda_{\gamma\alpha}(^{32}\text{S})} \right) \left(\frac{\rho Y_\alpha \lambda_{\alpha\gamma}(^{32}\text{S})}{\lambda_{\gamma\alpha}(^{36}\text{Ar})} \right) \left(\frac{\rho Y_\alpha \lambda_{\alpha\gamma}(^{36}\text{Ar})}{\lambda_{\gamma\alpha}(^{40}\text{Ca})} \right) \\ &= f(T, Q_{\alpha\gamma}) \rho^3 Y_\alpha^3 \quad \text{etc.} \end{aligned}$$

Within that one group, which contains ^{28}Si and the vast majority of the mass, one can evaluate any abundance relative to e.g., ^{28}Si

$$Y(^A Z) = C(^A Z, \rho, T_9) Y(^{28}\text{Si}) Y_\alpha^{\delta_\alpha} Y_p^{\delta_p} Y_n^{\delta_n}$$

where $\delta_\alpha = \text{largest integer} \leq \frac{Z-14}{2}$ *Need 6 parameters: Y_α, Y_p, Y_n and $Y(^{28}\text{Si})$ plus T and ρ , but ...*

$$\delta_n = N - 14 - 2\delta_\alpha$$

$$\delta_p = Z - 14 - 2\delta_\alpha$$

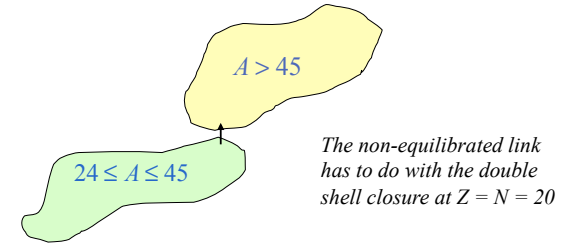
e.g., ^{35}Cl 17 protons; 20 neutrons

$$\delta_\alpha = 1 \quad \delta_p = 1 \quad \delta_n = 2$$

^{40}K 19 protons 21 neutrons

$$\delta_\alpha = 2 \quad \delta_p = 1 \quad \delta_n = 3$$

The situation at the end of oxygen burning is that there are two large QE groups coupled by non-equilibrated links near $A = 45$.



Early during silicon burning these two groups merge and the only remaining non-equilibrated reactions are for $A < 24$.

$$C(^A Z) = (\rho N_A)^{\delta_\alpha + \delta_p + \delta_n} C'(^A Z)$$

$$C'(^A Z) = (5.942 \times 10^{33} T_9^{3/2})^{-(\delta_\alpha + \delta_p + \delta_n)}$$

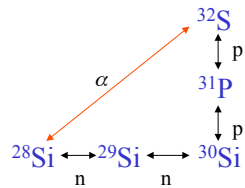
$$\frac{G(^A Z)}{G(^{28}\text{Si})} 2^{-(\delta_p + \delta_n)} \left(\frac{A}{28} \right)^{3/2} \left(\frac{1}{4} \right)^{3\delta_\alpha/2} \exp(Q / kT)$$

where

$$Q = BE(^A Z) - BE(^{28}\text{Si}) - \delta_\alpha BE(\alpha)$$

i.e., the energy required to dissociate the nucleus $^A Z$ into ^{28}Si and δ_α alpha particles. The binding energy of a neutron or proton is zero.

Moreover there exist loops like:



For constant ρ and T

$$Y(^{32}\text{S}) = \text{const} \cdot Y(^{30}\text{Si})Y_p^2$$

$$Y(^{30}\text{Si}) = \text{const}' \cdot Y(^{28}\text{Si})Y_n^2$$

$$Y(^{32}\text{S}) = \text{const}'' \cdot Y(^{28}\text{Si})Y_\alpha$$

$$Y_\alpha \propto \left(\frac{Y(^{32}\text{S})}{Y(^{28}\text{Si})} \right) \propto \left(\frac{Y(^{32}\text{S})}{Y(^{30}\text{Si})} \right) \left(\frac{Y(^{30}\text{Si})}{Y(^{28}\text{Si})} \right) \propto Y_p^2 Y_n^2$$

$$Y_\alpha = C_\alpha(T_9, \rho) Y_n^2 Y_p^2$$

$$C_\alpha = \frac{1}{2} (5.94 \times 10^{33})^{-3} \rho^3 T_9^{9/2} \exp(328.36 / T_9)$$

where $328.36 = \text{BE}(\alpha)/kT$

This reduces the number of independent variables to 5, but wait ...

The large QE cluster that includes nuclei from $A = 24$ through at least $A = 60$ contains most of the matter (^{20}Ne , ^{16}O , ^{12}C , and α are all small), so we have the additional two constraints

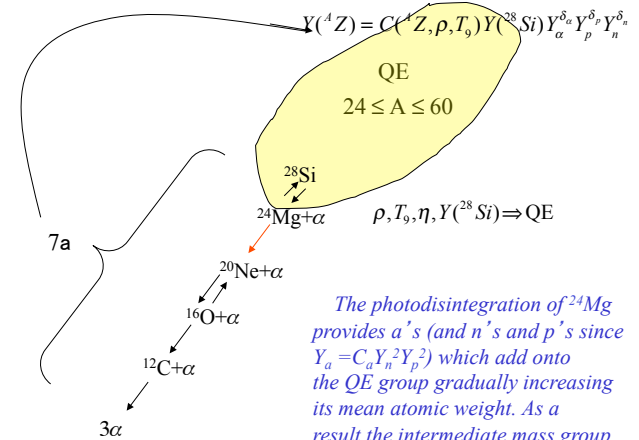
$$\sum_{60 \geq A \geq 24} A_i Y_i \approx 1 \quad \text{mass conservation}$$

$$\sum_{60 \geq A \geq 24} (N_i - Z_i) Y_i \approx \eta \quad \text{charge conservation}$$

The first equation can be used to eliminate one more unknown, say Y_p , and the second can be used to replace Y_n with an easier to use variable, η . *Thus 4 variables now specify the abundances of all nuclei heavier than magnesium.* These are

$$\rho, T_9, \eta, \text{ and } Y(^{28}\text{Si})$$

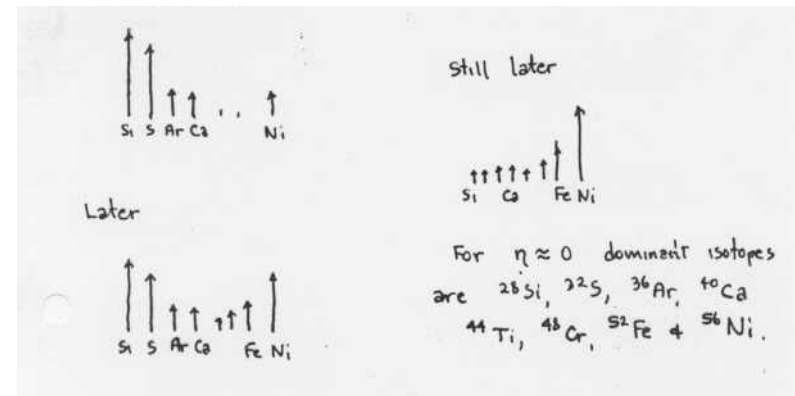
The cluster evolves at a rate given by $^{24}\text{Mg}(\gamma, \alpha)^{20}\text{Ne}$



The photodisintegration of ^{24}Mg provides α 's (and n 's and p 's since $Y_\alpha = C_\alpha Y_n^2 Y_p^2$) which add onto the QE group gradually increasing its mean atomic weight. As a result the intermediate mass group, Si-Ca gradually "melts" into the iron group.

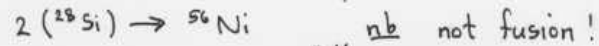
Nature of the burning:

Lighter species melt away while the iron group grows



Energetics and Approximations

Gross approximation



$$q_{\text{nuc}} = 9.65 \times 10^{17} \left[\frac{1}{2} (483.982) - 236.531 \right] / 28$$

$$= 1.9 \times 10^{17} \text{ erg g}^{-1} \quad (\text{not much})$$

But this assumes ${}^{28}\text{Si}$ burns to ${}^{56}\text{Ni}$ ($\eta = 0.002$ approximation)

This is all a bit misleading because, except explosively (later), silicon burning does not produce ${}^{56}\text{Ni}$. There has been a lot of electron capture during oxygen burning and more happens in silicon burning.

| | | |
|--|-----------------------|--------------------|
| E.g., Si ignition in a $15 M_{\odot}$ star | $\eta_c \approx 0.07$ | $Y_e \approx 0.46$ |
| Si depletion | $\eta_c \approx 0.13$ | $Y_e \approx 0.44$ |

Under these conditions silicon burning produces ${}^{54}\text{Fe}$, ${}^{56}\text{Fe}$, ${}^{58}\text{Fe}$ and other neutron rich species in the iron group.

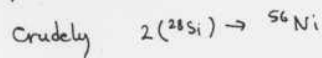
Suppose $\left(\frac{56}{30}\right) {}^{30}\text{Si} \rightarrow {}^{56}\text{Fe} \quad \text{i.e., } q_{\text{nuc}} = 9.65 \times 10^{17} \sum \frac{X_i}{A_i} BE(A_i)$

$$q_{\text{nuc}} = 9.65 \times 10^{17} \left[(492.26) / 56 - (255.62) / 30 \right]$$

$$= 2.6 \times 10^{17} \text{ erg g}^{-1} \text{ which is closer to correct for Si}$$

core burning than $1.9 \times 10^{17} \text{ erg g}^{-1}$

Energy Generation Rate



$$\epsilon_{\text{nuc}}^{\text{Si}} = (9.65 \times 10^{17}) (5.46) \frac{dY({}^{28}\text{Si})}{dt}$$

$$= 5.3 \times 10^{18} \frac{dY({}^{28}\text{Si})}{dt}$$

$$\frac{dY({}^{28}\text{Si})}{dt} = -2 Y({}^{24}\text{Mg}) \lambda_{\gamma\alpha}({}^{24}\text{Mg})$$

↑ one photodisintegrates
one goes to Ni

This is very like neon burning except that 7 alpha-particles are involved instead of one.

Reaction rates governing the rate at which silicon burns:

Generally speaking, the most critical reactions will be those connecting equilibrated nuclei with $A > 24$ (magnesium) with alpha-particles. The answer depends on temperature and neutron excess:

Most frequently, for η small, the critical slow link is ${}^{24}\text{Mg}(\gamma, \alpha){}^{20}\text{Ne}$

The reaction ${}^{20}\text{Ne}(\gamma, \alpha){}^{16}\text{O}$ has been in equilibrium with ${}^{16}\text{O}(\alpha, \gamma){}^{20}\text{Ne}$ ever since neon burning. At high temperatures and low Si-mass fractions, ${}^{20}\text{Ne}(\alpha, \gamma){}^{24}\text{Mg}$ equilibrates with ${}^{24}\text{Mg}(\gamma, \alpha){}^{20}\text{Ne}$ and ${}^{16}\text{O}(\gamma, \alpha){}^{12}\text{C}$ becomes the critical link.

However for the values of η actually appropriate to silicon burning in a massive stellar core, the critical rate is ${}^{26}\text{Mg}(p, \alpha){}^{23}\text{Na}(p, \alpha){}^{20}\text{Ne}$

$$Y(^{24}\text{Mg}) = Y(^{28}\text{Si}) \frac{\lambda_{\alpha} (^{28}\text{Si})}{\rho Y_{\alpha} \lambda_{\alpha} (^{24}\text{Mg})}$$

To get ^{24}Mg

$$= 9.87 \times 10^9 T_9^{3/2} \left(\frac{24X}{28}\right)^{3/2} e^{-11.605 Q_{\alpha} (^{24}\text{Mg})/T_9} \cdot \frac{Y(^{28}\text{Si})}{\rho Y_{\alpha}}$$

$$= 6.3 \times 10^{10} e^{-115.83} \frac{T_9^{3/2}}{\rho Y_{\alpha}} Y(^{28}\text{Si})$$

what is ρY_{α} ?

Assume $X(^{28}\text{Si}) + X(^{56}\text{Ni}) = 1$ let $X(^{28}\text{Si}) = f$

To get α

$$\frac{Y(^{56}\text{Ni})}{Y(^{28}\text{Si})} = \frac{1-f}{2f} = C(^{56}\text{Ni}, \rho, T_9) Y_{\alpha}^7$$

$$C(^{56}\text{Ni}, \rho, T_9) = (\rho N_A)^7 (5.94 \times 10^{33})^{-7} T_9^{-21/2} \left(\frac{56}{28}\right)^{3/2} \left(\frac{1}{4}\right)^{21/2} \exp [BE(^{56}\text{Ni}) - BE(^{28}\text{Si}) - 7 BE(\alpha)] / kT$$

$$\log_{10} C_{56} = -242.288 - \frac{21}{2} \log_{10} T_9 + 248.936/T_9 + 7 \log_{10} (\rho N_A)$$

a) solve for ρY_{α} in terms of f

b) substitute $Y(^{24}\text{Mg})$ as f^2 of $Y(^{28}\text{Si})$

c) express $\lambda_{\alpha} (^{24}\text{Mg}) \dots \lambda_{\alpha} (^{20}\text{Ne})$

get

$$\frac{dY(^{28}\text{Si})}{dt} = -Y(^{28}\text{Si}) Q(^{28}\text{Si})$$

$$Q(^{28}\text{Si}) = 9.55 \times 10^{10} T_9^{3/2} \left(\frac{2f}{1-f}\right)^{17} e^{-142.06/T_9} \lambda_{\alpha} (^{20}\text{Ne}) \approx 1$$

independent of ρ .

$$\epsilon_{\text{nuc}}^{Si} \approx 1.8 \times 10^{28} T_9^{3/2} X(^{28}\text{Si}) e^{-142.06/T_9} \lambda_{\alpha} (^{20}\text{Ne})$$

$$\lambda_{\alpha} (^{20}\text{Ne}) \approx 124 \left(\frac{T_9}{3.5}\right)^5$$

$$\epsilon_{\text{nuc}}^{Si} \propto T_9^{3.5} \propto T^{47} \quad 47 = 5 + 1.5 + \frac{142.06}{3.5}$$

Balanced Power

$$\epsilon_{\text{nuc}} \propto \rho^0 T^{47} \quad \epsilon_{\nu} \propto \rho^{-1} T^9$$

$$\frac{\langle \epsilon_{\text{nuc}} \rangle}{\langle \epsilon_{\nu} \rangle} = 1 \Rightarrow \epsilon_{\text{nuc}} = 13.1 \epsilon_{\nu} \quad \text{take } X(^{28}\text{Si}) = 0.5$$

| T_9 | ϵ_{nuc} | $\epsilon_{\nu} (\rho = 10^7)$ | $\epsilon_{\nu} (\rho = 10^8)$ |
|-------|-------------------------|--------------------------------|--------------------------------|
| 3.0 | 7.0(9) | 4.1(12) | 1.1(11) |
| 3.5 | 1.7(13) | 2.0(13) | 7.8(11) |
| 4.0 | 6.2(15) | 7.3(13) | 3.7(12) |

So burn Si at about $T_9 \approx 3.5$

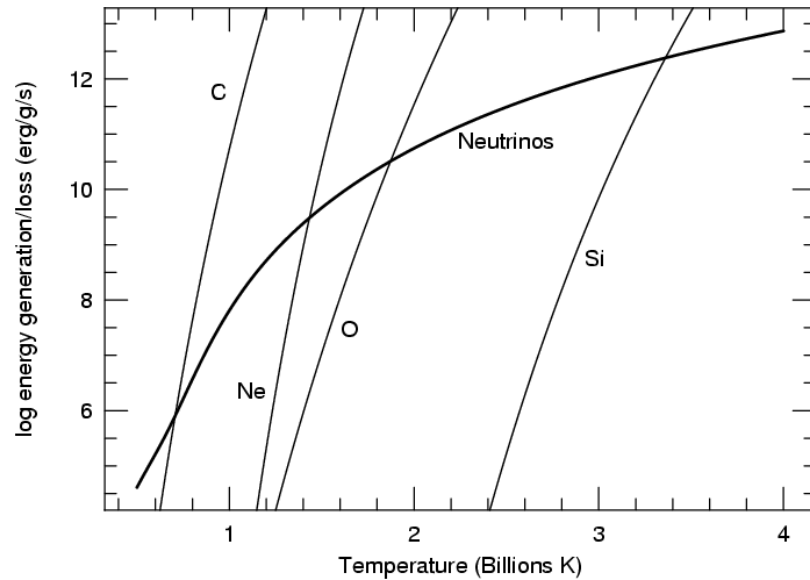
Lifetime

$$\frac{1}{\tau_{28}} = \frac{1}{128} \frac{dY_{28}}{dt}$$

$$= 9.5 \times 10^{10} (3.5)^{3/2} (124) e^{-142.06/3.5}$$

$$= 1.8 \times 10^{-4} \Rightarrow \tau = 1.5 \text{ hr.}$$

lengthened by convection to ~ 1 day



Nucleosynthesis

Basically, silicon burning in the star's core turns the products of oxygen burning (Si, S, Ar, Ca, etc.) into the most tightly bound nuclei (in the iron group) for a given neutron excess, η .

The silicon-burning nucleosynthesis that is ejected by a supernova is produced explosively, and has a different composition dominated by ^{56}Ni .

The products of silicon-core and shell burning in the core are both so neutron-rich (η so large) that they need to be left behind in a neutron star or black hole. However, even in that case, the composition and its evolution is critical to setting the stage for core collapse and the supernova explosion that follows.

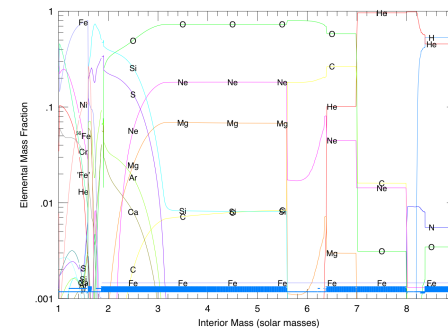
Silicon burning nucleosynthesis

Following Si-burning at the middle of a 25 solar mass star:

| | | |
|------------------|-------|---------------------------------------|
| ^{54}Fe | 0.487 | |
| ^{58}Ni | 0.147 | |
| ^{56}Fe | 0.141 | Neutron-rich nuclei in the iron peak. |
| ^{55}Fe | 0.071 | $Y_e = 0.4775$ |
| ^{57}Co | 0.044 | |

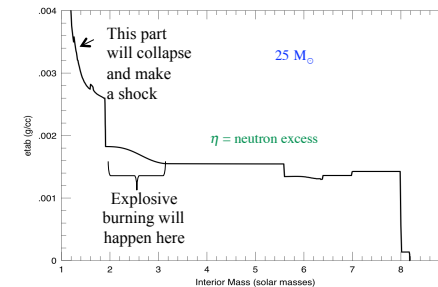
Following *explosive* Si-burning in a 25 solar mass supernova, interesting species produced at $Y_e = 0.498$ to 0.499.

| <u>product</u> | <u>parent</u> | |
|---------------------------|------------------------|---|
| ^{44}Ca | ^{44}Ti | ^{44}Ti and $^{56,57}\text{Ni}$ are important targets of γ -ray astronomy |
| $^{47,48,49}\text{Ti}$ | $^{48,49}\text{Cr}$ | |
| ^{51}V | ^{51}Cr | |
| ^{55}Mn | ^{55}Co | |
| $^{50,52,53}\text{Cr}$ | $^{52,53}\text{Fe}$ | |
| $^{54,56,57}\text{Fe}$ | $^{56,57}\text{Ni}$ | |
| ^{59}Co | ^{59}Cu | |
| $^{58,60,61,62}\text{Ni}$ | $^{60,61,62}\text{Zn}$ | |



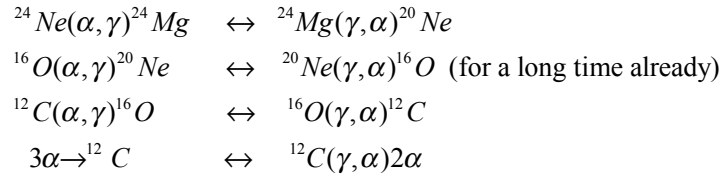
25 Solar Mass
Pop I
PreSupernova

(Woosley and Heger 2007)



Nuclear Statistical Equilibrium

As the silicon abundance tends towards zero (though it never becomes microscopically small), the unequilibrated reactions below $A = 24$ finally come into equilibrium



The 3α reaction is the last to equilibrate. Once this occurs, **every** isotope is in equilibrium with every other isotope by strong and electromagnetic reactions (but not by weak interactions)

Until the temperature becomes very high ($kT \gg 1 \text{ MeV}$)
 The most abundant nuclei are those with large binding energy per nucleon and "natural" values of η . For example,

$$\begin{aligned}
 \eta = 0 & \quad {}^{56}\text{Ni} \\
 0.037 & \quad {}^{54}\text{Fe} \\
 0.071 & \quad {}^{56}\text{Fe} \\
 & \text{etc.}
 \end{aligned}$$

In general, the abundance of an isotope peaks at its natural value for η . E. g.,

$$\eta({}^{56}\text{Fe}) = \frac{N-Z}{A} = \frac{30-26}{56} = 0.0714$$

In particular, $Y({}^{28}\text{Si}) = f(T, \rho) Y_\alpha^7$ with the result that now only **3 variables, ρ , T_9 , and η** now specify the abundances of everything

$$Y({}^AZ) = C({}^AZ, \rho, T_9) Y_n^N Y_p^Z$$

$$C({}^AZ, \rho, T_9) = (\rho N_A)^{A-1} C'({}^AZ, T_9)$$

$$C'({}^AZ, T_9) = \frac{G({}^AZ, T_9)}{2^A} \theta^{1-A} \exp[BE({}^AZ) / kT]$$

$$\theta = 5.943 \times 10^{33} T_9^{3/2}$$

$G({}^AZ, T_9)$ is the temperature-dependent partition function.

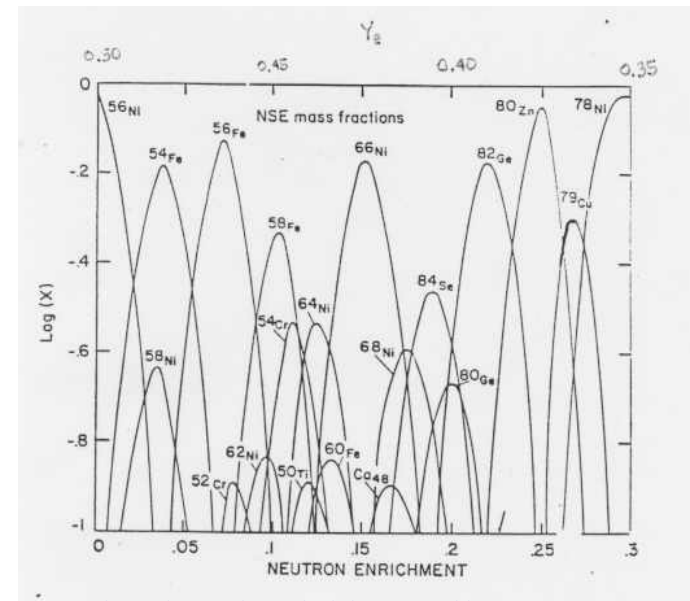
At low T

$$G({}^AZ, T_9) = \sum (2J_i + 1) e^{-E_i/kT}$$

At high T, though (see earlier discussion of nuclear level density)

$$G({}^AZ, T_9) \approx \frac{\pi}{6akT} e^{a(kT)}$$

$$a \approx \frac{A}{9} \text{ MeV}^{-1}$$



The resultant nucleosynthesis is most sensitive to η .

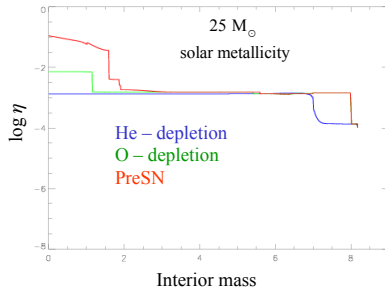
True Equilibrium

If the weak interactions were also to be balanced, (e.g., neutrino capture occurring as frequently on the daughter nucleus as electron capture on the parent), one would have a state of true equilibrium. Only two parameters, r and T , would specify the abundances of everything. The last time this occurred in the universe was for temperatures above 10 billion K in the Big Bang.

However, one can also have a *dynamic weak equilibrium* where neutrino emission balances anti-neutrino emission, i.e., when

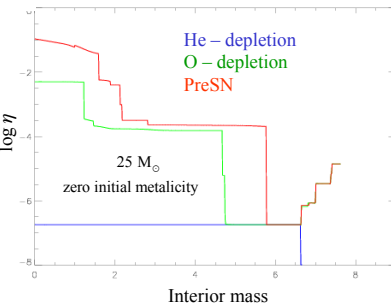
$$\frac{dY_e}{dt} = 0$$

This could occur, and for some stars does, when electron-capture balances beta-decay globally, but not on individual nuclei. The abundances would be set by ρ and T , but would also depend on the weak interaction rate set employed.



The distribution of neutron excess, η , within two stars of 25 solar masses (8 solar mass helium cores) is remarkably different. In the Pop I star, η is approximately 1.5×10^{-3} everywhere except in the inner core (destined to become a collapsed remnant)

In the Pop III ($Z = 0$) star the neutron excess is essentially zero at the end of helium burning (some primordial nitrogen was created) Outside of the core η is a few $\times 10^{-4}$, chiefly from weak interactions during carbon burning. Note some primary nitrogen production at the outer edge where convection has mixed ^{12}C and protons.



Weak Interactions

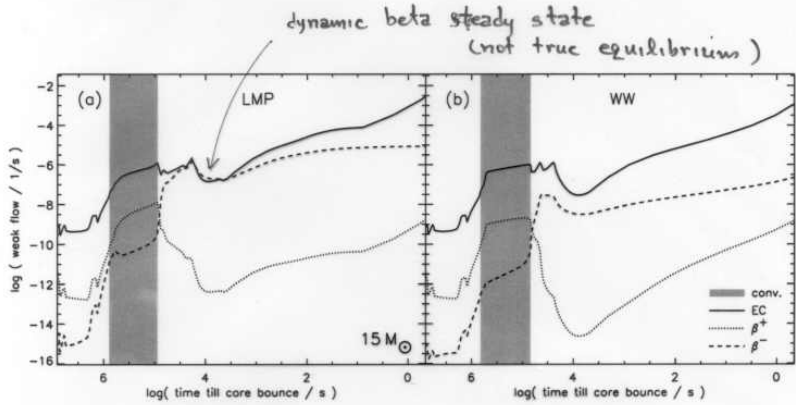
Electron capture, and at late times beta-decay, occur for a variety of isotopes whose identity depends on the star, the weak reaction rates employed, and the stage of evolution examined. During the late stages it is most sensitive to η , the neutron excess.

Aside from their nucleosynthetic implications, the weak interactions determine Y_e , which in turn affects the structure of the star. The most important isotopes are not generally the most abundant, but those that have some combination of significant abundance and favorable nuclear structure (especially Q-value) for weak decay.

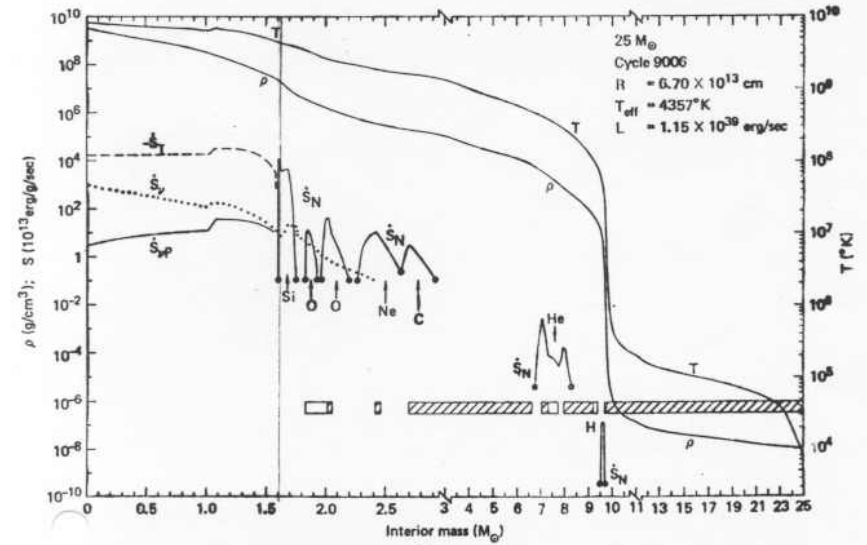
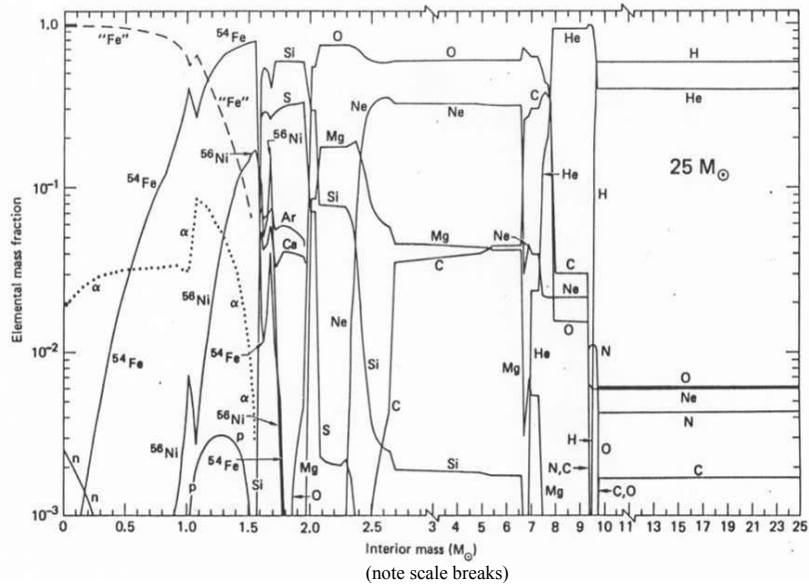
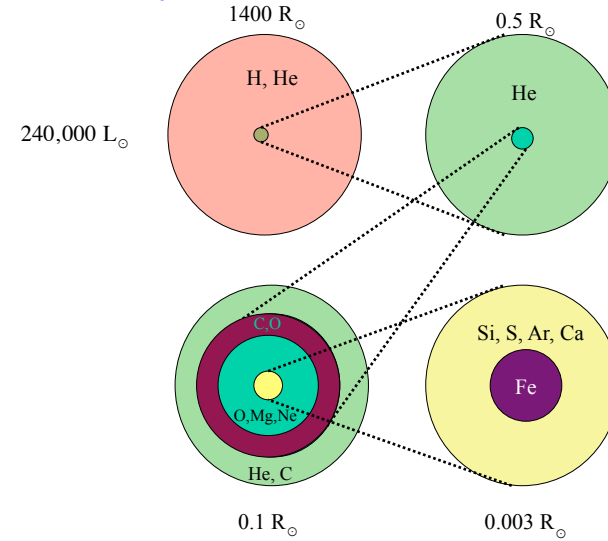
From silicon burning onwards these weak decays provide neutrino emission that competes with and ultimately dominates that from thermal processes (i.e., pair annihilation).

| | <u>O-depletion</u> | <u>O-shell</u> | <u>Si-ignition</u> | <u>Si-shell</u> |
|--------------------------|----------------------------------|---------------------------------|---------------------------------|---------------------|
| $T(10^9\text{K})$ | 2.26 | 1.90 | 2.86 | 3.39 |
| $\rho(\text{g cm}^{-3})$ | 1.2×10^7 | 2.8×10^7 | 1.1×10^8 | 4.5×10^7 |
| Y_e | 0.498 | 0.495 | 0.489 | 0.480 |
| e-capture | $^{35}\text{Cl}, ^{37}\text{Ar}$ | $^{35}\text{Cl}, ^{33}\text{S}$ | $^{33}\text{S}, ^{35}\text{Cl}$ | $^{54,55}\text{Fe}$ |
| β -decay | $^{32}\text{P}, ^{36}\text{Cl}$ | $^{32}\text{P}, ^{36}\text{Cl}$ | $^{32}\text{P}, ^{28}\text{Al}$ | $^{54,55}\text{Mn}$ |

| | <u>Si-depletion</u> | <u>Si-shell burn</u> | <u>Core contraction</u> | <u>PreSN</u> |
|--------------------------|----------------------------------|----------------------------------|----------------------------------|----------------------------------|
| $T(10^9\text{K})$ | 3.78 | 4.13 | 3.55 | 7.16 |
| $\rho(\text{g cm}^{-3})$ | 5.9×10^7 | 3.2×10^8 | 5.4×10^8 | 9.1×10^9 |
| Y_e | 0.467 | 0.449 | 0.445 | 0.432 |
| e-capture | $^{54,55}\text{Fe}$ | $^{57}\text{Fe}, ^{61}\text{Ni}$ | $^{57}\text{Fe}, ^{55}\text{Mn}$ | $^{65}\text{Ni}, ^{59}\text{Fe}$ |
| β -decay | $^{54}\text{Mn}, ^{53}\text{Cr}$ | $^{56}\text{Mn}, ^{52}\text{V}$ | $^{62}\text{Co}, ^{58}\text{Mn}$ | $^{64}\text{Co}, ^{58}\text{Mn}$ |



25 M_{\odot} Presupernova Star (typical for 9 - 130 M_{\odot})

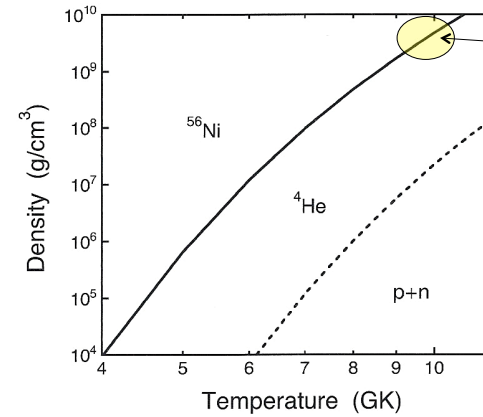


As silicon shells, typically one or at most two, burn out, the iron core grows in discontinuous spurts. It approaches instability.

Pressure is predominantly due to relativistic electrons. As they become increasingly relativistic, the structural adiabatic index of the iron core hovers precariously near 4/3. The presence of non-degenerate ions has a stabilizing influence, but the core is rapidly losing entropy to neutrinos making the concept of a Chandrasekhar Mass relevant.

In addition to neutrino losses there are also two other important instabilities:

- **Electron capture** – since pressure is dominantly from electrons, removing them reduces the pressure.
- **Photodisintegration** – which takes energy that might have provided pressure and uses it instead to pay a debt of negative nuclear energy generation.



Typical presupernova central conditions

The transition from the iron group to He and especially from He to $2n + 2p$ absorbs an enormous amount of energy.

Figure needs correcting for high temperature nuclear partition functions.

See also Clayton Fig 7-9 and discussion

Photodisintegration:

from page 12 of this lecture

$$Y_\alpha = C_\alpha(T_9, \rho) Y_n^2 Y_p^2$$

$$C_\alpha = \frac{1}{2} (5.94 \times 10^{33})^{-3} \rho^3 T_9^{9/2} \exp(328.36 / T_9)$$

where $328.36 = BE(\alpha)/kT \Rightarrow C_\alpha$ increases rapidly as $T \downarrow$

Setting $Y_\alpha = 1/8$ (i.e., $X = 1/2$ divided by 4) and $Y_p = Y_n = 1/4$ gives the line for the helium-nucleon transition on the previous page.

The Ni- α transition is given by Clayton problem 7-11 and eqn 7-22. It comes from solving the Saha equation for NSE (page 30 of these notes) for the case

$$X(^{56}\text{Ni}) = X_\alpha = 0.5 \Rightarrow Y(^{56}\text{Ni}) = 1/112; Y_\alpha = 1/8$$

$$Y(^{56}\text{Ni}) = C_{56}(T_9) \rho^{12} Y_\alpha^{13}$$

The photodisintegration of one gram of ^{56}Ni to one gram of α -particles absorbs:

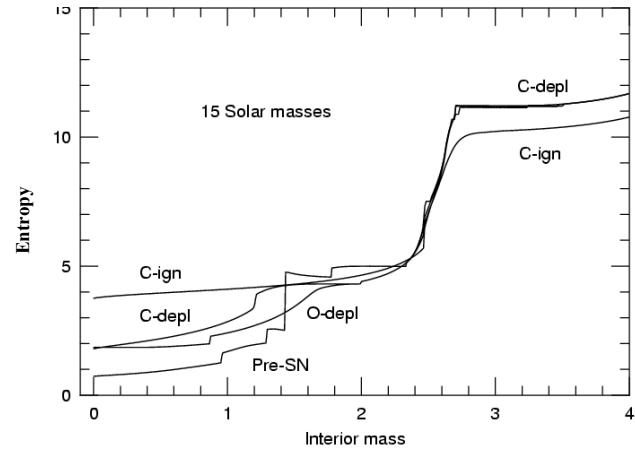
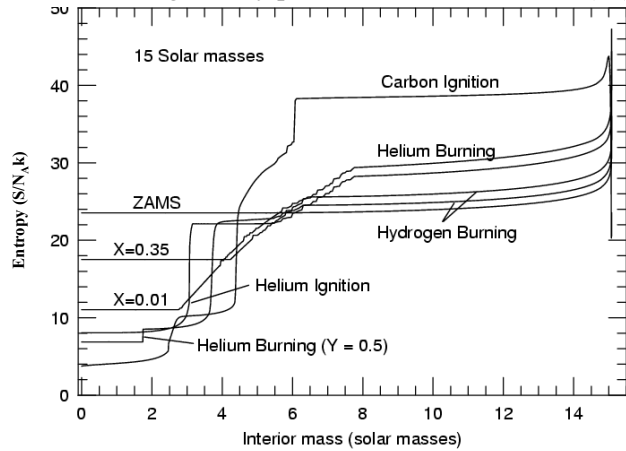
$$\begin{aligned} q_{nuc} &= 1.602 \times 10^{-6} N_A \sum (\delta Y_i)(BE_i) - q_v \text{ erg/gm} \\ &= 9.64 \times 10^{17} \left[\left(-\frac{1}{56}\right)(483.993) + \left(\frac{1}{4}\right)(28.296) \right] \\ &= -1.51 \times 10^{18} \text{ erg g}^{-1} \end{aligned}$$

Similarly, the photodisintegration of one gram of α 's to one gram of nucleons absorbs:

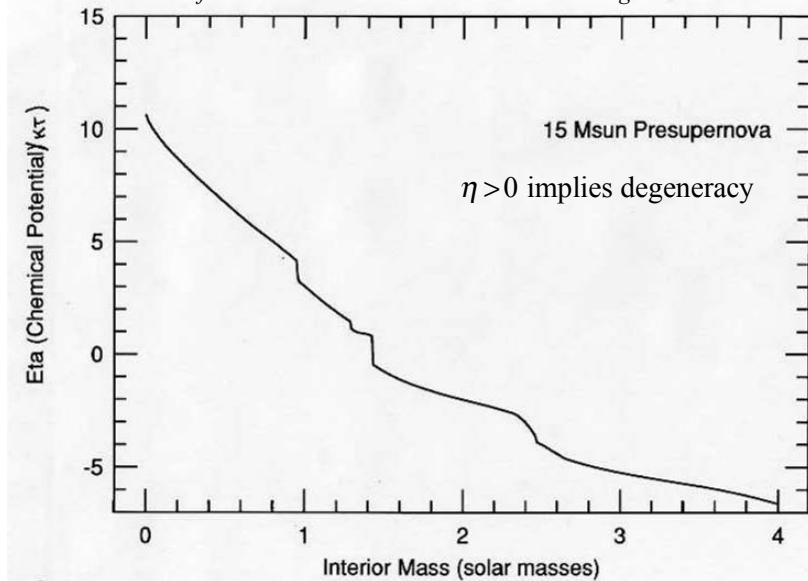
$$\begin{aligned} q_{nuc} &= 9.64 \times 10^{17} \left[-\left(\frac{1}{4}\right)(28.296) + 0 \right] \\ &= -6.82 \times 10^{18} \text{ erg g}^{-1} \end{aligned}$$

essentially undoing all the energy released by nuclear reactions since the zero age main sequence.

What about degeneracy? Can the core be supported by electron degeneracy pressure and form a stable (Fe) dwarf?



Because of increasing degeneracy the concept of a Chandrasekhar Mass for the iron core is relevant – but it must be generalized.



The Chandrasekhar Mass

Traditionally, for a fully relativistic, completely degenerate gas:

see lecture 1

$$\frac{P_c^3}{\rho_c^4} = \frac{K_{4/3}^3 \rho_c^4 Y_e^4}{\rho_c^4} = \frac{G^3 M^2}{20.745}$$

$$K_{4/3} = 1.2435 \times 10^{15} \text{ dyne cm}^{-2}$$

$$M_{Ch} = 5.83 Y_e^2$$

$$= 1.457 M_{\odot} \text{ at } Y_e = 0.50$$

This is often referred to loosely as 1.4 or even 1.44 solar masses

BUT

- | | <u>Effect on M_{Ch}</u> |
|---|--------------------------------------|
| 1) Y_e here is not 0.50 (Y_e is actually a function of radius or interior mass) | ↓ |
| 2) The electrons are not fully relativistic in the outer layers (γ is not 4/3 everywhere) | ↑ |
| 3) General relativity implies that gravity is stronger than classical and an infinite central density is not allowed (there exists a critical ρ for stability) | ↓ |
| 4) The gas is not ideal. Coulomb interactions reduce the pressure at high density | ↓ |
| 5) Finite temperature (entropy) corrections | ↑ |
| 6) Surface boundary pressure (if WD is inside a massive star) | ↓ |
| 7) Rotation | ↑ |

Coulomb Corrections

Three effects must be summed – electron-electron repulsion, ion-ion repulsion and electron ion attraction. Clayton p. 139 – 153 gives a simplified treatment and finds, over all, a decrement to the pressure (eq. 2-275)

$$\Delta P_{Coul} = -\frac{3}{10} \left(\frac{4\pi}{3} \right) Z^{2/3} e^2 n_e^{4/3}$$

Fortunately, the dependence of this correction on n_e is the same as relativistic degeneracy pressure. One can then just proceed to use a corrected

$$K_{4/3} = K_{4/3}^0 \left[1 - Z^{2/3} \frac{e^2}{\hbar c} \frac{2^{5/3}}{5} \left(\frac{3}{\pi} \right)^{1/3} \right]$$

$$= K_{4/3}^0 \left[1 - 4.56 \times 10^{-3} Z^{2/3} \right]$$

Effect on M_{Ch}

Relativistic corrections, both special and general, are treated by Shapiro and Teukolsky in *Black Holes, White Dwarfs, and Neutron Stars* pages 156ff. They find a critical density (entropy = 0).

$$\rho_c = 2.646 \times 10^{10} \left(\frac{0.50}{Y_e} \right)^2 \text{ gm cm}^{-3}$$

Above this density the white dwarf is (general relativistically) unstable to collapse. For $Y_e = 0.50$ this corresponds to a mass

$$M_{Ch} = 1.415 M_\odot$$

in general, the relativistic correction to the Newtonian value is

$$\frac{\Delta M}{M} = -9 \left[4 + 310 \left(\frac{0.50}{Y_e} \right)^{2/3} \right]^{-1}$$

$$= -2.87\% \quad Y_e = 0.50$$

$$= -2.67\% \quad Y_e = 0.45$$

where $K_{4/3}^0 = \frac{\hbar c}{4} (3\pi^2)^{1/3} N_A^{4/3}$

and $\frac{M_{Ch}}{M_{Ch}^0} = \left(\frac{K}{K^0} \right)^{3/2}$

hence

$$M_{Ch} = M_{Ch}^0 \left[1 - 0.0226 \left(\frac{Z}{6} \right)^{2/3} \right]$$

Putting the relativistic and Coulomb corrections together with the dependence on Y_e^2 one has

$$M_{Ch} = 1.38 M_\odot \quad \text{for } {}^{12}\text{C} \quad (Y_e = 0.50)$$

$$= 1.15 M_\odot \quad \text{for } {}^{56}\text{Fe} \quad (Y_e = \frac{26}{56} = 0.464)$$

$$= 1.08 M_\odot \quad \text{for Fe-core with } \langle Y_e \rangle \approx 0.45$$

So why are iron cores so big at collapse (1.3 - 2.0 M_\odot) and why do neutron stars have masses $\approx 1.4 M_\odot$?

Finite Entropy Corrections

Chandrasekhar (1938)
 Fowler & Hoyle (1960) p 573, eq. (17)
 Baron & Cooperstein, ApJ, 353, 597, (1990)

For $n=3$, $\gamma=4/3$, relativistic degeneracy

$$P_c = K_{4/3}^0 \left[1 + \frac{1}{x^2} \left(\frac{2\pi^2 k^2 T^2}{m_e^2 c^4} - 1 \right) + \dots \right] (\rho Y_e)^{4/3}$$

$$x = \frac{h}{mc} \left(\frac{3}{8\pi} n_e \right)^{1/3} = \frac{p_F}{m_e c} = 0.01009 (\rho Y_e)^{1/3}$$

(Clayton 2-48)

And since, in the appendix to these notes we show

$$s_e = \frac{\pi^2 k T Y_e}{\epsilon_F} \quad (\text{relativistic degeneracy})$$

one also has

$$M_{Ch} \approx M_{Ch}^0 \left(1 + \left(\frac{s_e}{\pi Y_e} \right)^2 + \dots \right)$$

The entropy of the radiation and ions also affects M_{Ch} , but much less.

This finite entropy correction is not important for isolated white dwarfs. They're too cold. But it is very important for understanding the final evolution of massive stars.

In particular, Baron & Cooperstein (1990) show that

$$P = P_o \left(1 + \frac{2}{3} \left(\frac{\pi k T}{\epsilon_F} \right)^2 + \dots \right)$$

$$\epsilon_F = p_F c = \left(\frac{3h^3 n_e}{8\pi} \right)^{1/3}$$

$$\epsilon_F = 1.11 (\rho_7 Y_e)^{1/3} \text{ MeV}$$

and since $M_{Ch} \propto K_{4/3}^{3/2}$ a first order expansion gives

$$M_{Ch} \approx M_{Ch}^0 \left(1 + \left(\frac{\pi k T}{\epsilon_F} \right)^2 \right)$$

Because of its finite entropy (i.e., because it is hot) the iron core develops a mass that, if it were cold, could not be supported by degeneracy pressure.

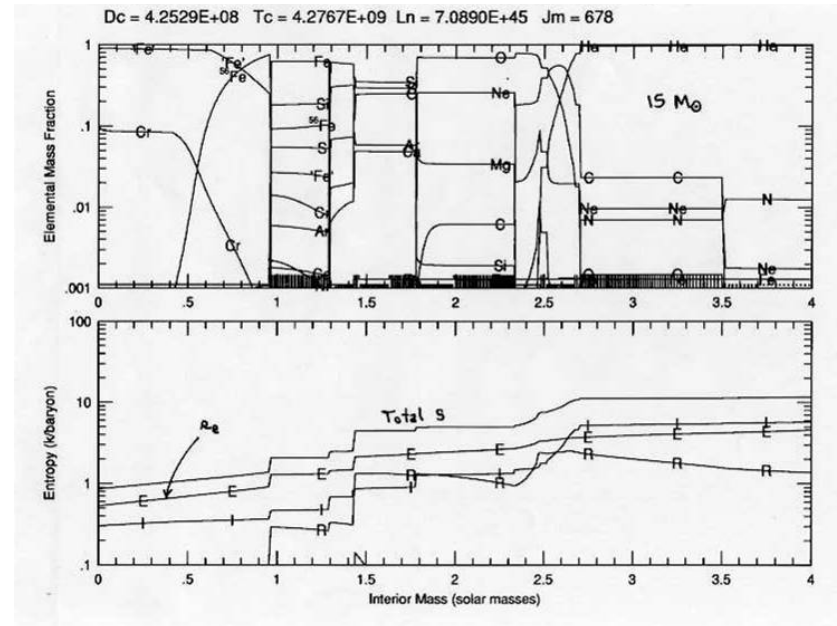
Because the core has no choice but to decrease its electronic entropy (by neutrino radiation, electron capture, and photodisintegration), and because its (hot) mass exceeds the Chandrasekhar mass, it must eventually collapse.

E.g., on the following pages are excerpts from the final day in the life of a $15 M_{\odot}$ star. During silicon shell burning, the electronic entropy ranges from 0.5 to 0.9 in the Fe core and is about 1.3 in the convective shell.

$$M_{\text{ch}} \approx 1.08 M_{\odot} \left(1 + \left(\frac{0.7}{\pi(0.45)} \right)^2 \right) = 1.34 M_{\odot} > 0.95 M_{\odot}$$

The Fe core plus Si shell is also stable because

$$M_{\text{ch}} \approx 1.08 M_{\odot} \left(1 + \left(\frac{1.0}{\pi(0.47)} \right)^2 \right) = 1.57 M_{\odot} > 1.3 M_{\odot}$$



But when Si burning in this shell is complete:

3) The Fe core is now $\sim 1.3 M_{\odot}$.

s_e central = 0.4

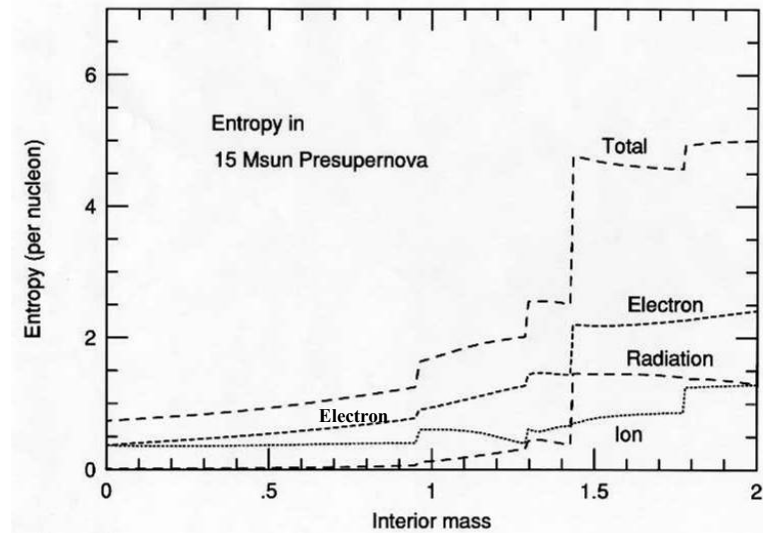
s_e at edge of Fe core = 1.1

average ≈ 0.7

M_{ch} now about $1.34 M_{\odot}$ (uncertain to at least a few times $0.01 M$)

Neutrino losses farther reduce s_e . So too do photodisintegration and electron capture as we shall see. And the boundary pressure of the overlying silicon shell is not entirely negligible.

The core collapses



The collapse begins on a Kelvin-Helmholz (neutrino) time scale and accelerates to a dynamic implosion as other instabilities are encountered.

Photodisintegration:

As the temperature and density rise, the star seeks a new fuel to burn, but instead encounters a phase transition in which the NSE distribution favors α – particles over bound nuclei. In fact, this transition never goes to completion owing to the large statistical weight afforded the excited states of the nuclei. But considerable energy is lost in a partial transformation. Recall for α 's.

$${}^{56}\text{Fe} \rightarrow 13\alpha + 4n$$

$$q_{nuc} \approx 9.65 \times 10^{17} \left(\frac{28.296}{4} - \frac{492.262}{56} \right) \Delta X_{56}$$

$q_{photo} = -1.7 \times 10^{18} \text{ erg g}^{-1}$

i.e., BE/A for a typical iron group nucleus

not really free neutrons. They stay locked inside bound nuclei that are progressively more neutron rich.

Electron capture

The pressure and entropy come mainly from electrons, but as the density increases, so does the Fermi energy, ϵ_F . The rise in ϵ_F means more electrons have enough energy to capture on nuclei turning protons to neutrons inside them. This reduces Y_e which in turn makes the pressure and entropy at a given density smaller.

$$\epsilon_F = 1.11 (\rho_7 Y_e)^{1/3} \text{ MeV}$$

By $2 \times 10^{10} \text{ g cm}^{-3}$, $\epsilon_F = 10 \text{ MeV}$ which is above the capture threshold for all but the most neutron-rich nuclei. There is also briefly a small abundance of free protons (up to 10^{-3} by mass) which captures electrons.

What happens?

As the density rises, so does the pressure (it never decreases at the middle), but so does gravity. The rise in pressure is not enough to maintain hydrostatic equilibrium, i.e., Γ remains slightly less than $4/3$. The collapse accelerates.

Photodisintegration decreases s_e because at constant total entropy (the collapse is almost adiabatic), s_i increases since 14 α -particles have more statistical weight than one nucleus. The increase in s_i comes at the expense of s_e .

But the star does not a) photodisintegrate to neutrons and protons; then b) capture electrons on free protons; and c) collapse to nuclear density as a free neutron gas as some texts naively describe.

Bound nuclei persist, then finally touch and melt into one gigantic nucleus with 10^{57} nucleons – the neutron star.

Y_e declines to about 0.37 before the core becomes opaque to neutrinos. (Y_e for an old cold neutron star is about 0.05; Y_e for the neutron star that bounces when a supernova occurs is about 0.29).

The effects of a) exceeding the Chandrasekhar mass, b) photodisintegration and c) electron capture operate together, not independently.

Appendix on Entropy

Zentralfriedhof
Vienna, Austria



$S = k \log W$

As discussed previously

1.3. ENTROPY

25

For a mixture of ions and radiation,

$$\epsilon = \frac{aT^4}{\rho} + \frac{3N_A \rho k T}{2\mu\rho}, \quad P = \frac{1}{3}aT^4 + \frac{\rho N_A k T}{\mu}$$

it follows (Clayton, page 120-121) that

$$S = \text{const} + \frac{N_A k}{\mu} \ln\left(\frac{T^{3/2}}{\rho}\right) + \frac{4aT^3}{3\rho} \quad \text{for ideal gas plus radiation}$$

Note that this implies, if T^3/ρ is a constant, that S will *decrease* with increasing T . The constant is both complicated and arbitrary. By convention

$$S_{\text{rad}} = \frac{4aT^3}{3\rho}$$

It is more common to normalize this to the gas constant, $N_A k$, to obtain the *entropy per baryon*

$$s_{\text{rad}} = \frac{4}{3} \frac{aT^3}{N_A k \rho}$$

The ions are rarely relativistic or degenerate. Any good thermodynamics text (e.g.,

$$(T^{3/2}/\rho)/(T^3/\rho) = T^{3/2}$$

dividing by k makes
 s dimensionless

Eg., just the radiation part

$$V \equiv 1/\rho$$

$$T dS = d\epsilon + P dV$$

$$\begin{aligned} T dS &= \frac{4aT^3}{\rho} dT + aT^4 \left(-\frac{1}{\rho^2} d\rho \right) + \frac{1}{3} aT^4 dV \\ &= \frac{4aT^3}{\rho} dT + aT^4 dV + \frac{1}{3} aT^4 dV \end{aligned}$$

$$dS = 4aT^2 V dT + \frac{4}{3} aT^3 dV$$

$$= d\left(\frac{4}{3} aT^3 V \right)$$

$$\text{So } S_{\text{rad}} = \left(\frac{4}{3} \frac{aT^3}{\rho} \right)$$

Reif p. 362) gives

$$s_{\text{ion}} = \frac{1}{\bar{A}} \left[\ln \left[\frac{g_0}{n} \left(\frac{2\pi M k T}{h^2} \right)^{3/2} \right] + \frac{5}{2} \right]$$

Reif
*Fundamentals of Statistical
and Thermal Physics*
McGraw Hill

where g_0 is the partition function (usually taken as 1), $n = \rho N_A / \bar{A}$, $\bar{A} = (\sum Y_i)^{-1}$, $M = \bar{A} / N_A$.

A similar expression holds for non-degenerate, non-relativistic electrons. The more general expression for electrons is more complicated (Cox and Guili 1966).

Cox and Guili
Principles of Stellar Structure
Second edition
A. Weiss et al
Cambridge Scientific Publishers

$$s_e = \frac{1}{\rho N_A} \left(\frac{P_e + E_e}{kT} - \eta n_e \right) \\ = \frac{P_e + E_e}{\rho N_A k T} - \eta Y_e$$

Cox and Guili
(24.76b)

where η is the electron chemical potential/kT

For a non-relativistic, non-degenerate electron gas,
Clayton 2-63 and 2-57 imply (for $\eta \ll 0$)

$$n_e = \frac{2(2\pi m_e k T)^{3/2}}{h^3} \left(e^\eta - \frac{e^{2\eta}}{2^{3/2}} + \dots \right)$$

which implies

$$-\eta \approx \ln \left[\frac{2(2\pi m_e k T)^{3/2}}{n_e h^3} \right]$$

which gives the ideal gas limit for electron entropy
(similar to ions but has Y_e and m_e)

$$\eta = \frac{\mu}{kT}$$

where μ , the chemical potential is defined by

$$T S = E + P V - \mu N \quad (\text{CG 10.20})$$

hence

$$S_e / V = (\epsilon_e \rho + P_e - \mu n_e) / T \\ S_e = \left(\epsilon_e + \frac{P_e}{\rho} - \frac{\mu n_e}{\rho} \right) / T$$

$$Y_e = \frac{n_e}{\rho N_A}$$

For an ideal gas

$$s_e = S_e / N_A k = \left(\frac{\epsilon_e \rho + P_e}{\rho N_A k T} \right) - \eta Y_e \\ = \left(\frac{5}{2} - \eta \right) Y_e \quad \eta \ll 0$$

$$\epsilon \rho + P = \left(\frac{3}{2} + 1 \right) \rho N_A k T$$

For $\eta \gg 1$ (great degeneracy)

$$n_e \approx \frac{8\pi}{c^3 h^3} (kT)^3 \frac{1}{3} \eta^3 \left[1 + \frac{\pi^2}{\eta^2} \right]$$

$$P_e \approx \frac{8\pi}{3c^3 h^3} (kT)^4 \frac{1}{4} \eta^4 \left[1 + \frac{2\pi^2}{\eta^2} + \frac{7\pi^4}{15\eta^4} \right] \quad (P_e \propto n_e^{4/3})$$

$$\frac{P_e}{n_e k T} \approx \frac{1}{4} \eta \left[\frac{1 + \frac{2\pi^2}{\eta^2}}{1 + \frac{\pi^2}{\eta^2}} \right] \gg 1$$

$\eta \gg \pi$ and keeping only terms in η^{-2}

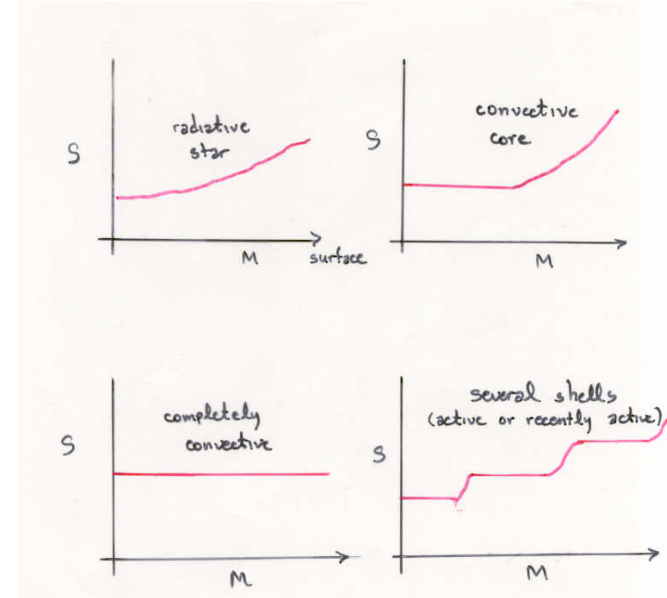
$$\frac{P_e}{n_e k T} \approx \frac{\eta}{4} \left[\left(1 + \frac{2\pi^2}{\eta^2} \right) \left(1 - \frac{\pi^2}{\eta^2} \right) \right] \\ \approx \frac{\eta}{4} \left(1 + \frac{2\pi^2}{\eta^2} - \frac{\pi^2}{\eta^2} \right) = \frac{\eta}{4} + \frac{\pi^2}{4\eta}$$

$$u_e = 3P_e \quad \eta = \frac{\mu}{kT} \quad \mu \approx \varepsilon_F$$

$$\begin{aligned} \frac{S_e}{V} &= (u_e + P_e - \mu n_e) / T \approx \frac{(4P_e - \mu n_e)}{T} \\ &\approx \frac{n_e kT \left(\eta + \frac{\pi^2}{\eta} \right) - \eta n_e kT}{T} = \frac{\pi^2 n_e k}{\eta} \\ &= \frac{\pi^2 \rho N_A Y_e k}{\eta} \end{aligned}$$

$$s_e = \frac{S_e}{\rho N_A k} \approx \frac{\pi^2 Y_e}{\eta}$$

$$s_e \approx \frac{\pi^2 k T Y_e}{\varepsilon_F}$$



3. Completely convective stars with negligible degeneracy will be polytropes of index $n = 3/2$. $\ln(T^{3/2}/\rho) = \text{constant}$ implies $T \propto \rho^{2/3}$ and $P \propto \rho^{5/3}$ and negligible radiation pressure (entropy)

4. Stars in which radiation completely dominates the entropy which are completely convective will be $n = 3$ polytropes. This is the case for extremely massive stars. $T^3/\rho = \text{constant}$ and $P \propto T^4$ implies $P \propto \rho^{4/3}$

5. Stars have negative heat capacity.

$$C = \frac{dQ}{dT} = T \frac{dS}{dT} = \frac{dS}{d \ln T}$$

For an ideal gas with negligible radiation,

$$\begin{aligned} S &\sim \left(\ln \frac{T^{3/2}}{\rho} \right) \frac{N_A k}{\mu} \\ \rho_c &\sim \frac{T_c^3}{M^2} \\ S &\sim \left(\ln \frac{M^2}{T_c^{3/2}} \right) \frac{N_A k}{\mu} \\ &= \left(2 \ln M - \frac{3}{2} \ln T_c \right) \frac{N_A k}{\mu} \\ C &\approx -\frac{3}{2} \frac{N_A k}{\mu} \end{aligned}$$

Thus the well known property of stars to get hotter as they radiate. It can be shown that the presence of radiation acts to inhibit this negative heat capacity

6. Stars of greater mass have higher entropy. Again $S \sim \ln T^{3/2}/\rho$, but T_c^3/ρ_c increases as M^2 . Thus as M rises so does S . This implies that more massive stars have higher entropy and a less centrally condensed

structure - less core convergence. This will be quite important.